MULTIFREQUENCY VLBI STUDIES
OF
COMPACT EXTRAGALACTIC JETS

by

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II  An Extremely Curved Relativistic Jet in PKS 2136+141

III  Coordinated Multiwavelength Observations of 3C 66A during the WEBT Campaign of 2003 – 2004

IV  Multifrequency VLBA Monitoring of 3C 273 during the INTEGRAL Campaign in 2003 – I. Kinematics of the Parsec Scale Jet from 43 GHz Data
T. Savolainen, K. Wiik, E. Valtaoja, and M. Tornikoski

V  Multifrequency VLBA Monitoring of 3C 273 during the INTEGRAL Campaign in 2003 – II. Extraction of the VLBI Component Spectra
T. Savolainen, K. Wiik, E. Valtaoja, and M. Tornikoski
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Abstract

In this doctoral thesis, multifrequency very long baseline interferometry (VLBI) observations of compact relativistic jets in active galactic nuclei are presented and analysed. Particularly, spectra of emission features in the parsec scale jet of the nearby quasar 3C 273 are measured by using simultaneous VLBI data at six frequencies from 5 to 86 GHz. In addition to this, VLBI monitoring data are combined with light curves from single-dish telescopes as well as with X-ray/γ-ray data from satellite observations.

The results obtained in this thesis give observational support to several such properties of compact jets that are usually assumed in the current standard model of AGN but seldom demonstrated observationally. First of all, a clear correlation between total flux density flares at high radio frequencies and ejections of superluminal components is found for a sample of 27 γ-bright blazars. This is in accordance with predictions of the shock-in-jet models. Typically, the VLBI core brightens during the flare and a new moving component is seen in the jet after the flare has peaked, indicating that much of the energy dissipation happens close to the VLBI core. Secondly, frequency dependent angular sizes found for some superluminal components in 3C 273 can be readily understood in the context of shock models. It is also shown that the flat radio spectrum of 3C 273 is composed of a number of synchrotron emitting features, each becoming self-absorbed at progressively lower frequency as they move out along the jet.

The measured spectra and sizes of the superluminal components are used to calculate the magnetic field density and the electron energy distribution normalisation factor in the parsec scale jet of 3C 273, independent of any equipartition assumption. The core shows magnetic field density $B \sim 1 \text{ G}$, and it seems that the decay of $B$ is roughly inversely proportional to the distance from the core. Unless the core is unbeamed, its magnetic energy density dominates over that of the relativistic electrons. Significant gradients across the jet in both velocity and in magnetic field density are found ~ 1.5 mas downstream of the core in 3C 273. On the northern side of the jet, the bulk Lorentz factor is ~ 7 and the magnetic field density is ~ $10^{-3}$ G, while on the southern side $\Gamma \sim 17$ and $B \sim 10^{-1}$ G. Hence, larger bulk velocity corresponds to larger $B$ and vice versa.

Multifrequency VLBI observations led also to a surprising discovery of an extremely curved jet in the quasar PKS 2136+141. The jet turns 210° in the plane of the sky, which is, to the author’s knowledge, the largest change ever observed in the position angle of an astrophysical jet. A model is proposed where the bending is due to a helical Kelvin-Helmholtz normal mode driven by a periodic perturbation at the base of the jet.
“There is a theory which states that if ever anybody discovers exactly what the Universe is for and why it is here, it will instantly disappear and be replaced by something even more bizarre and inexplicable. There is another theory which states that this has already happened.”

from ‘The Hitchhiker’s Guide to the Galaxy’ by Douglas Adams

Galaxies are the basic building blocks of the Universe – they are gravitationally bound systems containing typically $10^7 - 10^{12}$ stars, interstellar gas and dust, and (probably) dark matter. The vast majority of them are labelled as ‘normal’ galaxies whose radiative output is the combined shine of their billions of stars. Scattered among normal galaxies are intriguing systems showing the same morphological properties but much greater luminosity than their normal cousins of a comparable size. The radiative power of the nuclei of these active galaxies can outshine their stellar population by up to a thousand-fold and they are the most powerful long-lasting bodies in the Universe. Besides being highly luminous, the active galaxies are often also strongly variable sources, showing brightness variations down to time scales of hours. Active galaxies, or active galactic nuclei (AGN) as they are also commonly referred to, have been a subject of vigorous study for over four decades, and we believe that their primary energy source is the gravitational accretion process on a supermassive black hole. However, details of this process and much of the related physics still remain poorly understood.

One of the spectacular phenomena related to AGNs are the highly relativistic jets of plasma ejected from the most extreme objects. The jets can reach from the immediate vicinity of the supermassive black hole out into the intergalactic medium and distances of hundreds of kiloparsecs. Their radiation is due to non-thermal processes implying that particles in the jet are accelerated to extremely high energies – for example, it is well established that the radio-infrared-optical part of the jet spectrum is incoherent synchrotron radiation. In the past 25 years, considerable progress has been made in explaining spectral properties and flux density variability of radio bright quasars and BL Lac objects – together with basic kinematic properties of their parsec scale jets – in terms of shocks propagating in an underlying relativistic flow (e.g. Blandford & Königl 1979, Marscher & Gear 1985, Hughes et al. 1985, Türler et al. 2000). Although simple analytical shock models have been quite successful in explaining the radio observations, new high quality observational data together with state-of-the-art relativistic magneto-hydrodynamical simulations of the jets suggest that the jet physics is more complicated, and the observed
properties arise from a complex mixture of e.g. bulk and phase motions, viewing angle selection effects, and non-linear interactions between the perturbations and the underlying jet (Agudo et al. 2001, Gómez et al. 2001, Aloy et al. 2003, Hardee et al. 2005).

Besides radio and optical emission, many quasars and BL Lac objects emit a significant part of their energy in X-rays and $\gamma$-rays. This high energy component is commonly interpreted as inverse Compton scattering of soft target photons off the same relativistic electrons that emit the synchrotron radiation, but the origin of the target photons and the actual location of the emission region are matters of much controversy. Also, little is known about the launching mechanism of the jets, although magnetic fields are suspected to be involved. As a matter of fact, even the matter content of the jets is still unknown (normal or pair plasma) as is their dominant energy carrier (cold plasma, relativistic electrons or Poynting flux) and their total kinetic flux. We do not know where and how the jets are collimated, exactly how the electrons in the jet gain their non-thermal energy distribution, where in the jet and through which mechanism is the high energy radiation emitted, how is the behaviour of the jet connected to activity in the central engine, or how the jets remain stable up to distances of hundreds of kiloparsecs. There are also things related to shock models that are poorly understood – like the lack of a simple model for the exponentially rising part of the radio flare. Answering these questions requires advancement in all the fronts of investigation: observational, theoretical and computational. In this dissertation the observational path is pursued.

**Very long baseline interferometry (VLBI)** is a technique for combining radio telescopes separated by distances up to the size of the Earth (and more with antennas in space) into a single instrument, which provides the most accurate astronomical images in terms of angular resolution – a beam size of 50 microarcseconds is achieved with global observations at 3 mm wavelength (Bååth et al. 1992, Rantakyrö et al. 1998). Since VLBI requires targets with high brightness temperatures, the parsec scale jets in the cores of active galaxies – usually referred to as compact jets – are perfect sources to be observed with this technique. Already early VLBI observations revealed that the emission features in the jets move at apparently superluminal speeds (first result by Whitney et al. 1971, followed by an immediate confirmation by e.g. Cohen et al. 1971). This was a direct evidence for the existence of bulk relativistic outflow in AGN.

Before the advent of the Very Long Baseline Array (VLBA; Zensus et al. 1995) – a dedicated, U.S.-based VLBI array with excellent monitoring capabilities, frequency agility and good suitability for polarimetry – most VLBI studies of compact extragalactic jets concentrated on imaging the sources in total intensity at a given frequency band in order either to do morphological surveys or to monitor motions of the emission features in the jet (for a comprehensive review, see Zensus 1997). Morphological and kinematical data are interesting, but they alone give us a limited view of the physics of the jets. In order to study the matter and fields in the jets, and to understand the physical processes involved, more information is needed.

With the VLBI technique, polarisation in the parsec scale jets can also be measured, and this is an excellent probe of the degree of ordering of the magnetic field and the
mean direction of the field. Another important piece of information is available from multifrequency VLBI observations, since these allow one to study the shape of the radio spectrum at different parts of the jet, which in turn probes the magnetic field density and the relativistic particle energy density in the given region – or reveals thermal plasma through free-free absorption. While a number of polarisation monitoring campaigns have been conducted during the last ten years with the VLBA (for example, Lister & Homan 2005, and Jorstad et al. 2005), multifrequency VLBI studies targeting to observe the continuum spectrum of the jets have been less numerous. However, VLBA’s capability to make practically simultaneous multifrequency observations allows one to make such spectral studies and to obtain important information on the physical conditions within regions in the innermost few parsecs of AGNs. This has been demonstrated in the studies of free-free absorption in nearby AGNs by Walker et al. (2000), Marr et al. (2001) and Vermeulen et al. (2003). One aim of the work presented in this thesis was to use the multifrequency approach in the VLBI observations of blazar jets, and to measure the radio spectrum along the parsec scale jet of nearby quasar 3C 273.

For studying the relativistic jets, blazars – a violently variable subclass of AGN, grouping together flat-spectrum radio loud quasars and BL Lac objects – are the most convenient sources, since their highly relativistic jets are pointed towards us causing the jet emission to dominate over the unbeamed emission from the central engine. Because blazars radiate over the whole electromagnetic spectrum from radio to TeV γ-rays, it is important that these sources are also studied over a wide range of wavebands. A promising way of exploration in the future is to combine the (polarimetric) VLBI imaging data with multiwaveband variability monitoring data covering as many frequencies as possible from radio to γ-rays (see e.g. Marscher 2006a). Such an approach should ultimately allow us to connect different events in the light curves with features in the VLBI images, to place the emission regions at different frequencies relative to each other, and to construct crude emission maps over the whole spectrum. The second aim of the work presented in this thesis was to combine VLBI images with light curves from single-dish telescopes, as well as with X-ray/γ-ray data from satellite observations.

The thesis consists of a general introductory part and five research articles (hereafter Papers I–V). In Paper I, VLBI monitoring data of a sample of γ-bright AGNs is compared with the longterm flux density curves at 22 and 37 GHz, and a clear connection between the ejection events in the jet and the mm-wave flares is established. Paper II shows how the multifrequency approach in VLBI imaging can sometimes provide surprising results just by displaying the source in different angular scales. A striking example of this is our discovery of a breathtaking 210° turn of the jet in a distant quasar PKS 2136+141. First results from a large multiwaveband campaign of 3C 66A, containing also multifrequency VLBI data, are presented in Paper III.

Papers IV and V discuss results from a multifrequency VLBA monitoring of 3C 273, which was carried out in conjunction with X-ray and γ-ray observations made with XMM-Newton and INTEGRAL satellites in 2003. In Paper IV, complex component motions in the parsec scale jet of 3C 273 are studied, and the jet Doppler factor is accurately
measured. We extract spectra of 16 jet emission features from the multifrequency VLBI
data in Paper V. The spectra are of high quality and have a frequency coverage from 5
to 86 GHz, which allows us to measure the spectral turnover of several jet features and
to calculate important physical parameters – including e.g. the magnetic field density
and the predicted amount of high energy radiation emitted through synchrotron self-
Compton (SSC) process. The latter is compared with hard X-ray flux density from the
satellite observations.
“Some things are just fundamentally not easy.”

– Chris Carilli (from professor Karl Menten’s web page of Famous Last Words)

At the diffraction limit, the angular resolution of a telescope is approximately $\frac{\lambda}{D}$ where $D$ is the diameter of the aperture and $\lambda$ is the wavelength of observation. Since radio wavelengths are some $10^3$ to $10^7$ times longer than optical wavelengths, the angular resolution achieved by any single-aperture radio telescope is very crude by optical standards. Together with a number of ingenious technological innovations the desire to overcome this handicap drove the development of radio interferometry in the 1950s and 1960s. The angular resolution of an interferometer depends on the separation of the interferometer elements, so-called baseline, not on the aperture size of an individual telescope. Therefore, by increasing the distance between the interferometer elements, the angular resolution of the instrument can be improved.

As explained later, an interferometer is basically a spatial filter and a single baseline can measure only one spatial frequency of the source brightness distribution, i.e. one angular scale. The early two-element interferometers were designed to measure angular diameters of celestial objects by progressively increasing the baseline until coherence was lost and interference pattern disappeared. A technique called aperture synthesis, the simultaneous use of several interferometer baselines of an array of pairwise connected telescopes, made it possible to observe a large number of spatial frequencies and form images of the source brightness distribution (Ryle 1957; Scott et al. 1961). The use of the rotation of the Earth further increases the effective number of baselines (Ryle & Neville 1962; Ryle 1962), and these techniques have led to construction of radio telescope arrays such as VLA, MERLIN, and the Westerbork Synthesis Radio Telescope, which have been very successful in charting the radio sky and making detailed images of radio sources at angular resolutions comparable to optical telescopes. For a good historical review of the development of high-resolution imaging in radio astronomy, the reader is referred to Kellermann & Moran (2001).

The radio interferometers in the 1960s were arrays of locally distributed antennas that were directly connected by wires or by radio links. Already in the mid-1960s, the radio-linked interferometer measurements had shown directly that there exists a source population with angular sizes smaller than 0.05 arcsec (Adgie et al. 1965), and observations of interplanetary scintillation indicated a number of objects that would be as small as a few milliarcseconds (Cohen 1965). The measured spectra of these compact radio
sources showed signs of synchrotron self-absorption and suggested that their sizes could be as small as 1 mas (Slysh 1963; Williams 1963). The discovery of rapid variability in flat spectrum radio sources (Sholomitskii 1965; Dent 1965) implied that the linear sizes of the emission regions are less than one light-year. If the redshifts of these sources had cosmological origin, their angular sizes would be below 1 mas. The baselines needed to resolve these compact sources were thousands of kilometres, a lot longer than could be realised with a connected-element array.

The solution was to record the observations on tape, together with time marks, independently at each antenna, and to form the interference fringes after the actual observation. This technique required very high stability of the local frequency standard as well as tape recorders capable of a high speed, broad band recording. The first successful experiments with the new very long baseline interferometry technique were carried out in the late 1960s and the first fringes on a baseline of several thousand kilometres were observed in 1967 (Broten et al. 1967). Since then the progress has been fast, and it has been intimately linked to both the developments in receiver and recorder technology, and to the rapidly increasing computational power. Also the development of sophisticated algorithms for the data processing have played a significant part.

Today, the field has reached some level of maturity, and there exist several VLBI facilities such as the European VLBI Network (EVN), the NRAO’s Very Long Baseline Array (VLBA) in the United States, the Japanese VLBI Exploration of Radio Astrometry (VERA), the Australian Long Baseline Array, and the Global Millimetre-VLBI Array (GMVA). Also the observing techniques in VLBI have become more user-friendly – for instance, total intensity imaging observations with the VLBA are made with comparable ease to those done with other modern astronomical instruments. However, there are still some difficult aspects and pitfalls in VLBI observations, especially when more advanced techniques such as polarimetry, spectrometry or phase-referencing are used. Also, the analysis and interpretation of the VLBI data involves some problematic issues.

People outside the radio astronomy or interferometry community are often unfamiliar with concepts related to VLBI or interferometry in general. Therefore, I will shortly present the central ideas of astronomical interferometry, and give a quick introduction to VLBI observations and data processing. The following treatment concentrates mostly on simple imaging observations in total intensity, and more difficult techniques are either totally omitted or just briefly mentioned – except for a section in the end of the chapter that discusses multifrequency observations. Good references for VLBI technique are books by Thompson et al. (1986) and Zensus et al. (1995). A recommended reference in VLBI polarimetry is Leppänen (1995).

2.1 Basics of interferometry

In the following, basic concepts of the interferometry are introduced, and the spatio-temporal coherence function is explained. To avoid handling of tensor functions, for
a moment we ignore the vector nature of electromagnetic radiation and treat it as if it were a scalar field. This basically means that we leave out polarisation phenomena from our treatment. Also, we consider all fields to be complex quantities although in nature the true field is, of course, described by the real part of the complex field quantity. No proofs for the theorems are given here since those can be found in the excellent textbook *Principles of Optics* by Born & Wolf (1980). A concise presentation of the central ideas of interferometry can also be found in Clark (1998).

### 2.1.1 Coherence in a stationary electric field

Let us consider electromagnetic radiation with an effective spectral width $\Delta \nu$ ("bandwidth") coming from a distant source with a finite extent. For simplicity, we assume that the radiation propagates in a vacuum and therefore, we may ignore all the amplitude and phase distortions introduced by the intervening medium. According to Fourier’s theorem, over a finite time interval such a wave field can be expressed as a sum of strictly monochromatic wave trains. While in a monochromatic wave field the amplitude of vibrations at any point $P$ is constant and the phase is changing linearly, in a wave field produced by a real source the amplitude and phase undergo irregular fluctuations. The rapidity of these fluctuations depends primarily on $\Delta \nu$. The complex amplitude of the wave field remains reasonably constant during a time interval which is short compared to a *coherence time* $\Delta t \sim 1/\Delta \nu$, i.e. as long as the change of the relative phase of any two Fourier components is much less than $2\pi$. For such time intervals, the general wave field with a bandwidth $\Delta \nu$ behaves like a monochromatic wave with the mean frequency $\bar{\nu}$.

Let us next consider the disturbances at two points $P_1$ and $P_2$ in a quasi-monochromatic wave field described above. Points $P_1$ and $P_2$ are assumed to be at a large distance from the source of the radiation. When $P_1$ and $P_2$ are sufficiently close to each other, it is expected that the fluctuations of the wave field at these two points are not independent. If the distance between $P_1$ and $P_2$ is so short that the difference in the path lengths between the source and these two points is small compared to the mean wavelength $\bar{\lambda}$, the fluctuations at $P_1$ and $P_2$ will be essentially the same. Also, some correlation between the fluctuations will be present even at larger separations as long as the path difference does not exceed a *coherence length* $c\Delta t \sim \bar{\lambda}^2/\Delta \lambda$. It is expected that the degree of this correlation is intimately related to the contrast of the interference pattern (i.e. fringes) in the Young’s double slit experiment. In a coherent case (e.g. $P_1$ and $P_2$ both receive very narrow bandwidth light from the same point source) the correlation is high and the fringes are sharp, while in an incoherent case (e.g. $P_1$ and $P_2$ receive light from two different sources) the degree of correlation is zero and the interference pattern disappears. Normally, we have a situation which is somewhere between these two extremes and the field fluctuations are called partially coherent.
Coherence function

The degree of correlation between the vibrations measured at different points $P_1$ and $P_2$ in the electric field is expected to describe the field properties. The exact way how the degree of correlation relates to the field properties is contained in two elegant theorems: Wiener–Khinchin theorem and van Cittert–Zernike theorem.

First, we define the spatio-temporal coherence function (also known as mutual intensity in optics) of the electric field $E(P, t)$ between points $P_1$ and $P_2$ at times $t_1$ and $t_2$:

$$V(P_1, t_1, P_2, t_2) = \langle E(P_1, t_1)E^*(P_2, t_2) \rangle \quad (2.1)$$

where the asterisk marks the complex conjugate of the field and the brackets denote an average over a time interval, which is long compared to the mean period of oscillation of the quasi-monochromatic field. We have two important special cases of $V(P_1, t_1, P_2, t_2)$, namely $P_1 = P_2$ and $t_1 = t_2$.

Wiener–Khinchin theorem

If the coherence function is measured at a single point of a stationary electric field (such a field is expected from e.g. an astronomical source whose properties do not vary significantly between the measurements),

$$V(P_1, t_1, P_1, t_1 + \tau) = V(P_1, t_1, P_1, t_1) \equiv V_\tau(\tau) \quad (2.2)$$

i.e. the degree of coherence depends solely on the delay between the measurements. Physically, the temporal coherence function $V_\tau(\tau)$ describes the degree of correlation between fields along a given wave train and it is exactly what is measured in laboratory with the Michelson interferometer where light from a single source is split, then propagated along two paths of different lengths, and finally recombined and observed.

The temporal coherence function is very useful because of the Wiener–Khinchin theorem of stationary random processes. Its optical analogue states that the normalised value of $V_\tau(\tau)$ equals to the normalised Fourier transform of the spectral density $G(\nu)$:

$$\frac{V_\tau(\tau)}{V_\tau(0)} = \frac{\int_0^{+\infty} G(\nu)e^{-i2\pi\nu\tau} d\nu}{\int_0^{+\infty} G(\nu) d\nu} \quad (2.3)$$

Now, according to this theorem it is possible to unambiguously recover the spectrum of the source from measurements of $V_\tau(\tau)$ without use of any dispersing element. The Fourier transform spectrometers are the instruments to do this.

Van Cittert–Zernike theorem

Let us next consider a measurement of a quasi-monochromatic wave field at two separate points $P_1$ and $P_2$. If the time delay, $\tau \ll 1/\Delta \nu$, is very short compared to the coherence time, $V(P_1, t_1, P_2, t_1 + \tau)$ does not differ appreciably from the spatial coherence
function \( V(P_1, t_1, P_2, t_1) \equiv V_s(r) \). Here, \( r \) is the vector separation between the points \( P_1 \) and \( P_2 \). From astronomer’s point of view, the interest in \( V_s(r) \) arises from the van Cittert–Zernike theorem, which states that for a spatially incoherent source \( \sigma \) in the far field, the normalised value of \( V_s(r) \) is equal to the normalised Fourier transform of the source brightness distribution, \( I(s) \):

\[
\frac{V_s(r)}{|V_s(0)|} = \frac{\int \sigma I(s) e^{-\imath 2\pi v_s \cdot r/c} ds}{\int \sigma I(s) ds},
\]

(2.4)

where \( s \) is a direction vector pointing to the source.

If we assume that the observed source subtends only a small portion of the celestial sphere, we may simplify equation 2.4:

\[
\frac{V_s(u, v)}{|V_s(0, 0)|} = \frac{\int \int I(l, m) e^{-\imath 2\pi v(u l + v m)/c} dl dm}{\int \int I(l, m) dl dm},
\]

(2.5)

where angular coordinates \((l, m)\) are measured with respect to a direction \( s_0 = (0, 0, 1) \), which is called the phase tracking centre in the sky. In this coordinate system, \( u \) and \( v \) are the baseline coordinates perpendicular to \( s_0 \) and measured in wavelengths. Since the Fourier transform is invertible, it is possible to recover the sky brightness distribution from a suitable number of \( V_s(u, v) \) measurements:

\[
I(l, m) = \int \int V_s(u, v) e^{\imath 2\pi v(u l + v m)/c} du dv,
\]

(2.6)

where the scaling constant \((\int \int I(l, m) dl dm)/|V_s(0, 0)|\) is absorbed in \( V_s(u, v) \). This is the basis of astronomical interferometry.

How is the spatial coherence function \( V_s(u, v) \) related to the interference pattern observed in the classic Young’s double slit experiment? The visibility of the fringes \( \mathcal{V}(u, v) \) in Young’s experiment is defined as the contrast between light and dark stripes:

\[
\mathcal{V}(u, v) = \frac{I_{\text{max}} - I_{\text{min}}}{I_{\text{max}} + I_{\text{min}}}
\]

(2.7)

where \( I_{\text{max}} \) and \( I_{\text{min}} \) are the maximum and minimum intensities of the fringe pattern, respectively. It turns out that \( \mathcal{V}(u, v) \) is exactly the normalised amplitude of the complex spatial coherence function:

\[
\mathcal{V}(u, v) = \left| \frac{V_s(u, v)}{|V_s(0, 0)|} \right|.
\]

(2.8)

On the other hand, the phase of \( V_s(u, v) \) turns out to be proportional to the displacement of the quasi-monochromatic fringes relative to fringes that would be formed with monochromatic and co-phasal illumination of the slit openings in the direction parallel to lines joining the openings. Basically, the phase of \( V_s(u, v) \) tells how far a point source is from the phase tracking centre. Since there is a close bearing with the spatial coherence function and the fringe visibility in Young’s experiment, \( V_s(u, v) \) is commonly called the complex visibility.
2.1 BASICS OF INTERFEROMETRY

2.1.2 Measuring the spatial coherence function

Sparse sampling of the \((u, v)\) plane

Radio (and optical) interferometers are devices to measure the spatial coherence function \(V(u,v)\) (from now on, we drop out the subscript \(s\)). In practice, \(V(u,v)\) cannot be measured everywhere, but is sampled at certain points of the \((u,v)\) plane, and we can describe the sampling by a sampling function \(S(u,v)\), which is unity at points where \(V(u,v)\) is measured and zero elsewhere (see Fig. 2.1 for examples of the \((u,v)\) coverage at the VLBA). Now, the source brightness distribution derived from interferometric measurements with sampling function \(S(u,v)\) is

\[
I^D(l,m) = \int \int V(u,v)S(u,v)e^{i2\pi\nu(ul+vm)/c} \, du \, dv, \quad (2.9)
\]

which is often called the **dirty image**\(^1\). \(I^D(l,m)\) is a convolution between the true brightness distribution \(I(l,m)\) and the point spread function of the interferometer

\[
B(l,m) = \int \int S(u,v)e^{i2\pi\nu(ul+vm)/c} \, du \, dv, \quad (2.10)
\]

which is also known as the **synthesised beam**. It is therefore clear that in order to find out the true source brightness distribution, we need to deconvolve the dirty image. Methods for this are introduced in section 2.3.

\(^1\)Notice that we have again absorbed the scaling constant \((\int \int I(l,m) \, dl \, dm)/|V(0,0)|\) in \(V(u,v)\).
Since a measurement of $V$ in one point of the $(u,v)$ plane gives us one Fourier component of the source brightness distribution, one can also think the interferometer as a spatial filter, which passes through only some of the spatial frequencies of the Fourier transform of the source brightness distribution. Therefore, the angular size scales observable by an interferometer depend on the range of baseline lengths: the shortest baseline defines the largest observable structure size and the longest baseline defines the smallest observable structure, i.e. angular resolution$^2$.

**Primary beam pattern**

In the real world, the interferometer elements are for example radio telescopes, which have some directional sensitivity pattern (primary beam) $A(l,m)$. Hence, in practice, the measured complex visibility is

$$V(u,v) = \int \int A(l,m) I(l,m) e^{-i2\pi\nu(ul+vm)/c} \, dl \, dm.$$  \hspace{1cm} (2.11)

As Clark (1998) points out, $A(l,m)$ is not only nuissance. In fact, since $A(l,m)$ falls quickly to zero except in the vicinity of the pointing centre, the assumption of a small source used earlier in the text is valid in most practical cases.

**A note about spectroscopy**

In the above discussion, we have all the time considered quasi-monochromatic wave fields. The complex visibility function $V(u,v)$ given above relates to the source brightness distribution at a narrow frequency band and it really should be noted as $V(u,v,\nu)$. In an interferometer, calculating $V(u,v,\nu)$ can be realised with filter banks in the early, linear, stages of the signal path. However, with the current technology, it is easier to insert additional multipliers in the digital correlator to calculate the spatial correlation as a function of delay $\tau$. Since adding a delay simply means adding a phase shift of $\tau \nu$ to each quasi-monochromatic component $V(u,v,\nu)$,

$$V(u,v,\tau) = \int V(u,v,\nu) e^{-i2\pi\nu\tau} \, d\nu.$$  \hspace{1cm} (2.12)

This is a Fourier transform, and it can be inverted to obtain $V(u,v,\nu)^3$. The reader should note that it can be expected already from equation 2.3 that $V(u,v,\tau)$ is related to

---

$^2$Conventional telescopes can also be treated as devices for measuring the spatial coherence function -- they just sample all the points of the $(u,v)$ plane within their aperture, and do the Fourier-transform automatically in their image plane. The angular resolution is again defined by the longest baseline, i.e. the aperture size.

$^3$This is how the conventional lag-correlator works. There is also another correlator type -- so-called FX correlator, which is used e.g. at the VLBA. In an FX correlator, the delay and phase compensated input signals from each station are directly Fourier transformed to “station spectra”, which are then cross-multiplied (and averaged) pairwise to obtain the cross-power spectrum for each baseline. See e.g. Romney (1998) for a review of the correlator types.
the source spectrum. Modern correlators produce baseline cross-power spectra at their output, which are then averaged over frequency for continuum applications. Therefore, an interferometer is naturally also a spectrometer.

2.2 VLBI in practice

In VLBI, there is no direct link between the interferometer elements. The signal from the source is recorded at each antenna and transported to the correlator where the fringes are formed after the observing session. The correlator reproduces and resynchronises the recorded samples recreating in a single room the situation that existed across the world at the time of observation. In principle however, the observer does not have to care much about the operation of a VLBI system: the correlated visibilities are delivered to him just as in the case of a connected-element interferometer. Therefore, the details of the VLBI signal path from the antenna to the correlator and the intricacies of the correlation process are glossed over here. What the observer needs to care about are the unstable phases of the uncalibrated VLBI data and the phase gradients with respect to frequency and time.

2.2.1 About phases

The main sources of unstable phases in VLBI are the non-ideal frequency standards of the antennas, imperfections of the correlator model, and variations in the refractive index of the atmosphere. In the early years of the VLBI experiments, the quality of the frequency standard used to be the factor limiting the coherence time, but since then the technology has advanced so that nowadays the atmosphere is the main source of phase noise.

Since the correlator needs to calculate cross-correlation for signals that correspond to the same arriving wavefront, the signals must be corrected for the delay due to the antennas having different path lengths to the source, and for the fringe rate, which is due to antennas moving at different speeds along the direction to the source, causing different Doppler shifts. These corrections are derived from the correlator model, which is a geometrical model of the positions of the antennas in 3-D space. The model takes into account some very subtle geodetic effects (see e.g. Table 22.1 in Walker 1998). Errors in the correlator model – which can be due to e.g. inaccurate source or antenna positions, station clock offsets, errors in the clock rates, errors in the Earth model, or errors in the atmosphere model – leave residual delay and rate into the data causing phase slopes over frequency and time. In an imaging experiment, it is desirable to average the data in frequency and in time in order to reduce the amount of data and, most importantly, to allow self-calibration to be done over longer time intervals and wider bandwidths, which increases the sensitivity. Now, if there are phase slopes present, the averaging

\[ \text{An interested reader is again referred to Thompson et al. (1986) and Zensus et al. (1995).} \]
will degrade the visibility amplitudes and introduce non-closing errors, which cannot be described as a product of antenna gains and thus cannot be handled with self-calibration techniques (see section 2.3.2). Therefore, one needs to estimate and remove the residual delay and rate from the data before averaging. This is called fringe fitting and it is discussed in the next section.

At low frequencies, the refractivity of the ionosphere introduces a phase shift, which is inversely proportional to frequency. This phase shift depends on the electron content of the ionosphere, which varies on both diurnal and minutes timescales. At microwave and at millimetre wavelengths, the troposphere introduces a phase shift, which is proportional to frequency (i.e. the tropospheric delay is frequency independent). It is due to the refractivity caused by oxygen, nitrogen, and water vapour. The effect of the atmosphere on the phases can be divided into two parts: the dry component causes a systematic variation of the delay and the variable wet component causes phase noise. The large excess delays caused by the dry atmosphere, which can be over 200 cm, are normally removed by the correlator model. The rapidly changing wet component, on the other hand, is much more difficult to deal with, and it usually dominates the phase errors in the VLBI data at frequencies $\gtrsim 5$ GHz (see Carilli et al. 1998 for a detailed discussion of the tropospheric effects). This phase noise can be reduced with self-calibration in the case of a target source that is strong enough to be detected within the coherence time (see section 2.3.2). For weaker sources, phase-referencing to a strong calibrator, located near the target source, can be used\(^5\).

2.2.2 VLBI calibration path

The basic task of calibration can be stated as follows: to a rather good accuracy, we may assume that the effects introduced on $V$ by the measurement device and by the atmosphere are station-dependent, and that the measured complex visibility $V'_{ij}(t, \nu)$ between the interferometer elements $i$ and $j$ at frequency $\nu$ and time $t$ is

$$V'_{ij}(t, \nu) = G_i(t, \nu)G^*_j(t, \nu)V_{ij}(t, \nu)$$  \hspace{1cm} (2.13)

where the complex quantity $G_i(t, \nu)$ is the element gain. There is also a baseline-based gain term, which, however, is usually small if the delay and rate errors have been corrected sufficiently before data averaging. Therefore, the calibration of $V'$ boils down to finding out $G_i(t, \nu)$ for each antenna.

In the following, we will briefly overview the calibration steps, which are carried out before the final self-calibration (which is discussed in the next section in the context of imaging). These calibration steps include a priori amplitude calibration, corrections for residual delays and rates, and bandpass calibration. All of them can be done in the NRAO Astronomical Image Processing System (AIPS; Greisen 1988; Bridle & Greisen

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\(^5\)In the case of phase-referencing, the calibrator data are self-calibrated.
1994) – a powerful data reduction package for radio interferometry. Detailed instructions for using the data reduction tasks in the AIPS are given in the AIPS Cookbook\(^6\).

**Amplitude calibration**

The correlator output are raw correlation coefficients \( \rho \), which need to be converted to the correlated flux densities

\[
S_{12}^c = \rho_{12} \cdot b \cdot 2k \cdot 10^{26} \sqrt{\frac{T_{\text{sys},1} T_{\text{sys},2}}{A_{\text{eff},1} A_{\text{eff},2}}} \quad [\text{Jy}],
\]

where \( b \) is a correlator specific scaling factor including also corrections for coarse digitisation, \( k \) is the Boltzmann constant, \( T_{\text{sys},i} \) is the system temperature of the \( i \)th antenna, and \( A_{\text{eff},i} \) is the effective aperture of the \( i \)th antenna. System temperatures at the stations are measured during the observation e.g. by injecting a low level, broadband noise signal of known strength in the receiver front end at some duty cycle. The effective aperture of the antenna, \( A_{\text{eff},i} \), is a product of the geometric area of the antenna and the elevation dependent aperture efficiency, which is constituted of several factors, such as surface accuracy, spillover, illumination, subreflector blockage, radiation efficiency, pointing errors etc. The effective aperture as a function of direction in the sky – so-called gain curve – is provided by calibration measurements carried out by the observatory staff at times different from VLBI observations. The application of measured system temperatures and gain curves on the correlation coefficients from the correlator is called \textit{a priori} amplitude calibration, and it can be done with the AIPS task APACAL.

At frequencies higher than about 15 GHz, it is necessary to take into account the atmospheric absorption, which is mostly due to the resonance lines of the oxygen and water vapour. Since we may assume the total opacity of the atmosphere to be approximately proportional to \( \sec z \), where \( z \) is the zenith angle, correcting for the opacity in the VLBI data can be done if the system temperature measurements cover a sufficient elevation range and the effective atmospheric temperature \( T_{\text{atm}} \) is known. The latter can be estimated from the maximum ground level temperature of the day, and at the VLBA this information is provided by a weather station at the antenna site (Leppänen 1993). From \( T_{\text{sys}}, T_{\text{atm}}, \) and elevation dependent spillover temperature \( T_{\text{spill}} \), the receiver temperature \( T_{\text{rec}} \) and the zenith opacity \( \tau_0 \) can be obtained by fitting. After \( T_{\text{rec}} \) (which can be assumed to be constant during the observation) is known, the atmospheric attenuation can be calculated from

\[
L \equiv e^\tau = \frac{T_{\text{atm}}}{T_{\text{atm}} - T_{\text{sky}}},
\]

where

\[
T_{\text{sky}} = T_{\text{sys}} - T_{\text{rec}} - T_{\text{spill}}.
\]

\(^6\)http://www.aoc.nrao.edu/aips/cook.html
The opacity correction can be calculated in the APCAL along with the a priori amplitude calibration. However, when fitting for $T_{\text{rec}}$, APCAL assumes that the zenith opacity is constant, which is generally not true at high frequencies. This, together with short period disturbances like rain showers, make the least-squares fit sometimes difficult, and it is advisable to examine carefully the fit results before applying the opacity correction determined by the APCAL. It sometimes helps to split the observation into two or more parts in order to find a period of reasonably constant zenith opacity.

Correcting the amplitudes for possible errors in the sampler threshold levels can be also considered as a part of a priori amplitude calibration, and it is done with the AIPS task ACCOR. At this point, it is also a good idea to look through the data and flag any obviously bad visibilities, which may have gone unnoticed by the automatic flagging system. Flagging the clearly bad data at this point may help fringe fitting later.

**Single-band delay**

In uncalibrated VLBI data, there will be a significant residual delay due to clock offsets at the stations. Furthermore, the delays will be different for different basebands, since the signals go through different electronics and are affected for example by cable stretching. Also, there will be phase offsets between the basebands. Removing the large single-band delay, and aligning the phases across the basebands can be done in two ways. At the VLBA antennas, there exists a so-called pulse-cal system, which injects a string of sharp pulses at one microsecond intervals in the receiver front end. This generates a spectrum consisting of a series of lines at harmonics of 1 MHz, which can be used to calibrate variations in the delay and phase due to the instrumental effects. The AIPS task PCCOR uses the pulse-cal tones together with a sort of a fringe fit on a short reference scan to calculate the instrumental phase and delay offsets. The calibrator scan is needed to solve for $2\pi$ ambiguities.

If pulse-cal tones are for some reason unavailable, one can determine the single band delays and phase offsets by fringe fitting a short scan of a strong calibrator source and zeroing the fringe rate solutions. The delays obtained for this scan are then applied for all the data. The obvious weakness of this approach compared to using pulse-cal tones is that it does not take into account possible time dependent variations of the instrumental delays. At the VLBA, however, such time variability does not seem to be a serious problem.

**Fringe fitting**

At this point, there will still be time dependent residual delays and rates in the data due to errors in the correlator model. Estimating these delay and rate residuals is referred to as fringe fitting. A fringe fit is basically a self-calibration (see section 2.3.2) that includes, in addition to phase gains, the first derivatives of the phase gains.

Usually, the fringe fitting programmes, such as the AIPS task FRING, first Fourier
transform the data from time-frequency domain to delay-rate domain. The peak amplitude point in this space provides a starting guess for the actual least-squares solution. The fringe fit can be done for each baseline separately or it can be station-based (so-called global fringe fit), which allows weaker sources to be detected. There are a lot of intricate details in the fringe fitting, and the observer can adjust a large number of parameters in the fringe fitting programmes like FRING. It is beyond the scope of this introductory text to go into these details, and therefore, reader is referred to Cotton (1995) for an in-depth discussion of fringe fitting.

**Other possible calibration steps**

Due to imperfect analog filter shapes at the stations, there are variations in amplitude and phase across the individual baseband channels. These can be removed by using a self-calibration-like technique on strong calibration sources with the AIPS task BPASS. The baseband correction is essential for spectroscopic observations, and it may slightly improve the dynamic range also in continuum observations. For polarimetric VLBI observations, the parallactic angle rotation due to alt-azimuth mount of the antennas needs to be removed (the AIPS procedure VLBAPANG), and cross polarised single-band delays (and phase offsets between the basebands) need to be corrected. This can be done with e.g. the AIPS procedure CRSFRING (Leppänen 1995). The polarisation calibration also includes determining the instrumental polarisation (i.e. so-called leakage or D-terms) and the phase offset between the right and left handed channels (which causes a rotation of the polarisation angle in the sky).

After *a priori* amplitude calibration and removing the residual delays and rates, the data can be averaged in frequency and in time. At this point, we have usually used the AIPS task SPLIT to average the data over the spectral channels and to write it out so that the Caltech DIFMAP programme (Shepherd 1997) can be used in interactive editing, imaging and self-calibration. Averaging in time is done in DIFMAP. Before averaging, the phases are first self-calibrated (see section 2.3.2) to increase the coherence. The averaging time must be kept short enough so that the coherence time is not exceeded and that a time average smearing does not start to limit the undistorted field of view.

### 2.3 Imaging

There are two things which prevent us from just simply applying an inverse Fourier transform to the measured complex visibilities in order to get an image of the source (equation 2.6): a sparsely sampled \((u, v)\) plane and residual amplitude and phase errors still left in the visibility data. Modern VLBI imaging solves these problems by using an iterative method combining *deconvolution*, which applies *a priori* information of the sky brightness distribution to interpolate the visibilities at the points in the \((u, v)\) plane that were not sampled, and *self-calibration*, which tries to correct the antenna gains by
using trial visibilities calculated from a source model. Practice has taught that this iterative imaging technique, which both produces the deconvolved image and does the final calibration simultaneously, can produce a faithful representation of the sky brightness distribution. However, it is vital to understand the dangers that are involved in the imaging process: a wrongly driven deconvolution-self-calibration loop can easily add artefacts in the final image.

2.3.1 Deconvolution

According to equation 2.9, the dirty image is a convolution of the true sky brightness distribution with the dirty beam given in equation 2.10:

$$I^D = B * I. \quad (2.17)$$

Since $B$ is composed of a finite number of Fourier components, there are functions $Z(l, m)$ so that

$$B * Z = 0. \quad (2.18)$$

Now, if $I^S$ is a solution to the convolution equation, so is $I^S + \alpha Z$ for any $\alpha$, meaning that equation 2.17 is not invertible by any linear convolution. Therefore, we always need extra information to constrain the unobserved visibility data. Luckily, quite simple a priori assumptions are sufficient for this. Commonly used assumptions are e.g. the finite source size, positivity of the true sky brightness, smoothness of the sky, and that the sky is assumed to be representable by a small number of point sources. A deconvolution algorithm uses this extra information to solve equation 2.17 for an estimate of the true sky brightness distribution.

The most widely used deconvolution algorithm in VLBI is CLEAN, invented by Jan Högbom in the early 70’s (Högbom 1974) and further developed by Clark (1980) and by Cotton & Schwab (Schwab 1984). CLEAN is based on the assumption that the sky brightness distribution can be represented by a small number of point sources, and a simple iterative approach is devised to find the positions and flux densities of these point sources. The basic CLEAN starts by locating the peak of the dirty image. It then subtracts from the dirty image, at the position of the peak, the dirty beam multiplied by the peak strength and a downscaling factor ($\leq 1$), called loop gain. The position and flux density of the subtracted point source is recorded, and the algorithm is repeated from the beginning. Iteration is continued until some stopping condition is achieved – this can be e.g. the first negative $\delta$-component, number of iterations, or an adequate noise level in the residual map. Finally, the accumulated point source model is convolved with a CLEAN beam (normally an elliptical Gaussian fitted to the central lobe of the dirty beam) to suppress the higher spatial frequencies, which are usually not well extrapolated by the CLEAN algorithm. The residuals from the last round of CLEAN are commonly added to the CLEAN image in order to give an indication of the true noise level. The a priori information of the finite source size can be used in CLEAN by constraining the algorithm to a restricted area in the image plane (so-called CLEAN windows).
Theoretical understanding of CLEAN is relatively poor. Although the convergence conditions are known (Schwarz 1978), nobody has so far managed to do an error analysis of CLEAN. Such an analysis would be important since the CLEAN solutions are not unique: slightly different final images can be produced from the same visibility data just by changing the control parameters of the algorithm. Also, there are instabilities in CLEAN (such as the well-known striping with extended sources) since the convergence conditions may be violated by numerical errors and by gridding, and the algorithm will diverge eventually. However, with most real sources, CLEAN seems to work well – the observer just needs to keep in mind the possible pitfalls.

The next most popular deconvolution algorithm is the Maximum Entropy Method (MEM; see e.g. Cornwell & Evans 1985; Narayan & Nityananda 1986 and references therein). There is no consensus on the justification for MEM. Following Cornwell et al. (1998) we define the entropy as something, which when maximised, produces a positive image and minimises the variance of the pixel values. There are many possible forms of this type of entropy, but the one often used is:

\[ \mathcal{H} = - \sum_k I_k \ln \frac{I_k}{M_k^e}, \]

(2.19)

where \( I_k \) is the intensity of the pixel \( k \) and \( M_k \) is a default image that represents the optimum solution in the absence of data. \( M_k \) can simply be a flat image. A MEM algorithm maximises \( \mathcal{H} \) while keeping the fit, \( \chi^2 \), of the predicted visibility to that observed close to its expected value. MEM is generally better than CLEAN for large and complex sources. On the other hand, there is a known problem with MEM in situations when a point source lies on background level of emission – in these cases MEM fails to fully remove the sidelobes. This can be, however, cured by subtracting the point source prior to deconvolution.

Other non-linear deconvolution algorithms exist beside CLEAN and MEM, but they are currently much less used. Of these algorithms, the non-negative least squares (NNLS) optimisation applied to directly solving the parameters of the point source model is especially interesting (Briggs 1995). With increasing computing power, such an approach has become feasible, and the quality of the images deconvolved with NNLS is excellent.

Finally, it should be mentioned that in VLBI imaging, weighting of the \((u, v)\) data prior the Fourier transforming of the visibilities to the dirty image has a significant impact on the final image. However, we will not cover the different weighting schemes here, but rather refer to the detailed discussion by Briggs (1995). More information about practical VLBI imaging can be found in Walker (1995).

### 2.3.2 Self-calibration

Deconvolution handles (or at least tries to handle) the \((u, v)\) plane sampling problem in VLBI, but there is still another factor limiting the image quality – namely the residual
amplitude and phase errors of the measured visibilities. Phase errors are especially severe, and in the early days of VLBI, they were thought to prevent any imaging of VLBI data. The calibration method normally used in the connected element arrays – i.e. using an unresolved calibrator of a known flux density to calibrate the visibilities of the programme source – is not possible in VLBI since there are only few sources of sufficient strength that are unresolved in milliarcsecond scales. Therefore, the calibration sources would be generally far from the programme source causing the atmospheric phase shift and elevation dependent antenna gain to differ. Also, the station clocks in VLBI are not running synchronously.

Eventually, it was realised that even uncalibrated phases imposed strong constraints on the source structure if the antenna dependence of any phase offset was maintained. Here, a rather old discovery was used: Roger Jennison found out already in the 1950s that if the phases of three baselines $\theta_{ij}$ are summed up, all the antenna-based phase errors are cancelled (Jennison 1958):

$$\Phi_{123} = \theta_{12}^M + \theta_{23}^M + \theta_{31}^M = \theta_{12} + (\phi_1 - \phi_2) + \theta_{23} + (\phi_2 - \phi_3) + \theta_{31} + (\phi_3 - \phi_1)$$

$$= \theta_{12} + \theta_{23} + \theta_{31},$$

(2.20)

where $\theta_{ij}$ is the phase due to the source on the baseline $i-j$ and $\phi_i$ is the phase error of the station $i$. This closure phase is a good observable, and it allows VLBI imaging in the presence of phase errors.

There is also a closing quantity for the visibility amplitude:

$$\Gamma_{ijkl} = \frac{g_i g_j A_{ij} \cdot g_k g_l A_{kl}}{g_i g_k A_{ik} \cdot g_j g_l A_{jl}} = \frac{A_{ij} A_{kl}}{A_{ik} A_{jl}},$$

(2.21)

where $A_{ij}$ is the visibility amplitude due to the source on the baseline $i-j$ and $g_i$ is the gain error of the antenna $i$.

Unfortunately, closure quantities cannot be used to produce images directly. Instead, they have been devised in imaging mainly in two ways: hybrid mapping or self-calibration, the latter being mostly used nowadays. Hybrid mapping (Readhead & Wilkinson 1978) starts by choosing an initial model and transforming it to trial visibilities. Observed visibility amplitudes are kept, and new visibility phases are selected by modifying the model visibility phases to be consistent with the observed closure phases. The new visibilities are then transformed to a dirty image, which is deconvolved and used as a starting model in the next iteration round. Problem with this algorithm is that the resulting error distribution of the phases is non-Gaussian, which makes subsequent analysis difficult. Also, antenna based a priori information cannot be easily incorporated in hybrid mapping.

In self-calibration, the complex gains of the stations are altered (see e.g. Cornwell & Fomalont 1998). This clearly conserves the closure quantities, imposing constraints
to the method. In a self-calibration algorithm, a model image is first chosen and transformed to model visibilities $V_{ij}^{\text{model}}$. Corrections to element gains $G_i$ at time $t_k$ are solved by minimising

$$M = \sum_k \sum_{i,j \neq j} w_{ik}(t_k) \left| V_{ij}(t_k) - G_i(t_k)G_j^*(t_k)V_{ij}^{\text{model}} \right|^2,$$  

(2.22)

where $w_{ik}(t_k)$ is the weight depending on the variance of the observed visibility $V_{ij}(t_k)$. A new dirty image is produced from the visibilities that are corrected with the new antenna gains solved from equation 2.22. This image is then deconvolved and used as a new model image in the next round of iteration.

Convergence of the deconvolution-self-calibration loop depends on how the algorithm is driven. It is important that amplitudes are not self-calibrated before the model is already quite good. Otherwise, artefacts due to an erroneous model can get frozen in the gains and the convergence is stopped. Such a situation easily occurs if there is a lot of emission missing from the model at short $(u, v)$ spacings due to a shallow CLEAN. Also, it is advisable to start amplitude self-calibration with a long solution interval, which is shortened as the model improves. The phases are naturally solved over time intervals that are short compared to the coherence time from the beginning. As was noted in section 2.2.1, baseline based errors do not close – violating the basic assumption behind self-calibration. Therefore, they can be devastating to the VLBI imaging. Luckily, such errors are usually small – especially with the VLBA data.

### 2.3.3 Model fitting

There are several situations where it is useful to inspect the VLBI data directly in the visibility domain. Such situations include e.g. interpretation of very sparsely sampled or poorly calibrated data, which cannot be reliably imaged. In these cases, direct examination of the visibilities can still give some information of the source structure. Errors are also easier to locate in the visibility data than in the image because Fourier transform spreads a localised error in the $(u, v)$ plane over the whole image.

Model fitting is usually better for a quantitative analysis of the source properties. This is due to several reasons. The final images from VLBI observations contain a wealth of information, which, however, can be difficult to parameterise. Representing the source by a small number of e.g. Gaussian components is often advantageous in order to get a convenient number of parameters for the quantitative analysis. Also, it may be difficult to determine from the images whether apparent changes are due to real changes in the source structure, or just to differences in the $(u, v)$ sampling and imaging parameters.

One can, of course, fit the image plane with functions (e.g. Gaussians), but there are problems with this approach. Firstly, possible deconvolution errors enter the image plane model fitting, and secondly, the noise properties of the image after non-linear
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Figure 2.2: An example of using closure phases in comparing two models for the core of PKS 0420-014. The figure shows closure phases from triangles consisting of long (left panels) and short (right panels) baselines from a series of observations. There are two core-region models presented: the dashed line represents a single, potentially extended component, and the solid line represents two point-like components. The fourth epoch (1997 July 31) is for comparison, no component close to the core is expected at this epoch – and none is seen.

deconvolution, like CLEAN, are poorly understood. On the contrary, if the data are the real and imaginary part of the visibilities, the errors are Gaussian. Another advantage of the \((u, v)\) plane fitting over the image plane fitting is that former is also a deconvolution process, which allows component sizes and separations significantly smaller than the beam to be measured if the signal-to-noise ratio is high enough.

Model fitting works best with a rather simple source, which can be represented by a collection of well-separated, discrete components. Complex sources showing a lot of extended emission are problematic for model fitting, but it is notable that even in these
cases the images produced from a small number of model components convolved with the restoring beam (with residuals added) are remarkably similar to the \textit{CLEAN} images from the same data (Lobanov & Zensus 1999; Homan et al. 2001; Kadler et al. 2004; Paper IV).

Model fitting can be also used in checking and adjusting the amplitude calibration since, although there are not many unresolved calibrator sources in VLBI, in many cases the calibrators can be represented rather accurately by a simple one- or two-component model. Also, model fitting can provide a starting model that is better than a point source for the deconvolution-self-calibration loop.

Model fitting involves designing the model, choosing a figure-of-merit function, and then adjusting the model parameters to minimise this function. In practice, the currently used model fitting programmes are based on non-linear least squares algorithms minimising $\chi^2$ function. This implicitly assumes that the model is actually a good fit to the data, the errors are uncorrelated, known and Gaussian, and there are no systematic calibration errors. The least squares algorithms can have problems with finding the global minimum and in some cases with the slow convergence. However, perhaps the most problematic aspect is that the model fitting does not have a unique solution. Therefore it is important that one does not introduce more parameters in the model than are clearly needed by the data. A \textit{CLEAN} image of the source, if available, and closure phases can provide additional constraints on choosing the model.

One can sometimes do simple hypothesis testing with the model fitting by inspecting the closure phases. An example is given in Fig. 2.3.3 where two models for the core of PKS 0420-014 are compared. It is clear that the model composed of two point-like components gives a much better fit to the closure phases from long baselines than a single, perhaps extended, component. As is expected, both models fit short baselines equally well. From this, it can be concluded that there are two components in the core of PKS 0420-014 (see Paper I). The power of this method is that the closure phases are not affected by any calibration problems (unless there are severe non-closing errors).

### 2.3.4 Error analysis in model fitting

Knowing the uncertainties in the model parameters is important for any subsequent analysis like determining the proper motions or measuring the component flux density variability. Unfortunately, estimating the real ($1\sigma$) measurement errors for the parameters of the model fitted to VLBI data is a notoriously difficult task. Errors taken from the diagonal of the $\chi^2$ covariance matrix are almost guaranteed to be overoptimistic since they assume the perfectly calibrated visibility data and uncorrelated model parameters. Neither of these assumptions is generally true: it may be possible to adjust the antenna gains to support a slightly different model, and it is common that the parameters of the adjacent components depend on each other. Still another complicating factor is that after

\[7\text{All the model fitting presented in Papers I-V is done with the MODELFIT task in DIFMAP. This programme directly fits the real and imaginary part of the visibilities using the Levenberg-Marquardt method.} \]
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self-calibration is applied, the data are not independent anymore, and therefore, the true number of degrees of freedom in the data is difficult to calculate.

The first two problems – calibration errors and interdependency of the model parameters – can be solved by a method first described by Tzioumis et al. (1989), where the value of the model parameter under scrutiny is adjusted by a small amount and fixed. The other parameters of the model are then solved by model fitting (which can also involve self-calibration of the element gains) and the results are inspected. The cycle is repeated until a clear discrepancy between the data and the model is achieved. DIFWRAP is a programme that can do this iteration easily (Lovell 2000). There is, however, a problem also with this approach. Since the number of degrees of freedom is unknown, there is no statistical limit for determining when the particular model does not fit the data anymore, and one must rely on highly subjective visual inspection of the data. In Paper IV, we discuss the error determination with DIFWRAP at length, and the positional errors derived from DIFWRAP analysis are compared with errors estimated from the variance about the best-fit proper motion model. The errors are comparable, and they agree well with the often quoted “rule-of-thumb”, $1/5$ of the beam size. This gives some confidence in the positional errors determined by DIFWRAP despite the problematic visual inspection.

However, a more robust model fitting method, which provides an accurate estimate of the true ($1\sigma$) measurement errors, would be highly desirable since it would allow firm statistical methods to be used in analysing the source properties. One possible approach in the future could be Bayesian analysis (see e.g. D’Agostini 2003). Bayesian methods are good for choosing the model that best describes the data with the given a priori information. Especially promising for VLBI data is the ability of Bayesian methods to determine and marginalise over a number of unknown parameters like the atmospheric or instrumental phase fluctuations. The first method for producing radio interferometry images that, given a certain prior, are optimal in the Bayesian sense of maximum posterior probability density, has been recently published by Sutton & Wandelt (2006).

2.4 Multifrequency VLBI

Studying the radiation from celestial sources as a function of frequency – were it a detailed scrutiny of the spectral lines or a broadband study of the overall spectral energy distribution – has been for over a century the most powerful tool in our quest to understand astronomical objects. It is only natural to employ this tool also on the study of compact extragalactic jets with the VLBI. For instance, the broadband radio spectra of emission features in a parsec scale jet are important observables; in the framework of incoherent, random pitch-angle synchrotron radiation model (e.g. Marscher 1987), the frequency and flux density of the spectral turnover together with the emission region’s angular size provide means to derive the physical conditions in the jet – such as magnetic field density and energy distribution of the radiating electron population. In Paper V, we have derived these quantities for the archetypical quasar 3C 273.
Before the commissioning of the VLBA, multifrequency VLBI studies of compact jets were difficult and rare (see, however, pioneering studies by Cotton et al. 1980, Bartel et al. 1984, Marscher & Broderick 1985, and Marscher 1988), since it was not usually possible to observe the source simultaneously at several frequencies (apart from the dual-frequency S/X band system for geodetic VLBI) and the variability of the compact jets made the comparison of observations taken at different epochs difficult. With the frequency agility of the VLBA, it is now possible to make practically simultaneous measurements at several different frequencies, and this has made the studies of the continuum spectrum in the parsec scale jets feasible. The VLBA also provides array homogeneity and accurate a priori flux density calibration procedure, which further help the calibration of the spectrum. However, multifrequency VLBI studies are still time-consuming, and extraction of the broad band spectrum from a multifrequency VLBI data set is a difficult task with a number of problems like uneven \((u, v)\) plane coverage at different frequencies, flux density calibration, and image alignment. Despite the difficulties, we believe that this task is worth doing, and moreover we believe that the multifrequency (polarimetric) VLBI monitoring will be the way of the future in studying the AGN jets.

### 2.4.1 Spectral extraction by model transfer

There are basically two possible – and somewhat complementary – approaches to extracting spectral information from multifrequency VLBI maps. The first is to form spectral index images or, if data at more than two frequencies are available, maps of synchrotron turnover frequency (Lobanov 1998). The second approach is to use model-fit components, which provide a convenient number of spectra for a detailed scrutiny.

An advantage of using the model components in spectral extraction is that by applying a priori information of the source structure from the higher frequency data, it is possible to fit the low frequency data with components having sizes and mutual separations significantly smaller than the beam size at the given frequency. Hence, we are able to measure the spectra for much smaller features than is possible by spectral index mapping\(^8\). The highest spatial frequency reliably reached by this extrapolation naturally depends on the signal-to-noise ratio of the data and on the complexity of the source structure. In Paper V, where we apply this so-called model transfer method to the parsec scale jet of 3C 273, we limit the extrapolation to spatial frequencies that are below the \(uv\)-radius corresponding to \(\sim 1/5\) of the beamsize at the given frequency. The procedure used in the case of 3C 273 can be summarised as follows:

1) First, a simple model consisting of two-dimensional Gaussian components is formed at 43 GHz, which is the frequency that provides the best angular resolution together with a good signal-to-noise ratio.

\(^8\)The problem with spectral index imaging is that CLEAN gives poor results when extrapolating to the high spatial frequencies, and hence super-resolved images cannot be used in constructing spectral index maps.
2) The 43 GHz model is transferred to other frequencies. The relative positions of the components are kept fixed, and the models are aligned by assuming that optically thin components have frequency independent positions, i.e. that their brightness centroids are co-spatial at all frequencies.

3) Model-fitting is run at each frequency allowing only component sizes and flux densities to vary. If any two components have a separation smaller than \( \sim 1/5 \) of the beam size at the given frequency, these components are replaced by a single component. This is to avoid extrapolating to spatial frequencies that are not well constrained by the data (see section 2.3.4).

4) The sizes of the components are plotted against frequency and fitted with either a constant value or with a power-law after removing clear outliers. Here we try to remove any bad size measurement and to estimate possible frequency dependence of the component sizes. In fact, frequency dependent sizes are found for four components in the jet of 3C 273, and such a frequency stratification is likely a sign of radiative losses producing a gradient in the electron energy distribution across the component (Paper V).

5) The angular sizes of the components are fixed to values derived from the power-law fit (or to a constant value) and the model-fitting is run again with component flux density as the only variable. This gives the final spectra.

As can be seen from Fig. 11 in Paper V, the method produces good quality spectra and, for some components, it allows also measurements in the self-absorbed part of the spectrum. Using \textit{a priori} information of the relative component positions and assuming smoothly varying component sizes reduces the number of free parameters in the final model and makes the out-of-band extrapolation of spatial frequencies possible. However, there are also problems with this approach; the transferred source model is not unique (see section 2.3.3), the results depend a lot on the correct alignment of the model between the frequencies, and true \( 1\sigma \) uncertainties in the final spectra are hard to determine.

### 2.4.2 Image alignment

Unless the observations are carried out using the phase-referencing technique, there will be no absolute source position information in the images after self-calibration is applied. Therefore, one cannot use co-ordinates to register the images at different frequencies. One way to align the images is to use the position of the compact core. This, however, is likely a poor choice since at normal VLBI frequencies the core is roughly the point in the flow where the jet becomes optically thin, and hence, its location depends on the observing frequency.

A better option for aligning the images without phase-referencing is to use bright components in the optically thin part of the jet. If one can assume that there are no significant variations of the spectral index across a component, it can be used to register the images. For example, in Paper V we simply use the brightness centroids of two bright and well-defined components to align the models at different frequencies. A more accu-
rate alignment can be obtained by cross-correlating a portion of a pair of images (which have been convolved to the same resolution), and finding a set of position offsets that maximise the correlation for all the pairs; an accuracy of 5µas is achieved by Walker et al. (2000) for 3C 84. Yet another way to align the images is to find relative positions that minimise the spectral index gradients, but this approach is tedious and not as accurate as cross-correlation.

2.4.3 Flux density calibration

The absolute flux density calibration at different frequencies is naturally important for the quality of the final spectra. The \textit{a priori} amplitude calibration of the VLBA is known to be better than 10% at frequencies below 15 GHz. At 22, 43, and especially at 86 GHz, the high and variable atmospheric opacity and pointing errors (mainly at 86 GHz) start to cause problems, and it is advisable to observe a calibrator source to check and, if needed, to correct the flux density scale. The calibrator, which must be a compact source with little or no extended emission, is observed with a single-dish telescope or with a connected-element array like VLA. If the source is variable (as they usually are), it must be observed either simultaneously with the VLBI session or monitored frequently enough to allow reliable interpolation. The measured total flux density is then compared with the extrapolated zero baseline flux density from the VLBI data. Any known large-scale emission that will be resolved out in the VLBI scale must be subtracted from the total flux density before comparison.

We find in Paper V that the \textit{a priori} amplitude calibration system of the VLBA is actually more accurate than the nominal value of 10%: the flux density of the calibrator source 3C 279 in Paper V is correct to $\sim$ 5% at all frequencies from 5 to 43 GHz. However, there is a large discrepancy between the single-dish flux density observed by the Swedish-ESO Submillimetre Telescope (SEST) and the VLBA flux density at 86 GHz, the VLBA flux density being only $\sim$ 60% of the SEST measurement. This is likely due to a problem in amplitude calibration of the VLBA data (like poor opacity correction) – or there have been pointing problems during the observation.

2.4.4 Uneven sampling of the \((u, v)\) plane

As was emphasised in section 2.1.2, an interferometer is a spatial filter, and the size scales it probes in the sky are determined by the range of baseline lengths and by the observation frequency. For VLBI networks, which cannot be scaled like some reconfigurable arrays (such as VLA), this means that the observations at different frequencies show different parts of the source. Increasing the observing frequency is like zooming in the source. There are both positive and negative sides to it: it is good that we can increase the angular resolution of our observations by increasing our observing frequency without having to increase the baseline length. On the other hand, the mismatch between \((u, v)\) coverages at different frequencies poses a potentially serious problem for
the extraction of the spectral information from multifrequency VLBI observations.

Imaging the source at several different angular scales by multifrequency VLBI is sometimes very useful, and it can yield surprises as was the case with a distant quasar PKS 2136+141 (Paper II). This quasar is a high frequency peaker candidate (Tornikoski et al. 2001), and in 2001 we observed it with the VLBA at six frequencies in order to study the origin of the inverted spectrum. The source turned out to have a core-jet morphology, and it also held a surprise: when PKS 2136+141 was seen at different frequencies, its jet direction changed significantly. The jet position angle was rotated by \(210^\circ\) in the plane of the sky between 2.3 and 43 GHz (see Fig. 2 in Paper II). This is, to our knowledge, the largest ever observed change in the jet position angle. Hence, in the case of PKS 2136+141, the multifrequency VLBI observations revealed something very important about the jet structure – something single frequency studies could not have shown.

The uneven spatial samplings of the \((u, v)\) plane at different frequencies result in changes in the corresponding interferometer beams, and can (and in practice will) lead to severe artifacts appearing in spectral index images if convolving beams smaller than that of the lower frequency synthesised beam are used. However, simply convolving images at different frequencies with a common beam corresponding to the interferometer beam size at the lower frequency does not guarantee that errors in the spectral images will not be made (Swain 1999). The normal practice in making spectral index maps is either to use \((u, v)\) plane tapering or to flag the data in order to produce matching \((u, v)\) ranges at both frequencies. Lobanov (1998) has studied the errors in the spectral index maps caused by uneven \((u, v)\) plane sampling. His simulated data suggest that the pixel SNR is the main factor determining the errors and, for sufficiently high SNR data \((\gtrsim 7)\), the fractional errors do not exceed \(\sim 10\%\). The SNR here is measured with respect to the self-calibration noise, which is taken to be equal to the largest negative pixel in the map.

In Paper V, we have estimated the effect of the uneven \((u, v)\) plane coverage for the model transfer method by using simulated data. The average fractional error in the component flux density due to the differences in \((u, v)\) plane sampling, model fitting algorithm, and added noise is \(\approx 13\%\). The component angular size seems to be more uncertain quantity with an average fractional error larger than 20\%. The size of the simulated data set used in Paper V was rather small, and therefore, more extensive simulations are required to accurately characterise the errors resulting from uneven \((u, v)\) plane sampling. The above figures are, however, indicative.
CHAPTER 3

Compact jets in AGN

“The Universe is full of magical things, patiently waiting for our wits to grow sharper.”

– Eden Phillpotts

In this and the next chapter, some central theoretical and observational results regarding the compact extragalactic jets will be reviewed, and a short introduction to AGNs in general is given. This is by no means a comprehensive review – such would be beyond the scope of this thesis – but rather it is intended to provide the necessary background for understanding the results presented in the Papers I–V.

3.1 A short introduction to active galactic nuclei

In chapter 1, we defined AGNs as galaxies with unusually high radiative output from their nuclei. However, apart from this common property, the AGN class is like a zoo – there is a colourful bunch of subclasses, or species, having vastly differing properties. In the following, we will look at a few most prominent categories and briefly describe their observational properties.

3.1.1 Seyfert Galaxies

In 1940’s Carl Seyfert discovered a class of (mostly spiral) galaxies having a very bright star-like nucleus and prominent emission lines, which were either unusually broad or arised from atypically highly ionised elements (Seyfert 1943). The Seyfert galaxies – as they are now called – were the first AGN class discovered. Seyfert’s original classification was later modified by Khachikian & Weedman (1974) who divided Seyferts into two subclasses known as Seyfert 1 and Seyfert 2 galaxies. In Seyfert 1 galaxies the permitted emission lines are broad (1000 – 10 000 km s⁻¹) and forbidden lines are narrow (a few hundred km s⁻¹), while Seyfert 2 galaxies show only narrow lines. A simple interpretation of these line properties is that in Seyfert 2s both the permitted and forbidden lines originate from the same region, whilst in Seyfert 1s the permitted and forbidden lines are formed in different regions. The broad line region (BLR) is believed to lie close to the central engine (< 1 pc) so that the high velocities inferred from the linewidths can be explained. The narrow line region (NLR), on the other hand, is thought to be located ~ 10 – 1000 pc from the central engine.
There are no broad emission lines in Seyfert 2s, but is this due to the absence of the BLR – or is the BLR hidden in Seyfert 2s? This is one of the key points in the AGN unification scheme, and we will return to it shortly. Since the work of Khachikian & Weedman, the classification of Seyferts has been further modified. Now there exist such additional categories as Seyferts 1.5, 1.8 and 1.9 (Osterbrock 1981). This shows how the improvements in the observational capabilities (spectroscopy in this case) tend to make the classifications more complex as the sources can be studied in more detail.

Seyfert galaxies make up a few percent of all bright galaxies, and there seem to be at least a few times as many Seyfert 2s as Seyfert 1s. Only few Seyferts are strong radio sources.

3.1.2 Radio Galaxies

Extragalactic radio sources can be divided into core-dominated radio sources in which radio emission from a compact structure in the host galaxy dominates over the extended structure, and to lobe-dominated radio sources which show powerful radio emission from double lobes lying symmetrically around the host galaxy. Core-dominated sources will be discussed later in connection to the blazar phenomenon. The lobe-dominated sources are associated with giant elliptical galaxies, but their radio structure extends well beyond the confines of a galaxy. The largest double radio sources have sizes of hundreds of kiloparsecs or even a few megaparsecs, which makes them the largest single entities in the Universe.

The radio galaxies radiate synchrotron emission implying that the lobes contain a vast amount of energy in form of relativistic electrons and magnetic field. In order to explain the double-lobed morphology and the fact that radio lobes had not yet radiated away all their energy although they must be at least several million years old, Scheuer (1974) and Blandford & Rees (1974) suggested that the lobes are continuously fed by jets ejected from the central galaxy. This view was later confirmed by high-sensitivity, high-resolution interferometric imaging. For example, in NGC 6251 the jet can be followed from the parsec scales (VLBI images) all the way to kiloparsec scale (VLA) and to the radio lobes (WSRT; see figure in Bridle & Perley 1984). This is important, since in the parsec scales, high-resolution VLBI imaging reveals motions indicating that there indeed is outflow. Generally a compact core-jet component is seen in the middle of the double radio sources in VLBI images.

Fanaroff & Riley (1974) noticed that there is a distinct change in the properties of lobe-dominated radio sources at a flux level of $5 \times 10^{25}$ W at 178 MHz. The lower luminosity sources, called Fanaroff-Riley Class I (FRI), show edge-darkened, steep-spectrum radio lobes that are connected to the central galaxy by continuous, smooth, double-sided jets. The jets in FRIs are often distorted with bends, and their luminosity decreases as a function of distance from the central galaxy. Fanaroff-Riley Class II (FRII) sources have larger luminosity, their lobes are often edge-brightened with distinctive “hot spots” at their outermost regions, and their jets are usually one-sided and more linear than in
FRIs. The FRII jets also show a knotty structure instead of a smooth jet. It should be noted that although the jets in FRII sources often look less striking than in FRIs (due to

Figure 3.1: Top: An example of FRI radio source, 3C 31 observed with the VLA. Bottom: An example of FRII radio source, Cygnus A. The high-resolution, high-dynamic range image is made with the VLA. (Images courtesy of NRAO/AUI)
the higher luminosities of FRII lobes), they are in fact usually more luminous. The jet speed in FRII sources is likely to be higher than in FRIs. See Fig. 3.1 for examples of FRI and FRII sources.

Besides the classes described above, there are extended radio sources that show distorted and complex morphologies such as narrow-angle tailed (NAT) and wide-angle tailed (WAT) sources. These sources can actually tell us a lot of the properties of radio emitting gas, since the observed distortions arise from interaction of the jet with intracluster medium (at least in the case of NATs) whose characteristics can be inferred from X-ray observations.

### 3.1.3 Quasars

Quasars are the true giants of the AGN zoo: they show luminosities up to overwhelming $10^{13}\, L_\odot$, often together with very broad emission lines, and they have been observed from $z \sim 0.1$ up to $z \sim 6$ (Fan et al. 2003). These enigmatic sources were first discovered in 1963. Lunar occultation observations made by radio astronomer Cyril Hazard had yielded an accurate position for a strong radio source 3C 273 in late 1962. The radio source appeared to coincide with a 13th magnitude “star”. The identification was communicated to Maarten Schmidt at Mount Palomar Observatory and he obtained a spectrum of 3C 273. To the amazement of Schmidt, the spectrum showed strong and broad Balmer lines, which were redshifted by 16% (Schmidt 1963; Greenstein & Schmidt 1964). The source was extragalactic and the Hubble law yielded a distance of almost one gigaparsec meaning that 3C 273, a 13th magnitude source, was by far the most luminous object observed by that time.

After the first quasars were discovered, it was soon found out that these extremely luminous sources are also variable, and moreover, variable at timescales less than a year, implying the size of the emitting region smaller than a light-year. It was clear that ordinary stellar processes could not produce such large power from such a small volume for a time long enough to explain the existence of quasars. Theoretical search for a suitable power source soon pointed towards supermassive black holes (SMBHs) as the only viable engine to power the quasars.

Many quasars are spectroscopically similar to the most luminous Seyfert 1 galaxies as emission line widths up to 10000 km s$^{-1}$ are observed in both classes. Indeed, nowadays quasars and Seyferts are considered as members of the same class but with differing luminosities. The dividing line is set at $M_V = -23.5$, but this is a rather arbitrary choice.

About 10% of all quasars are radio-loud while the rest are radio weak compared to their optical luminosity. The radio emission is associated with a jet of relativistic plasma radiating synchrotron emission with a power-law spectrum from radio to optical. The radio-loud quasars can be further divided according to their spectra (steep spectrum lobe-dominated or flat-spectrum core-dominated) or their optical polarisation (high polarisation quasars with polarisation $> 3\%$ or low polarisation quasars with polarisation
3.1. A SHORT INTRODUCTION TO ACTIVE GALACTIC NUCLEI

<3%; see Moore & Stockman 1984; Valtaoja 1996). The radio structure is often similar to Fanaroff-Riley II radio galaxies, albeit with much larger brightness ratio between the compact jet and the lobes. The radio-loud quasars reside almost exclusively in bright elliptical galaxies (Dunlop et al. 2003; Floyd et al. 2004 and references therein) while for the radio-quiet quasars, high resolution HST imaging has shown that host galaxies of nearly all types exist, although the most luminous radio-quiet quasars tend to almost always reside in ellipticals (Bahcall et al. 1997; Dunlop et al. 2003). The hosts in both cases are often distorted.

The quasar luminosity function shows a significant cosmological evolution with a strong peak at $z \sim 2$, and there seems to be a connection to the evolution of galaxies. Since quasars have been observed at redshifts exceeding 6, it is possible that they have also played a role in re-ionising the early Universe.

3.1.4 BL Lac objects, FSRQs and blazar phenomenon

BL Lac objects are yet another AGN class, named after the prototypical source BL Lacertae. Classical BL Lac objects show compact, core-dominated radio emission, high polarisation and strong and rapid variability at all frequencies. The characteristic property is the absence of emission lines\(^1\). Optical and X-ray variability is sometimes detected in timescales of hours (Wagner & Witzel 1995). The luminosity range of BL Lac objects is large, although it is difficult to estimate the intrinsic luminosity because the radiation from BL Lac objects is strongly beamed (see section 3.2.2). In general, BL Lac objects have low redshifts ($< 0.2$), but there are also some sources at larger distances.

BL Lac objects have been mainly found through radio and X-ray observations, and it was initially thought that radio selected – or classical – BL Lac objects (RBL) and X-ray selected BL Lac objects (XBL) are two distinct classes, since RBLs seemed to be more luminous in radio and optical, more variable and more core-dominated than XBLs (e.g. Perlman & Stocke 1993; Jannuzzi et al. 1994). Later, it was realised that the explaining factor is the synchrotron cut-off frequency, and BL Lac objects were divided to low-frequency peaking (LBL) and high-frequency peaking (HBL) sources (Padovani & Giommi 1995). Also intermediate sources (IBL) have been found, and recently, Nieppola et al. (2006) showed that the BL Lac population is continuous in the frequency of the synchrotron peak of the spectral energy distribution (SED) and there is no bimodality. They also showed that there is no correlation between the peak frequency and the luminosity of the source for BL Lac objects, contrary to a so-called “blazar sequence” hypothesis (Fossati et al. 1998).

Some BL Lac objects have strikingly similar radiation properties with the flat-spectrum radio quasars (FSRQs). They emit radiation over the whole electromagnetic spectrum from radio to $\gamma$-rays, they are variable at all frequencies, and they show compact radio structure in the VLBI images with clear evidence of superluminal motion in most cases (see section 3.2.2). All together, these objects are commonly referred to as

\(^1\)It should be noted that with modern instruments, weak emission lines have been found later in many BL Lac objects – including BL Lacertae itself.
“blazars” (a term originally proposed by Ed Spiegel in his after-dinner speech at a BL Lac conference in 1978). Blazars, however, are not a physically meaningful class of objects, since FSRQs and BL Lacs have different parent populations, different jet velocities, average luminosities etc. (Valtaoja 1996). The term rather refers to a phenomenon, namely that of a relativistic jet of plasma pointing nearly towards us. The flux density of the emission from the jet is amplified greatly due to relativistic beaming, and it dominates over the other emission components giving birth to the distinctive radiation properties of blazars, which are discussed in more detail in chapter 4.

Blazars are naturally interesting sources on their own right because they show violent variability at energies up to TeV $\gamma$-rays, but also because they allow studies of the jet in the vicinity of the central engine. Their radiation is strongly dominated by the non-thermal emission from the innermost portions of the relativistic jet and that is why they are good targets for the studies of jet formation.

3.1.5 Basic model and unification

The above-described classification scheme of AGNs is already confusing, and it does not even discuss the whole plethora of subclasses within each category. It is thus not surprising that much effort has been devoted to unifying different AGN classes on a basis of simple physical ideas. And indeed, some insight has been gained.

In the basic model of an AGN, we have a supermassive black hole in the centre of a galaxy, and it is a gravitational accretion process on the black hole that powers the AGN. The gravitational field of the black hole attracts gas clouds in the central parsec. The clouds will go through a number of collisions, lose kinetic energy, and eventually take up a form of a disk, where matter is flowing inwards and angular momentum is transferred outwards by viscous or magnetic drag. The viscous heating makes the innermost part of the disk an extremely hot source of thermal emission. This is the classical picture of an accretion disk (Shakura & Sunyaev 1973). Actually, besides the standard thin disk, there are a number of other models for the accretion flow in AGNs such as slim disk, ion supported or radiation supported thick disks, advection-dominated accretion flows etc. (see e.g. chapter 7 in Krolik 1999), but discussion of these models is beyond the scope of this text. There must also be some sort of a hot corona of energetic electrons which scatter photons from accretion disk up to hard X-ray energies.

The basic concept in the unification models is the orientation of the AGN with respect to the observer (Antonucci 1993; Urry & Padovani 1995). As was mentioned in the context of Seyfert galaxies, around the supermassive black hole and the accretion disk there are a large number of highly ionised gas clouds that have velocities up to 10 000 km s$^{-1}$. This broad line region extends up to 1 parsec from the central engine. In the standard unified model the difference between Seyfert 1 and Seyfert 2 galaxies is explained to be due to an optically and geometrically thick molecular torus which surrounds the central engine and blocks the view to the BLR in Seyfert 2s while in Seyfert 1s the torus is more face-on towards us and the BLR is seen. The narrow-line region is
further away from the central engine than the BLR, and therefore, torus cannot hide it. The main evidence for this unification comes from the detection of polarised (Antonucci & Miller 1985) or infra-red (Reunanen et al. 2003) broad lines in Seyfert 2s, large scale [OIII] emission cones (Mulchaey et al. 1996), and from direct interferometric observations of a warm dust structure in mid-infrared (Jaffe et al. 2004) and in radio (Klöckner et al. 2003). Also, similar to Seyfert 2s, a population of Type 2 quasars that do not show a BLR has been found. The same unification scheme containing the obscuring torus and an orientation effect applies to them as well.

The abovementioned unification does not address the second major division of AGNs, namely the radio-loud / radio-quiet dichotomy. It now seems clear that the difference between radio-loud and radio-quiet AGNs is that the former have a jet of synchrotron emitting relativistic plasma while the latter do not. But what is the fundamental parameter of the AGN that determines whether there will be a powerful jet or not? This is currently an open issue, and possible candidates are for example mass and spin of the black hole, the accretion rate and the properties of the host galaxy.

For radio-loud AGNs, another orientation-based unification scheme has been proposed. The idea is that FRI radio galaxies and BL Lacs on the one hand, and FRIIs and quasars on the other hand, belong to the same class of objects, and it is the angle between the jet and our line-of-sight which determines whether we will see for example a classic FRII double radio source or a quasar (Antonucci & Ulvestad 1985; Barthel 1989). The idea is that relativistic effects, which are dependent on the viewing angle, will greatly enhance the observed flux density, shorten the timescales, blueshift the spectrum, and increase the apparent speed of the jets that point towards us. At large angles the observer sees an ordinary radio galaxy, either FRI or FRII. Then, depending on the type of the radio galaxy, as the angle is decreased ($\lesssim 45^\circ$), the observer sees a quasar (FRII) or a BL Lac object (FRI). When the jet points almost directly towards the observer, the relativistically beamed emission from the jet dominates all other sources of radiation and a blazar is seen.

3.2 Compact jets

3.2.1 Radiative processes

There are two, rather exotic, radiative processes that are important in extragalactic jets: synchrotron radiation and inverse Compton scattering.

**Synchrotron radiation**

Synchrotron radiation is a non-thermal emission mechanism, which was first discovered in man-made particle accelerators – hence its name. Charged particles accelerated in a magnetic field emit electromagnetic radiation. For nonrelativistic particles, the emission is cyclotron radiation at the particle gyrofrequency. In the case of a highly relativistic
charged particle circling around in a magnetic field, the emission will become beamed in the direction of motion (aberration) and an observer will see radiation from the particle only when its emission cone sweeps through his line-of-sight. The resulting spectrum is much more complex than in the nonrelativistic case and it can extend to many times the gyrofrequency. For a detailed discussion and derivation of the relevant formulas for synchrotron radiation, see e.g. Pacholczyk (1970), and chapter 6 in Rybicki & Lightman (1979).

The synchrotron radiation from an ensemble of relativistic electrons with a power-law energy distribution has a distinct shape. At optically thin frequencies, the emitted flux density is

$$S_\nu \propto (B \sin \eta)^{1-\alpha} \left( \frac{\nu}{2c_e} \right)^\alpha,$$

(3.1)

where $\nu$ is emission frequency, $B$ is magnetic field density, $\eta$ is pitch angle between the magnetic field and line-of-sight, $c_e = 3e/4\pi m_e$, with $e$ and $m_e$ being the charge and mass of the electron, respectively. The optically thin spectral index is $\alpha = -(s - 1)/2$, where $s$ is the power-law index of the electron energy distribution; $N(E) = N_0 E^{-s}$. The resulting spectrum has a power-law shape – except at the low and high frequency ends of the spectrum. A synchrotron source will become self-absorbed at low frequencies and below the turnover frequency $\nu_m$, the flux density is

$$S_\nu \propto \nu^{5/2},$$

(3.2)

in a homogeneous source. On the other hand, radiative losses cause the electron energy distribution to steepen at high frequencies, yielding a corresponding steepening also in the emitted synchrotron spectrum above the break frequency, $\nu_b$. Finally, there is a high-energy cut-off in the spectrum defined by the highest energy electrons in the source. As discussed in Paper V, it is possible to derive the strength of the magnetic field, as well as the normalisation factor of the electron energy distribution from observations of the spectral turnover and size of a synchrotron source. An observation of the break frequency $\nu_b$ can tell about the age of the emitting electrons; the highest energy electrons radiate at the highest frequencies and, thus, lose their energy fastest.

Another characteristic property of synchrotron radiation, besides its spectrum, is the high degree of linear polarisation. For particles of a power-law energy distribution in a uniform magnetic field, the degree of linear polarisation is

$$\Pi = \frac{s + 1}{s + \frac{7}{3}}.$$

(3.3)

at optically thin part of the spectrum, resulting in $\Pi > 70\%$ for typical values of $s$. The observed degree of polarisation in extragalactic compact jets is usually much lower, only a few percent, implying disordered magnetic fields. However, Lister & Homan (2005) have recently measured fractional polarisation of $> 50\%$ in some parsec scale jets.
Inverse Compton scattering

Radiation can be scattered by free charges. Whenever the moving charge (e.g., electron) has sufficient kinetic energy compared to the photon, net energy can be transferred from the electron to the photon resulting in an increase in the frequency of the photon. This process is called inverse Compton scattering. Again, for a detailed discussion and derivation of the relevant formulas, see Pacholczyk (1970), and chapter 7 in Rybicki & Lightman (1979).

When a photon of energy $h\nu \ll m_e c^2$ (i.e., the scattering is in the Thompson regime\(^2\)) scatters from a relativistic electron, which has a Lorentz factor\(^3\) $\gamma$, its frequency is on average boosted by a factor of $\approx \gamma^2$. The inverse Compton power emitted by the electron is

$$P_{\text{IC}} = \frac{4}{3} \sigma_T c \gamma^2 \beta^2 U_{\text{ph}}, \quad (3.4)$$

where $\sigma_T$ is the Thompson cross section, $c$ is the speed of light, and $U_{\text{ph}}$ is the energy density of the target photon field. This can be compared with the synchrotron power emitted by an electron:

$$P_{\text{synch}} = \frac{4}{3} \sigma_T c \gamma^2 \beta^2 U_B, \quad (3.5)$$

where $U_B$ is the energy density of the magnetic field. Therefore,

$$\frac{P_{\text{synch}}}{P_{\text{IC}}} = \frac{U_B}{U_{\text{ph}}} \quad (3.6)$$

that is, the radiation losses due to synchrotron and inverse Compton mechanisms are in the same ratio as the magnetic field energy density is to the photon field energy density in the source.

3.2.2 Structure of a jet

The two classical morphologies of the compact radio structure in AGNs are the asymmetric “core-jet” and the compact symmetric objects (CSOs), which resemble miniature FRII sources. Here, we are interested in the first group. The typical “core-jet” source has an unresolved or barely resolved flat-spectrum core, which is presumed to be stationary, and from which a one-sided, steep-spectrum jet is ejected (for examples, see figures in Papers III, IV, and V). The jets can appear as short or long, straight or curved – even “wigging”, and smooth or knotted. For examples of different parsec scale morphologies of compact extragalactic sources, the reader is referred to VLBI surveys. The Caltech-Jodrell Bank Survey of bright flat-spectrum radio sources is a complete flux limited sample of 293 sources, which have been observed with the global VLBI and VLBA.

\(^2\)If the photon energy is comparable to the electron rest mass energy, recoil becomes important, and the scattering cross section is reduced to the so-called Klein-Nishina cross section. The maximum attainable photon energy in scattering is $\approx \gamma m_e c^2$, and the pair production threshold can be exceeded.

\(^3\) $\gamma = 1 / \sqrt{1 - \beta^2}$ where $\beta$ is the speed of the electron in units of light speed.
at 6 cm (Taylor et al. 1994). A higher resolution survey is the VLBA 2 cm Survey tar-
gated to monitor the structural evolution of over 130 sources (Kellermann et al. (1998);  
Zensus et al. 2002; Kellermann et al. 2004).

**Relativistic effects**

The bright knots in the jet usually move at apparently superluminal speeds, but also sta-
tionary or subluminal features have been observed (Jorstad et al. 2001a; Kellermann 
et al. 2004). The apparent superluminal motion is a result of a simple geometric effect 
occuring when relativistic motion at velocity $\beta = v/c$ is observed in a small angle, $\theta$, be-
tween the direction of motion and the observer’s line-of-sight. The observed transverse 
velocity of an emitting feature is then

$$\beta_{\text{app}} = \frac{\beta \sin \theta}{1 - \beta \cos \theta},$$

which can significantly exceed $c$. The apparent superluminal speeds (which are a direct 
evidence for relativistic motion) in compact extragalactic sources were predicted already 
in 1966 by Martin Rees, and first observed with the pioneering VLBI experiments in 
early 1970’s (Whitney et al. 1971; Cohen et al. 1971).

There are also other relativistic effects affecting the appearance of the compact jets. 
For a source approaching the observer at velocity $\beta$, the relativistic Doppler shift will 
affect the frequency of the radiation by

$$\nu = \delta \nu',$$

where $\delta$ is the relativistic *Doppler factor:*

$$\delta = [\Gamma (1 - \beta \cos \theta)]^{-1},$$

where $\Gamma$ denotes the bulk Lorentz factor$^4$ $\Gamma = 1/\sqrt{1 - \beta^2}$. The primed quantities refer 
to the rest frame of the source and unprimed to the observer’s frame. All the involved 
time scales will also be shortened in the rest frame of the observer:

$$t = \delta^{-1} t'.$$

The abovementioned transformations of frequency and time, together with relativistic 
beaming of the radiation in the direction of motion of the source will result in the fol-
lowing transformation for the flux density:

$$S_\nu = \delta^{x-\alpha} S'_\nu,$$

$^4$Here we distinguish between the Lorentz factor of an individual particle, $\gamma$, and the Lorentz factor of 
the bulk flow, $\Gamma$. 
where \( \alpha \) is the spectral index of the emission, \( x = 2 \) for a continuous flow, and \( x = 3 \) for a point source or a “blob” of finite extent. In addition to this enhancement of observed flux density from an approaching jet, there will be corresponding dimming of the receding counter-jet. The flux density ratio between the two is

\[
\frac{S_j}{S_{cj}} = \left( \frac{1 + \beta \cos \theta}{1 - \beta \cos \theta} \right)^{2-\alpha}.
\]

(3.12)

As is evident from this equation, intrinsically two-sided relativistic jets will appear as one-sided when observed in a small angle to the line-of-sight.

Relativistic aberration of light will also affect the appearance of structures in the jets. Let us, for example, consider a thin emission feature propagating in a relativistic jet. Unless the jet is observed exactly end-on, the feature will be rotated in the observer’s frame. If \( \theta = \sin^{-1}(\Gamma^{-1}) \), we will see the feature from the side.

**Jet speed**

The observed superluminal motion of the knots places a constraint on the true velocity of the plasma in the jet. If the observed pattern speed corresponds to the speed of the flowing plasma, the jet Lorentz factor is \( \Gamma \geq \sqrt{1 + \beta_{app}^2} \). Hence, from VLBI observations we have immediately a lower limit for the jet speed. If the flux density variability timescale of the flares at high radio frequencies is controlled by the light travel-time across the emitting region (as it seems to be, see section 4.1.1), variability data (either total flux density or VLBI) can be used to estimate the Doppler factor of the emitting feature (Lähteenmäki & Valtaoja 1999; Jorstad et al. 2005; Paper IV). From the Doppler factor and the apparent speed of the superluminal components, it is possible to directly calculate the bulk Lorentz factor:

\[
\Gamma = \frac{\beta_{app}^2 + \delta^2 + 1}{2\delta},
\]

(3.13)

and the viewing angle of the jet:

\[
\theta = \arctan \left( \frac{2\beta_{app}}{\beta_{app}^2 + \delta^2 - 1} \right).
\]

(3.14)

However, the pattern speed does not necessarily correspond to the flow speed. Best example of this are the standing features seen in many jets (Jorstad et al. 2001a). They can occur e.g. due to a change in the direction of the flow (jet brightens where it is flowing more directly towards our line-of-sight) or due to standing shock waves (like recollimation shocks). Although the feature in these cases is not moving, the emitting plasma in them still flows at high speed. On the other hand, if the moving pattern is a shock, it can travel at a higher speed than the energised plasma overtaken by it. Fortunately, the difference in this case is only modest: for extreme ultrarelativistic plasma,
\[ \Gamma_{\text{plasma}} = 1.06 \Gamma_{\text{shock}} \left(1 - \beta_{\text{shock}}/3 \right) \sim 0.7 \Gamma_{\text{shock}} \] (Marscher 2006b). If the jet sound speed is less than \( c/\sqrt{3} \) of the ultrarelativistic case, the difference between \( \Gamma_{\text{plasma}} \) and \( \Gamma_{\text{shock}} \) is smaller. What is measured from the VLBI data is typically the motion of the centroid of the emission feature, which is roughly the mean of the shock and plasma velocities. Hence, the observed motion of superluminal components is only slightly faster than the bulk motion of the energised plasma.

The fastest observed superluminal speeds are over 40 \( c \), implying \( \Gamma \gtrsim 40 \) (Jorstad et al. 2001a; 2005). These are observed in \( \gamma \)-bright quasars, which generally show fast jet speeds and high Doppler factors (Jorstad et al. 2001a; Kellermann et al. 2004; Lähteenmäki & Valtaoja 2003). Typical radio-loud AGNs, however, do not show jet velocities that high; due to their large brightness enhancement by relativistic beaming, the sources with high speed jets overpopulate the flux limited samples, while the majority of the jets have \( \Gamma < 10 \) (Lister & Marscher 1997).

It has been suggested that significant velocity gradients could exist across a jet so that central portion of the jet is flowing much faster than the outer layer (e.g. Sol et al. 1989). There is observational evidence supporting this so-called “spine-sheath” structure in the kiloparsec scale jets of FRI radio galaxies (Laing et al. 1999), where lower speed in the layer is thought to arise from entrainment of the external medium. Recently Giroletti et al. (2004) have observed limb brightening in the parsec scale jet of the TeV blazar Mrk 501, which they interpret as a spine-sheath structure with \( \Gamma = 15 \) for the spine and \( \Gamma \sim 3 \) for the layer. The presence of velocity structure already at very small scale suggests that it may originate in the base of the jet instead of gradual entraining of surrounding material. In Papers IV and V, we find a velocity gradient across the parsec scale jet in 3C 273. There the bulk Lorentz factor changes from \( \Gamma = 6.6 \pm 1.4 \) on the northern side of the jet to \( \Gamma = 18 \pm 8 \) on the southern side.

Ghisellini et al. (2005) propose structured jets for TeV blazars in order to explain the bright and rapidly variable TeV emission with less extreme physical parameters than what are needed in one-zone SSC models. In the extreme versions of the structured jet model, the spine is an ultrafast jet of pair plasma, while the sheath is a slower velocity flow of normal plasma. As Marscher (2006b) points out, there is, however, a problem with the parent population of such jets since only a small fraction of blazar jets have Lorentz factors exceeding 10.

**VLBI core**

VLBI core is the compact, flat-spectrum, stationary component at one end of the jet in the VLBI image of a typical blazar. As already mentioned in section 2.4.2, the feature seen as a radio core in the low frequency VLBI maps (\( \lesssim 20 \) GHz) is the section of the jet where the optical depth is close to unity at the frequency of observation. Thus its location is frequency dependent (e.g. Lobanov 1998). However, the parsec scale jet cannot stay self-similar all the way to its launching site since even in the flat-spectrum quasars the synchrotron spectrum steepens somewhere between a few tens of GHz and
3.2. COMPACT JETS

\[ \sim 1000 \text{ GHz and becomes optically thin (Impey & Neugebauer 1988; Marscher 1995).} \]

The radio cores are in the order of 0.1 mas at frequencies of a few tens of GHz, resulting in a jet width of a few times \( 10^{16} \text{ cm} \) at \( \nu_m \sim 1000 \text{ GHz} \) for a source at redshift of 0.5 (Marscher 2006b). This is still hundreds of gravitational radii even for black hole mass of \( 10^9 \text{ M}_\odot \), and moreover, it is just the width of the jet. For an opening angle of a few degrees, the distance from the central engine can be as high as one parsec, i.e. the jet has its brightest point well downstream of its launching site. There are two models that can explain this. In the first one, the core is a structure of standing, conical recollimation shocks that accelerate particles locally and compress the magnetic field component that is parallel to the shock. Such recollimation shocks can occur if there is a drop sudden enough in the external gas pressure and the resulting compression waves steepen into shocks (Daly & Marscher 1988; Gómez et al. 1995). The second possible explanation is that the jet accelerates out to parsec scales and the core is the point where the flow reaches its terminal velocity (e.g. Vlahakis & Königl 2004; see also critique by Sikora et al. 2005).

**Brightness temperature**

An interesting observational quantity of the compact radio sources is their brightness temperature, which is the temperature that a black body is needed to have in order to emit the observed intensity at the given frequency. For a VLBI feature having a Gaussian surface brightness profile, the brightness temperature at the observing frequency \( \nu \) (in GHz) is

\[ T_{b,\text{VLBI}} = 1.22 \times 10^{12} \frac{S_{\text{VLBI}}(1 + z)}{a_{\text{min}}a_{\text{max}}\nu^2}, \]  

where \( S_{\text{VLBI}} \) is the flux density of the component in Jy, \( z \) is redshift, and \( a_{\text{min}} \) and \( a_{\text{max}} \) are the sizes of its minor and major axes in milliarcseconds, respectively. Brightness temperature can also be calculated from the flux density variability data by assuming that the variability timescale equals the light-crossing time across the source (for the relevant formula, see Lähteenmäki & Valtaoja 1999 or Paper IV).

Brightness temperature is a good diagnostic of the radiative processes working in compact radio sources. First of all, in a stationary synchrotron source, \( T_b \) has a strong upper limit at \( \sim \ 10^{12} \delta/(1 + z) \) K. This is due to catastrophic cooling of the emitting electrons by inverse Compton scattering in sources with higher \( T_b \) (Kellermann & Pauliny-Toth 1969). Since brightness temperatures exceeding \( 10^{12} \) K are observed (especially from variability data), the emission from compact extragalactic sources must be Doppler-boosted.

By applying a requirement of equipartition between the magnetic field energy density and the radiating particle energy density, which corresponds roughly to requiring minimum energy in the source, Readhead (1994) derived another limit for the brightness temperature in the source frame, \( T_{b,\text{eq}} \sim 5 \times 10^{10} - 10^{11} \) K. Based on the analysis of
several sets of observational data, Lähteenmäki et al. (1999) argued that the intrinsic brightness temperature in radio-loud AGN does not exceed this value.

**Bends and instabilities**

Curved trajectories of individual knots are observed commonly (Kellermann et al. 2004; Jorstad et al. 2005; Paper II), and the jets can sometimes exhibit very sharp bends. In reality, however, the jets do not go through 90° degree bending; instead, the projection effects play a role here: in a jet pointing close to our line-of-sight, even very small changes in the jet direction are exaggerated to a large degree.

In the introduction of Paper II, we list different mechanisms capable of producing the observed curved structures. These include (deflecting) interaction between the jet and a cloud of interstellar matter, precession of the jet base, and fluid instabilities, which include Kelvin-Helmholtz and, where the electrical currents are high, also the current-driven instability. The helical Kelvin-Helmholtz fundamental mode, for example, is able to displace the ridge line of the flow and produce helical jets, as observed in Paper II. Observations of these helical structures can allow derivation of sound speeds in the jet and external medium, as well as an independent measurement of the flow speed (e.g. Hardee et al. 2005). The problem here is that disentangling the effects of shocks and instability modes may not be easy since the observed jet properties can be a complex mixture of bulk and phase motions, viewing angle selection effects, and non-linear interactions between the perturbations and the underlying jet (Aloy et al. 2003).

**3.2.3 Launching of the jets**

**Matter and energy content of the jet**

The composition of the plasma in jets – i.e. whether it is pair plasma, electron-proton plasma or a combination of both – is one of the crucial questions regarding the jet physics since it is intimately linked to the jet formation process, its acceleration and other dynamics, production of high energy radiation, etc. Currently, we do not know how large is the fraction of pair plasma in jets; various studies using indirect methods to measure it give contradictory results. For example, Celotti & Fabian (1993) favour heavy jets on the basis of bulk kinetic energy and number flux of emitting particles estimated from VLBI data, while Reynolds et al. (1996) – also using VLBI data – argue for pair plasma in the nearby radio galaxy M87. Sikora & Madejski (2000), on the other hand, find that X-ray observations of optically violently variable quasars support a model in which $e^+ - e^-$ pairs are more numerous, but protons still dominate the dynamics.

Energy content of the jet is another interesting but poorly known quantity. Studies of the radio lobes of FRII galaxies give an average kinetic power of the jet, $\langle L_{\text{jet}} \rangle \gtrsim 10^{44.0\pm2}$ erg s$^{-1}$, but measuring $L_{\text{kin}}$ in compact jets is difficult. The energy density of the emitting particles and the magnetic field can be obtained from the radio data by two ways. One is to assume an equipartition between the two; this corresponds to assuming
the minimum energy condition (see the previous section about the brightness temperature). The other one is to use multifrequency VLBI data to measure the turnover frequency and flux density of the emitting region, as well as its size, optically thin spectral index, and Doppler factor. From these one can calculate the magnetic field density and the electron energy distribution normalisation factor without any assumption of equipartition. This is done in Paper V for quasar 3C 273. The electron number density depends strongly on the low frequency cutoff of their energy distribution, and hence, it cannot be reliably measured. This is unfortunate, since knowing the number density of the electrons would give us the total kinetic luminosity of the particles in the jet as a function of the ratio of cold protons to pairs.

**Acceleration and collimation by magnetic fields**

It seems nowadays very likely that the launching mechanism of relativistic jets involves differentially twisting magnetic fields that are tied either to the accretion disk or to the ergosphere of a Kerr black hole (e.g. Meier et al. 2000). General relativistic magnetohydrodynamic simulations indicate that the magnetic winds from the inner accretion disk can only reach velocities of $\Gamma \sim 3$, and currently the most promising way to produce highly collimated, high $\Gamma$ jets from AGNs is the Blandford–Znajek mechanism, which is electromagnetic extraction of the rotational energy of a spinning black hole (Blandford & Znajek 1977; see e.g. McKinney 2006 and references therein for results of GRMHD simulations).

In the magnetic jet launching, the jet will accelerate along the polar axis due to a gradient in the magnetic pressure as long as it stays Poynting flux dominated. According to Vlahakis & Königl (2004), $\sim 1/2$ of the magnetic energy is converted to kinetic energy of the flow. Collimation is provided by magnetic hoop stress due to the toroidal field component. The magnetic launching model therefore predicts that the jet starts as Poynting-flux dominated and with a helical magnetic field geometry. It is not, however, clear how far out the jet continues to accelerate, and whether the helical magnetic field is still present in the parsec scales (see Vlahakis & Königl 2004 and the critique by Sikora et al. 2005). Recently, Jorstad et al. (2005) have found that the intrinsic opening angle of the jet is inversely proportional to its Lorentz factor. This result can be explained, if there is extended acceleration of the jet flow, either by gas dynamics (original jet model by Blandford & Rees 1974) or by magnetic forces. There is also some evidence for helical magnetic fields in BL Lac objects (Gabuzda et al. 2004) and in the quasar 3C 273 (Asada et al. 2002; Zavala & Taylor 2005; Attridge et al. 2005). However, Sikora et al. (2005) consider several observational constraints and they conclude that while the acceleration is likely to take at least $10^5$ gravitational radii, the jet is proton rest mass dominated already at the parsec scale.
Highly beamed non-thermal emission from a relativistic outflow pointing nearly towards our line-of-sight dominates the observed spectral energy distribution (SED) of blazars, which makes them good targets for jet studies. The blazar SED typically exhibits two broad bumps in a $(\log(\nu), \log(\nu S_\nu))$ plot – one peaking between $10^{13}$ and $10^{17}$ Hz and the other between $10^{21}$ and $10^{24}$ Hz (Fossati et al. 1998; see Fig. 4.1 for an example). The low-frequency spectral component is due to synchrotron emission from an ensemble of relativistic electrons moving in a magnetic field. This is shown, not only by the spectral shape, but also by the polarisation and variability data, which all point towards synchrotron mechanism. The high-frequency component is usually interpreted as a result of inverse Compton (IC) scattering of soft photons off the same relativistic electrons that produce the synchrotron bump.

A characteristic property of blazars is a strong and rapid flux variability across the whole electromagnetic spectrum. The variability, which was a major riddle in itself after its discovery, is now an important tool in understanding the workings of the relativistic jets. In the following, we will discuss the variability of the synchrotron and IC components, their connection to each other, and the relation between flux density variability and structural variability in parsec scale jets.

**4.1 Synchrotron component**

**4.1.1 Observational properties**

The characteristics of synchrotron spectrum are a self-absorption turnover occurring somewhere in the radio to mm region and a power-law at optically thin frequencies, typically with a spectral index of $-0.7$, declining through IR and optical. The power-law usually steepens above a break frequency, $\nu_b$, before an exponential cut-off. The break frequency corresponds to the low-frequency peak in a $\log(\nu S_\nu)$ vs. $\log(\nu)$ plot, and the steepening is caused by radiative cooling of the highest energy electrons. In some blazars the radio to millimetre spectrum is flatter than what is expected for optically thin synchrotron emission. The apparently flat spectrum is, however, actually composed of several synchrotron components each becoming self-absorbed at a progressively lower frequency as they move down the jet (Marscher 1988; Paper V).

The flux density curves of blazars at centimetre and millimetre wavelengths show outbursts at random intervals typically lasting from months to years, depending on the frequency; the fastest flares at high radio frequencies can even have timescales as short
as weeks. The outbursts between cm and mm wavelengths are well correlated, and the flares at high frequencies precede those at low frequencies (e.g. Stevens et al. 1994; see Fig. 4.2). Also, variations in far-IR and radio are at least sometimes correlated, as are, to some extent, the millimetre waves and the optical (e.g. Robson et al. 1993; Tornikoski et al. 1994; Hanski et al. 2002). We note, however, that there are also optical flares for which no radio counterpart can be found. The variability timescale of the synchrotron flares decreases and their amplitude increases with increasing frequency. As shown in Paper I (see also references therein), the ejections of superluminal components are connected to the flares at high radio frequencies. Typically, the VLBI core brightens during the flare and a new moving component is seen in the jet after the flare has peaked.

The long term monitoring of AGNs at high radio frequencies, carried out at Metsähovi Radio Observatory since 1980, has shown that synchrotron flares are typically nearly symmetric, having (quasi-)exponential rise and decay with a sharp peak in between (Teräsranta & Valtaoja 1994; Valtaoja et al. 1999). The decay timescale is slightly longer than the rise timescale. The sharply peaking flares are very problematic for the models of blazar variability. In the case of the shock-in-jet model for the flux density...
variability (see below), the symmetric flux density curves of the flares indicate that the variability is controlled by the light-crossing time across the emitting region (Sokolov et al. 2004). However, such flares should have flat peaks in general – not sharp. Currently, we are not aware of any model that is able to naturally reproduce the observed, sharply peaked flare shape. Chiaberge & Ghisellini (1999) stress that the symmetric flares without plateau strongly constrain the injection and the cooling timescales: according to them, either the timescale of relativistic particle injection must be close to and the cooling timescale shorter than the light-crossing time, or there must be repeated injections within the light-crossing time, in order to explain the shape of the light curves.

4.1.2 Shock models

The first models trying to explain the radio variability of AGNs contained a spherical, adiabatically expanding cloud of plasma (“a blob”), which emitted incoherent synchrotron radiation from electrons with a power-law energy distribution (Shklosky 1965; van der Laan 1966; Pauliny-Toth & Kellermann 1966). These models, however, failed to explain the observed variability amplitudes at different frequencies, and they also predicted too long time lags at low radio frequencies.

Figure 4.2: Multiwaveband light curves of 3C 273. (The figure is from Stevens et al. 1994)
Nowadays, radio to infrared flares in blazars are usually explained by the shock-in-jet model (Blandford & Königl 1979; Marscher & Gear 1985; Hughes et al. 1985), which postulates that relativistic shock waves are responsible for the particle acceleration needed for the non-thermal emission, and that they appear as superluminal components in the VLBI maps (Paper I). Relativistic shocks travelling down the jet may result from variability of the physical conditions in the region where the jet is formed. If the quasi-steady process that generates the jet is occasionally disturbed by instabilities occurring in the accretion disk, the jet could be momentarily saturated by a pile of highly energetic plasma, which is much denser than the plasma in the steady flow downstream. If this high pressure plasma propagates in the jet rest frame with speed exceeding the signal speed of the medium (typically sound speed in non-magnetic medium or Alfvén speed in magnetised plasma), the speed and density changes are not communicated to the matter ahead of the disturbance, and a shock wave is formed in front of the dense material. A typical shock structure is a pair shock: the undisturbed, slower moving material ahead of the disturbance is swept up, compressed and heated by a forward shock, while the information about the shocked plasma is communicated to the pushing pile of dense upstream material by a compression wave that can steepen into a reverse shock. A contact discontinuity between the forward and reverse shock separates these two shells of plasma. The reverse shock moves backward in the frame of the forward shock but forward in the observer’s frame.

A shock front compresses the plasma, which causes an increase in the density, an adiabatic increase in the internal energy, amplification of the magnetic field component parallel to the shock front, and particle acceleration. Energisation by diffusive shock acceleration, which is a variant of the first-order Fermi mechanism, is usually thought to be behind the high electron energies required by the observed synchrotron emission. In the diffusive shock acceleration, a particle gains energy by subsequent shock crossings. The particle is scattered off magnetic irregularities, which are carried by the plasma flow that converges at the shock. The scattering-centre rest frames at different sides of the shock move at substantially different speeds, and, consequently, the particle crossing the shock gains a boost in its scattering-centre frame energy that is proportional to the velocity difference of the scattering centres. One such encounter does not lead to a very substantial energy gain in a non-relativistic shock, but the particle can be scattered back upstream where it is caught up by the shock again and a second crossing takes place. Repeated crossings can thus lead to a very large increase in the particle energy. On the other hand, in the case of a relativistic shock even one interaction with the shock front can produce a large gain in the particle energy.

In order to explain the large flare of 1983 in the quasar 3C 273, Marscher & Gear (1985) developed a model in which electrons are accelerated in a thin shock front, and then cooled by radiative and expansion losses as they advect behind the front, forming a gradient in the maximum electron energy across the shocked region i.e. the highest frequency radiation can be emitted only within a thin sheet behind the shock front, and the thickness of the sheet increases as the frequency decreases. This frequency stratifi-
CHAPTER 4. MULTIWAVEBAND Variability of BLAZARS

cation nicely explains why a flare caused by a shock can be spread over many decades in frequency, yet the variability timescale is generally much shorter at the high frequencies than at the low frequencies. The shock front is assumed to move in a smooth jet flow that is relativistic, adiabatic, and conically expanding. There are three stages of the shock evolution in the Marscher & Gear model according to the dominant cooling mechanism of the electrons: (1) the Compton scattering loss phase, (2) the synchrotron loss phase, and (3) the adiabatic expansion loss phase. During the first, Compton dominated phase, the synchrotron self-absorption turnover frequency, $\nu_{m}$, decreases a little while the turnover flux density, $S_{m}$, is increasing; this behaviour is in accordance with the observed spectral evolution of the 1983 flare of 3C 273, which is very difficult to explain with the standard model of adiabatically expanding source (Marscher & Gear 1985; see, however, also critique by Björnsson & Aslaksen 2000). During the synchrotron loss dominated second phase, $\nu_{m}$ continues to decrease while $S_{m}$ changes only a little. Once $\nu_{m}$ reaches the frequency at which the emitting electrons have lifetimes larger than the source crossing-time, the adiabatic losses become dominant and both $\nu_{m}$ and $S_{m}$ begin to decrease.

The Marscher & Gear model successfully explains the spectral evolution of the large flare of 1983 in 3C 273, and later multiwavelength studies of single outbursts in 3C 279, PKS 0420-014, 3C 345, and 3C 273 have provided additional support to the model (Litchfield et al. 1995; Stevens et al. 1995; 1996; 1998). The basic shape and the overall evolution of the flares at radio wavelengths also agree with the predictions of the Marscher & Gear model (e.g. Valtaoja et al. 1988; 1992), as do the coincident ejections of superluminal components (Paper I). Likewise, the frequency dependent angular sizes of the emission features in the parsec scale jet of 3C 273 are in overall agreement with the Marscher & Gear model (Paper V). Türler et al. (1999; 2000) have generalised the original model, and used it to fit thirteen longterm radio-submillimetre light curves of 3C 273 with distinct outbursts. Their model is able to reproduce the flare shapes at different frequencies from radio to millimetre rather well, thus giving support to the Marscher & Gear model. They also observe the predicted flattening of the optically thin spectral index from the rising to declining phase of the flare (Türler et al. 1999).

There are, however, also problems with the simple analytical model of Marscher & Gear. For example, it does not reproduce well the sharp peaks of the flares observed at high frequencies and it tends to predict too low flux densities for large flares at high frequencies (Türler et al. 2000; Lindfors et al. 2006). The rising part of the flare is another problem; in the original model it was assumed that a reduction in the synchrotron self-Compton losses cause the flare to rise, but this has been criticised by Björnsson & Aslaksen (2000) who consider one of the assumptions made by Marscher & Gear invalid.

Another shock model was developed by Hughes et al. (1985; 1989a; 1989b; 1991) who applied the fundamental idea of a shocked jet to explaining both total flux density and polarisation variations at low radio frequencies. Their numerical model is based on a piston-driven shock, which is formed inside the optically thick part of the jet and is thus first hidden. The flux density begins to rise when the shock emerges from behind the
4.2. HIGH ENERGY COMPONENT

τ = 1 surface in the jet, and reaches its maximum when the entire shock is visible. The decaying part of the flare is explained by adiabatic expansion of the shocked material. The model explains well the radio light curves and polarisation variability at low radio frequencies, but it does not incorporate the radiative energy losses, which are important at higher frequencies.

Modern versions of the shock-in-jet model include e.g. an application of the “internal shock” scenario of gamma-ray bursts to blazar variability (Spada et al. 2001). In this model, two shells of plasma, moving at different speeds, collide resulting in a pair of shocks. The formation of moving shells is attributed to intermittent variability of the central engine. A slightly different approach was adopted by Sokolov et al. (2004). In their model, a flare is triggered by a collision between a moving shock and a system of standing shocks that form the stationary core (see section 3.2.2). Sokolov et al. model both the synchrotron and the SSC components, and they take into account all relevant light-travel delay effects in calculating the light curves. If the electron cooling timescale is shorter than the light-crossing timescale across the emitting region, light-travel effects become important in determining the source appearance. For example, the observer will see radiation from the far side of the jet delayed with respect to that produced at the near side. Therefore, at any given time the observer detects radiation from various locations emitted at different times and characterised by particle distributions of different ages.

4.2 High energy component

4.2.1 Observational properties

One of the most important results in AGN studies during the 1990s was the detection of roughly 60 blazars with the EGRET instrument on board the Compton Gamma Ray Observatory at energies > 100 MeV (Mukherjee et al. 1997; Hartman et al. 1999). In many cases the γ-ray energy flux is actually dominant over the energy flux at lower frequencies – at least during outbursts. All the EGRET-detected AGNs are radio-loud and many of them exhibit variations in the γ-rays at timescales of days to months (Mukherjee et al. 1997). The variability of the high-frequency peak of the SED tends to be stronger than that of the synchrotron peak. The γ-ray emission must be relativistically beamed similarly to the synchrotron radiation since otherwise the photon-photon absorption would quench the high energy radiation.

In high-polarisation quasars, the millimetre wave variability is correlated with the γ-rays: the high γ-ray states occur during the rising part or the peak of the millimetre wave flare i.e. the millimetre wave flux density starts to rise before the γ-event (Lähteenmäki & Valtaoja 2003, and references therein). VLBI monitoring of EGRET blazars has shown that their apparent speeds are on average considerably higher than in the general class of compact radio sources (Jorstad et al. 2001a; Kellermann et al. 2004). The high γ-ray states are also statistically associated with ejections of superluminal components – the γ-flare occuring on average two months later than the zero-separation epoch of the
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superluminal component (Jorstad et al. 2001b).

Interesting correlations between the X-ray and lower frequency variability, with reverse time-lags, has been found for a few well-observed sources. McHardy et al. (1999) report correlated infrared and X-ray flaring in 3C 273 with infrared leading X-rays by 0.75 ± 0.25 days. In 3C 279 X-rays are particularly well correlated with optical emission, with the optical leading by 15 ± 15 days, in PKS 1510-089 there is a correlation between X-rays and high radio frequencies – radio preceding X-rays by 6 days, and in BL Lac the longterm X-ray and optical light curves are strongly correlated with zero lag (Marscher 2004). The ejections of superluminal components tend to occur near the epochs of X-ray flares in PKS 1510-089 and 3C 279 (Marscher 2004).

A small number of nearby low luminosity, high frequency peaking BL Lacs have been detected with atmospheric Čerenkov telescopes at TeV energies (see e.g. Krawczynski 2004 for a list of sources with corresponding references). These objects show very rapid outbursts at TeV γ-rays, with variability timescales as short as 30 minutes (Gaidos et al. 1996). The short variability timescale implies a small emitting region and this, together with the requirement that TeV γ-rays are able to escape from their production site, places a lower limit to the Doppler factor, δ ≳ 9. The variability at TeV energies is correlated with X-rays at least in two sources, Mrk 421 and Mrk 501 (Takahashi et al. 1996; Sambruna et al. 2000).

4.2.2 Models

The high energy part of the blazar SED is usually interpreted as inverse Compton (IC) scattering of soft photons from the relativistic leptons in the jet\(^1\), but there is still little agreement over a number of essential features of the production of X-rays and γ-rays. First of all, there are several possible seed photon populations for the IC scattering. In the synchrotron self-Compton model, the seed photons are synchrotron photons produced by the same relativistic electrons that scatter them later (e.g. Maraschi et al. 1992; Bloom & Marscher 1996). The obvious advantage of this model is that SSC scattering must happen at some level in a relativistic jet. In the external Compton (EC) models, the relativistic electrons interact with the diffuse ambient radiation field, which can be provided by e.g. the accretion disk (Dermer & Schlickeiser 1993), the broad line region clouds (Sikora et al. 1994), or the molecular torus around the central engine (Sikora et al. 1994; Blazejowski et al. 2000).

Secondly, the location of the high energy emission site, as well as the mechanism triggering the flaring are open issues. In the external Compton models, the emission region is usually approximated as a homogeneous blob travelling in the jet (Sikora et al.

\(^1\)Although the leptonic inverse Compton models are currently the most favoured alternative as an explanation of the high energy emission from blazars, there are also other alternatives. In hadronic models, the energy of relativistic protons is converted to high energy radiation by such processes as direct synchrotron emission of protons, photomeson production, photon-proton pair production or nuclear collisions (see e.g. Mannheim 1993).
4.2. HIGH ENERGY COMPONENT

it is natural to assume that also the high energy emission originates in the shocks, which
are the likely cause of the synchrotron flares. The correlated variations between the
synchrotron and the IC component discussed in the previous section strongly suggest
that both emission components originate from the same region or contiguous regions.
The bright synchrotron flares usually correspond to an increase in the flux density of the
VLBI core, followed by an ejection of a new superluminal component (Paper I), which is
well in agreement with the shock models. Therefore, it seems likely that also the X-ray
and γ-ray flares are linked to the formation of a new shock component in the jet, and that
the high energy emission probably originates from a site close to the VLBI core. This
poses a serious problem for those EC models in which the seed photons are emitted by
the accretion disk or by the broad line region clouds. At the distance of the VLBI core
from the central engine, the photon field provided by the accretion disk or the BLR is far
too weak to significantly contribute to the high energy emission. The reverse time-lags
observed for 3C 273, 3C 279, and PKS 1510-089 also support the SSC scenario for the
X-rays if light-travel time effects are important (Marscher 2004; Sokolov et al. 2004).

The SSC model describes quite well the X-rays from FSRQs, but fails to produce
the strongly peaked MeV-GeV emission (e.g. Maraschi et al. 1994; Lindfors et al.
2005), and therefore, EC models are usually invoked to explain the Compton peak in
these sources. However, as just pointed out, the correlated γ-ray and millimetre wave
variations mean that the Comptonised photons cannot come from the accretion disk or
from the BLR clouds. This still leaves the hot dust in the circumnuclear torus as a viable
alternative for the source of EC seed photons (Blazejowski et al. 2000). Contrary to
FSRQs, the TeV γ-ray emission from BL Lacs is well modelled by simple single-zone
SSC, although high Doppler factors up to 50 are required (e.g. Krawczynski et al. 2001).

Although SSC emission from relativistic shocks seems a good model for at least X-
ray emission in blazars, incorporating the SSC losses in the shock models in a fully con-
sistent way is difficult. Typically, seed photons for IC scattering are all assumed to have
the same frequency (so-called δ-function approximation), but this only true for strongly
peaked seed photon spectrum, like a blackbody. In the SSC case, all seed photons that
can scatter up to a given frequency contribute equally, i.e. abundant low-frequency pho-
ton are scattered by rare high-energy electrons, mid-frequency photons are scattered by
mid-energy electrons, and rare high-frequency photons are scattered by abundant low-
energy electrons, all producing roughly equal amount of Compton scattered photons of
a given frequency (McHardy et al. 1999). Worse still, light-travel time effects add an-
other complexity. Namely, the SSC losses depend on the retarded radiation field seen by
electrons at each location of the source, and the radiation field depends on the history
of the electron energy distribution, which depends on the history of energy losses, and
so on. Thus, the problem is highly non-linear and non-local in space and time (see e.g.

\(^2\)However, according to Sokolov & Marscher (2005), similar reverse time-lags are, in principle, also
possible in the EC scenario, but it should be possible to distinct between the two by observing the X-ray
spectral index variability during a flare.
Figure 4.3: Cartoon of the physical structure and emission regions of a radio-loud AGN by Marscher (2005). The length scale is logarithmic beyond $10R_S$, where $R_S$ is the Schwarzschild radius.

Sokolov et al. 2004).

Finally, in Fig. 4.3, there is a cartoon depicting the likely locations of emission regions at different wavelengths in a radio-loud quasar. The sketch is by Alan Marscher, and some parts of it are still rather speculative; for example, it is not clear whether the emission from the ambient jet between the black hole and the radio core is visible. Nevertheless, it collects together the current view of a blazar emission structure.
CHAPTER 5

Summary of the papers

5.1 Relationship between VLBI components and total flux density flares at mm-wavelengths

Paper I: Connections between Millimetre Continuum Variations and VLBI Structure in 27 AGN
by T. Savolainen, K. Wiik, E. Valtaoja, S. G. Jorstad & A. P. Marscher

As discussed in the previous chapter, the shocked jet models have been quite successful in explaining the flux variability of blazars at centimetre and millimetre wavelengths. In this framework, the moving knots in the VLBI maps are interpreted as shocks propagating down the jet. However, conclusive evidence of a link between radio variability and VLBI components has remained scarce.

In this paper, the connection between moving jet components and large flux density flares at high radio frequencies is studied by comparing multi-epoch VLBA images at 22 and 43 GHz with 22 and 37 GHz flux density curves from Metsähovi Radio Observatory quasar monitoring programme for a sample of 27 γ-bright blazars. A clear link between the two is found: for essentially every new prominent ($S > 0.1$ Jy) VLBI component emerging in the jet, there is a coincident flare in the total flux density curve. The extrapolated ejection time of a VLBI component is on average 0.2 yr after the beginning of the flare, and the flux evolution of the component mimics that of the flare.

Usually the new moving VLBI component is first observed after the total flux density flare has peaked and is already decaying. What is seen flaring in the VLBI maps, is the compact core. However, it is shown in Paper I that in most cases, this is probably a resolution effect: there is a new component, but it starts to decay already within a distance from the radio core that is comparable to the typical VLBA beam size at 43 GHz ($\sim 0.15$ mas). If the moving components are plasma energised by a shock (which seems plausible in most cases), the shocks must develop much more quickly, and closer to the core, than previously thought. Also, the results imply that the highest energy electrons in the knot lose their energy rapidly beyond the core region – probably due to strong radiative losses near the core. It seems that in situ particle acceleration is weak beyond the core region. Since there are also a few core flares for which no new moving component is observed, it is possible that in some cases the radiative losses are so strong that the electrons are cooled and the emission dies out before the knot has travelled far enough so that it could be resolved from the core.
5.2 Highly curved jet structures

**Paper II: An Extremely Curved Relativistic Jet in PKS 2136+141**


Detailed studies of sources showing curved jet structures are important since the observed properties of the bend can constrain several physical parameters – like sound speed in the jet and in the external medium (Hardee 2003; 2005). In Paper II, the discovery of an extremely curved relativistic jet in a radio-loud quasar PKS 2136+141 ($z = 2.427$) is reported. Our multifrequency VLBA data show a 210° turn, which is, to our knowledge, the largest ever observed change in the position angle of a jet. The jet has a spiral-like trajectory, which is likely a sign of an intrinsic helical geometry. The jet kinematics are derived from over eight years of VLBA monitoring of the source at 15 GHz and a sign of apparent acceleration along the jet is found. The acceleration and changes in the apparent opening angle of the jet allow us to constrain the angle between our line-of-sight and the jet.

A precessing ballistic jet – like in the case of a well-known Galactic source SS 433 – cannot produce the observed properties of PKS 2136+141. The deflections of the jet from clouds of interstellar matter are also unlikely. Instead, we propose a model where the bending is due to a helical Kelvin-Helmholtz normal mode that is driven by a periodic perturbation at the base of the jet. The model gives a very good fit to the observed jet trajectory and to the derived viewing angles. The origin of the periodic perturbation at the base of the jet could be e.g. orbital motion of a pair of supermassive black holes.

5.3 Multiwavelength campaign and VLBI monitoring of blazar 3C 66A

**Paper III: Coordinated Multiwavelength Observations of 3C 66A during the WEBT Campaign of 2003 – 2004**

by M. Böttcher, J. Harvey, M. Joshi, M. Villata, C. M. Raiteri, D. Bramel, R. Mukherjee, T. Savolainen, W. Cui, G. Fossati, I. A. Smith et al.

3C 66A ($z = 0.444$) is a low-frequency peaked BL Lac object, which has been suggested as a promising candidate for detection by the new generation of atmospheric Čerenkov telescopes. As a matter of fact, multiple detections of 3C 66A at TeV energies with the GT-48 Telescope have been reported by Neshpor et al. (1998; 2000), but so far no other group has been able to confirm these results. 3C 66A has also been previously detected at $E > 100$ MeV by the EGRET instrument on board the *Compton Gamma Ray Observatory*, and it has been a target of numerous observations at radio, IR, optical and X-rays (see e.g. Takalo et al. 1996 for an intensive optical monitoring). However, the
5.3. MULTIWAVELENGTH CAMPAIGN AND VLBI MONITORING OF BLAZAR 3C 66A

SED of the object and the correlations between the different wavebands have remained rather poorly understood since the multiwavelength observations have not been simultaneous and the source is known to exhibit large-amplitude variations on timescales down to $\sim 1$ week. To correct this, an intensive multiwavelength campaign was organised to observe 3C 66A from July 2003 to April 2004. The source was observed at radio, near-IR and optical by the Whole Earth Blazar Telescope (WEBT) collaboration, in X-rays by the RXTE satellite, and in TeV $\gamma$-rays by STACEE and by the Whipple Telescope of the VERITAS collaboration. The multiwaveband monitoring was complemented by multifrequency polarimetric imaging with the VLBA. The VLBA monitoring consisted of 9 observation epochs, three of which were during the core campaign.

In optical, the source was in an active state with its brightness gradually increasing throughout the core campaign and culminating in February 2004. On the top of the overall brightening trend, several major outbursts of $\Delta R \sim 0.3 - 0.5$ occurred on timescales of $\sim 10$ days. Also, evidence for intraday microvariability on timescales down to $\sim 2$ hours was found. Interestingly, large optical flares seem to exhibit spectral hysteresis with the $B - R$ hardness peaking several days prior to the peak of the $B$- and $R$-band flux densities, i.e. the optical spectrum softens already before the flare reaches its maximum. This might be explained within the framework of the shocked jet models as the evolution of the synchrotron break frequency, $\nu_b$, during the flare (see Figs. 7–10 in Sokolov et al. 2004). It can be seen from Fig. 15 in Paper III that $\nu_b$ of 3C 66A lies in the optical-UV region, and, on the other hand, radiative losses will probably shift $\nu_b$ towards lower frequencies as the flare progresses. If $\nu_b$ is above the $B$ band at the onset of the flare, the $B - R$ colour index will start as hard and soften after $\nu_b$ drops below $B$ band; at this point the flare can still be rising if the emitting volume is increasing, and a delay between the $B - R$ minimum and the flux density maximum is observed.

In Paper III, we present VLBA data from the core campaign period. The 22 and 43 GHz Stokes $I$ images from September 2003 to January 2004 reveal a rather smooth jet with no clearly discernible knots except for brightening at $\sim 2$ mas from the core where the jet is bending. Superluminal motion is seen for only one component, but the monitoring period is rather short; analysis of all the nine epochs will be needed to reliably study the kinematics (Savolainen et al., in prep.). An interesting result is that the brightness temperature as a function of distance along the jet, $r$, can be well fitted with a simple power-law: $T_b \propto r^{-2}$. As is shown in the paper, this, together with an equipartition assumption, suggests a magnetic field decay $B \propto D^{-1.3}$, indicating a predominantly perpendicular magnetic field orientation. Here $D$ is the cross-sectional diameter of the jet that is found to behave as $D \propto r^{0.6}$, i.e. the jet is not free within the first $2 - 3$ mas from the core. This suggests that some collimation process still works at the parsec scale. Using a totally different method, Türler et al. (2000) and Lindfors et al. (2006) have found evidence for similar collimation in 3C 273 and 3C 279, respectively. In their cases, $D \propto r^{0.8}$. 

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5.4 Multifrequency VLBA monitoring of 3C 273

**Paper IV:** Multifrequency VLBA Monitoring of 3C 273 during the INTEGRAL Campaign in 2003 – I. Kinematics of the Parsec Scale Jet from 43 GHz Data
by T. Savolainen, K. Wiik, E. Valtaoja & M. Tornikoski

**Paper V:** Multifrequency VLBA Monitoring of 3C 273 during the INTEGRAL Campaign in 2003 – II. Extraction of the VLBI Component Spectra
by T. Savolainen, K. Wiik, E. Valtaoja & M. Tornikoski

In 2003–2004 the archetypical quasar 3C 273 \((z = 0.158); \) see Courvoisier 1998 for a comprehensive review of the source properties\) was the target of an X-ray and γ-ray observing campaign with the XMM-Newton and INTEGRAL satellites (Courvoisier 2003). The satellite observations were supported by ground based monitoring at radio, millimetre and optical. Also a polarimetric multifrequency VLBA campaign was carried out in order to complement the multiwaveband monitoring data with imaging of the parsec scale jet. The aim of the campaign was to study the kinematics, polarisation, spectra and spectral evolution of the emission features in the parsec scale jet down to size scale of \(\sim 0.1\) mas, which corresponds to \(\sim 0.3\) pc at the distance of 3C 273. To achieve these goals, the source was observed with the VLBA five times during 2003. At every epoch, dual polarisation was recorded at six frequencies (5, 8.4, 15, 22, 43, and 86 GHz).

Paper IV reports the kinematic analysis of the 43 GHz VLBA data from all the five epochs, as well as discusses model fitting and estimation of the model parameters’ uncertainties in detail. The jet structure is found to be “wiggling”, a morphology also reported in several earlier studies (see references in Paper IV). It has been proposed by Abraham & Romero (1999) that the observed variation in the ejection angle of the components could be explained by a model consisting of a simple ballistic jet with a precessing nozzle. We compare our kinematic data from the year 2003 with the model of Abraham & Romero and find that the observed variations in the ejection angle and component velocity are much faster than predicted by the model. In fact, the change in the ejection angle is so fast that it is difficult to see how a relativistic jet could change its direction so abruptly. It also seems that the components are unlikely to be ballistic. However, changes in the average jet direction are, in principle, compatible with the precession model.

The kinematical analysis results in apparent component speeds ranging from \(4.6h^{-1}c\) to \(13.0h^{-1}c\). As can be seen from Fig. 10 in Paper IV, a small total flux density flare at 37 GHz occurred in January 2003, and this flare seems to be connected to the component C2, which was ejected in 2003.0 and shows a flux density evolution similar to the flare. We note that this is in accordance with the general behaviour of millimetre flares.

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\(^1\)We note that Papers IV and V discuss only part of the data gathered during the VLBA campaign; polarisation images and the evolution of the component spectra will be studied in forthcoming papers.
5.4. MULTIFREQUENCY VLBA MONITORING OF 3C 273

in blazars as reported in Paper I. Using the flux density variability timescale measured from the total flux density curve and from the flux density curves of VLBI components, we estimate the Doppler factors of five moving knots\(^2\). From the apparent speed and the Doppler factor, Lorentz factor and viewing angle for these components can be calculated. The average angle between the milliarcsecond scale jet and our line-of-sight is \(9^\circ\). The Lorentz factor of the newly ejected components is \(\sim 10\).

In Paper V, spectra of 16 jet features are extracted from the first epoch multifrequency VLBA data by applying a model transfer method, which uses a priori information of the source structure from the higher frequencies\(^3\). The presented component spectra are of good quality and they can be well fitted with the standard self-absorbed synchrotron spectrum from a homogeneous source. It is demonstrated how the flat radio spectrum in 3C 273 between 1 and 100 GHz, as observed by the single-dish radio telescopes, is composed of a number of synchrotron emitting features, each becoming self-absorbed at progressively lower frequencies as they move out along the jet. This is something that is usually assumed in the standard model of quasar jets, but seldom demonstrated observationally. In Fig. 5.1, we have plotted the component turnover frequency as a function of its size. It can be seen from the figure that the turnover frequency, \(\nu_m\), is inversely proportional to the angular size of the component at the turnover, \(a(\nu_m)\) – exactly what is expected for a flat spectrum source. We also find evidence for frequency dependent angular sizes above the peak frequency in the case of four components. This can be readily explained in the framework of the shocked jet models: the emitting electrons are accelerated in a thin shock front, and as soon as the electrons leave the acceleration region, their energy starts to decrease due to radiative cooling, establishing a gradient in the electron energy distribution across the source (Marscher 1987; Sokolov et al. 2004).

Adopting the standard incoherent synchrotron theory and assuming uniform and spherical emission sources, we use the measured spectra and component sizes to calculate the magnetic field density, \(B\), the electron energy distribution normalisation factor, \(N_0\), energy densities of the relativistic electrons and the magnetic field, and the anticipated amount of synchrotron self-Compton radiation for the emission features in the jet. The errors are estimated by Monte Carlo simulations. Despite the highly non-linear dependence of the abovementioned physical parameters on the component size and the synchrotron peak frequency, we are able to measure e.g. the magnetic field density with only a factor of two uncertainty. What makes this important, is that these results are independent of the equipartition assumption. The core shows magnetic field density of

\(^2\)The measured variability timescale, and consequently, the Doppler factor, the Lorentz factor and the viewing angle of component B2 are erroneous in Paper IV. See footnote 4 in Paper V for correct values.

\(^3\)The source model used in the spectral extraction is simpler than the one used for kinematic analysis in Paper IV; for example, the components in Paper V are circular instead of elliptical. Therefore, for the same component, there can be a difference in its size and flux density between these two papers. This, however, does not appreciably affect the shape of the spectrum – or the brightness temperature of the component, which is a measure of the surface brightness – and we can reliably calculate parameters like magnetic field density or anticipated SSC flux density since these mostly depend on the brightness temperature at the synchrotron peak frequency.
Figure 5.1: Turnover frequency of VLBI components in the parsec scale jet of 3C 273 as a function of their size at the turnover. The solid line shows a power-law fit to the data. The best fit is obtained for a power-law index of $-1.1$.

$\sim 1 \text{ G}$, and it seems that the decay of $B$ is roughly inversely proportional to the distance from the core. Unless the core is unbeamed, its magnetic energy density dominates over that of the relativistic electrons.

An interesting transverse structure is found at a distance of $1 - 2 \text{ mas}$ from the core. There are both a significant velocity (Paper IV) and magnetic field (Paper V) gradients across the jet. On the northern side of the jet, component B2 has bulk Lorentz factor $\Gamma = 6.6 \pm 1.4$, while on the southern side, components B3 and B4 show $\Gamma = 17 \pm 7$ and $\Gamma = 18 \pm 8$, respectively. The magnetic field density is $\sim 10^{-3} \text{ G}$ for B2 and $\sim 10^{-1} \text{ G}$ for B3 and B4. Hence, the larger bulk velocity corresponds to larger $B$ and vice versa. This can be understood if emission features are shocks and the larger $\Gamma$ causes larger compression ratio, which enhances the magnetic field more effectively. Another possible explanation is that we are seeing a spine/sheath structure, where $B$ decreases from the high on-axis value in the fast spine to the lower value in the slow sheath.

In 2003–2004, 3C 273 was in a very weak state at both high energies and millimetre waves (Courvoisier et al. 2003; Türler et al. 2006). This – together with correlated
IR and X-ray variations (with IR leading X-rays) observed by McHardy et al. (1999) — supports the synchrotron self-Compton origin of the jet X-ray emission in 3C 273. We estimate the anticipated first-order SSC flux density for the different jet features by using a formula given in Marscher (1987), and compare these predictions with the observed 100 keV flux density from the INTEGRAL pointing of January 2003. We find that, in general, the radio components in the parsec scale jet of 3C 273 were weak SSC emitters at the time of our VLBA observation; their estimated flux densities are several orders of magnitude below the observed value. We can identify three components that could have been significant sources of SSC X-rays:

1) The mm-wave core could produce the observed 100 keV flux density, but only if it were unbeamed. This, however, seems unlikely since the VLBI observations generally suggest beaming.

2) The VLBA and INTEGRAL data are not exactly simultaneous: there is a delay of over one month between the observations. As was already mentioned earlier, a small total flux density flare at 37 GHz occurred in January 2003 and this flare is connected to the ejection of component C2, which already shows a decaying flux density when first observed with the VLBA on February 28. Hence, at the time of the INTEGRAL pointing, C2 was closer to the core and it probably had higher brightness temperature than what is measured in late February, and consequently, also higher SSC flux density – possibly a significant percentage of the observed 100 keV flux density. Unfortunately, there is not enough data to quantify this.

3) The only strong SSC source in Fig. 16 in Paper V is the northern component B2, which could, in principle, produce all the hard X-ray emission observed in the 2003 quiescent state of 3C 273. B2 is an anomalous emission feature and the high calculated SSC flux density is due to its low magnetic field density. It is surprising – and suspicious – that one single component at a de-projected distance of ~ 25 pc from the core could emit all the observed hard X-rays. There are, however, large uncertainties involved in determining $S_{SSC}(100 \text{ keV})$, and one cannot therefore draw strong conclusions about B2’s contribution to the X-ray flux density from a single observation.

In the shock models, X-rays and $\gamma$-rays are expected to be produced by a new moving shock at or near the mm-wave core, and it is thus unfortunate that we do not have VLBA data from January 2003 since it would have been very interesting to calculate the SSC flux density of the core and component C2 at the peak of the small flare at that time.
Bibliography

Bridle, A. H., & Greisen, E. W. 1994, The NRAO AIPS Project – A Summary, AIPS Memo 87, NRAO
BIBLIOGRAPHY

Dent, W. 1965, Science, 148, 1458
Hirabayashi, H. 2000, PASJ, 52, 997
Krawczynski, H. 2004, NewAR, 48, 367
Leppänen, K. J. 1993, VLBA Scientific Memo No. 1
Leppänen, K. J. 1995, 22 GHz polarimetric imaging with the Very Long Baseline Array, Metsähovi Radio Research Station Report A20, Helsinki University of Technology, Espoo
Marscher, A. P. 2005, Mem. S. A. It., 76, 13
Rees, M. J. 1966, Nature, 211, 468
Schmidt, M. 1963, Nature, 197, 1040
Shklovsky, I. S. 1965, Nature, 206, 176
Sholomitskii, G. 1965, Astron. Zh., 42, 673
Slysh, V. I. 1963, Nature, 199, 682
van der Laan, H. 1966, Nature, 211, 1131
Williams, P. J. S. 1963, Nature, 200, 56

Abbreviations of journal names

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Original Publications
Connections between millimetre continuum variations and VLBI structure in 27 AGN

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Abstract. We compare total flux density variations in 27 γ-ray blazars with structural changes in their parsec-scale jets using multi-epoch VLBA observations at 22 and 43 GHz together with data from the Metsähovi quasar monitoring program at 22 and 37 GHz. There is a clear connection between total flux density outbursts and VLBI components emerging into the jet. For essentially every new moving VLBI component, there is a coincident total flux density flare, with evolution similar to that of the component. Furthermore, extrapolated ejection times of the new VLBI components correspond to the beginnings of associated flares. Our results suggest that it is possible to explain all the radio variations as shocks propagating down the jet. A large fraction of the shocks grow and decay within the innermost few tenths of a milliarcsecond and therefore we see them only as “core flares” in the VLBI images. However, with present data we cannot exclude the possibility that the core itself also brightens (and thus contributes to the flare) as a shock passes through it.

Key words. BL Lacertae objects: general – galaxies: active – galaxies: jets – quasars: general – radio continuum: galaxies – techniques: interferometric

1. Introduction

Blazars are an interesting and violent subclass of active galactic nuclei (AGN), grouping together (although somewhat artificially from a physical point of view) radio-loud quasars and BL Lacertae objects. These sources have in common flat cm-wave radio spectrum, high and variable polarization, and pronounced variability of the flux density at all frequencies. The superluminal motion observed in these sources together with brightness temperatures in excess of the $10^{12}$ K inverse Compton limit (Kellermann & Pauliny-Toth 1968) indicate highly beamed emission from relativistic jets oriented towards the line of sight of the observer. Using modern-day very long baseline interferometry (VLBI) techniques we can resolve the jet-like structures in blazars on angular scales down to ~0.1 milliarcsec (mas).

Relativistic jets also offer an explanation for radio-to-infrared variability of blazars. Marscher & Gear (1985) studied the strong 1983 outburst in 3C 273 and managed to fit the flaring spectra with self-absorbed synchrotron emission. They explained successfully the time-evolution of the flare as being due to a shock wave propagating in the relativistic jet. The Marscher & Gear model (hereafter MG-model) has three stages of shock evolution based on the dominant cooling mechanisms of the electrons: 1) the Compton scattering loss phase, 2) the synchrotron radiation loss phase and 3) the adiabatic expansion loss phase. MG-model is a simple, analytical model, which describes well the general behaviour of the radio outbursts in AGN (but see the critique of Björnsson & Aslak 2000). The model was generalized by Marscher et al. (1992) to include the effects of bending in jets and turbulence on the light curves.

Hughes et al. (1985, 1989a, 1989b, 1991) proposed a similar shock model based on a numerical code simulating a piston-driven shock. Their model successfully explains the lower frequency variability, but it does not incorporate radiative energy losses of the electrons, which are important at high frequencies and in the earliest stages of the shock evolution. Valtaoja et al. (1992b) presented a generalized shock model describing qualitatively the three stages of the shock evolution (growth, plateau and decay) without going into details, thus providing a framework for comparison between the theory and observations. Total flux density (TFD) monitoring campaigns, which provide nearly fully sampled flux curves at radio wavelengths, and VLBI images, which allow us to map the parsec-scale
structure of the blazar jets, are the two main observational tools for constraining theoretical models.

In VLBI observations of blazars, bright knots of emission referred to as “components” are seen. These components line up to form jet-like features appearing in various forms from very straight to heavily bent structures. The so-called “core” is the point where the jet becomes visible. The core is presumed stationary (see Bartel et al. 1986), but most of the other VLBI components move outward in the jet at apparent superluminal speeds. However, in some sources there are also stationary components other than the core. These may be due, for example, to interactions between the jet and the surrounding interstellar medium.

According to the shocked jet models, moving components in the VLBI maps are interpreted as shock propagating down the jet. However, there has been a dearth of conclusive evidence linking VLBI components with radio flux variations; only a relatively small number of individual sources have been investigated thus far. One of the first studies linking TFD variations with moving knots in the VLBI maps was carried out by Mutel et al. (1990). They found that each of four major TFD outbursts of BL Lac between 1980 and 1988 can be associated with the emergence of a new superluminal component. Abraham et al. (1996) estimated the ejection times of seven VLBI components in 3C 273 and noticed that all ejections were related to increases in the single-dish flux density at frequencies higher than 22 GHz. Türrler et al. (1999) also studied 3C 273 by decomposing multi-frequency light curves into a series of self-similar flares. They found good correspondence between the ejection times of the VLBI components and the beginning times of the flares. Krichbaum et al. (1998) have reported a correlation between mm-VLBI component ejections and local minima in the 90 GHz total flux density curve of PKS 0528+134. For 3C 345, which is one of the best observed sources with VLBI at 22 GHz, Valtaoja et al. (1999) were able to associate VLBI components with individual millimetre flares. Similar correlations were also found for PKS 0420–014 (Britzen et al. 2000) and for 3C 279 (Wehrle et al. 2001). In our study, we compare for the first time two large data sets: multi-epoch VLBA images of 42 blazars (Jorstad et al. 2001a) detected at 0.1–3 GeV by EGRET and TFD data from the mm-wave Metsähovi Radio Observatory quasar monitoring program. A description of our data is given in Sect. 2. Our aim is to establish connections between TFD variations and structural changes in the jets. The results from the analysis, as we will show in Sect. 3, strongly support the shocked jet model.

The VLBI core is the dominant component in almost all the cases studied. The core region was usually also highly variable, being responsible for most of the observed TFD variability in these sources. Variations in the VLBI core flux are reported in the literature quite often (see, e.g., the results of the recent VLBA monitoring of 3C 279 by Wehrle et al. 2001). Since the core is usually assumed to be the apex of the jet, the implicit assumption is that a core flare results from a change in the jet flow parameters. However, according to our study, these variations are rather related to moving VLBI components that blend with the radio core. This will be discussed in Sect. 4.

2. The data

Our data set consists of observations from two separate campaigns, namely the VLBA monitoring of EGRET-detected blazars by Jorstad et al. (2001a) and the Metsähovi Radio Observatory quasar monitoring program. Jorstad et al. (2001a) monitored a sample of 42 γ-bright blazars at 22 and 43 GHz with VLBA between 1993 and 1997. For 27 of these sources (see Table 1), there were enough TFD variation data available from Metsähovi monitoring (Teräsranta et al. 1998) to reliably identify large outbursts in the flux curves.

Selection criteria for sources observed by Jorstad et al. (2001a) with the VLBA were: (1) detection by EGRET ($E > 100$ MeV; Hartman et al. 1999); (2) flux density at 37 GHz $\geq 1$ Jy; and (3) declination ($\delta$) $\geq$ $-30^\circ$. These criteria give a sample containing roughly 60% of the known γ-ray blazars (Hartman et al. 1999). The VLBA observations were made at high radio frequencies giving $\approx 0.1$–$0.3$ mas resolution and $\approx 10$ mas map size. High resolution allows us to study the inner parts of the jet and possibly see how the shock formation is connected to the flaring behaviour of these sources. However, as we discuss below, VLBA maps at 43 GHz in many cases still have insufficient resolution to separate the new shock in the jet from the core before the millimetre flare is over.

Since its beginning in 1980, the Metsähovi quasar monitoring program has been the most comprehensive such program at high radio frequencies. The Metsähovi sample contains 157 individual sources including about 100 of the brightest radio-loud AGN in the Northern hemisphere (declination $\geq -10^\circ$), which are observed at 22, 37 and 87 GHz (see Teräsranta et al. 1998 for details). The sample also includes the Northern 2 Jy catalogue of flat spectrum sources (Valtaoja et al. 1992a) fulfilling the following criteria: $\delta \geq 0^\circ$, $\alpha_{2000}(2.7–5$ GHz) $\geq -0.5$ ($S \propto \nu^\alpha$), with $\alpha$ taken from the catalogue of Kühr et al. (1981), and $S_{max}(22$ GHz) $> 2$ Jy. Of the 27 sources in our study, 13 belong to this 2 Jy catalogue. Of the 14 sources that do not belong, 5 have declination below $0^\circ$ and the rest were fainter than 2 Jy at 22 GHz prior to 1992. In our study, we used 22 and 37 GHz Metsähovi data from 1990 to 1998 together with 22 and 43 GHz VLBA maps. Comparing 37 GHz TFD-data with 43 GHz VLBA maps is justified by the typically flat spectra of our sources in the millimetre region.

Our sample of 27 sources with good VLBI and TFD data consists of 12 high optical polarization quasars (HPQs), 7 low optical polarization quasars (LPQs), 7 BL Lacertae objects (BLOs) and one object classified as a radio galaxy (GAL). The percentage of each class of radio-loud AGN is given in Table 2 for our sample, for the 2 Jy sample (Valtaoja et al. 1992a), and for a sample containing the EGRET blazar identifications that have a high probability of being correct (Mattox et al. 2001). As one can see, our sample is very similar to the γ-ray blazars as well as to the 2 Jy sample representing the brightest radio-loud AGN. [The one radio galaxy in our sample, 0446+112, is a less certain EGRET identification, which is the reason why it is not included in the list by Mattox et al. (2001).] This supports the notion that the results presented in this paper are applicable to all γ-ray blazars and, to some extent, to radio-loud AGN in general.
3. Connections between millimetre flux curves and VLBI components

In order to compare TFD and VLBI events in our data, we decompose the total flux density variations in the Metsähovi 22/37 GHz flux curves into exponential flares of the form

$$\Delta S(t) = \begin{cases} S_{\text{max}}e^{(t-t_{\text{max}})/\tau}, & t < t_{\text{max}} \\ S_{\text{max}}e^{-(t-t_{\text{max}})/1.3\tau}, & t > t_{\text{max}} \end{cases}$$

(1)

Here $S_{\text{max}}$ is the maximum amplitude of the flare, $t_{\text{max}}$ is the epoch of the flare maximum and $\tau$ is the flare rise timescale. It has been shown earlier that all TFD variations can be modelled to surprising accuracy with a small number of flares consisting of an exponential rise, sharp peak and exponential decay superposed on a constant baseline flux (Valtaoja et al. 1999). This decomposition (described in more detail by Valtaoja et al. 1999) helps us to identify and isolate individual events as well as to estimate amplitudes and timescales of the outbursts.

Next we plotted the TFD decompositions and the flux variations of the VLBI components for each source. Two examples illustrate the results of this comparison: 1633+382 (4C 38.41) and PKS 2230+114 (CTA 102) are shown in Figs. 1 and 2. The component identifications can be found in Jorstad et al. (2001a). Even at first glance, it is evident that there is a clear connection between the millimetre continuum variations and the VLBI component fluxes. Whenever there are enough VLBA observations, the summed flux curve of the VLBI components is similar to the continuum flux curve; only the amplitude of the former is ~90% that of the latter. This is expected with the missing 10% of the flux in the VLBI maps probably just due to the insensitivity of high-frequency VLBI to diffuse emission. There is a slight time shift between the 37 GHz TFD curves and the 43 GHz VLBI component flux curves. This is understandable according to the shock models, since the maximum amplitude of the flare moves from high frequencies to lower frequencies as the shock evolves.

A much more interesting result is that for every superluminal ejection seen in the VLBA data, the TFD decomposition shows a coinciding flare. We examine ejections having zero epochs after the year 1990. For most of our sources, Metsähovi TFD monitoring is rather sparse before this and therefore not suitable for our comparison. We exclude two ejections because of large gaps in the Metsähovi flux curve at their zero epochs (the observation gap in 1994). We require that there be at least three observations of the ejected component and that the

<table>
<thead>
<tr>
<th>Source</th>
<th>Other desig.</th>
<th>Class</th>
<th>$z$</th>
<th>Epochs</th>
<th>22 GHz</th>
<th>43 GHz</th>
<th>$N$</th>
</tr>
</thead>
<tbody>
<tr>
<td>0202+149</td>
<td>HPQ</td>
<td>0.833</td>
<td>1995–97</td>
<td>–</td>
<td>+</td>
<td>4</td>
<td></td>
</tr>
<tr>
<td>0219+428</td>
<td>3C 66A</td>
<td>BLO</td>
<td>0.444</td>
<td>1995–97</td>
<td>+</td>
<td>+</td>
<td>7</td>
</tr>
<tr>
<td>0234+285</td>
<td>CTD 20</td>
<td>HPQ</td>
<td>1.207</td>
<td>1995–97</td>
<td>+</td>
<td>+</td>
<td>4</td>
</tr>
<tr>
<td>0235+164</td>
<td>AO 235+164</td>
<td>BLO</td>
<td>0.94</td>
<td>1995–96</td>
<td>–</td>
<td>+</td>
<td>6</td>
</tr>
<tr>
<td>0420–041</td>
<td>OA 129</td>
<td>HPQ</td>
<td>0.915</td>
<td>1995–97</td>
<td>+</td>
<td>+</td>
<td>8</td>
</tr>
<tr>
<td>0446+112</td>
<td>GAR</td>
<td>BLO</td>
<td>1.207</td>
<td>1995–97</td>
<td>+</td>
<td>+</td>
<td>4</td>
</tr>
<tr>
<td>0448–020</td>
<td>HPQ</td>
<td>2.82</td>
<td>1995–97</td>
<td>+</td>
<td>+</td>
<td>5</td>
<td></td>
</tr>
<tr>
<td>0528+134</td>
<td>LPQ</td>
<td>2.07</td>
<td>1994–97</td>
<td>+</td>
<td>+</td>
<td>11</td>
<td></td>
</tr>
<tr>
<td>0716+714</td>
<td>BLO</td>
<td>&gt;0.2</td>
<td>1995–97</td>
<td>+</td>
<td>–</td>
<td>9</td>
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</tr>
<tr>
<td>0804+499</td>
<td>OJ 508</td>
<td>HPQ</td>
<td>1.43</td>
<td>1996–97</td>
<td>+</td>
<td>+</td>
<td>3</td>
</tr>
<tr>
<td>0827+243</td>
<td>LPQ</td>
<td>2.046</td>
<td>1995–97</td>
<td>+</td>
<td>+</td>
<td>6</td>
<td></td>
</tr>
<tr>
<td>0836+710</td>
<td>HPQ</td>
<td>2.17</td>
<td>1995–97</td>
<td>+</td>
<td>+</td>
<td>7</td>
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</tr>
<tr>
<td>0851+202</td>
<td>OJ 287</td>
<td>BLO</td>
<td>0.306</td>
<td>1995–96</td>
<td>+</td>
<td>–</td>
<td>7</td>
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<tr>
<td>0954+658</td>
<td>BLO</td>
<td>0.367</td>
<td>1995–96</td>
<td>+</td>
<td>–</td>
<td>3</td>
<td></td>
</tr>
<tr>
<td>1101+384</td>
<td>Mkn 421</td>
<td>BLO</td>
<td>0.031</td>
<td>1995–97</td>
<td>+</td>
<td>–</td>
<td>8</td>
</tr>
<tr>
<td>1156+295</td>
<td>4C 29.45</td>
<td>HPQ</td>
<td>0.729</td>
<td>1995–97</td>
<td>+</td>
<td>+</td>
<td>5</td>
</tr>
<tr>
<td>1219+285</td>
<td>ON 231</td>
<td>BLO</td>
<td>0.102</td>
<td>1995–97</td>
<td>+</td>
<td>–</td>
<td>3</td>
</tr>
<tr>
<td>1222+216</td>
<td>4C 21.35</td>
<td>LPQ</td>
<td>0.435</td>
<td>1996–97</td>
<td>+</td>
<td>–</td>
<td>2</td>
</tr>
<tr>
<td>1226+023</td>
<td>3C 273</td>
<td>LPQ</td>
<td>0.158</td>
<td>1993–95</td>
<td>+</td>
<td>+</td>
<td>5</td>
</tr>
<tr>
<td>1253–055</td>
<td>3C 279</td>
<td>HPQ</td>
<td>0.538</td>
<td>1993–97</td>
<td>+</td>
<td>+</td>
<td>10</td>
</tr>
<tr>
<td>1510–089</td>
<td>HPQ</td>
<td>0.361</td>
<td>1995–97</td>
<td>–</td>
<td>+</td>
<td>5</td>
<td></td>
</tr>
<tr>
<td>1606+106</td>
<td>LPQ</td>
<td>1.226</td>
<td>1994–97</td>
<td>+</td>
<td>+</td>
<td>5</td>
<td></td>
</tr>
<tr>
<td>1611+343</td>
<td>DA 406</td>
<td>LPQ</td>
<td>1.401</td>
<td>1994–97</td>
<td>+</td>
<td>+</td>
<td>11</td>
</tr>
<tr>
<td>1633+382</td>
<td>4C 38.41</td>
<td>LPQ</td>
<td>1.814</td>
<td>1994–96</td>
<td>+</td>
<td>–</td>
<td>6</td>
</tr>
<tr>
<td>1741–038</td>
<td>HPQ</td>
<td>1.054</td>
<td>1995–97</td>
<td>+</td>
<td>–</td>
<td>2</td>
<td></td>
</tr>
<tr>
<td>2230+114</td>
<td>CTA 102</td>
<td>HPQ</td>
<td>1.037</td>
<td>1995–97</td>
<td>+</td>
<td>–</td>
<td>7</td>
</tr>
<tr>
<td>2251+158</td>
<td>3C 454.3</td>
<td>HPQ</td>
<td>0.859</td>
<td>1995–96</td>
<td>+</td>
<td>+</td>
<td>12</td>
</tr>
</tbody>
</table>

Table 1. List of sources in our sample. Here $z$ is the redshift and $N$ is the total number of VLBA observations.

<table>
<thead>
<tr>
<th>Class</th>
<th>Our sample</th>
<th>2 Jy sample</th>
<th>EGRET blazars</th>
</tr>
</thead>
<tbody>
<tr>
<td>HPQ</td>
<td>44%</td>
<td>28%</td>
<td>37%</td>
</tr>
<tr>
<td>LPQ</td>
<td>26%</td>
<td>32%</td>
<td>30%</td>
</tr>
<tr>
<td>BLO</td>
<td>26%</td>
<td>28%</td>
<td>26%</td>
</tr>
<tr>
<td>GAL</td>
<td>4%</td>
<td>8%</td>
<td>0%</td>
</tr>
<tr>
<td>N/A</td>
<td>0%</td>
<td>4%</td>
<td>7%</td>
</tr>
</tbody>
</table>

Table 2. The percentage of each class of the radio-loud AGN in three different samples. [HPQ = high optical polarization quasar, LPQ = low optical polarization quasar, BLO = BL Lacertae object, GAL = radio galaxy and N/A = No exact classification available].
observed flux density of the component be greater than 0.1 Jy (the approximate noise level of Metsähovi observations) at some time. In our data, there are 29 ejections of VLBI components fulfilling the above criteria (see Table 3). The TFD flares corresponding to these 29 ejections are identified by comparing the component ejection times with the beginning times of the TFD flares, as well as by comparing the light curves of the VLBI components with those of the decomposed TFD flares.

We define the beginning of an exponential TFD flare as $t_{0,\text{TFD}} = t_{\text{max}} - \tau$, where $\tau$ is the variability timescale (e-folding time). This definition gives the point where $S(t_{0,\text{TFD}}) = \frac{1}{2}S_{\text{max}}$. While there is no mathematical sense in defining the beginning of an exponential function, in reality there must be a starting point to a flare. We could instead estimate the beginning of the flare as the previous local minimum of the flux curve ($t_{\text{lm}}$). If we compare $t_{\text{lm}}$ to $t_{0,\text{TFD}}$ for an outburst that starts just after a local minimum, we see that the average time difference between the two is 0.0 years with a standard deviation of 0.4 years (see Fig. 3). Therefore, the average values of $t_{0,\text{TFD}}$ and $t_{\text{lm}}$ are the same. When two or more closely spaced outbursts blend together, the local minimum is no longer a good indicator of the start of the flare. In such a case the local minimum is near the peak rather than the beginning of the later flare. Hence, $t_{0,\text{TFD}}$ is a more reliable and practical starting point to a flare.

We compare the extrapolated ejection epochs of the superluminal knots (from Jorstad et al. 2001a) with the beginning times of the TFD flares. In 28 of the 29 cases we find a TFD flare that occurred within 0.5 yr of the ejection epoch. The only exception is component E2+B1 of 3C 279, for which $t_{0,\text{TFD}}$ is not very well determined.

The frequency of large TFD flares ($\Delta S > 0.3 \cdot S_{\text{quiescent}}$) estimated from the Metsähovi data is 1 per 1.6 years. On the other hand, the frequency of observed superluminal ejections is approximately 1 per 2.3 years. Using these values we calculate the probability that a superluminal ejection could occur by random chance within a time interval $dt$ before or after the
beginning of the TFD flare. The results are given in Table 4. For every applied $\Delta t$ range the expected number of random occurrences is clearly much lower than the observed number of coincidences. The probability that 28 out of 29 ejections would be observed to occur randomly within 0.5 yr of the beginnings of TFD flares is $<10^{-7}$. Hence, the correspondence between $t_{0,\text{VLBI}}$ and $t_{0,\text{TFD}}$ is real at a very high level of significance.

We therefore find that, at high radio frequencies, the start of a TFD flare precedes the arrival of a new superluminal knot at the position of the brightness centroid of the core of the jet. However, we do not have enough VLBI data to say if the converse is true, i.e., whether there is a new VLBI component for every TFD flare. When we see the new VLBI component for the first time, the flux of the TFD flare is usually already decreasing. This behaviour is analysed in Sect. 4, in which we discuss the so-called core flares.

The mean time difference between the zero epoch of the VLBI components and the beginning of the TFD flares $\Delta t = t_{0,\text{VLBI}} - t_{0,\text{TFD}}$ is $(0.19 \pm 0.04)$ yr (ignoring component E2+B1 in 3C 279). The extrapolated ejection time of a VLBI component is therefore $\sim 0.2$ yr after the beginning of the associated TFD flare, on average. This may indicate that the proper motion of a typical VLBI knot accelerates during the early stages in the component’s evolution. On the other hand, we note that $t_{0,\text{VLBI}}$ is the moment when the component is coincident with the brightness centroid of the core. In this case, the TFD flare might begin when the disturbance that creates the shock first hits the inner edge of the core, which would occur before $t_{0,\text{VLBI}}$.

The fluxes of the VLBI components and the decomposed TFD flares are correlated. In Fig. 4 we plot the VLBI component fluxes vs. the decomposed TFD flare fluxes at the VLBI epochs (from our exponential-flare model fits). The Spearman correlation coefficient of this graph is $r_S = 0.817$; the probability that $r_S$ would be this high from uncorrelated data is $<10^{-14}$. The linear Pearson correlation coefficient $r_P = 0.76$, which corresponds to a probability of $<10^{-12}$ that the correlation is by chance. Furthermore, in 59% of the cases the fluxes differ by less than a factor of two. There is therefore a clear connection between new VLBI components in the jet and mm-wave flares in γ-ray blazars.
Table 3. The zero epochs of the VLBI components ($t_{0,\text{VLBI}}$) determined by Jorstad et al. (2001a) and the start times of the corresponding TFD flares ($t_{0,\text{TFD}}$). The last column gives the time difference $\Delta t = t_{0,\text{VLBI}} - t_{0,\text{TFD}}$. The designation of the components follows Jorstad et al. (2001a).

<table>
<thead>
<tr>
<th>Source</th>
<th>Comp.</th>
<th>$t_{0,\text{VLBI}}$</th>
<th>$t_{0,\text{TFD}}$</th>
<th>$\Delta t$ [yr]</th>
</tr>
</thead>
<tbody>
<tr>
<td>0202+149</td>
<td>B</td>
<td>1994.8 ± 0.1</td>
<td>1994.7</td>
<td>0.1</td>
</tr>
<tr>
<td>0219+428</td>
<td>B</td>
<td>1995.7 ± 0.4</td>
<td>1995.7</td>
<td>0.0</td>
</tr>
<tr>
<td>0235+164</td>
<td>B</td>
<td>1995.4 ± 0.1</td>
<td>1994.9</td>
<td>0.5</td>
</tr>
<tr>
<td>0420–014</td>
<td>B</td>
<td>1995.3 ± 0.1</td>
<td>1994.9</td>
<td>0.3</td>
</tr>
<tr>
<td>0458–020</td>
<td>B</td>
<td>1994.0 ± 0.1</td>
<td>1993.6</td>
<td>0.4</td>
</tr>
<tr>
<td>0528+134</td>
<td>B</td>
<td>1995.5 ± 0.1</td>
<td>1995.5</td>
<td>0.0</td>
</tr>
<tr>
<td>0827+243</td>
<td>B</td>
<td>1994.7 ± 0.1</td>
<td>1994.6</td>
<td>0.1</td>
</tr>
<tr>
<td>0851+202</td>
<td>B</td>
<td>≈ 1996.6</td>
<td>1996.6</td>
<td>0.0</td>
</tr>
<tr>
<td>1156+295</td>
<td>B</td>
<td>1995.6 ± 0.1</td>
<td>1995.6</td>
<td>0.0</td>
</tr>
<tr>
<td>1226+023</td>
<td>B</td>
<td>1993.4 ± 0.3</td>
<td>1993.4</td>
<td>0.0</td>
</tr>
<tr>
<td>1253–055</td>
<td>B</td>
<td>1992.4 ± 0.3</td>
<td>1991.9</td>
<td>0.5</td>
</tr>
<tr>
<td>1510–089</td>
<td>B</td>
<td>1995.3 ± 0.1</td>
<td>1995.1</td>
<td>0.2</td>
</tr>
<tr>
<td>1633+382</td>
<td>B</td>
<td>1990.4 ± 0.3</td>
<td>1990.7</td>
<td>0.3</td>
</tr>
<tr>
<td>2230+114</td>
<td>B</td>
<td>1995.5 ± 0.1</td>
<td>1995.5</td>
<td>0.2</td>
</tr>
<tr>
<td>2251+158</td>
<td>B</td>
<td>1994.8 ± 0.1</td>
<td>1994.7</td>
<td>0.1</td>
</tr>
</tbody>
</table>

Table 4. The probability $P$ of coincidence of superluminal ejections and the start of TFD flares, for both the random and observed case ($P_{\text{Observed}} = N_{\text{Associated}}/N_{\text{Total}}$). The interval $\Delta t$ is the maximum time difference between the ejection of the knot and the beginning of a TFD flare for them to be considered associated. $\rho$ is the probability that the observed number of associations could occur by random chance.

<table>
<thead>
<tr>
<th>$\Delta t$ [yr]</th>
<th>$P_{\text{Random}}$</th>
<th>$P_{\text{Observed}}$</th>
<th>$\rho$</th>
</tr>
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<tr>
<td>0.5</td>
<td>47%</td>
<td>97%</td>
<td>&lt; 10$^{-7}$</td>
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<tr>
<td>0.4</td>
<td>40%</td>
<td>90%</td>
<td>&lt; 10$^{-7}$</td>
</tr>
<tr>
<td>0.3</td>
<td>32%</td>
<td>76%</td>
<td>&lt; 10$^{-8}$</td>
</tr>
<tr>
<td>0.2</td>
<td>22%</td>
<td>59%</td>
<td>&lt; 10$^{-4}$</td>
</tr>
<tr>
<td>0.1</td>
<td>12%</td>
<td>41%</td>
<td>&lt; 10$^{-5}$</td>
</tr>
</tbody>
</table>

4. VLBI core flares

In our sample, the core is the brightest VLBI component in 152 of the 165 VLBA images. It usually contains 50–90% of the total flux density of the source. Typically it is the core region that becomes brighter and fainter during the TFD flares. Similar behaviour has been noted previously in, for example, 3C 279 (Wehrle et al. 2001). This pattern is completely different from that observed in 3C 345, in which the fluxes of the components peak well downstream of the core (Valtaoja et al. 1999). In our sample there is one source, PKS 0202+149, with a component (B) whose flux clearly peaks downstream. However, this component is not very strong (0.3 Jy at maximum) and it still peaks quite close to the core – at a separation of only 0.36 mas. There are also a few other sources in our sample with a component that might peak downstream, but we do not have sufficiently well-sampled observations to confirm these. In any case, these features are considerably weaker than the flux level involved in most of the core flares.

These apparent VLBI core flares are therefore largely responsible for the TFD variability in our sources. The question is: Are these core flares due to true changes in the radio core (changes in bulk Lorentz factor, electron energy spectrum, etc.) or are they associated with the birth of new moving VLBI components that blend with the core at the resolution of the images?
We define a core flare to be an increase in the VLBI core-region flux that is (1) >30% within one year and (2) greater than the noise level of the TFD curve. There were altogether 24 core flux variations satisfying our conditions. When analysing these VLBI core flares and coincident TFD variations, we noticed that, in 11 cases, after the core flare and an associated TFD outburst had peaked and were already fading, a new VLBI component appeared in the jet (see Table 5). Moreover, the flux of this new component was decreasing as well. Of 13 cases not showing a new component after the core flare, 6 simply did not have enough data because the core flare occurred during the last observing epoch. Thus, in 11 of the 18 cases (61%) the observed increase in VLBI core flux is consistent with being caused by the appearance of a new moving component, which brightens and then starts to decay within ≈0.15 mas (the typical VLBA beamsize at 43 GHz) of the actual radio core. This represents a high percentage if we consider the sparsity of the VLBI time coverage, systematic uncertainties in VLBI model fitting, and the problems related to component identification across epochs.

We explore the above interpretation further by reconstructing the core flares under the assumption that only new components contribute to the variations in flux while the brightness of the actual core remains constant. We then decompose the total flux into a constant baseline flux from the core plus a new shock component with variable flux. We choose the value of the baseline flux density to be the lowest observed level (either the core component flux or the weakest TFD flux, whichever is lower). Two examples of our reconstructions are presented in Figs. 5 and 6 (PKS 0420−014 and PKS 0528+134).

In order to test the above scenario, we estimate the separation between the true core and the new VLBI component during core flares by assuming that the proper motions remain constant from the birth of a knot to its last appearance on the images of Jorstad et al. (2001a). We find five cases (PKS 0420−014: 1995 Apr. 22, 1995 Jun. 21, 1995 Aug. 04; PKS 0528+134: 1995 Apr. 22, and 1156+295: 1996 May 04) in which the expected separation during a VLBA observation was larger than about one-third of the beam size. If the source is not very complicated, i.e., if it can be described using a small number of circular-Gaussian components, partially resolved compact structure is revealed by non-zero closure phases on the longest baselines. Closure phases are constructed to cancel out antenna-based calibration errors, hence they indicate the presence of non-point components in a model-independent fashion.

Partially resolved structure in the core region can take on two basic forms: (1) a slightly extended core with no other components or (2) one or more non-core components within one beamwidth of the brightness centroid of the core. To compare these two hypotheses we plot closure phases from triangles consisting of both long and short baselines (Figs. 7 and 8). We superimpose two best-fit models corresponding to the above hypotheses. It is evident – and also expected – that these models differ significantly only on long baselines. It can be clearly seen that the two-component core models provide much better fits to the data than do the single-component models. Also, in Fig. 7 the progressively increasing absolute value of the closure phase over the three successive epochs can be explained by a moving component close to the core.

We therefore find evidence for new components in the vicinity of the VLBI core during major TFD flares, but the brightness of these moving knots has usually decayed considerably by the time the features become distinct from the core on VLBI maps. This implies that a majority – perhaps all – of the strongest TFD flares are associated with bright knots that fade within ~0.1−0.3 mas of the core. (The only source in our data set that has substantial variations in flux outside the core region is 3C 454.3, where stationary feature C has brightness and variability comparable to the core.) If the moving components are shock waves, the shocks must therefore develop much more quickly, and closer to the core, than previously thought.

The bright features found in jets of classical superluminal radio sources well downstream of the core are mainly lower-frequency phenomena, although in some objects (e.g., 3C 279; see Wehrle et al. 2001) an occasional prominent component remains bright long after it detaches from the core on the images. The implication is that the critical synchrotron frequency of the highest energy electrons in the knot decays rapidly beyond the core region. This is probably the combination of (1) radiative energy losses near the core, (2) decay of the magnetic field and electron energies from downstream expansion of the jet, and (3) weak in situ particle acceleration outside the core region. These conditions have important implications for models of shocks and other structures in relativistic jets.

It is possible that all of the variations in flux in the VLBI core region during these flares actually occur in new moving knots. However, our observations are also consistent with the possibility that a disturbance both creates a new shock wave and brightens the core itself as it passes through the core region. There may be some evidence supporting this scenario, since in two sources, 3C 279 and 1510−089, the new VLBI component is observed during (rather than only after) a major core flare. Based on the present data, we cannot determine which scenario is correct. The model in which the TFD variability is completely due to shocks is the simpler and requires fewer assumptions. On the other hand, it would be surprising (and revealing in terms of models for the core; see Daly & Marscher 1988) if the disturbance in velocity and/or energy density that creates the shock did not also affect the core (see, e.g., the hydrodynamical simulations of Gómez et al. 1997).

We calculate the distance from the core where a new component reaches its maximum luminosity:

\[ l = \frac{\Delta t_{\text{obs}}}{\Gamma(1 + \frac{z}{c})}, \]

where \( l \) denotes the distance from the core along the jet, \( \Gamma \) is the bulk Lorentz factor, \( D \) is the Doppler factor, \( \Delta t_{\text{obs}} \) is the observed timescale for a knot to reach maximum luminosity, \( z \) is the redshift, \( \beta = v/c \approx 1 \), \( v \) is the bulk velocity of the component and \( c \) is the velocity of light. We derive the Doppler factor \( D \) from the variability data using the procedure described in Lähteenmäki & Valtaoja (1999). We then determine the Lorentz factor \( \Gamma \) from this value of \( D \) together with the apparent superluminal speed of the associated VLBI component (Jorstad et al. 2001a). The rise time \( \Delta t_{\text{obs}} \) was determined as the difference between the epoch of the TFD flare peak and the extrapolated zero epoch of the associated VLBI component,
hence $\Delta t_{\text{obs}} = t_{\text{max,TFD}} - t_{0,\text{VLBI}}$. The mean observed timescale for shock evolution $\Delta t_{\text{obs}}$ is $140_{-115}^{+135}$ days, which gives an average value for $l$ of $6 \pm 3$ pc along the jet at 43/37 GHz for our sample. A timescale of 140 days corresponds on average to a movement of $\approx 0.11$ mas in our sample (the average proper motion for components near the core is $\mu \approx 0.3$ mas yr$^{-1}$), so the components usually reach their maximum brightness well within the beamsizes of the VLBA.

5. Conclusions

Our comparisons between VLBI and TFD data show that there is a clear connection between TFD outbursts of blazars and structural changes in their jets. For every new VLBI component emerging into the jet, there is a coincident TFD flare, and also the extrapolated zero epoch of the VLBI component agrees well with the beginning of the TFD flare. At a later stage, there is a clear correspondence between the flux density of the VLBI component and that of the modelled exponential TFD flare. We conclude that one can – for most events – use the beginning of a major TFD flare as an indicator of the time of ejection of a VLBI component. The most pronounced TFD outbursts seem to occur within the innermost few tenths of a milliarcsecond of the core, which is comparable to the maximum resolution of present-day VLBI. The flares are associated with the production of new superluminal components, commonly interpreted as shocks. However, with the present observations, we cannot determine definitively how much of the flare is contained in the shock and how much is due to changes in the flux of the core itself as it reacts to the disturbance that creates the shock wave.

Due to insufficient resolution even at 43 GHz, we usually see only “core flares” in our sources. An example of the effects of limited resolution is BLO 1749+096. Despite violent variability in TFD, this object is virtually unresolved even with a beamsize of 0.6 mas along the jet in global 22 GHz VLBI observations (Wiik et al. 2001). Only global 86 GHz VLBI, with a 0.22 mas beam, has been able to resolve the jet (Lobanov et al. 2000).
Fig. 6. A reconstruction of the “core flares” in quasar PKS 0528+134. That corresponding to the TFD flare in 1995.5 and component B3 of Jorstad et al. (2001a) can be reconstructed easily (dotted line), but for the flare in 1995.9 (corresponding to component B4 of Jorstad et al. 2001a) we do not have enough VLBI observations to determine whether its behaviour is similar.

Table 5. A list of VLBI core flares with a new moving VLBI component appearing after the flare. The time span of the core flare is $t_{CF}$, $\Delta S_{CF}$ is the lower limit for the observed amplitude of the core flare, $t_{NC}$ is the time when the new moving component is first detected, $S_{NC}$ is its flux density, and $R$ is its distance from the core at this time.

<table>
<thead>
<tr>
<th>Source</th>
<th>$\nu$ [GHz]</th>
<th>$t_{CF}$</th>
<th>$\Delta S_{CF}$ [Jy]</th>
<th>$t_{NC}$</th>
<th>$S_{NC}$ [Jy]</th>
<th>$R$ [mas]</th>
</tr>
</thead>
<tbody>
<tr>
<td>0219+428</td>
<td>43</td>
<td>1995.47–1996.60</td>
<td>0.45</td>
<td>1996.60</td>
<td>0.22</td>
<td>0.51</td>
</tr>
<tr>
<td>0420−014</td>
<td>43</td>
<td>1995.31–1995.59</td>
<td>1.04</td>
<td>1996.34</td>
<td>0.09</td>
<td>0.21</td>
</tr>
<tr>
<td>0458−020</td>
<td>43</td>
<td>1995.31–1995.47</td>
<td>0.40</td>
<td>1995.47</td>
<td>0.60</td>
<td>0.20</td>
</tr>
<tr>
<td>0528+134</td>
<td>22</td>
<td>1994.65–1995.01</td>
<td>2.30</td>
<td>1995.01</td>
<td>1.69</td>
<td>0.19</td>
</tr>
<tr>
<td>0851+202</td>
<td>43</td>
<td>1995.47–1996.34</td>
<td>1.91</td>
<td>1996.34</td>
<td>0.28</td>
<td>0.36</td>
</tr>
<tr>
<td>1156+295</td>
<td>22</td>
<td>1996.34–1996.60</td>
<td>0.31</td>
<td>1996.60</td>
<td>0.34</td>
<td>0.20</td>
</tr>
<tr>
<td>1222+216</td>
<td>22</td>
<td>1996.90–1997.58</td>
<td>0.40</td>
<td>1997.58</td>
<td>0.09</td>
<td>0.45</td>
</tr>
<tr>
<td>1633+382</td>
<td>22</td>
<td>1994.76–1995.79</td>
<td>1.04</td>
<td>1995.79</td>
<td>0.60</td>
<td>0.13</td>
</tr>
<tr>
<td>2230+114#</td>
<td>43</td>
<td>1996.60–1998.48</td>
<td>5.00</td>
<td>1998.48</td>
<td>2.14</td>
<td>0.11</td>
</tr>
<tr>
<td>2251+158</td>
<td>43</td>
<td>1995.01–1995.31</td>
<td>4.92</td>
<td>1995.31</td>
<td>4.43</td>
<td>0.06</td>
</tr>
</tbody>
</table>

* Includes 1998 VLBA data from Rantakyrö et al. (2002).

The electron energy spectrum. The essential difference between these two models is that, in the jet-parameter model, luminosity changes in the underlying flow cause the observed flux variability, whereas in the shock model the flares represent an evolving shock wave. The core itself can also become brighter if the shock wave disturbs it when passing through the nozzle of the jet. However, this does not change our suggestion that there is a shock associated with every large TFD flare. This has important implications for models of $\gamma$-ray emission from blazars (see, e.g., Valtaoja et al. 2002; Lähteenmäki & Valtaoja 2002; Jorstad et al. 2001b).
Fig. 7. Closure phases from triangles consisting of long (left panels) and short (right panels) baselines from a series of observations of PKS 0420−014. Two core-region models are superimposed: dashed curve: a single, perhaps extended component; solid curve: a point-like core plus a second point-like component. The fourth epoch (1997 July 31) is for comparison only: no component close to the core is expected from the TFD light curves nor indicated by the VLBA data.

Given the possible blending of a new shock component with the core on VLBI images, one has to be very careful when drawing conclusions based on the behaviour of the core. For example, Lobanov & Zensus (1999) have studied the spectral evolution of the jet components in quasar 3C 345. They find two time intervals during which they could not reproduce the observed spectral changes (frequency of the spectral maximum was rising while the flux density at this frequency was constant or even decreasing slightly) of the core component in terms of the relativistic shock model. However, their spectral fits included only a single component in the core region. According to our results, there is likely to be a second, variable component blending with the core that can cause the observed spectral changes. We suggest that it may be possible to explain the behaviour of the core in a way which is consistent with the relativistic shock model.

The results presented in this article indicate that VLBI observations at 90 GHz are needed in order to study new moving
components during the stage when they are still growing, and still greater resolution is required to see the actual formation of the shock. This type of study would benefit greatly from future space interferometry missions such as the proposed VSOP-2 and iARISE missions. In addition to finer resolution, we also need better sampled VLBI component flux curves, which can be obtained from intensive VLBI monitoring.

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References
Valtaoja, E., Savolainen, T., Wiik, K., & Lähteenmäki, A. 2002, PASA, 19, 117

Fig. 8. Closure phases from observations of 0528+134 and 1156+295. Continued from Fig. 7.
We report the discovery of an extremely curved jet in the radio-loud quasar PKS 2136+141. Multifrequency VLBA images show a bending jet making a turnaround of 210° in the sky plane, to our knowledge the largest ever observed change in an astrophysical jet’s P.A. Images taken at six different frequencies, from 2.3 to 43 GHz, reveal a spiral-like trajectory, which is likely a sign of an intrinsic helical geometry. A space-VLBI image, taken with HALCA at 5 GHz and having comparable resolution to our ground-based 15 GHz data, confirms that the bend is a frequency-independent structure. Eight years of VLBA monitoring data at 15 GHz show knots in the jet clearly deviating from ballistic motion, which suggests that the bending may be caused by a growing helical Kelvin-Helmholtz normal mode. The jet appearance suggests a helical wave at a frequency well below the “resonant” frequency of the jet, indicating that the wave is driven by a periodic perturbation at the base of the jet. We fit the observed structure in the source with a helical twist and find that a simple isothermal model with a constant wave speed and wavelength gives a good fit. The measured apparent velocities indicate some degree of acceleration along the jet, which, together with an observed change in the jet’s apparent half-opening angle, allows estimation of the changes in the angle between the local jet direction and our line of sight. We suggest that the jet in PKS 2136+141 is distorted by a helical Kelvin-Helmholtz normal mode externally driven into the jet (e.g., by precession) and that our line of sight falls within the opening angle of the helix cone.

Subject headings: galaxies: jets — quasars: individual (PKS 2136+141) — radio continuum: galaxies

Online material: color figure

1. INTRODUCTION

A significant fraction of extragalactic jets show some degree of bending—from slightly curved jets up to a complete turnaround of almost 180°. Recently, in their large study of jet kinematics of radio-loud active galactic nuclei, Kellermann et al. (2004) measured vector velocities for 60 bright jet features (also called components). They found that approximately one-third of these components show a significant nonradial motion; i.e., the direction of their velocity vector differs by at least 3° from the mean structural position angle of the jet. If these observed velocities trace the underlying jet flow, their result indicates that bends in jet direction are very common.

For core-dominated radio sources with high optical polarization, there is a well-known bimodal distribution of the angles between jets on parsec and kiloparsec scales, with a main peak of misalignment angles around 0° and a secondary peak around 90° (Pearson & Readhead 1988; Lister et al. 2001). However, a large-angle misalignment exceeding 120° is rare (Wilkinson et al. 1986; Tingay et al. 1998; Lister et al. 2001). Up to today, the largest observed ∆P.A. is 177° in the gamma-ray blazar PKS 1510−089, which shows a jet bending almost directly across our line of sight (Homan et al. 2002). Since core-dominated radio sources have jets oriented close to our line of sight, all intrinsic variations in the jet trajectories are exaggerated in projection—often to a large degree. This implies that rather small intrinsic bends can manifest themselves as large-angle misalignments between the jet axes observed on parsec and kiloparsec scales, or as high as ~90° turns in the VLBI images.

Observations of relativistic jets on parsec scales provide evidence that active galactic nucleus (AGN) jets can exhibit “wigglng” structures (e.g., 4C 73.18 [Roos et al. 1993]; 3C 345 [Zensus et al. 1995]; 3C 273 [Lobanov & Zensus 2001]; 3C 120 [Hardee et al. 2005]) reminiscent of helically twisted patterns. It has been proposed that Doppler boosting, together with parsec-scale jets traveling in helical paths, could explain the excess of sources showing 90° misalignment angle between parsec- and kiloparsec-scale jets without invoking an uncomfortable 90° intrinsic curvature (Conway & Murphy 1993). As the number of sources showing apparently helical structures has grown, the helical jet models have become increasingly popular also as an explanation for the (quasi-) periodic flux variations in AGNs (Abraham & Romero 1999; Ostorero et al. 2004). However, the mechanism producing an apparently helical shape of the jet is unclear—as are the explanations for more modest observed bends.
The “corkscrew” structure of the jet in the well-known Galactic source SS 433 is successfully explained by ballistic motion of material ejected from a precessing jet nozzle (Stirling et al. 2002), and a similar model has also been suggested for several extragalactic jets showing “wiggling” (e.g., 3C 273 [Abraham & Romero 1999]; BL Lac [Stirling et al. 2003]; OJ 287 [Tateyama & Kingham 2004]). In these models, the jet precession is either due to the Lense-Thirring effect in a case of misalignment between the angular momenta of the accretion disk and a Kerr black hole (see Caproni et al. 2004 and references therein) or due to a binary black hole system where a secondary black hole tidally induces the precession.

Contrary to the above-mentioned cases, Lister et al. (2003) report that although the powerful radio source 4C +12.50 exhibits a jet ridgeline highly reminiscent of that in SS 433, it is most likely due to streaming instead of ballistic motion. Streaming helical motion arises naturally from spatial stability analysis of relativistic jets, since the jets are unstable against growing Kelvin-Helmholz (K-H) normal modes (Hardee 1987). Provided there is a suitable perturbation mechanism present in the inner part of the jet, the distortion waves propagating down the jet can displace the whole jet (helical fundamental mode) or produce helically twisted patterns on the jet surface (fluting modes). If the jet carries a large-scale electric current (so-called Poynting flux–dominated jets), it is in addition unstable against magnetic kink instability, which could also produce observed “wiggling” structures (Nakamura & Meier 2004).

Not all jets with observed bends exhibit wiggling structure, and many of the observed changes in the jet direction can be explained without invoking helical motions. Proposed explanations for curving jets include ram pressure due to winds in the intracluster medium, a density gradient in a transition to the intergalactic medium, and deflections by massive clouds in the interstellar medium. Most likely, different mechanisms work in different sources. It would be valuable to be able to reliably identify the reason for bending in individual sources, since the observed properties of the bend—correctly interpreted—can constrain several physical parameters of the jet and the external medium (see, e.g., Hardee [2003] for the case of K-H instabilities).

In this paper, we present Very Long Baseline Array (VLBA) images from a dedicated multifrequency observation and from the VLBA 2 cm Survey4 (Kellermann et al. 1998) showing that PKS 2136+141 (OX 161), a radio-loud quasar at a moderately high redshift of 2.427, has a parsec-scale jet, which appears to bend over 180° on the plane of the sky, being—to our knowledge—the largest ever observed change in the position angle of an astrophysical jet (other sources showing very pronounced changes in the jet direction include, e.g., PKS 1510–089 [Homan et al. 2002], 1803+784 [Britzen et al. 1999], and NRAO 150 [Agudo et al. 2006]). In Very Large Array images, PKS 2136+141 is a compact source, showing no extended emission on arcsecond scales (Murphy et al. 1993). Both 5 GHz (Fomalont et al. 2000) and 15 GHz (Kellermann et al. 1998) VLBA observations reveal a core-dominated source with a short, slightly bending jet.

Originally, Tornikoski et al. (2001) identified PKS 2136+141 as a candidate gigahertz-peaked—spectrum source (GPS) having a slightly inverted spectrum up to 8–10 GHz in the intermediate-to-quiescent state and a clearly inverted spectrum (α > 0.5) during outbursts. The high turnover frequency reported in Tornikoski et al. (2001) even puts PKS 2136+141 in the class of high-frequency peakers (HFPs). Although Torniainen et al. (2005) have recently classified the source as a flat-spectrum radio source having a convex spectrum only during outbursts, the simultaneous continuum spectra from RATAN-600 (Kovalev et al. 1999; S. Trushkin 2005, private communication) do show a convex shape also in the intermediate-to-quiescent state, albeit with a slightly lower peak frequency. The source is variable at radio frequencies, showing a factor of ∼3 variation in centimeter-wavelength flux curves with a characteristic timescale of ∼5–6 yr (see Fig. 1). The last strong outburst started around 1998 and peaked in late 2002 and in early 2004 at 14.5 and 8 GHz, respectively. This indicates that both observing programs, our multifrequency observations in 2001 and 15 GHz VLBA monitoring during 1995–2004, caught the source during a major flare.

The paper is organized as follows: the multifrequency VLBA data demonstrating the 210° bend of the jet, together with a space-VLBI observation from the HALCA satellite (Highly Advanced Laboratory for Communications and Astronomy), are presented in §2. In §3, we present a kinematic analysis, derived from over 8 yr of the VLBA 2 cm Survey monitoring data at 15 GHz, indicating nonballistic motion of the jet components. In addition, changes in βapp and in the apparent half-opening angle of the jet are investigated. In §4, possible reasons for observed bending are discussed and a helical streaming model explaining the observed structure is presented. Conclusions are summarized in §5.

Throughout the paper we use a contemporary cosmology with H0 = 71 km s⁻¹ Mpc⁻¹, ΩM = 0.27, and ΩΛ = 0.73. For this cosmology and a redshift of 2.427, an angular distance of 1 mas transforms to 8.2 pc and a proper motion of 0.1 mas yr⁻¹ to an apparent speed of 9.2 c. We choose the positive spectral index convention, Sν ∝ ν⁻α.

2. MULTIFREQUENCY OBSERVATIONS

In 2001 May we made multifrequency polarimetric VLBI observations of four HFP quasars, including PKS 2136+141, using the VLBA. Observations were split into a high-frequency part (15, 22, and 43 GHz), which was observed on May 12, and a
low-frequency part (2.3, 5, and 8.4 GHz) observed on May 14. Dual polarization was recorded at all frequencies.

2.1. Reduction of the VLBA Data

The data were correlated with the VLBA correlator in Socorro and were postprocessed at Tuorla Observatory using the NRAO’s Astronomical Image Processing System, AIPS (Bridle & Greisen 1994, Greisen 1988) and the Caltech DIFMAP package (Shepherd 1997). Standard methods for VLBI data reduction and imaging were used. An a priori amplitude calibration was performed using measured system temperatures and gain curves. For the high-frequency data (15–43 GHz), a correction for atmospheric opacity was applied. After the removal of a parallactic angle phase, a single-band delay and phase offsets were calculated manually by fringe-fitting a short scan of data of a bright source. We did manual phase calibration instead of using pulse-cal tones because there were unexpected jumps in the phases of the pulse-cal tones during the observations. Global fringe-fitting was performed, and the delay difference between right- and left-hand systems was removed (for the purpose of future polarization studies). Bandpass corrections were determined and applied before averaging across the channels, after which the data were imported into DIFMAP.

In DIFMAP the data were first phase self-calibrated using a point-source model and then averaged in time. We performed data editing in a station-based manner and ran several iterations of CLEAN and phase self-calibration in Stokes I. After a reasonable fit to the closure phases was obtained, we also performed amplitude self-calibration, first with a solution interval corresponding to the whole observation length. The solution interval was gradually shortened as the model improved by further cleaning. Final images were produced with the Perl library FITSPLOT.

We have checked the absolute flux calibration by comparing the extrapolated zero baseline flux density of our compact calibrator source 1749+096 at 5, 8.4, and 15 GHz to the single-dish measurements made at the University of Michigan Radio Astronomy Observatory (UMRAO) and at 22 and 43 GHz to the fluxes from Metsähovi Radio Observatory’s quasar monitoring program at 22 and 37 GHz, respectively (Terasranta et al. 2004). The flux densities agree to 5% at 8.4, 15, 22, and 37/43 GHz and to 8% at 5 GHz, which is better than the expected nominal accuracy of 10% for the a priori amplitude calibration. Being unable to make a flux check for the 2.3 GHz data, we conservatively estimate it to have an absolute flux calibration accurate to 10%.

In order to estimate the parameters of the emission regions in the jet, we model-fitted to the self-calibrated \((u, v)\) data in DIFMAP. The data were fitted with a combination of elliptical and circular Gaussian components, and we sought to obtain the best possible fit to the visibilities and to the closure phases. Several starting points were tried in order to avoid a local minimum fit. We note that since the source structure is complex, the models

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Fig. 2.—Total intensity VLBA observations of PKS 2136+141 from 2001 May. The compiled figure shows images of the source at all six observed frequencies (2.3, 5, 8.4, 15, 22, and 43 GHz). The letter “C” in the top left panel marks the location of the core. The size and orientation of the beam are shown in the lower left corner of each image. Peak intensities and contour levels are given in Table 1.
are not unique but rather show one consistent parameterization of the data. Based on the experiences in error estimation reported by several groups, we assume uncertainties in component flux \(\sim 5\%\), in position \(\sim \frac{1}{2}\) of the beam size (or of the component size if it is larger than the beam), and in size \(\sim 10\%\) (see, e.g., Jorstad et al. [2005] and Savolainen et al. [2006] for recent discussions on the model fitting errors). Although Jorstad et al. (2005) use larger positional uncertainties for weak knots having flux densities below 50 mJy, we also use \(\sim \frac{1}{2}\) of the beam size (or of the component size) for these components. Bigger uncertainties would result in such a large ratio of the individual errors to the scatter of the component positions (about the best-fit polynomial describing the component motion) that it would be statistically unlikely (see § 3.1.2). A detailed description of the polarization data reduction and imaging in Stokes \(Q\) and \(U\), together with polarization images, will appear in T. Savolainen et al. (2006, in preparation).

2.2. Reduction of the HALCA data

The 5 GHz space-VLBI observations of PKS 2136+141 were carried out as a part of the VLBI Space Observatory Programme (VSOP) survey program (Hirabayashi 2000; R. Dodson et al. 2006, in preparation) on 1998 May 28. In addition to the \textit{HALCA} satellite, the array consisted of Arecibo, Haardesthoek, and the Green Bank (140 foot) telescopes. The data were correlated using the Penticton correlator and reduced using the same standard methods as for the ground-based images. Due to the small aperture of the \textit{HALCA} satellite, its weight was increased to 10 to persuade the fitting algorithm to take better account of the long space baselines.

The dynamic range of the resulting image is rather small. This is due to a component that had a strong effect only on a single baseline between Arecibo and Green Bank. Because the \((u, v)\) coverage is very limited, all attempts to model this component diverged. By manually changing the model to follow the visibility amplitudes in this scan, we estimate that the position of this component is about 7 mas at a P.A. of 100\(^\circ\). Because this component could not be formally included in the model, the residuals are rather strong and thus the imaging noise is high.

2.3. Source Structure from the Multifrequency Data

Figure 2 displays images of PKS 2136+141 at all six observed frequencies. In the images at 15, 22, and 43 GHz, uniformly weighted \((u, v)\)-grids are employed in order to achieve the best possible resolution, whereas normal weighting is used in the low-frequency maps to highlight diffuse, low surface brightness emission. The restoring beam sizes, peak intensities, off-source rms noise, and contour levels of the images are given in Table 1.

The multifrequency images strikingly reveal a jet that gradually bends 210\(^\circ\) with its structural P.A. turning clockwise from \(-27\^\circ\) at 43 GHz to +123\(^\circ\) at 2.3 GHz. We identify the bright and the most compact model component lying in the southeast end of the jet in the 15–43 GHz images as the core and mark it with a letter “C”. The identification is confirmed by a self-absorbed spectrum of the component, as is shown below. In the range of about 0.4–1.0 mas from the core, the jet turns over 90\(^\circ\), which is visible in 15–43 GHz images, and in the image taken at 8.4 GHz, the jet direction continues to turn clockwise \(\sim 50\^\circ\) at about 3.5 mas from the core. There is also evident bending in the 5 and 8.4 GHz images: a curve of \(\sim 70\^\circ\) takes place at about 6 mas south of the core. It is not totally clear whether the trajectory of the jet is composed of a few distinct bends or whether it is a continuous helix. However, the gradual clockwise turn and the apparent spiral-like appearance of the jet in the multifrequency images are highly reminiscent of a helical trajectory. In addition, it seems unlikely that the jet goes through at least three consecutive deflections with all of them having the same sense of rotation in the plane of the sky.

The 5 GHz space-VLBI image of PKS 2136+141 from 1998 May 28, i.e., 3 yr before the multifrequency VLBA observations, shows a rather compact core-jet structure with an extended emission to the northwest (Fig. 3). The jet then takes a sharp 90\(^\circ\) bend to the southwest within about 1 mas from the core. This 5 GHz image shows a very similar curved structure near the core that can be seen in our ground-based 15 GHz image with matching resolution. Hence, the observed large-angle bending between the images taken at different frequencies cannot be attributed to frequency-dependent opacity effects.

PKS 2136+141 is rather compact at all frequencies, with a maximum jet extent of approximately 15 mas, corresponding to \(\sim 120\) pc at the source distance. The two brightest model components are the core and a newly ejected component I5 (see § 3 for component identifications), which is located at \(-0.25\) mas from the core—near the beginning of the first strong bend in the jet. Figure 4 shows the 8–43 GHz spectra of these two components. At 2.3 and 5 GHz the angular resolution is too poor to separate the core from I5, and hence, corresponding flux values are omitted in the figure. The core has a synchrotron peak frequency of \(\sim 20\) GHz, while the newly ejected component I5 shows an optically thin synchrotron spectrum with a spectral index \(\alpha = -0.6\) and a peak frequency below 8 GHz. The component I5 is brighter than the core at every frequency except at 22 GHz, where almost equal flux densities are measured. An optically thick spectrum of the core at frequencies below 22 GHz indicates that I5 is also brighter than the core at 2.3 and 5 GHz, although the spectra

<table>
<thead>
<tr>
<th>Frequency (GHz)</th>
<th>(\theta_{b,\text{maj}}) (mas)</th>
<th>(\theta_{b,\text{min}}) (mas)</th>
<th>P.A. (deg)</th>
<th>rms Noise (mJy beam(^{-1}))</th>
<th>Peak Intensity (mJy beam(^{-1}))</th>
<th>Contour (c_0)* (mJy beam(^{-1}))</th>
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<tr>
<td>2..................</td>
<td>6.12</td>
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</tr>
<tr>
<td>5(^b)...........</td>
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</tr>
<tr>
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</tr>
<tr>
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<td>-16.1</td>
<td>1.5</td>
<td>512</td>
<td>4.5</td>
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</table>

* Contour levels are represented by a geometric series \(c_0(1, \ldots, 2^n)\), where \(c_0\) is the lowest contour level indicated in the table (3 times rms noise).

\(^b\) HALCA image.
cannot be measured at those frequencies. The self-absorbed spectrum shown in Figure 4 confirms that we have correctly identified the core.

3. VLBA 2 cm SURVEY AND MOJAVE DATA

Seven observations of PKS 2136+141 have been performed as part of the VLBA 2 cm Survey (Kellermann et al. 1998; Zensus et al. 2002) between 1995 and 2004, of which the last one (epoch 2004.27)\(^6\) was part of the follow-up program MOJAVE (Monitoring of Jets in Active Galactic Nuclei with VLBA Experiments; Lister & Homan 2005). Details on the survey observing strategy, the observing procedures, and the data reduction can be found in the mentioned publications.

CLEAN VLBI images of PKS 2136+141 have been produced by applying standard self-calibration procedures with DIFMAP for all seven epochs. The calibrated visibility data were fitted in the ($u$, $v$)-domain with two-dimensional elliptical Gaussian components (see Fig. 5 and Table 2). Because of the complicated source structure, no adequate model representation of the source could be established with a smaller number of components than shown in Figure 5. Moreover, the components are found to be located close by along the curved inner part of the jet, in many cases separated from each other by considerably less than one beam size. Special care had to be taken to model this complicated structure consistently for all epochs.

3.1. Source Kinematics

A crucial question raised by the apparent helical shape of the jet in Figure 2 is whether it represents streaming motion or whether the helix is due to a precessing jet nozzle ejecting material that moves along ballistic trajectories. We have plotted positions of all model-fit components from the multifrequency data and from 8 yr of the VLBA 2 cm Survey monitoring data in Figure 6. The open circles, which correspond to the multiepoch data from the VLBA 2 cm Survey, form a dense and strongly curved region near the base of the jet, a ∼2 mas long continuous and slightly curved section 3.5 mas southwest of the core, and a few small isolated groups. This subtle finding alone, without invoking any component identification scenarios, demonstrates that ballistic-motion models are unlikely to yield a meaningful representation of the jet kinematics in PKS 2136+141. Brightening at certain points of the jet can be either due to an increased Doppler factor (if a section of the jet bends toward our line of sight) or due to an impulsive particle acceleration in a standing shock wave (e.g., forming in the bend). There seems to be a zone of avoidance in the 15 GHz data between the base of the jet and the large western group, further supporting this idea.

We have analyzed the jet kinematics using source models derived from the VLBA 2 cm Survey monitoring data. In a complicated source like PKS 2136+141 it is often difficult to identify components across epochs with confidence. We have based our identifications on the most consistent trajectories and on the flux density evolution of the components. We restrict our analysis to a full kinematical model for the inner 2 mas of the jet, since beyond that distance a fully self-consistent model could not be established due to the lower surface brightness and high complexity in the outer region.

3.1.1. Component Trajectories and Flux Evolution

Figure 7 shows the trajectories of the six compact components (I1–I6) that we have identified in the inner 2 mas of the jet. The components travel toward the southwest and their trajectories do not extrapolate back to the core, implying that the components travel along a curved path and that the jet is nonballistic. The components fade below the detection limit within ∼1.6 mas from the core, and we cannot follow them further.

We present the flux density evolution of the six inner-jet components in Figure 8. The overall picture is that the fluxes decrease as the components move forward along their path. They reach about 5–15 mJy by the time they have traveled to a distance of ∼1.5 mas from the core, after which they are not seen anymore. The knot I3 is the brightest component of the source at the first two epochs, and its flux density evolution matches well with the strong 1993 flare visible in the 14.5 GHz UMRAO flux density curve (see Fig. 1). The ejections of components I1 and I2 (see Table 4) take place during the rising phase of the 1993 outburst, and component I3 is ejected just before the outburst reaches its maximum.
The core has a steadily rising flux density until 2001.20, when the flux suddenly drops and two new components, I4 and I5, appear. At the next two epochs, the core flux increases again, and at the last epoch (2004.27) there is a drop accompanied by the appearance of a new component, I6. The brightening of the core during our monitoring and the ejections of components I4, I5, and I6 correspond to a strong total flux density flare peaking in late 2002 at 14.5 GHz (again, see the total flux density curve in Fig. 1).

The above-mentioned flux density evolution of the components is self-consistent and is well in accordance with the general behavior of flat-spectrum radio quasars during strong total flux density flares (Savolainen et al. 2002); i.e., a compact VLBI core is mostly responsible for the rising part of the observed flares in single-dish flux curves, and a new component appears in the jet during or after the flare peaks accompanied by a simultaneous decrease in the core flux density. The fact that the overall flux density evolution of the identified components in PKS 2136+141 seems to obey the common behavior identified for a number of other sources supports our kinematical model. An intriguing detail is the consecutive ejection of not one but three new components in connection with a single, strong total flux density flare. This may indicate that there is some substructure in the flares; i.e., the outbursts in 1993 and 2002 could be composed of smaller flares. On the other hand, it could also mean that a single strong event in the total flux density curve is able to produce complicated structural changes in the jet, e.g., forward and reverse shocks could both be visible, or there could be trailing shocks forming in the wake of the main perturbation as theoretically predicted by Agudo et al. (2001) and recently observed in several objects, most prominently in 3C 111 (Jorstad et al. 2005; Kadler 2005; M. Kadler et al. 2006, in preparation).

![VLBA images of PKS 2136+141 at 15 GHz. The figure includes observations from the VLBA 2 cm Survey, from the MOJAVE Survey, and a 15 GHz image from our multifrequency data set observed on 2001.36. In all images, a uniformly weighted (u,v)-grid is used. The Gaussian components fitted to the visibility data are shown as ellipses overlaid on each image. The size and orientation of the beam are shown in the lower left corner of each image. Peak intensities and contour levels are given in Table 2.](image)

**Fig. 5.**

**TABLE 2**

<table>
<thead>
<tr>
<th>Epoch (yr)</th>
<th>$\Theta_{b,\text{maj}}$ (mas)</th>
<th>$\Theta_{b,\text{min}}$ (mas)</th>
<th>P.A. (deg)</th>
<th>rms Noise (mJy beam$^{-1}$)</th>
<th>Peak Intensity (mJy beam$^{-1}$)</th>
<th>Contour $c_0^*$ (mJy beam$^{-1}$)</th>
</tr>
</thead>
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<td>1.01</td>
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<td>11.5</td>
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<td>1254</td>
<td>2.7</td>
</tr>
<tr>
<td>1996.82</td>
<td>0.92</td>
<td>0.41</td>
<td>-1.6</td>
<td>0.8</td>
<td>633</td>
<td>2.4</td>
</tr>
<tr>
<td>1998.18</td>
<td>1.31</td>
<td>0.61</td>
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<td>1.4</td>
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<td>4.2</td>
</tr>
<tr>
<td>1999.55</td>
<td>1.45</td>
<td>0.49</td>
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<td>1429</td>
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</tr>
<tr>
<td>2001.20</td>
<td>1.14</td>
<td>0.52</td>
<td>-3.1</td>
<td>0.4</td>
<td>1153</td>
<td>1.2</td>
</tr>
<tr>
<td>2001.36</td>
<td>0.90</td>
<td>0.50</td>
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<td>0.5</td>
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</tr>
<tr>
<td>2002.90</td>
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<td>2004.27</td>
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<td>0.3</td>
<td>1745</td>
<td>0.9</td>
</tr>
</tbody>
</table>

$^*$ Contour levels are represented by a geometric series $c_0(1, \ldots, 2^n)$, where $c_0$ is the lowest contour level indicated in the table (3 times rms noise).
3.1.2. Apparent Velocities and Acceleration

Identification of the model fit components across the epochs allows us to estimate the velocities of the components and to search for possible acceleration or deceleration. Since the components follow a highly curved path, we choose to measure the distance traveled along the jet by using an average ridgeline, which is shown in Figure 6 (dashed line). The ridgeline is determined by fitting a smooth function to the positions of model components within the innermost 2 \times 1.5 mas, and it also displays an approximate ridgeline of the jet.

Figure 9 displays the motion of the components along the jet. To measure the component velocities and to estimate the ejection epochs, we have made standard linear least-squares fits to the motions:

\[ l(t_i) = a_0 + a_1(t_i - t_{\text{mid}}), \]

where \( l \) is the traveled distance along the jet, \( t_i \) is the epoch of observation \( (i = 1, \ldots, N) \), and \( t_{\text{mid}} = (t_1 + t_N)/2 \). In order to search for a possible acceleration or a deceleration, we have also fitted components I2 and I3, which are observed over most epochs, with second-order polynomials:

\[ l(t_i) = a_0 + a_1(t_i - t_{\text{mid}}) + a_2(t_i - t_{\text{mid}})^2. \]

To assess the goodness of the fits, we have calculated a relative \( \chi^2 \) value for each fit. The \( \chi^2 \) values are relative, not absolute,
since we do not strictly know the real 1 σ measurement errors of the component positions but have instead used $1/2$ of the beam size projected onto the ridgeline. Although the $1/2$ of the beam size estimate for the errors is rather well justified in the literature, it is likely to be conservative. Thus, the tabulated values of the $\chi^2$ distribution do not provide a good statistical test of the goodness of the fits in our case. However, we assume that the beam sizes give a good estimate of the relative errors between the epochs, and hence, only one scaling factor (which is close to unity) for the positional errors (and for the $\chi^2$ values) remains unknown. With this assumption, the relative $\chi^2$ values can be used to compare the goodness of the fits against each other. The parameters of the fitted polynomials are given in Table 3, together with relative reduced $\chi^2$ values and the number of degrees of freedom, $\nu$, for each fit.

![Component motion along the jet ridgeline showing I1 (open circles), I2 (filled squares), I4 (crosses), I5 (filled triangles), and I6 (open square). The solid lines represent linear fits to the data, and the dashed line represents a second-order fit.](image)

We have gathered the average angular velocities along the jet, $\langle \dot{\mu} \rangle = \alpha_i$, the average apparent speeds, $\langle \beta_{\text{app}} \rangle$, and the epochs of zero separation, $T_0$, for each component in Table 4. For the second-order fits, the angular acceleration along the jet, $\dot{\mu} = 2\dot{\alpha}$, is also reported. The proper motions range from 0.09 to 0.27 mas yr$^{-1}$, corresponding to the apparent superluminal velocities of 8.7c–25.1c.

While component I2 is well fitted by a straight line, the linear proper motion model does not seem to adequately represent the motion of component I3, as can be seen from Figure 9. On the other hand, a second-order polynomial gives an acceptable fit to I3, and its relative $\chi^2$ value divided by the number of degrees of freedom is smaller by a factor of 3.5 than the value for the first-order polynomial. However, the small number of data points and the uncertainty about the real (1 σ) errors in the component positions make it difficult to give the statistical significance of the deviation from the constant-speed model. We have used two approaches to assess the significance. The most robust and most straightforward test is to use a statistic $F_{\alpha,\nu}$, which has an $F$-distribution with $(\nu_1, \nu_2)$ degrees of freedom and which is independent of the unknown scaling factor in the positional errors. This statistic can be used to test whether the squared residuals of the linear model are significantly larger than those of the accelerating model. As already mentioned, for the first- and second-order polynomials in the case of I3, $F_{5,5} = 3.5$, which implies a difference in the residuals only at the confidence level of $\alpha = 0.10$; i.e., according to the $F$-test, the statistical significance of the difference in the residuals is only marginal. Another approach to this problem is to try to estimate the real 1 σ uncertainties of the component positions by applying a method described by Homan et al. (2001). They estimated the uncertainties of the fitted parameters of their proper-motion models by using the variance about the best-fit model, and as a by-product they obtained an upper bound estimate of the component position uncertainty. Since the relative $\chi^2$ values divided by the degrees of freedom are significantly less than unity for components I2 and I5 (the situation of component I4 is discussed below), we suspect that the real 1 σ positional uncertainties are in fact smaller than $1/2$ of the beam size. If the beam sizes give a good estimate of the relative errors between the epochs, we can try to estimate the uniform scaling factor for the positional uncertainties by requiring that $\chi^2 \approx \nu$ for the linear proper motion models of I2, I4, and I5. The sum $\sum_i (\chi^2_i/k - \nu_i)^2$, ($i = 12, 14, 15$) is minimized for the positional error scaling factor $k = 0.77$; i.e., the 1 σ positional uncertainties are $0.77 \times \chi^2/k$ of the beam size $=0.08$–0.15 mas. If these (1 σ) uncertainties are assumed, the linear fit to the motion of component I3 has $\chi^2 = 15.2$ with 6 degrees of freedom, implying that the speed of I3 is not constant at a significance level of $\alpha = 0.025$.

Component I4 seems to show motion that indicates acceleration similar to I3, but since there are only four data points for I4, we have not fitted it with a second-order polynomial, which would leave only 1 degree of freedom for the fit. By looking at Figure 9, one can see that I4 should already have been present in the jet at epoch 1999.55, but it cannot be unambiguously inserted into the source model; i.e., adding one more component to the model does not improve the fit but rather makes the final fit highly dependent on the chosen initial model parameters. One possible explanation is that the proper motion for I4 is faster than...
indicated by the best-fit line in Figure 9, and it is not actually present in the jet at 1999.55. A second possibility is that I4 is a trailing shock released in the wake of the primary superluminal component I3 instead of being ejected from the core (Agudo et al. 2001; Gómez et al. 2001). This alternative is not a very plausible one, since according to Agudo et al. (2001) trailing shocks should have a significantly slower speed than the main perturbation. A third possible explanation for the nondetection of I4 at 1999.55 is that either I4 and the core or I4 and I3 are so close to one another at 1999.55 that they do not appear as separate components in our data.

As discussed above, we cannot firmly establish that component I3 accelerates along the jet ridgeline, although it seems probable. However, in addition, the other components (apart from I4, which itself might show acceleration) appear to be systematically faster the farther away from the core they lie. We have averaged the velocities of the individual components over traveled distance bins of 1.0 and 0.5 mas. In averaging, we have taken into account the standard errors given in Table 4 and the fact that for a given component in a given bin, the weight has to be decreased if the first or last observation of the component falls into that bin. This is because we naturally do not have any knowledge about the component speed before or after our observations. For component I3 we have used the quadratic fit instead of the linear, since according to Agudo et al. (2001) trailing shocks should have a significantly slower speed than the main perturbation. A linear least-squares fit to the radial distances of I3 in the first bin to 1.0–1.5 mas. For the division in two bins, the change in the average component speed before or after our observations. For component I3 we have used the quadratic fit instead of the linear, since according to Agudo et al. (2001) trailing shocks should have a significantly slower speed than the main perturbation.

Kellermann et al. (2004) parameterized the 1995–2001 VLBA 2 cm Survey data of PKS 2136+141 by fitting Gaussian components to the brightest features of each epoch in the image plane and derived a speed of $\beta_{\text{app}} = 1.8 \pm 1.4$, which is much lower than our estimates. Basically, their approach traces component I3 over the time when its proper motion was slow (see Fig. 9), and due to the curved path of I3, the speed measured from the changes in its radial distance from the core is also lower than its true angular velocity. A linear least-squares fit to the radial distances of I3 in our data over the epochs 1995–2001 yields $\beta_{\text{app}} = 2.6$, which agrees with the speed given in Kellermann et al. (2004) within the errors.

### 3.2. Apparent Half-Opening Angles and Jet Inclination

The angle between the local jet direction and our line of sight also affects the apparent opening angle of the jet. Assuming that the jet has a conical structure and a filling factor of unity, we have estimated the local value of the apparent half-opening angle for component $i$ as $\Psi_{\text{app}}^i = \tan^{-1}(d_i/l_i)$, where $l_i$ is the distance of component $i$ from the core measured along the jet ridgeline and $d_i = a_i/2 + \Delta r_i$, with $a_i$ being the size of the elliptical component projected on the line perpendicular to the jet and $\Delta r_i$ being the normal distance of the component from the jet ridgeline.

Figure 11 shows the apparent half-opening angle for components 11–16 from 15 GHz data and for unnamed components from 22 and 43 GHz data as a function of traveled distance along the ridgeline. The components with zero axial ratio have been excluded from the figure. Naturally, $\Psi_{\text{app}}$ is very uncertain for a single component, but Figure 11 clearly shows that there is a change in the average $\Psi_{\text{app}}$ between the first ~0.5 mas and 0.8–2.0 mas, and the change seems to correspond to the first large bend in the jet. Within the first 0.5 mas, the average apparent half-opening angle is $\sim 25^\circ$, and farther down the jet it decreases to $\sim 10^\circ$. A natural explanation for this effect is that the viewing angle $\theta$ increases in the first apparent bend, since $\Psi_{\text{app}} \approx \Psi_{\text{in}}/\sin \theta$ if $\theta$ and the intrinsic half-opening angle $\Psi_{\text{in}}$ are both small. There is a caveat, however. Namely, the large value of $\Psi_{\text{app}}$ near the core could be due to a resolution effect: if the positional uncertainties for the components near the core were $\sim \frac{1}{4}$ of the beam size, $\Delta r$ of these components could be $\approx 0.1–0.2$ mas.
at 15 GHz merely due to the scatter, and this would result in \( \langle \Psi_{\text{app}} \rangle = 25^\circ \). However, there are two reasons why we think the observed change in \( \Psi_{\text{app}} \) is a true effect and due to a changing \( \theta \).

First, the model fit components from 22 and 43 GHz data show values for \( \Psi_{\text{app}} \) near the core similar to those of the components from 15 GHz data, although they have a factor of 1.5–2 smaller positional errors. Second, the components near the core are bright, having fluxes comparable to the core, which makes their positional uncertainties smaller than the adopted \( \pm 10' \) of the beam size (see also the previous section for a discussion about the positional uncertainties). According to Jorstad et al. (2005) bright (flux of the knot \( >100 \) rms noise level) and compact (size \( <0.1 \) mas) features in the jet have positional uncertainties \( \sim 0.01 \) mas at 7 mm. Thus, we regard the observed change in \( \Psi_{\text{app}} \) as a genuine effect.

The \( \beta_{\text{app}} \) of the jet seems to increase after \( \sim 0.5 \) mas from the core (see Fig. 10) while the \( \Psi_{\text{app}} \) decreases. This can be understood if the angle between the local jet direction and our line of sight within the first \( \sim 0.5 \) mas is smaller than the angle \( \theta_{\text{SL}} = \sin^{-1}(1/\Gamma) \), which maximizes \( \beta_{\text{app}} \). After \( 0.5 \) mas \( \theta \) increases toward the maximal superluminal angle, increasing \( \beta_{\text{app}} \) and decreasing \( \Psi_{\text{app}} \). Assuming that the largest velocity in Figure 10, \( \beta_{\text{app}} \approx 17c \), is close to the maximum apparent velocity (we do not consider the velocity of 11, 25.1c, very reliable, since it is based on only two data points, and hence, we do not use it in our estimation of the maximum \( \beta_{\text{app}} \)), we can estimate that the jet Lorentz factor \( \Gamma \sim 20 \). Now, if we assume a constant Lorentz factor \( \Gamma = 20 \), we can fit for the values of the viewing angle and the intrinsic jet half-opening angle, giving the observed superluminal speeds and half-opening angles before and after the bend. For \( \langle \beta_{\text{app}} \rangle \approx 7c \), \( \langle \Psi_{\text{app}} \rangle \approx 25^\circ \), within the first \( 0.5 \) mas from the core and \( \langle \beta_{\text{app}} \rangle \approx 17c, \langle \Psi_{\text{app}} \rangle \approx 10^\circ \), after the bend, we obtain the following best-fit values: \( \theta(0–0.5 \) mas) = \( 0.6^\circ \), \( \theta(1.0–2.0 \) mas) = \( 1.5^\circ \), and \( \Psi_{\text{int}} = 0.26^\circ \). According to this result, the jet bends away from our line of sight by \( \approx0.9^\circ \) after the first \( 0.5 \) mas from the core.

Naturally, it is possible that the components also exhibit real acceleration along the jet in addition to the apparent acceleration due to the projection effect. Unfortunately, the observations do not allow us to decide between these cases. Applying Occam’s razor, we consider a constant jet speed in the following discussion.

4. DISCUSSION

4.1. Possible Reasons for Bending

In this section we discuss the mechanisms capable of producing the curved structure of PKS 2136+141. In principle, there are several possible scenarios for the observed bending in relativistic jets, but we can rule out some of them in this particular case on the grounds of our analyzed data.

Projection effects play an important role in this case, making the intrinsic bending angle much less than the observed one. If the angle between the jet and our line of sight is as small as the analysis in the previous section indicates, it is possible that, for example, within \( \sim 0.5–1.0 \) mas from the core where the apparent bend is \( \sim90^\circ \), the intrinsic bending angle is only \( \sim1–2^\circ \). This implies that the formation of internal shocks in the bend, and their possibly destructive effect on the collimation of the jet, is not of great concern here. Mendoza & Longair (2002) calculate an upper limit to the bending angle of a jet in order not to create a shock wave at the end of the curvature. Their result for relativistic jets is \( \sim50^\circ \), which leaves our estimated intrinsic bending angle for PKS 2136+141 well within the limits. Therefore, we do not consider internal shocks formed by bending to restrict the possible explanations for the curvature in this source.

One of the scenarios that we can rule out is a precessing jet in which the components ejected at different times in different directions move ballistically and form an apparently curved locus. Such models have been used to describe oscillating “nozzles” observed in some BL Lac sources (Stirling et al. 2003; Tateyama & Kingham 2004). In the case of PKS 2136+141, we can reject the precessing ballistic jet hypothesis, since the individual components do not follow ballistic trajectories but rather exhibit streaming motion along a curved path (see §3.1.1). However, precession of the jet inlet may still be behind the observed structure if it serves as an initial perturbation driving a helical Kelvin-Helmholtz normal mode (see § 4.2).

Homan et al. (2002) explained the large misalignment between the parsec- and kiloparsec-scale jets of PKS 1510–089 with a scenario in which the jet is bent after it departs the host galaxy, either by a density gradient in the transition region or by ram pressure due to the winds in the intracluster medium. In PKS 2136+141 the bending starts within 0.5 mas from the core, and because \( \theta = 0.6^\circ = 0.01 \) mas for the inner part of the jet (the error range here refers to an uncertainty introduced by errors in \( \langle \beta_{\text{app}} \rangle \); see §§ 3.1.2 and 3.2), the corresponding deprojected distance is smaller than 0.8 kpc. The whole bending visible in Figure 2 takes place within about 15 mas from the core. Given the viewing angles estimated in § 3.2, it is fair to say that the deprojected length of the jet is—probably significantly—less than \( \sim12 \) kpc, and the whole observed bending takes place within that distance from the core. There are no observations of the host galaxy of PKS 2136+141, but the deprojected lengths estimated above can be compared with typical scale lengths of the elliptical hosts of radio-loud quasars (RLQs), which are reported by several groups. In their near-infrared study, Taylor et al. (1996) found low-redshift (z \( \sim0.2 \)) RLQ hosts to have very large half-light scale lengths \( R_{1/2} \), from 14.0 to 106.8 kpc with an average of \( \sim30 \) kpc. Other studies have yielded smaller values: e.g., Floyd et al. (2004) used Hubble Space Telescope WFPC2 data to study hosts of 17 quasars at z \( \sim0.4 \), finding \( R_{1/2} = 10.2 \pm 1.8 \) kpc for RLQs, and Kotilainen et al. (1998) reported \( R_{1/2} = 13 \pm 7 \) kpc for 12 flat-spectrum radio quasars up to z \( =1.0 \) in their near-infrared study. In the framework of hierarchical models of galaxy formation, the hosts of quasars at high redshift are also expected to be more compact than their low-redshift counterparts. Falomo et al. (2005) have recently managed to resolve a quasar host at z \( =2.555 \), and they report an effective radius of 7.5 ± 3 kpc. As the deprojected distance \( \sim0.8 \) kpc indicates, the bending of the jet in PKS 2136+141 starts well within the host galaxy, and the typical scale lengths listed above suggest that most of the curved jet is located within the host, although it is possible that part of it lies in the outskirts of the host galaxy and is susceptible to the density gradient in the transition region. However, the bending clearly starts inside the galaxy and thus requires some other cause.

A collision inside the host galaxy, between the jet and a cloud of interstellar matter, can change the direction of the flow. For example, Homan et al. (2003) have found component C4 in 3C 279 to change its trajectory by \( 26^\circ \) in the plane of the sky, and they suggest this to be due to a collimation event resulting from an interaction of the component with the boundary between the outflow and the interstellar medium. However, the mere fact that the observed \( \Delta P.A. \) is larger than \( 180^\circ \) in our case constrains the collision scenario, since it excludes the situations where the jet is bent by a single deflection, e.g., from a single massive cloud. In principle, the observed structure may be due to several consecutive
4.2. Helical Streaming Model

Relativistic jets are known to be Kelvin-Helmholtz unstable, and they can naturally develop helical distortions if an initial seed perturbation is present at the jet origin. The helical K-H fundamental mode is capable of displacing the entire jet and consequently producing large-scale helical structures where the plasma streams along a curved path (Hardee 1987). The initial perturbations can have a random spectrum and be due to, for example, jet-cloud interaction, or they can originate from a periodic variation in the flow direction of the central engine (precession or orbital motion). These perturbations can trigger pinch, helical, or higher order normal modes propagating down the jet, and the appearance of this structure depends on the original wave frequency and amplitude, as well as subsequent growth or damping of the modes. Since the individual components in PKS 2136+141 do not follow ballistic trajectories but rather seem to stream along a helical path, we consider a helical K-H fundamental mode to be a possible explanation for the curious appearance of this source. For simplicity, in the following discussion we limit ourselves to a purely hydrodynamic case and do not consider the effect of magnetic fields for the growth of K-H modes, nor do we discuss the current-driven instabilities, which can produce helical patterns in Poynting flux-dominated jets (Nakamura & Meier 2004).

The appearance of the jet in PKS 2136+141 already suggests some properties of the wave. The fact that components are observed to follow a nearly stationary helix implies that the wave frequency of the helical twist, \( \omega \), is much below the “resonant” (or maximally unstable) frequency \( \omega' \), which corresponds to the frequency of the fastest growing helical wave (Hardee 1987). Such a low-frequency, long-wavelength helical twist suggests that the wave is driven by a periodic perturbation at the jet base. If the central source induced white-noise–like perturbations, we would expect to see structures corresponding to the fastest growing frequency, i.e., \( \omega' \) (Hardee et al. 1994).

If isotropical jet expansion without gradients in the jet speed or in the ratio of the jet and external medium sound speeds is assumed, the wave speed, \( \beta_{\infty} \) (in units of \( c \)), in the low-frequency limit remains constant as the jet expands (Hardee 2003). Since the intrinsic wavelength, \( \lambda \), for a given \( \omega \) varies proportional to the wave speed, it can also be assumed to remain constant along the jet (as long as \( \omega \ll \omega' \)). Applying this assumption, we have fitted a simple helical twist to the observed positions of the VLBI components. The helical twist is specified in cylindrical coordinates with \( z \) along the axis of helix, by an amplitude \( A \) in the radial direction, and by a phase angle \( \phi \) given by

\[
\phi = 2\pi h \int_{z_1}^{z} \frac{dz}{\lambda(z)} + \phi_1, \tag{3}
\]

where \( h \) is the handedness of the helix (−1 for right-handed and +1 for left-handed) and \( \phi_1 = \phi(z_1) \). With constant \( \lambda \), the integral in equation (3) becomes trivial and \( \phi \propto z \). The amplitude growth is assumed to be conical: \( A = z \tan \psi_k \), where \( \psi_k \) is the opening angle of the helix cone. To describe the orientation of the helix in the sky we use two angles: \( i \) is the angle between the axis of the helix and our line of sight and \( \chi \) is the position angle of the axis of the helix projected on the plane of the sky. There are five parameters in the model altogether, since the handedness can be fixed to \( h = +1 \) by simply looking at the images.

A computer program was written to fit the positions of the VLBI components with different models expected to explain the shape of the jet and the kinematics of individual components. In the program, the cumulative sum of perpendicular projected image-plane distances between a three-dimensional model and the component centroids from the VLBI observations was minimized using a differential evolution (Storn & Price 1997) algorithm. This evolutionary algorithm was chosen because of its good performance with nonlinear and multimodal problems. All the observed components (and all frequencies) were used in a single model fit because the components seem to follow a common path, i.e., the helical structure appears to persist for longer than the component propagation times. In addition to component positions on the plane of the sky, the model was constrained by the viewing angles determined in \( \S \) 3.2. The angle \( \theta \) between the normal to the jet’s cross section and our line of sight was evaluated along the model, and it was required to be compatible with the viewing angles determined in \( \S \) 3.2. The following limits for \( \theta \) were used: at 0.3 mas from the core \( \theta \) is between 0°4 and 0°8, and at 1.5 mas from the core \( \theta \) is between 1°3 and 1°7. If the model exceeded these limits, the cost function in the fitting algorithm was multiplied with a second-order function normalized to the width of the allowed range.

The best-fit trajectory of the model is presented in Figure 12, and Figure 13 shows the angle between the local jet direction and our line of sight, the Doppler factor, and the apparent velocity as a function of distance along the jet, which have been calculated by assuming \( \Gamma = 20 \). As is clear from these figures, the simple helical twist with constant wavelength along the jet gives a good fit to the data. The best-fit parameters of the model are listed in Table 5. In the best-fit model, we are looking straight into the cone of the helix (\( i = 0°3 \)), which has a small half-opening angle of 1°0, meaning that the orientation of the helix is a very lucky coincidence. The fitted helical wavelength of the perturbation is 776 mas, corresponding to 6.4 kpc. This still needs to be
corrected for a combined effect of the geometry and the possibly relativistic wave speed. The true intrinsic helical wavelength is given by $\lambda_{\text{int}} = \lambda_{\text{mod}}(1 - \beta_w \cos \varphi)$. Some constraints on $\beta_w$ can be derived from the fact that we do not see any systematic change in the helical trajectory during 8.7 yr; i.e., the change of the trajectory is less than the components’ positional uncertainty in our observations. This gives an upper limit on the apparent wave speed, $\beta_{\text{app}} < 1.1c$. As we look into the helix cone, the appropriate viewing angle for the wave motion is between $\psi_c - i$ and $\psi_c + i$, which gives $\beta_w < 0.989$ and $\lambda_{\text{int}} > 0.01\lambda_{\text{mod}}$. These limits are purely due to the uncertainty in the positions of the VLBI components ($\sim \frac{1}{2}$ of the beam size), and it is likely that the true wave speed is much slower and the intrinsic wavelength is closer to $\lambda_{\text{mod}}$. Further observations, e.g., within the MOJAVE program, should give a tighter constraint for $\beta_w$. In the low-frequency limit, the wave speed is

$$\beta_w = \frac{\Gamma^2\eta}{1 + \Gamma^2\eta} \beta_j, \tag{4}$$

where $\beta_j = \sqrt{1 - \Gamma^{-2}}$ is the flow velocity and $\eta = (a_j/a_j)^2$, with $a_j$ and $a_j$ being the sound speeds in the jet and in the external medium, respectively (Hardee 2003). Assuming $\Gamma = 20$ and applying the upper limit of $\beta_w$, we get a limit on the ratio of sound speeds: $a_j/a_j < 0.5$.

We would like to have an estimate of the wave frequency of the observed helical twist, since it could possibly tell us about the origin of the periodic perturbation. Unfortunately, we do not know the wave speed and hence cannot calculate the frequency. The limiting case with $\beta_w = 0.989$ considered above yields $\omega \sim 1.4 \times 10^{-10}$ Hz, corresponding to a period of $\sim 200$ yr, which is likely much below the actual period. However, some example values can be calculated for different combinations of $a_j$ and $a_j$. Let us first assume an ultrarelativistic jet with $a_j = c/\sqrt{3}$ and $\Gamma = 20$. Now, using equation (4) we can calculate $\beta_j$ and consequently $\omega$ for different values of $a_j$. For example, Conway & Wrobel (1995) use $a_j \approx 400$ km s$^{-1}$ in their study of Mrk 501 on the basis of X-ray observations and theoretical modeling of giant elliptical galaxies, but this value refers to an average sound speed in the central regions of an elliptical galaxy. Around the relativistic jet, there may be a hot wind or cocoon, where the sound speed is much higher, being a significant fraction of the light speed. For instance, Hardee et al. (2005) estimate that the sound speed immediately outside the jet in the radio galaxy 3C 120 is $a_j \gtrsim 0.1c$. For these two values, $a_j = 0.001c$ and $0.1c$, we get $\omega \sim 2 \times 10^{-15}$ and $2 \times 10^{-13}$ Hz, respectively. The corresponding driving periods are $P_d \approx 10^7$ and $10^5$ yr. If $a_j$ is less than $c/\sqrt{3}$ of the ultrarelativistic case, the frequencies will be higher and corresponding periods shorter.

The helical streaming model presented above describes only one normal mode, and it does not explain why the components in the 15 GHz monitoring data cluster in groups, i.e., why there seem to be certain parts in the jet where the flow becomes visible (see § 3.1). However, this might be explained if there are other, short-wavelength, instability modes present in the jet in addition to the externally driven mode. Our current data set does not allow us to test this hypothesis, but some hints of another instability mode may be present in Figure 12, where the components look like they are “oscillating” about the best-fit trajectory.

5. CONCLUSIONS

We have presented multifrequency VLBI data revealing a strongly curved jet in the gigahertz-peaked—spectrum quasar PKS 2136+141. The observations show a 210° change of the jet position angle, which is, to our knowledge, the largest ever observed $\Delta$PA. in an astrophysical jet. The jet appearance is highly reminiscent of a helix with the axis of the helix cone oriented toward our line of sight.

Eight years of monitoring with the VLBA at 15 GHz shows several components moving down the jet with clearly nonballistic trajectories, which excludes the precessing ballistic jet model from the list of possible explanations for the helical structure in PKS 2136+141. Instead, the individual components are streaming along a curved trajectory. The estimated ejection epochs of the components are coincident with two major total flux density outbursts in the 1990s, with three components being associated with both outbursts. This may suggest that a single flare event is
associated with complicated structural changes in the jet, possibly involving multiple shocks.

Most of the components have apparent velocities in the range of 8c–17c. One component even shows 25c, but this high speed is based on only two data points, and therefore it is unreliable. We find evidence suggesting that $\beta_{\text{app}}$ increases after the first 0.5 mas from the core, a distance corresponding to a strong bend. Since the apparent jet opening angle changes at this first 0.5 mas to $\approx$0.6 within the first 0.5 mas to $\approx$1.5 at distances between 1 and 2 mas.

We fit the observed jet trajectory with a model describing a jet displaced by a helical fundamental mode of Kelvin-Helmholtz instability. Our observations suggest that the wave is at the low-frequency limit (relative to the “resonant” frequency of the jet) and has a nearly constant wave speed and wavelength along the jet. This favors a periodic perturbation driven into the jet. The source of the perturbation could be, e.g., jet precession or orbital motion of a supermassive binary black hole. Our present data do not allow us to reliably calculate the driving period of the K-H wave, and hence we cannot further discuss the perturbation’s origin. Follow-up observations of the jet at angular scales larger than 15 mas, as well as further monitoring of the jet kinematics, will probably shed light on this question. In the best-fit model, the helix lies on the surface of the cone with a half-opening angle of 10°, and the angle between our line of sight and the axis of the cone is only 0.3°; i.e., we are looking right into the helix cone.

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REFERENCES

Hirabayashi, H. 2000, PASJ, 52, 997
Storn, R., & Price, K. 1997, J. Global Optimization, 11, 341
COORDINATED MULTIWAVELENGTH OBSERVATION OF 3C 66A DURING THE WEBT CAMPAIGN OF 2003–2004


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ABSTRACT

The BL Lac object 3C 66A was the target of an extensive multiwavelength monitoring campaign from 2003 July through 2004 April (with a core campaign from 2003 September to 2003 December) involving observations throughout the electromagnetic spectrum. Radio, infrared, and optical observations were carried out by the WEBT-ENIGMA collaboration. At higher energies, 3C 66A was observed in X-rays (RXTE) and at very high energy (VHE) in γ-rays (STACEE, VERITAS). In addition, the source has been observed with the VLBA at nine epochs throughout the period 2003 September to 2004 December, including three epochs contemporaneous with the core campaign. A gradual brightening of the source over the course of the campaign was observed at all optical frequencies, culminating in a very bright maximum around 2004 February 18. The WEBT campaign revealed microvariability with flux changes of ~5% on timescales as short as ~2 hr. The source was in a relatively bright state, with several bright flares on timescales of several days. The spectral energy distribution (SED) indicates a $i$F$_i$ peak in the optical regime. A weak trend of optical spectral hysteresis with a trend of spectral softening throughout both the rising and decaying phases has been found. On longer timescales, there appears to be a weak indication of a positive hardness-intensity correlation for low optical fluxes, which does not persist at higher flux levels. The 3–10 keV X-ray flux of 3C 66A during the core campaign was historically high and its spectrum very soft, indicating that the low-frequency component of the broadband SED extends beyond ~10 keV. No significant X-ray flux and/or spectral variability was detected. STACEE and Whipple observations provided upper flux limits at >150 and >390 GeV, respectively. The 22 and 43 GHz data from the three VLBA epochs made between 2003 September and 2004 January indicate a rather weak trend of optical spectral hysteresis with a trend of spectral softening throughout both the rising and decaying periods. The radio brightness profile suggests a magnetic field decay $E$ and, thus, a predominantly perpendicular magnetic field orientation.

Subject headings: BL Lacertae objects: individual (3C 66A) — galaxies: active — gamma rays: theory — radiation mechanisms: nonthermal

Online material: color figure

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1. INTRODUCTION

BL Lac objects and flat-spectrum radio quasars (FSRQs) are active galactic nuclei (AGNs) commonly unified in the class of blazars. They exhibit some of the most violent high-energy phenomena observed in AGNs to date. They are characterized by nonthermal continuum spectra, a high degree of linear polarization in the optical, rapid variability at all wavelengths, radio jets with individual components often exhibiting apparent supraluminal motion, and—at least episodically—a significant portion of the bolometric flux emitted in $\gamma$-rays. Forty-six blazars have been detected and identified with high confidence in high-energy ($>100$ MeV) $\gamma$-rays by the EGRET instrument on board the Compton Gamma Ray Observatory (Hartman et al. 1999; Mattiox et al. 2001). The properties of BL Lac objects and blazar-type FSRQs are generally very similar, except that BL Lac objects usually show only weak emission or absorption lines (with equivalent width in the rest-frame of the host galaxy of $<5$ Å), if any. In 3C 66A ($=0219+428 = \text{NRAO 102} = 4C 42.07$), a weak Mg ii emission line has been detected by Miller et al. (1978). This led to the determination of its redshift at $z = 0.444$, which was later confirmed by the detection of a weak Ly$\alpha$ line in the International Ultraviolet Explorer spectrum of 3C 66A (Lanzetta et al. 1993). However, as recently pointed out by Bramel et al. (2005), these redshift determinations are actually still quite uncertain. In this paper we do base our analysis on a redshift value of $z = 0.444$, but we remind the reader that some results of the physical interpretation should be considered as tentative pending a more solid redshift determination.

In the framework of relativistic jet models, the low-frequency (radio–optical/UV) emission from blazars is interpreted as synchrotron emission from nonthermal electrons in a relativistic jet. The high-frequency (X-ray to $\gamma$-ray) emission could be produced either via Compton upscattering of low-frequency radiation by the same electrons responsible for the synchrotron emission (leptonic jet models; for a recent review, see, e.g., Böttcher et al. 2002), or by hadronic processes initiated by relativistic protons co-accelerated with the electrons (hadronic models; for a recent discussion, see, e.g., Mücke & Protheroe 2001; Mücke et al. 2003).

To date, six blazars have been detected at very high energies ($>300$ GeV) with ground-based air Cerenkov detectors (Punch et al. 1992; Quinn et al. 1996; Catanesi et al. 1998; Chadwick et al. 1999; Aharonian et al. 2002; Horan et al. 2002; Holder et al. 2003). All of these belong to the subclass of high-frequency peaked BL Lac objects (HBLs). The field of extragalactic GeV–TeV astronomy is currently one of the most exciting research areas in astrophysics, as the steadily improving flux sensitivities of the new generation of air Cerenkov telescope arrays and their decreasing energy thresholds (for a recent review see, e.g., Weejes et al. 2002) provide a growing potential to extend their extragalactic source list toward intermediate- and even low-frequency peaked BL Lac objects (LBLs) with lower $\nu F_\nu$ peak frequencies in their broadband spectral energy distributions (SEDs). Detection of such objects at energies $\sim 40–100$ GeV might provide an opportunity to probe the intrinsic high-energy cutoff of their SEDs since at those energies, $\gamma\gamma$ absorption due to the intergalactic infrared background is still expected to be small (e.g., de Jager & Stecker 2002).

3C 66A has been suggested as a promising candidate for detection by the new generation of atmospheric Cerenkov telescope facilities such as STACEE (Solar Tower Air Cerenkov Effect Experiment) or VERITAS (Very Energetic Radiation Imaging Telescope Array System; e.g., Costamante & Ghisellini 2002). Neshpor et al. (1998, 2000) have actually reported multiple detections of the source with the GT-48 telescope of the Crimean Astrophysical Observatory, but those have not been confirmed by any other group so far (see, e.g., Horan et al. 2004). 3C 66A is classified as a low-frequency peaked BL Lac object (LBL), a class also commonly referred to as radio-selected BL Lac objects. Its low-frequency spectral component typically peaks at IR–UV wavelengths, while the high-frequency component seems to peak in the multi-MeV–GeV energy range. Since its optical identification by Wills & Wills (1974), 3C 66A has been the target of many radio, IR, optical, X-ray, and $\gamma$-ray observations, although it is not as regularly monitored at radio frequencies as many other blazars due to problems with source confusion with the nearby radio galaxy 3C 66B (6.5 from 3C 66A), in particular at lower (4.8 and 8 GHz) frequencies (Aller et al. 1994; Takalo et al. 1996).
The long-term variability of 3C 66A at near-IR (J, H, and K bands) and optical (UBVRI) wavelengths has recently been compiled and analyzed by Fan & Lin (1999 and 2000, respectively). Variability at those wavelengths is typically characterized by variations over \( \lesssim 1.5 \text{ mag} \) on timescales ranging from \( \sim 1 \text{ week} \) to several years. A positive correlation between the \( B - R \) color (spectral hardness) and the \( R \) magnitude has been found by Vagnetti et al. (2003). An intensive monitoring effort by Takalo et al. (1996) revealed evidence for rapid microvariability, including a decline \( \sim 0.2 \text{ mag} \) within \( \sim 6 \text{ hr} \). Microvariability had previously been detected in 3C 66A by Carini & Miller (1991) and De Diego et al. (1997), while other, similar efforts did not reveal such evidence (e.g., Miller & McGinsey 1978; Takalo et al. 1992; Xie et al. 1992). Lainela et al. (1999) also report on a 65 day periodicity of the source in its optically bright state, which has so far not been confirmed in any other analysis.

3C 66A is generally observed as a point source, with no indication of the host galaxy. The host galaxy of 3C 66A was marginally resolved by Wurtz et al. (1996). They found \( R_{\text{gamm}} = 19.0 \text{ mag} \) for the host galaxy; the Hubble type could not be determined.

In X-rays, the source has been previously detected by EXOSAT (European X-Ray Observatory Satellite; Sambruna et al. 1994), Einstein (Worrall & Wilkes 1990), ROSAT (Röntgensatellit; Fossati et al. 1998), BeppoSAX (Perri et al. 2003), and XMM-Newton (Croston et al. 2003). It shows large-amplitude soft X-ray variability among these various epochs of observation, with flux levels at 1 keV ranging from \( \sim 0.4 \text{ to } \sim 5 \mu \text{Jy} \) and generally steep (energy index \( \alpha > 1 \)) soft X-ray spectra below 1 keV. 3C 66A has also been detected in \( > 100 \text{ MeV} \) \( \gamma \)-rays by EGRET on several occasions, with flux levels up to \( F_{>100 \text{ MeV}} = (25.3 \pm 5.8) \times 10^{-9} \text{ photons cm}^{-2} \text{ s}^{-1} \) (Hartman et al. 1999).

Superluminal motion of individual radio components of the jet has been detected by Jorstad et al. (2001). While the identification of radio knots across different observing epochs is not unique, Jorstad et al. (2001) favor an interpretation implying superluminal motions of up to \( \beta_{\text{app}} \sim 19 \text{ h}^{-1} \sim 27 \). This would imply a lower limit on the bulk Lorentz factor of the radio-emitting regions of \( \Gamma \gtrsim 27 \). However, theoretical modeling of the nonsimultaneous SED of 3C 66A (Ghisellini et al. 1998) suggests a bulk Lorentz factor of the emitting region close to the core—where the \( \gamma \)-ray emission is commonly believed to be produced—of \( \Gamma \sim 14 \), more typical of the values obtained for other blazars as well.

In spite of the considerable amount of observational effort spent on 3C 66A (see, e.g., Takalo et al. 1996 for an extensive optical monitoring campaign on this source), its multiwavelength SED and correlated broadband spectral variability behavior are still surprisingly poorly understood, given its possible VHE \( \gamma \)-ray source candidacy. The object has never been studied in a dedicated multiwavelength campaign during the lifetime of EGRET. There have been few attempts of coordinated multiwavelength observations. For example, Worrall et al. (1984) present quasi-simultaneous radio, IR, optical, and UV observations of 3C 66A in 1983, but observations at different wavelength bands were still separated by up to \( \sim 2 \text{ weeks} \), and no simultaneous higher energy data were available. This is clearly inadequate to seriously constrain realistic, physical emission models, given the established large-amplitude variability on similar timescales and the importance of the high-energy emission in the broadband SED of the source.

For this reason, we organized an intensive multiwavelength campaign to observe 3C 66A from 2003 July through 2004 April, focusing on a core campaign from 2003 September to 2003 December. In § 2 we describe the observations and data analysis and present light curves in the various frequency bands. Spectral variability patterns are discussed in § 3, and the results of our search for interband cross-correlations and time lags are presented in § 4. Simultaneous broadband SEDs of 3C 66A at various optical brightness levels are presented in § 5. In § 6 we use our results to derive estimates of physical parameters, independent of the details of any specific model. We summarize in § 7. In a companion paper (M. Joshi & M. Böttcher 2005, in preparation), we will use time-dependent leptonic models to fit the spectra and variability patterns found in this campaign and make specific predictions concerning potentially observable X-ray spectral variability patterns and \( \gamma \)-ray emission.

Throughout this paper we refer to \( \alpha \) as the energy spectral index, \( F_v(Jy) \propto \nu^{-\alpha} \). A cosmology with \( \Omega_m = 0.3, \Omega \Lambda = 0.7, \) and \( H_0 = 70 \text{ km s}^{-1} \text{ Mpc}^{-1} \) is used. In this cosmology, and using the redshift of \( z = 0.444 \), the luminosity distance of 3C 66A is \( d_L = 2.46 \text{ Gpc} \).

2. OBSERVATIONS, DATA REDUCTION, AND LIGHT CURVES

3C 66A was observed in a coordinated multiwavelength campaign at radio, near-IR, optical (by the WEBT-ENIGMA collaboration), X-ray, and VHE \( \gamma \)-ray energies during a core campaign period from 2003 September to 2003 December. The object is being continuously monitored at radio and optical wavelengths by several ongoing projects, and we present data in those wavelength regimes starting in 2003 July. During the core campaign, we found that the source was gradually brightening, indicating increasing activity throughout the campaign period. For this reason, the WEBT campaign was extended, and we continued to achieve good time coverage until early 2004 March, and individual observatories still contributed data through 2004 April. The overall time line of the campaign, along with the measured long-term light curves at radio, infrared, optical, and X-ray frequencies, is illustrated in Figure 1. Table 1 lists all participating observatories that contributed data to this campaign. In this section we will describe the individual observations in the various frequency ranges and outline the data reduction and analysis.

2.1. Radio Observations

At radio frequencies, the object was monitored using the University of Michigan Radio Astronomy Observatory (UMRAO) 26 m telescope, at 4.8, 8, and 14.5 GHz; the 14 m Metsähovi Radio Telescope of the Helsinki University of Technology, at 22 and 37 GHz; and the 576 m ring telescope (RATAN-600) of the Russian Academy of Sciences, at 2.3, 3.9, 7.7, 11, and 22 GHz.

At the UMRAO, the source was monitored in the course of the ongoing long-term blazar monitoring program. The data were analyzed following the standard procedure described in Aller et al. (1985). As mentioned above, the sampling at the lower frequencies (4.5 and 8 GHz) was rather poor (about once every 1–2 weeks) and some individual errors were rather large due to source confusion problems with 3C 66B. At 14.5 GHz, a slightly better sampling of \( \sim 1–2 \text{ observations per week} \), at least until the end of 2003, was achieved with relative flux errors of typically a few percent. The resulting 14.5 GHz light curve is included in Figure 1. It seems to indicate generally moderate variability (\( \Delta F/F \lesssim 25\% \)) on timescales of \( \lesssim 1 \text{ week} \), although a discrete autocorrelation analysis (Edelson & Krolik 1988) does
not reveal any significant structure, primarily due to the insufficient sampling.

At 22 and 37 GHz, 3C 66A has been monitored using the 14 m radio telescope of the Metsähovi Radio Observatory of the Helsinki University of Technology. The data have been reduced with the standard procedure described in Teräsranta et al. (1998). The 22 and 37 GHz light curves reveal moderate-amplitude ($\Delta F/F \lesssim 30\%$), erratic variability that is clearly undersampled by the available data. The discrete autocorrelation functions indicate a timescale of a few days for this short-term variability.

We point out that the observed erratic variability may, at least in part, be a consequence of interstellar scintillation. At the Galactic coordinates of the source, the transition frequency, where the interstellar scattering strength (a measure of the phase change induced by interstellar scattering) becomes unity, is $\sim 7$ GHz (Walker 1998). At higher frequencies, as considered here, scattering will occur in the weak scattering regime, with only small phase changes through interstellar scattering. In this regime, we find fractional rms variability amplitudes for a point source due to interstellar scintillation of 0.36 and 0.09 at frequencies of 14.5 and 37 GHz, respectively, and the respective variability timescales due to interstellar scintillation are 1.4 and 0.9 hr. Consequently, in particular at lower frequencies ($\lesssim 20$ GHz), a substantial fraction of the observed variability may well be due to interstellar scintillation.

The RATAN-600 has monitored 3C 66A, performing 17 observations from mid-1997 through 2003 October. The last two of these observations (October 11 and 14) coincided with our core campaign and provided additional frequency coverage at 2.3, 3.9, 7.7, 11, and 22 GHz. Those data were analyzed as described in detail in Kovalev et al. (1999), and the time average of the resulting fluxes from the two observations in 2003 October are included in the SED shown in Figure 15.

A total of nine observing epochs using the VLBA have been approved to accompany our multiwavelength campaign on 3C 66A. In this paper we describe details of the data analysis and results of the first (2003.78; VLBA observing code BS133A), the second (2003.83; BS133B), and the fourth (2004.08; BS133E) epochs, concentrating on the results at 22 and 43 GHz. The third epoch (2003.96) suffered from radio frequency interference, and its reduction is in progress. Results of the complete set of all nine VLBA observing epochs, including all six frequencies (2.3, 5, 8.4, 22, 43, and 86 GHz) together with polarization data, will be presented in a separate paper (T. Savolainen et al. 2005, in preparation).

The data were correlated on the VLBA correlator and were postprocessed with the NRAO Astronomical Image Processing System, AIPS (Greisen 1988), and the Caltech DIFMAP package (Shepherd 1997). Standard methods for VLBI data reduction and imaging were used. A priori amplitude calibration was performed using measured system temperatures and gain curves in the AIPS task APCAL. At this point, a correction for atmospheric opacity was also applied. After removal of the parallactic angle phase, single-band delay and phase offsets were calculated manually by fringe-fitting a short scan of data of a bright calibration source (0420–014). We did manual phase calibration instead of using pulse-cal tones, because there were unexpected jumps in the phases of the pulse-cal during the observations. Global fringe-fitting was performed with the AIPS task FRING using Los Alamos (LA) as a reference antenna. The AIPS procedure CRSFRING was used to remove the delay difference between right- and left-hand systems (this is needed for polarization calibration). A bandpass correction was determined and applied before averaging across the channels, after which the data were imported into DIFMAP.

In DIFMAP, the data were first phase self-calibrated using a point source model and then averaged in time. We performed data editing in a station-based manner and ran several iterations of CLEAN and phase self-calibration in Stokes I. After a reasonable fit to the closure phases was obtained, we also performed amplitude self-calibration, first with a solution interval corresponding to the whole observation length. The solution interval was then gradually shortened as the model improved by further CLEANing. Final images were produced using the Perl library FITSplot. For final images, normal weighting of the $uv$-data was used in order to reveal the extended, low surface brightness emission.

We have checked the amplitude calibration by comparing the extrapolated zero baseline flux density of the compact source 0420–014 to the single-dish flux measurements at Metsähovi. The fluxes are comparable: at 22 GHz the VLBI flux is on average 5% lower than the single-dish flux, and the 43 GHz flux is on average 8% lower than the nearly simultaneous single-dish measurements at 37 GHz. The small amount of missing flux is probably due to some extended emission resolved out in the high-frequency VLBI images. The integrated VLBI flux of 3C 66A is about 30% smaller than the corresponding single-dish value at both 22 and 43 GHz, but since the source has notable kiloparsec-scale structure evident in the VLA images (Price et al. 1993; Taylor et al. 1996), it is most likely that the missing flux comes from this kiloparsec-scale jet, which is resolved out in our VLBI images. Thus, we conclude that the

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44 The VLBA is a facility of the National Radio Astronomy Observatory (NRAO). The NRAO is a facility of the National Science Foundation, operated under cooperative agreement by Associated Universities, Inc.

45 See http://personal.denison.edu/~homand.
In order to estimate the parameters of the emission regions in the jet, we did model fitting to the self-calibrated visibilities in DIFMAP. The data were fitted with circular Gaussian model components, and we sought to obtain the best possible fit to the visibilities and to the closure phases. Several starting points were tried in order to avoid a local minimum fit. The results of this fitting procedure are included in Figures 2 and 3. In all the lower frequencies.

Figure 4 shows the separation of the model components from the core, at each epoch. Components A–C are weak compared to the core (see Table 2), which renders the estimation of errors in the flux at this point could plausibly be due to increased Doppler boosting caused by the bending of the jet toward our line of sight. Region A shows weak extended emission at 8 mas from the core at 22 GHz. This emission is more pronounced at lower frequencies.

### Table 1

<table>
<thead>
<tr>
<th>Observatory</th>
<th>Specifications</th>
<th>Frequency/Filters/Energy Range</th>
<th>$N_{\text{obs}}$</th>
</tr>
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<tbody>
<tr>
<td><strong>Radio Observatories</strong></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>UMRAO, Michigan, USA</td>
<td>26 m</td>
<td>4.8, 8, 14.5 GHz</td>
<td>46</td>
</tr>
<tr>
<td>Metsahovi, Finland</td>
<td>14 m</td>
<td>22, 37 GHz</td>
<td>103</td>
</tr>
<tr>
<td>RATAN-600, Russia</td>
<td>576 m (ring)</td>
<td>2.3, 3.9, 7.7, 11, 22 GHz</td>
<td>17</td>
</tr>
<tr>
<td>VLBA</td>
<td>10 x 25 m</td>
<td>2.3, 5, 8.4, 22, 43, 86 GHz</td>
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<td><strong>Infrared Observatories</strong></td>
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<td></td>
</tr>
<tr>
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<td>1.1 m</td>
<td>$J, H, K$</td>
<td>171</td>
</tr>
<tr>
<td>Mount Abu, India (MIRO)</td>
<td>1.2 m (NICMOS3)</td>
<td>$J, H, K'$</td>
<td>79</td>
</tr>
<tr>
<td>Roque (NOT), Canary Islands</td>
<td>2.56 m</td>
<td>$J, H, K$</td>
<td>15</td>
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<tr>
<td><strong>Optical Observatories</strong></td>
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<td></td>
<td></td>
</tr>
<tr>
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<td>$R$</td>
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<tr>
<td>Armenzano, Italy</td>
<td>40 cm</td>
<td>$B, V, R, I$</td>
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</tr>
<tr>
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<td>60 cm</td>
<td>$R$</td>
<td>12</td>
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<td>Boltwood, Canada</td>
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<td>Catania, Italy</td>
<td>91 cm</td>
<td>$U, B, V$</td>
<td>835</td>
</tr>
<tr>
<td>Crimean Astron. Obs., Ukraine</td>
<td>70 cm ST-7</td>
<td>$B, V, R, I$</td>
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<tr>
<td>Heidelberg, Germany</td>
<td>70 cm</td>
<td>$B, R, I$</td>
<td>8</td>
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<tr>
<td>Kitt Peak (MDM), Arizona, USA</td>
<td>130 cm</td>
<td>$B, V, R, I$</td>
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<tr>
<td>Michael Adrian Obs., Germany</td>
<td>120 cm</td>
<td>$R$</td>
<td>30</td>
</tr>
<tr>
<td>Mount Lemmon, South Korea</td>
<td>100 cm</td>
<td>$B, V, R, I$</td>
<td>399</td>
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<tr>
<td>Mount Maidanak, Uzbekistan</td>
<td>150 cm AZT-22</td>
<td>$B, R$</td>
<td>1208</td>
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<tr>
<td>Nyrölä, Finland</td>
<td>40 cm SCT</td>
<td>$R$</td>
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<tr>
<td>Osaka, Japan</td>
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<td>$R$</td>
<td>1167</td>
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<td>Perugia, Italy</td>
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<td>$V, R, I$</td>
<td>140</td>
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<td>Roque (KVA), Canary Islands</td>
<td>35 cm</td>
<td>$B, V, R$</td>
<td>653</td>
</tr>
<tr>
<td>Roque (NOT), Canary Islands</td>
<td>256 cm</td>
<td>$U, B, V, R, I$</td>
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<td>Sabadell, Spain</td>
<td>50 cm</td>
<td>$B, R$</td>
<td>4</td>
</tr>
<tr>
<td>San Pedro Martir, Mexico</td>
<td>150 cm</td>
<td>$B, V, R, I$</td>
<td>185</td>
</tr>
<tr>
<td>Shanghai, China</td>
<td>156 cm</td>
<td>$V, R$</td>
<td>36</td>
</tr>
<tr>
<td>Skinakas, Crete</td>
<td>130 cm</td>
<td>$B, V, R, I$</td>
<td>156</td>
</tr>
<tr>
<td>Sobaeksan, South Korea</td>
<td>61 cm</td>
<td>$B, V, R, I$</td>
<td>133</td>
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<tr>
<td>St. Louis, Missouri, USA</td>
<td>35 cm</td>
<td>$B, R$</td>
<td>16</td>
</tr>
<tr>
<td>Torino, Italy</td>
<td>105 cm REOSC</td>
<td>$B, V, R, I$</td>
<td>227</td>
</tr>
<tr>
<td>Tuorla, Finland</td>
<td>103 cm</td>
<td>$B, V, R$</td>
<td>1032</td>
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<td><strong>X-Ray Observatory</strong></td>
<td></td>
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<tr>
<td>RXTE</td>
<td></td>
<td>3–25 keV</td>
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<tr>
<td><strong>γ-Ray Observatories</strong></td>
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<td></td>
<td></td>
</tr>
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<td>STACEE</td>
<td>Solar-Tower Cerenkov Array</td>
<td>&gt;100 GeV</td>
<td>85</td>
</tr>
<tr>
<td>VERITAS</td>
<td>Whipple 10 m</td>
<td>&gt;390 GeV</td>
<td>31</td>
</tr>
</tbody>
</table>

accuracy of the amplitude calibration is better than 10% at both frequencies.

In order to estimate the parameters of the emission regions in the jet, we did model fitting to the self-calibrated visibilities in DIFMAP. The data were fitted with circular Gaussian model components, and we sought to obtain the best possible fit to the visibilities and to the closure phases. Several starting points were tried in order to avoid a local minimum fit. The results of this fitting procedure are included in Figures 2 and 3. In all the model fits, the core of the VLBI jet is the northernmost and also the brightest component. Beyond the core, we have divided the jet into three regions named A, B and C. Closest to the core is region C, which consists of three components with decreasing surface brightness as a function of the distance from the core. Region B, also made of three components, shows a clear bending of the jet together with rebrightening. The observed increase in the flux at this point could plausibly be due to increased Doppler boosting caused by the bending of the jet toward our line of sight. Region A shows weak extended emission at 8 mas from the core at 22 GHz. This emission is more pronounced at lower frequencies.

Figure 4 shows the separation of the model components from the core, at each epoch. Components A–C are weak compared to the core (see Table 2), which renders the estimation of errors in the flux at this point could plausibly be due to increased Doppler boosting caused by the bending of the jet toward our line of sight. Region A shows weak extended emission at 8 mas from the core at 22 GHz. This emission is more pronounced at lower frequencies.

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size of the knot—whichever is larger. These estimates are based on the experience with other sources (T. Savolainen et al. 2005, in preparation; K. Wiik et al. 2005, in preparation), and they follow the results by Jorstad et al. (2005), where uncertainties of model fit parameters for a large number of observations are estimated. Positional errors of similar size are also reported by Homan et al. (2001), who have estimated the uncertainties from the variance of the component position about the best-fit polynomial.

There is a $\sim 0.1$ mas shift of the components C1–C3 between the two frequencies, with the 22 GHz model components appearing further downstream. This cannot be explained by an opacity effect, since in that case the components would have shifted in the opposite direction: at higher frequency we would

![Diagram](image1)

**Fig. 2.**—22 GHz VLBA maps of 3C 66A during the three epochs in 2003–2004 contemporaneous with our campaign. Overlaid are the fit results from decomposing the structure into a sum of Gaussian components.

![Diagram](image2)

**Fig. 3.**—43 GHz VLBA maps of 3C 66A during the three epochs in 2003–2004 contemporaneous with our campaign. Overlaid are the fit results from decomposing the structure into a sum of Gaussian components.
expect to see emission from the region closer to the apex of the jet than at lower frequency, if the opacity effect is significant. This obviously is not the case in Figure 4, and moreover, for the components B2 and B3 the positions at 22 and 43 GHz are coincident. However, there is a simple explanation for the observed shift, if the brightness profile along the jet in region C is smooth. Namely, the flux in the outer part of region C decreases more steeply at 43 GHz than at 22 GHz (see Figs. 2 and 3). As region C is modelled by the same number of components at both frequencies, the model-fit procedure shifts the 22 GHz components downstream relative to 43 GHz components in order to represent the power law.

Figure 4 shows that, except for one component, all the components revealed by the analysis of the three epochs considered here are consistent with zero proper motion. Although a monitoring effort over a longer timescale will be necessary to settle this issue, our analysis shows slower component motion than that presented by Jorstad et al. (2001), where fast superluminal motions of several components with speeds ranging from 9 to 19 \(c\) were found. Note, however, that the data used by Jorstad et al. (2001) covered a much longer time frame and refers to different components and a rather different state of activity of 3C 66A. Moreover, there are also similarities between the data sets: Jorstad et al. (2001) reported a stationary component (their component “C”) at a distance of \(~0.5\) mas from the core, which could correspond to our component C2. New 7 mm VLBA monitoring data presented in a recent paper by Jorstad et al. (2005) suggest that there are two kinds of components in 3C 66A: fast and very weak components with apparent speeds \(>20\) c, and stronger components with moderate velocities of \(1.5–5\) c. The components C1–C3 in our analysis could be qualitatively similar to the latter. More solid conclusions concerning the question of superluminal motions of individual components might be possible after the final analysis of all nine epochs of the entire VLBA monitoring program of 3C 66A proposed in connection with this campaign (T. Savolainen et al. 2005, in preparation). Based on the analysis of the three epochs contemporaneous with the other observations of this campaign, one component, C1, shows superluminal motion of \((8.5 \pm 5.6)\) \(c\) for \(h = 0.7\).

The essentially zero proper motion of most of the components together with the observed trend of the brightness temperature versus distance along the jet (see § 6.2) suggests that in the case of 3C 66A the VLBA model-fit components might not correspond to actual “knots” in the jet, but a rather smooth flow with simple brightness profile (such as a power law) might better describe the jet during this campaign—at least in region C. More detailed parameter estimates will be extracted from these results in § 6.2.

### 2.2. Infrared Observations

In the context of the extensive WEBT campaign (see § 2.3), 3C 66A was also observed at near-IR wavelengths in the \(J, H,\) and \(K/K'\) bands at three observatories: the Campo Imperatore 1.1 m telescope of the Infrared Observatory of the Astronomical Observatory of Rome, Italy; the 1.2 m telescope (using the NICMOS3 HgCdTe IR array with 256×256 pixels, with each pixel corresponding to \(0.96\) on the sky) at the Mount Abu Infrared Observatory (MIRO) at Mount Abu, India; and the 2.56 m Nordic Optical Telescope (NOT) on Roque de los Muchachos on the Canary Island of La Palma. The primary data were analyzed using the same standard technique as the optical data (see § 2.3), including flat-field subtraction, extraction of instrumental magnitudes, calibration against comparison stars to obtain standard magnitudes, and dereddening. The sampling was generally not dense enough to allow an improvement of the data quality by rebinning. Unfortunately, the three observatories did not perform any observations on the same day, so that the cross-calibration between different instruments is problematic. We have therefore opted not to correct for possible systematic offsets beyond the calibration to standard magnitudes of well-calibrated comparison stars (see § 2.3).

The resulting \(H\)-band light curve is included in Figure 1, and a comparison of all three IR-band light curves (\(J, H,\) and \(K/K'\)) is shown in Figure 5. Generally, we find moderate variability of \(\Delta(J, H, K/K') \approx 0.2\) within \(~\)a few days. The various infrared bands trace each other very well, and there does not appear to be any significant time lags between the different bands. It is obvious that there are issues with the cross calibration between the three observatories. In particular, the Mount Abu observations (using the \(J, H,\) and \(K'\) filters) in mid-December of 2003 indicate a significant drop of the average level of near-IR flux. While there appears to be a similar flux drop around the same time in the radio light curves, the much more densely sampled optical light curves do not indicate a similar feature. However, it is obvious that the near-IR data around this time exhibit an unusually large scatter (by \(~\)0.5 mag on subhour timescales), they may be affected by calibration uncertainties. We note that rebinning of the IR data in time bins of up to 30 minutes did not improve the quality of the data because of the extreme nature of the apparent flux fluctuations. We also point out that the use of the \(K'\) filter at Mount Abu versus the \(K\) filter at the other two IR observatories could only explain offsets of less than the individual measurement errors, so this effect was neglected in our analysis.

### 2.3. Optical Observations

In the optical component of the extensive WEBT campaign, 24 observatories in 15 countries contributed 7611 individual photometric data points. The observing strategy and data analysis followed to a large extent the standard procedure described previously for a similar campaign on BL Lac in 2000 (Villata et al. 2002). For more information about WEBT campaigns, see...

It had been suggested that, optimally, observers perform photometric observations alternately in the $B$ and $R$ bands, and include complete ($U$)$BVRI$ sequences at the beginning and the end of each observing run. Exposure times should be chosen to obtain an optimal compromise between high-precision (instrumental errors less than $\sim 0.03$ mag for small telescopes and $\sim 0.01$ mag for larger ones) and high time resolution. If this precision requirement leads to gaps of 15–20 minutes in each light curve, we suggested carrying out observations in the $R$ band only.

Observers were asked to perform bias (dark) corrections, as well as flat-fielding on their frames, and obtain instrumental magnitudes, applying either aperture photometry (using IRAF or CCDPHOT) or Gaussian fitting for the source 3C 66A and the comparison stars 13, 14, 21, and 23 in the tables of González-Pérez et al.

### TABLE 2

<table>
<thead>
<tr>
<th>Frequency (GHz)</th>
<th>Component ID</th>
<th>Flux (mJy)</th>
<th>Core Distance (mas)</th>
<th>Position Angle (deg)</th>
<th>Diameter (mas)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Epoch 2003.78</td>
<td></td>
<td></td>
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<tr>
<td>22.................</td>
<td>Core 387</td>
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<td>0.0</td>
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<tr>
<td></td>
<td>C3 61</td>
<td>0.295</td>
<td>172.9</td>
<td>0.136</td>
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<tr>
<td></td>
<td>C2 36</td>
<td>0.658</td>
<td>168.2</td>
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<tr>
<td></td>
<td>C1 19</td>
<td>1.096</td>
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<td></td>
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<td></td>
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<td>A 37</td>
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(2001), where high-precision standard magnitudes for these stars have been published. This calibration has then been used to convert instrumental to standard photometric magnitudes for each data set. In the next step, unreliable data points (with large error bars at times when higher quality data points were available) were discarded. Our data did not provide evidence for significant variability on subhour timescales. Consequently, error bars on individual data sets could be further reduced by rebinning on timescales of typically 15–20 minutes.

Finally, there may be systematic offsets between different instruments and telescopes. Wherever our data sets contained sufficient independent measurements to clearly identify such offsets, individual data sets were corrected by applying appropriate correction factors. In the case of BL Lac and similar sources (e.g., Villata et al. 2002), such corrections need to be done on a night-by-night basis since changes in the seeing conditions affect the point-spread function and thus the precise amount of the host galaxy contamination. However, in the absence of a significant host galaxy contribution, these systematic effects should be purely instrumental in nature and should thus not depend on daily seeing conditions. Thus, we opted to introduce only global correction factors for entire single-instrument data sets on 3C 66A. This provided satisfactory results without obvious residual inter-instrumental inconsistencies. The resulting offsets are listed in Table 3.

In order to provide information on the intrinsic broadband spectral shape (and, in particular, a reliable extraction of $B-R$ color indices), the data were dereddened using the Galactic extinction coefficients of Schlegel et al. (1998), based on $A_B = 0.363$ mag and $E(B-V) = 0.084$ mag. As mentioned in the introduction, the $R$ magnitude of the host galaxy of 3C 66A is $\sim 19$ mag, so its contribution is negligible, and no host-galaxy correction was applied.

2.3.1. Light Curves

As a consequence of the observing strategy described above, the $R$- and $B$-band light curves are the most densely sampled ones, resulting in several well-sampled light curve segments over ~5–10 days each, with no major interruptions, except for a gap of a few hours due to the lack of coverage when 3C 66A would have been optimally observable from locations in the Pacific Ocean. The $R$-band light curve over the entire duration of the campaign is included in Figure 1 and compared to the light curves at all other optical bands in Figure 6. These figures illustrate that the object underwent a gradual brightening throughout the period 2003 July to 2004 February, reaching a maximum at $R \approx 13.4$ on 2004 February 18, followed by a sharp decline by $\Delta R \sim 0.4$ mag within ~15 days. On top of this overall brightening trend, several major outbursts by $\Delta R \sim 0.3$–0.5 mag on timescales of ~10 days occurred. The two most dramatic ones of these outbursts peaked on 2003 July 18 and November 1. Details of the November 1 outburst are displayed in Figures 7–9. We find evidence for intraday micro-variability of $\Delta R \sim 0.05$ mag on timescales down to ~2 hr. One example for such evidence is illustrated in Figure 10.
which shows the intranight data from the Torino Observatory for JD 2,452,955 (=2003 November 11). The figure presents the instrumental magnitudes of 3C 66A, comparison star A, and the difference of both. While the difference seems to be affected by seeing effects at the beginning and end of the observation (rising and falling in tandem with the instrumental magnitudes), there is evidence for intraday variability around JD 2,452,955.5, where there were only very minor changes in the atmospheric opacity.

Visual inspection of individual major outbursts suggests periods of more rapid decline than rise.

2.3.2. Periodicity Analysis

For several blazars, including 3C 66A, periodicities on various timescales, from several tens of days to several years, have been claimed (see, e.g., Rieger 2004 and references therein for a more complete discussion). In order to search for a possible periodicity in the optical $R$-band light curve from our WEBT observations, we performed a Fourier analysis on the light curve segments shown in more detail in Figs. 8 and 9.
campaign, we have performed a Fourier analysis of the R-band light curve after interpolation of the light curve over the (short) gaps in the available data. Such an analysis becomes unreliable for periods of more than $1/10$ of the length of the data segment. Consequently, our data can be used for a meaningful analysis on timescales of $\tau \lesssim 30$–50 days. No evidence for a periodicity in this range has been found. However, we note a sequence of large outbursts on 2003 July 18 (MJD 52,838), September 5 (MJD 52,887), November 1 (MJD 52,944), and December 28 (MJD 53,001), and 2004 February 18 (MJD 53,053), which are separated by intervals of 49, 57, 57, and 52 days, respectively. These intervals appear remarkably regular, and only slightly shorter than the 65 day period claimed by Lainela et al. (1999). The duration of our monitoring campaign is too short to do a meaningful analysis of the statistical significance of this quasi-periodicity. This question will be revisited in future work, on the basis of a larger data sample, including archival optical data.

2.4. X-Ray Observations

3C 66A was observed by the Rossi X-Ray Timing Explorer (RXTE) 26 times between 2003 September 19 and December 27, for a total observation time of approximately 200 ks. Analysis of the RXTE PCA data was carried out using faint-source procedures as described in the RXTE Cookbook. Standard selection criteria were used to remove disruptions from the South Atlantic Anomaly, bright Earth, instrument mispointing, and electron contamination. For the entire data set, only PCUs 0 and 2 were activated, but data from PCU 0 were discarded due to its missing propane veto. There is also a persistent problem in the background model of the PCA around the 4.78 keV xenon L edge, which causes a small anomaly in count rates near this energy. For faint sources like 3C 66A, this anomaly can be significant, so photons in the three energy bins surrounding this energy were discarded.

Photons from 3–10 keV were extracted from the data using only the top layer of PCU2. Extracting higher energy photons proved unreliable, as the source flux was close to the PCA source-confusion limit. A spectrum was compiled using data from all 26 source observations. The spectrum was fit to a simple power law, which resulted in an energy spectral index of $\alpha = 1.47 \pm 0.56$ and a flux $\Phi_{1\text{keV}} = (4.06 \pm 0.49) \times 10^{-7}$ photons cm$^{-2}$ s$^{-1}$ keV$^{-1}$, with $\chi^2/n = 13.6/14$. Galactic absorption was not included in the fit model, as the lack of data below 3 keV precludes a useful constraint on $N_H$. A plot of the spectrum and fit is shown in Figure 11.

The PCA is a nonimaging instrument with a FWHM response of $1.14$, so there is the potential that the signal found here is a sum over multiple sources. Most notably, the FR-1 galaxy 3C 66B lies only 6' from 3C 66A and is a known X-ray emitter. Although indistinguishable from 3C 66A with the PCA, its typical flux is

![Figure 9](image-url)  
**Fig. 9.**—Details of the R-band light curve during the decaying phase of the major outburst around 2003 November 1.

![Figure 10](image-url)  
**Fig. 10.**—Instrumental magnitudes of 3C 66A (top panel) and comparison star A (middle panel) of the University of Torino data of JD 2,452,955 (~2003 November 11) The difference (bottom panel) shows clear evidence for intra-night variability during the central portion of this observation (~JD 2,452,955.4 to 2,452,955.5).
at a level of $\sim 2\%$ of the measured PCA flux of 3C 66A (Croston et al. 2003), and it is thus not a significant contributor. In the 3EG catalog, there is also confusion of 3C 66A with a nearby pulsar, PSR J0218+4232. At a source separation of 58”, the PCA response for the pulsar is low for 3C 66A pointed observations, and thus it does not make a significant contribution.

The 3–10 keV photons were binned into 24 hr periods to render a light curve, which is included in Figure 1. No significant flux variations or transient events were seen. To check for spectral variations, the data set was divided into four roughly equal subperiods and power-law spectra were fitted to each subperiod. Fits to all four subperiods were consistent with the spectral parameters obtained from the entire data set, and all subperiods were consistent with each other. From this we conclude that there were no variations of either flux or spectral shape in the X-ray band during the observing period within the detection capabilities of RXTE.

2.5. Gamma-Ray Observations

At VHE $\gamma$-rays, 3C 66A was observed contemporaneously with our broadband campaign by STACEE and by the 10 m Whipple Telescope of the VERITAS collaboration.

STACEE took a total of 85 28 minute on-off pairs of data, totaling 33.7 hr of live time on-source. After data-quality cuts and padding, a total of 16.3 hr of on-source live time remained. A net on-source excess of 1134 events attributed to photons of energy $E > 100$ GeV was seen against a background of 231,742 events. The details of the STACEE data analysis are published in a separate paper (Bramel et al. 2005).

### 3. Spectral Variability

The source excess quoted above corresponds to a significance of 2.2 $\sigma$, which is insufficient to claim a detection, but can be used to establish firm upper flux limits. For various assumptions of an underlying intrinsic source power-law spectrum, those upper limits are listed in Table 4. The data were also binned in 24 hr segments. This revealed no significant excess on any given day and thus no evidence for a statistically significant transient event during the campaign period.

The VERITAS collaboration observed 3C 66A during the period 2003 September 27 to 2004 January 13, obtaining a total of 31 on-off pairs with typically $\sim 27.6$ minutes of on-source live time per pair. The data were analyzed following the standard Whipple data analysis procedure (see, e.g., Falcone et al. 2004). The log of the individual observations and 99.9% confidence upper limits, assuming an underlying Crab-like source spectrum with an underlying energy index of 2.49 (Hillas et al. 1998), are listed in Table 5. Combining all measurements results in a 99.9% confidence upper limit of $0.91 \times 10^{-11}$ ergs cm$^{-2}$ s$^{-1}$.

#### Table 4

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<th>$\Gamma$</th>
<th>$E_{\text{thr}}$ (GeV)</th>
<th>$dN/dE$ at $E_{\text{thr}}$ (photons m$^{-2}$ s$^{-1}$ GeV$^{-1}$)</th>
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<td>$5.23 \times 10^{-9}$</td>
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<td>2.5</td>
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<td>$9.39 \times 10^{-9}$</td>
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<tr>
<td>3.0</td>
<td>150</td>
<td>$2.26 \times 10^{-8}$</td>
</tr>
<tr>
<td>3.5</td>
<td>147</td>
<td>$3.10 \times 10^{-8}$</td>
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In this section we will describe spectral variability phenomena, i.e., the variability of spectral (and color) indices and their correlations with monochromatic source fluxes. As already mentioned in the previous section, no evidence for spectral variability in the X-ray regime was found. Consequently, we will concentrate here on the optical spectral variability as indicated by a change of...
the optical color. In particular, our observing strategy was optimized to obtain a good sampling of the $B/C_0$ color while the $R$-band flux is still rising at $R \approx 13.6$.

Based on our observation (see § 2.3) of no significant variability on timescale less than $\sim 1$ hr, we extracted $B - R$ color indices on measurements of $B$ and $R$ magnitudes separated by less than 20 minutes of each other. Figure 12 shows a comparison of the light curves of the $B$ and $R$ magnitudes and the $B - R$ color index. The figure reveals an interesting new result: while the $B$ and $R$ light curves are well correlated, with no significant detection of a time delay, maxima of the spectral hardness (minima of $B - R$) systematically precede $B$ and $R$ outbursts by a few days. In other words, the optical spectra harden at the onset of a major outburst and continue to soften as the flare progresses through both the rising and the decaying phase. However, Figure 12 also indicates that the apparent lead time of the $B - R$ hardness maximum varies within a range of a few days. This might be the reason that a cross-correlation analysis between the $B - R$ index and the $B$ or $R$ light curves does not reveal a strong peak. In fact, the DCF between those time series never exceeds values of $\sim 0.3$ at all.

Also very interesting and intriguing is the $B - R$ behavior of the source from around JD 2,452,955 to 2,452,995, where no major flux outbursts are observed, but several episodes of significant spectral hardening by $\Delta(B - R) \gtrsim 0.1$ (corresponding to spectral-index variations of $\Delta\alpha_{\text{opt}} \gtrsim 0.2$) are observed. In Figure 13 we have plotted the $B - R$ index versus $R$-band magnitude for the entire data set. The plot shows that there is a weak indication of a positive hardness-intensity correlation at low-flux states with $R \gtrsim 14.0$. At higher flux levels, no correlation is apparent, which might be a consequence of the result found above, that the $B - R$ hardness actually peaks during the rising phase of individual outbursts. This is in contrast with recent results of Vagnetti et al. (2003), who, on the basis of a smaller data set, have found a consistent trend of $B - R$ hardening with increasing $B$-band flux, independent of the actual flux value.

The optical spectral hysteresis visible in our data is further illustrated in Figure 14, where we have plotted the $B - R$ versus $R$ hardness-intensity diagram for the example of one individual outburst on 2003 November 1 (MJD 52,944). The top panel shows the hardness-intensity diagram for the entire flare, which is then split up into the rising and the decaying phases of the flare in the middle and lower panels, respectively. The figure illustrates that as the flare rises up to the peak $R$ brightness, the spectrum already begins to soften significantly as $R \lesssim 13.6$. Similar trends are found for other optical flares as well. Possible physical interpretations of this trend, along with detailed modeling of the SED and spectral variability will be presented in a separate paper (M. Joshi & M. Böttcher 2005, in preparation).

4. INTERBAND CROSS CORRELATIONS AND TIME LAGS

In this section we investigate cross-correlations between the measured light curves at different frequencies, within individual frequency bands as well as broadband correlations between different frequency bands. Because of relatively poor sampling of the radio and IR light curves, our searches for correlations...
between different radio and IR bands and between those wave-
length bands with the optical ones on short timescales com-
pared to the duration of our campaign did not return conclusive
results. Also, in agreement with our nondetection of a periodic-
ity on timescales of $\leq 50$ days, a discrete autocorrelation func-
tion (Edelson & Krolik 1988) analysis of individual optical light
curves did not return results beyond artificial “periodicities.”
Such artifacts can be attributed to the quasi-periodic, uneven
time coverage, which results in spurious detections of peri-
dicities at multiples of days and multiples of $\sim 0.5$ days (the
time delay between the peak coverage from the heavily contribut-
ing observatories in Europe and in East Asia). A discrete autocor-
relation function analysis of the $R$-band light curve also revealed
secondary peaks at $\pm 4$ days with a correlation coefficient close
to 1.

The cleanest results of a DCF analysis are obviously expected
for the most densely sampled light curves, which we obtained
in the $B$ and $R$ bands. We performed DCFs between the $R$-
and $B$-band light curves on a variety of timescales with a variety
of binning intervals, ranging from 15 minutes to 10 days. On in-
trady timescales, we find consistently a sharp peak at 0 delay
with all bin sizes we used, which indicates no evidence for a time
delay between the artificial variability patterns in the $R$ and $B$ band reaching
their peak fluxes in each band. The DCFs on short timescales are
dominated by the artificial 0.5 and 1 day periodicities mentioned
above. The DCF between $R$- and $B$-band magnitudes on time-
scales of several days seems to indicate a strong correlation at
$\tau \approx -4$ days, corresponding to a lead of the $B$ versus the $R$ band.
However, this might be a consequence of the probably artificial
4 day “periodicity” of the $R$-band light curve discussed in the
previous paragraph, which prevents us from making any claim
about the detection of a 4 day delay between the $R$- and $B$-band
light curves on the basis of our DCF analysis.

5. BROADBAND SPECTRAL ENERGY DISTRIBUTIONS

From the various flux measurements described in detail in the
previous sections, we can now compose contemporaneous SEDs of 3C 66A at various times during our campaign. Signifi-
cant variability was only detected in the radio, IR, and optical
bands. In those bands, we extracted SEDs for four epochs: dur-
ing two major outbursts around 2003 November 1 and December
28, during a minor outburst around 2003 October 1, and during
a rather quiescent state around 2003 November 11. Dereddened
optical and near-IR magnitudes were converted to fluxes using
the zero-point normalizations of Bessel et al. (1998). The result-
ing SEDs are plotted in Figure 15. Due to the relatively poor
sampling, radio and IR data were often not quite simultaneous
with the optical spectra, which were extracted near the peaks
of the individual outbursts mentioned above. In this case, the
closest near-IR and radio data points were chosen. This led to
time offsets between the optical and the radio data points of up
to $\sim 5$ days. Given the relatively long timescale and moderate
amplitude of variability at radio wavelengths, we are confident
that this did not introduce serious distortions of the low-frequency
SED.

In addition to the data taken during our 2003–2004 cam-
paign, we have included historical X-ray measurements to in-
dicate the degree of X-ray variability observed in this source
and the historical average GeV $\gamma$-ray flux measured by the EGRET
instrument on board the Compton Gamma-Ray Observatory from
five observations between 1991 November and 1995 September
(Hartman et al. 1999).

The shape of the time-averaged optical ($U$) $BVRI$ spectra, to-
gether with the very steep X-ray spectral index, indicates that

\[
\nu F_\nu \propto \nu^{-1.2} \quad \text{for } 0.1 \text{ keV} \leq \nu \leq 100 \text{ keV}
\]

the $\nu F_\nu$ peak of the synchrotron component of 3C 66A is typ-
ically located in the optical range. The shape of our best-fit
$RXTE$ spectrum provides strong evidence that the synchrotron
component extends far into the X-ray regime and intersects the
high-energy component in the time-averaged SED of 3C 66A at
$\geq 100$ GeV during our core campaign period. Due to the lack
of simultaneous GeV $\gamma$-ray coverage and of a firm detection at
$>100$ GeV, we cannot make a precise statement concerning the
level of $\gamma$-ray emission during our campaign. However, if the
historical EGRET flux is representative also for the time of our
campaign, then the total energy output in the low-frequency
and the high-frequency components of the SED of 3C 66A are
comparable, as is typical for the class of intermediate- and low-
frequency peaked BL Lac objects.

6. GENERIC PARAMETER ESTIMATES

In this section we discuss some general constraints on source
parameters. We first (§ 6.1) focus on parameter estimates using
the SED and optical intraday variability measurements, most
relevant to the innermost portion of the jet outflow, closest to the
central engine. In § 6.2 we use the results of our VLBA obser-
vations to estimate parameters of the relativistic flow on parsec
scales.

6.1. Parameters of the Inner Jet

From the minimum variability timescale of $\Delta t_{\text{min}} \sim 2$ hr, we
can estimate the size of the emitting region as

\[
R \approx c D \Delta t_{\text{min}}
\]

where $D = \Gamma(1 - \beta_1 \cos \theta_{\text{obs}})^{-1}$, where $\Gamma$ is the bulk Lorentz
factor of the emitting region, $\beta_1 c$ is the corresponding speed,
and $\theta_{\text{obs}}$ is the observing angle. This yields $R \approx 2.2 \times 10^{15} \delta_1 \text{ cm}$,
where $\delta = 10 \beta_1 \approx 15$ as an estimate from the limits on super-
luminal motion and from previous modeling efforts, as men-
tioned in the introduction.

An estimate of the comoving magnetic field can be found
by assuming that the dominant portion of the time-averaged
The synchrotron spectrum is emitted by a quasi-equilibrium power-law spectrum of electrons with \( N_e(\gamma) = n_0 V g \gamma^{-p} \) for \( \gamma_1 \leq \gamma \leq \gamma_2 \); here \( V g \) is the comoving blob volume. Based on the X-ray spectral index of \( \alpha \approx 1.5 \), we find a particle spectral index of \( p \approx 4 \). Since the synchrotron cooling timescale for X-ray-emitting electrons might be much shorter than the dynamical timescale (see eq. [3]), the X-rays are likely to be produced by a cooled electron distribution. In this case, the index \( p \approx 4 \) corresponds to a distribution of electrons injected into the emission region with an original index \( q \approx 3 \). The normalization constant \( n_0 \) is related to the magnetic field through an equipartition parameter \( e_B \equiv u_B/u_e \) (in the comoving frame). Note that this equipartition parameter only refers to the energy density of the electrons, not accounting for a (possibly dominant) energy content of a hadronic matter component in the jet. Under these assumptions, the magnetic field can be estimated as described, e.g., in Böttcher et al. (2003). Taking the \( \nu F_\nu \) peak synchrotron flux \( F_\nu \) at the dimensionless synchrotron peak energy \( \epsilon_{\text{sy}} \approx 5 \times 10^{-11} \) ergs cm\(^{-2}\) s\(^{-1}\) at \( \epsilon_{\text{sy}} \approx 5 \times 10^{-6} \) and \( R \approx 3.3 \times 10^{15} \) cm, we find

\[
B_{\text{so}} = 4.4 \epsilon^{-1}_B \epsilon^{2/7}_{\text{so}},
\]

which yields \( B_{\text{so}} \approx 2.9 \epsilon^{-7/3}_{\text{so}} \) G for \( \delta = 15 \).

We can further use this magnetic-field value to estimate the end points of the electron spectrum since the low-energy end, \( \gamma_1 \), might correspond to the \( \nu F_\nu \) peak of the synchrotron spectrum, and the synchrotron high-energy cutoff corresponds to \( \gamma_2 \). Generally, we find

\[
\gamma \approx 3.1 \times 10^3 \nu_{15}^{1/2} \left( \frac{\delta}{15} \right)^{-1/2} \left( \frac{B}{2.9 \text{ G}} \right)^{-1/2},
\]

where \( \nu_{15} \) is the characteristic synchrotron frequency in units of \( 10^{15} \) Hz. With our standard parameters, this yields \( \gamma_1 \approx 3.1 \times 10^3 \) and \( \gamma_2 \approx 1.5 \times 10^5 \) if the synchrotron cutoff occurs around 10 keV. We can also use this to estimate the synchrotron cooling timescale of electrons in the observer’s frame:

\[
\tau_{\text{cool, sy}}^{\text{obs}} \approx 2.8 \times 10^3 \left( \frac{\delta}{15} \right)^{-1/2} \left( \frac{B}{2.9 \text{ G}} \right)^{-3/2} \nu_{15}^{1/2} \text{ s}.
\]

For optical frequencies, this yields observed cooling timescales of the order of \( \sim 2 \) hr, in agreement with the observed minimum variability timescale. This, however, raises an important caveat. The observed minimum variability timescale may, in fact, be a reflection of the electron cooling timescale rather than the dynamical timescale, as we had assumed when choosing our estimate for the source size \( R \). For this reason, a more detailed future investigation of possible short-term variability at X-ray frequencies will be extremely important to resolve this issue. If X-ray variability on shorter timescales than \( \sim 2 \) hr is found, the emission region would then have to be more compact than the \( R \approx 3 \times 10^{15} \) cm that we had estimated here, and X-ray spectral hysteresis patterns could arise. In contrast, if the X-ray variability timescale is found to be consistent with the optical one, then it would have to be dominated by the dynamical timescale and thus be largely achromatic. Thus, in that case no significant X-ray spectral hysteresis would be expected.

6.2. Parameters of the Parsec-Scale Outflow

In our analysis of the 22 and 43 GHz VLBA maps of 3C 66A, we had found a rather smooth jet, with only one of the six Gaussian components (C1) showing evidence for superluminal motion of \( \beta_{\text{app}} = (8.5 \pm 5.6) h^{-1} \). If we consider the measured speed of the component C1 to be close to the maximum as observed under the superluminal angle given by \( \cos(\theta_{\text{SL}}) = \beta_{\text{G}} \), we can estimate the minimum Lorentz factor and maximum angle between the jet and our line of sight. With \( h = 0.7 \) and by taking the lower limit of the \( \beta_{\text{app}} \geq (8.5-5.6) h^{-1} = 2.9 h^{-1} \), we have a lower limit to the bulk Lorentz factor of \( \Gamma \geq 4.3 \) and an upper limit to the viewing angle of \( \theta_{\text{obs}} \leq 27.2 \). The large error in \( \beta_{\text{app}} \) results in very weak constraints on \( \Gamma \) and \( \theta_{\text{obs}} \), and they should be considered as very conservative estimates. Much more accurate values should be available after the analysis of all nine epochs of the VLBA monitoring program (T. Savolainen et al. 2005, in preparation).

From the measured fluxes at 22 and 43 GHz, we have calculated the brightness temperatures along the jet. The largest brightness temperatures are found for the core, yielding lower limits of \( T_{b \text{c}} \approx 10^{11} \text{ K} \) (see Table 6). Here we have to mention the caveat that the errors for the size of the core are very difficult to estimate, since our model-fitting yields very asymmetric probability density distributions for the core size. Essentially, the data are also consistent with a point-source core, so we can only give an upper limit to the core size. Thus, no upper limit to the core brightness temperature is available. Our calculated best-fit core brightness temperatures and corresponding lower limits are listed in Table 6 and can be compared with an equipartition brightness temperature (Readhead 1994) for 3C 66A of \( T_{b \text{c}, \text{sp}} = 6 \times 10^{10} \text{ K} \). If the equipartition assumption holds, the minimum Doppler factor is around 2. The same result was obtained by Lähteenmäki & Valtaoja (1999), who used variability arguments to calculate the Doppler factor. Because we do not have upper limits for the brightness temperature, we cannot calculate a more precise value for the Doppler factor. However, if the core brightness temperatures calculated from the best model-fit values are not too far from the truth, the core Doppler factor probably lies somewhere between 2 and 15, with the average best-fit value of 6.

We have also calculated the brightness temperatures for components other than the core at both 22 and 43 GHz. In Figure 16 the component brightness temperatures are plotted against the component distance from the core. The figure illustrates that the points form two well-defined power laws with a different variability timescale. This, however, raises an important caveat. The observed minimum variability timescale may, in fact, be a reflection of the electron cooling timescale rather than the dynamical timescale, as we had assumed when choosing our estimate for the source size \( R \). For this reason, a more detailed future investigation of possible short-term variability at X-ray frequencies will be extremely important to resolve this issue. If X-ray variability on shorter timescales than \( \sim 2 \) hr is found, the emission region would then have to be more compact than the \( R \approx 3 \times 10^{15} \) cm that we had estimated here, and X-ray spectral hysteresis patterns could arise. In contrast, if the X-ray variability timescale is found to be consistent with the optical one, then it would have to be dominated by the dynamical timescale and thus be largely achromatic. Thus, in that case no significant X-ray spectral hysteresis would be expected.

### Table 6

<table>
<thead>
<tr>
<th>Epoch</th>
<th>( T_{b \text{c}} ) (K)</th>
<th>( T_{b \text{c, min}} ) (K)</th>
</tr>
</thead>
<tbody>
<tr>
<td>2003.78</td>
<td>( 2.7 \times 10^{11} )</td>
<td>( 1.1 \times 10^{11} )</td>
</tr>
<tr>
<td>2003.83</td>
<td>( 2.0 \times 10^{11} )</td>
<td>( 1.0 \times 10^{11} )</td>
</tr>
<tr>
<td>2004.08</td>
<td>( 6.6 \times 10^{10} )</td>
<td>( 1.4 \times 10^{11} )</td>
</tr>
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</table>
wave caused by the interaction of the continuous, relativistic outflow with the observed kink.

The power law gradient of the brightness temperature along the jet suggests that 3C 66A exhibits a smooth parsec-scale jet without any prominent particle acceleration sites (shocks) other than the core and possibly the bend in region B. If we assume that the magnetic field \( B \), the electron density \( N \) and the cross-sectional diameter \( D \) of the jet can also be described by the power laws \( N \propto r^n \), \( B \propto r^b \), and \( D \propto r^d \), then the brightness temperature is expected to fall with \( r \) as \( T_b \propto r^a \) (with \( s < 0 \)). If optically thin synchrotron emission and a constant Lorentz factor of the emitting electrons are assumed, \( s = d + n + b(1 + \alpha) \), where \( \alpha \) is the spectral index (see, e.g., Kadler et al. 2004). If we know the values of \( s, d, \) and \( \alpha \), we can calculate the relationship between \( b \) and \( n \), or, by assuming equipartition, the actual values of \( b \) and \( N \). We know now that \( s = -2 \), and the average spectral index is \( \langle \alpha_{22-43} \rangle = 0.15 \).

We still need to find out the value of \( d \). The value \( d = 1 \) would correspond to a constant opening angle (conical jet), but our imaging results allow us to estimate a more precise value from the data by plotting the component size versus distance from the core (see Fig. 17). Surprisingly, up to 2.5 mas from the core, the jet is very tightly collimated. A power-law fit to the observed correlation between the lateral jet size and distance from the core, at 22 GHz (\( \times \) open symbols), for the three epochs: 2003.78 (circles), 2003.83 (squares), and 2004.08 (stars). The solid lines are our best-fit power laws to the brightness-temperature profiles; the vertical dashed lines indicate the jet regions B2 and B3, where a significant deviation from the power-law profile is found, coincident with an apparent kink in the jet flow direction.

jet (\( d = 0.6 \)), it decays as \( B \propto D^{-1.3} \), implying a magnetic field that is predominantly transverse to the jet axis. This is in accordance with other observational results on the magnetic field orientation in BL Lac objects in general (see, e.g., Gabuzda et al. 2004). A magnetic field being predominantly transverse to the jet agrees well with overall toroidal field configuration, which is expected in Poynting flux dominated jet models. Another possibility for creating predominantly transverse magnetic field is the enhancement of the transverse magnetic field component by a series of shocks, but this seems to be less likely in this case, since we do not observe bright knots in the jet, but a rather smooth flow.

7. SUMMARY

We have observed 3C 66A in a massive multiwavelength monitoring campaign from 2003 July to 2004 April. Monitoring observations were carried out at radio, infrared, optical, X-ray, and VHE \( \gamma \)-ray observations. The main observational results of our campaign are the following:

1. In the optical, several outbursts by \( \Delta m \sim 0.3-0.5 \) over timescales of several days were observed.
2. Optical intraday microvariability (\( \Delta m \leq 0.05 \)) on timescales of \( \Delta t_{\text{var}} \sim 2 \) hr was detected.
3. No clear evidence for periodicity was found, but a quasiregular sequence of several major outbursts separated by \( \sim 50-57 \) days was identified.
4. Large optical flares (on timescales of several days) seem to exhibit optical spectral hysteresis, with the \( B-R \) hardness peaking several days prior to the \( R \) and \( B \)-band flux peaks.
5. The 3–10 keV X-ray spectrum is best fitted with a single power law with energy index \( \alpha = 1.47 \pm 0.56 \), indicating that the transition between the synchrotron and the high-energy components occurs at photon energies of \( \geq 10 \) keV.
6. Radio VLBA monitoring observations reveal a rather smooth jet with no clearly discernible knots or hot spots except for moderate brightening in region B, where the jet is bending.
7. Decomposition of the VLBA radio structure into Gaussian components reveals superluminal motion only for one of six
components. Its apparent speed is \( v_{\text{app}} = (8.5 \pm 5.6) \text{\,hr}^{-1} = (12.1 \pm 8.0) \) for \( h = 0.7 \).

8. The radio brightness temperature profile along the jet, along with its observed geometry, \( D \propto r^{6.6} \), suggest a magnetic field decay \( B \propto r^{-1.3} \), indicating a predominantly perpendicular magnetic field orientation.

9. STACEE observations revealed a 2.2 \( \sigma \) excess, which provided strict upper limits at \( E_\gamma \gtrsim 150 \, \text{GeV} \). Additional VHE \( \gamma \)-ray limits at \( E_\gamma > 390 \, \text{GeV} \) resulted from simultaneous Whipple observations.

10. The broadband SED of 3C 66A during our campaign suggests that the synchrotron component peaks in the optical and extends far into the X-ray regime, out to at least \( \sim 10 \, \text{keV} \).

11. The following parameters of the synchrotron emission region near the core can be estimated:

\[
\delta \approx 15, \\
R \approx 3.3 \times 10^{15} \, \text{cm}, \\
B \approx 2.9 \, \text{eV}^{-1} \, \text{cm}^{-3}, \\
\gamma_1 \approx 3.1 \times 10^3, \\
\gamma_2 \approx 1.5 \times 10^4, \\
p \approx 4.
\]

No X-ray variability was detectable by RXTE on a \( \sim 1 \) day timescale. Future X-ray observations with more sensitive X-ray detectors will be important to probe for rapid X-ray variability and X-ray spectral hysteresis in order to put more stringent constraints on the source size, the nature of the variability mechanism, and the composition and energetics of the emitting plasma in the jet.

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REFERENCES

Carini, M. T., & Miller, H. R. 1991, BAAS, 23, 1420
——. 2004a, A&A, 421, 103
Multifrequency VLBA monitoring of 3C 273 during the INTEGRAL Campaign in 2003

I. Kinematics of the parsec scale jet from 43 GHz data

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ABSTRACT

In this first of a series of papers describing polarimetric multifrequency Very Long Baseline Array (VLBA) monitoring of 3C 273 during a simultaneous campaign with the INTEGRAL γ-ray satellite in 2003, we present 5 Stokes I images and source models at 7 mm. We show that a part of the inner jet (1–2 milliarcsec from the core) is resolved in a direction transverse to the flow, and we analyse the kinematics of the jet within the first 10 mas. Based on the VLBA data and simultaneous single-dish flux density monitoring, we determine an accurate value for the Doppler factor of the parsec scale jet, and using this value with observed proper motions, we calculate the Lorentz factors and the viewing angles for the emission components in the jet. Our data indicates a significant velocity gradient across the jet with the components travelling near the southern edge being faster than the components with more northern path. We discuss our observations in the light of jet precession model and growing plasma instabilities.

Key words. galaxies: active – galaxies: jets – galaxies: quasars: individual: 3C 273

1. Introduction

Although vigorously investigated for over forty years, the physics of active galactic nuclei (AGN) have proved to be far from easily understandable. A set of basic facts about the nature of the central engine being a supermassive black hole and the gravitational accretion process as the primary energy source have been established beyond reasonable doubt, but the details of the picture do not quite fit. The intricate physics of AGN is made difficult by the complex interplay between several emission mechanisms working simultaneously. For example, blazars, the class unifying flat-spectrum radio-loud quasars and BL Lac objects, contains sources that show significant radiation throughout the whole electromagnetic spectrum from radio to TeV γ-rays. The different emission mechanisms in AGNs include e.g. non-thermal synchrotron radiation from a relativistic jet, thermal emission from dust, optical-ultraviolet emission from the accretion disk, Comptonized hard X-rays from the accretion disk corona and γ-rays produced by inverse Compton scattering in the jet. Therefore, observing campaigns monitoring the sources simultaneously at as many wavelengths as possible are essential in studying AGNs.

Since the relativistic jet plays a crucial role in the objects belonging to the blazar class of AGNs, Very Long Baseline Interferometry (VLBI) observations imaging the parsec scale jet are an important addition to campaigns concentrating on this class. In particular, combined multifrequency VLBI monitoring and γ-ray measurements with satellite observatories may provide a solution to the question of the origin of the large amount of energy emitted in γ-rays by some blazars. By studying time lags between γ-ray flares and ejections of superluminal components into the VLBI jet, we should be able to establish whether the high energy emission comes from a region close to the central engine or whether it instead originates in the radio core or in the relativistic shocks – parsecs away from the black hole. In addition, multifrequency VLBI data yields spectra of the radio core and superluminal knots, which can be used to calculate the anticipated amount of the high energy emission due to the synchrotron self-Compton (SSC) process if an accurate value of the jet Doppler factor is known. Predicted SSC-flux
can be directly compared with simultaneous hard X-ray and γ-ray observations.

Due to its proximity, 3C 273 (\(z = 0.158\); Schmidt 1963) – the brightest quasar on the celestial sphere – is one of the most studied AGNs (see Courvoisier 1998, for a comprehensive review). Because 3C 273 exhibits all aspects of nuclear activity including a jet, a blue bump, superluminal motion, strong γ-ray emission and variability at all wavelengths, it was chosen as the target of an INTEGRAL campaign, where radiation above 1 keV was monitored in order to study the different high energy emission components (Courvoisier et al. 2003). INTEGRAL observed 3C 273 in January, June–July 2003, and January 2004 and the campaign was supported by ground based monitoring at lower frequencies carried out at several different observatories (Courvoisier et al. 2004). This article is the first in a series of papers describing a multifrequency polarimetric VLBI monitoring of 3C 273 using the Very Long Baseline Array (VLBA) in conjunction with the INTEGRAL campaign during 2003. The amount of data produced by such a VLBI campaign using six frequencies is huge. In order to divide the large volume of the gathered information into manageable chunks, we will treat the different aspects of the data (i.e. kinematics, spectra and polarisation) in different papers, and in the end we will attempt to draw a self-consistent picture taking into account the whole data set.

In the current paper we construct a template for the later analysis of the component spectra and polarisation, and for the comparison of the VLBI data with the INTEGRAL measurements. Since kinematics is important in all subsequent analyses, we start by studying the component motion in the jet within 10 mas from the radio core. The emphasis is on the components located close to the core (≤2 mas), since they are the most relevant for the combined analysis of the VLBI and INTEGRAL data. We use the 43 GHz data for the kinematical study, since they provide the best angular resolution with a good signal-to-noise ratio (SNR). The other frequencies are not discussed in this paper, since the positional errors at all other frequencies are significantly larger for the components we are interested in (i.e. those located near the core). Also, a kinematical analysis that takes into account multiple frequencies in a consistent way is very complicated using contemporary model fitting methods. However, the other frequencies are not ignored either; instead, we will check the compatibility of this paper’s results with the multifrequency data presented in the forthcoming papers treating the spectral information. In Paper II (Savolainen et al., in preparation) we will discuss the method to obtain the spectra of individual knots in the parsec scale jet, and present the spectra from the February 2003 observation. We will calculate the anticipated amount of SSC-radiation, and compare this with the quasi-simultaneous INTEGRAL data. In Papers III and IV, spectral evolution and polarisation properties of the superluminal components over the monitoring period will be presented. Paper V combines all the obtained information, to interpret it in a self-consistent way.

Throughout the paper, we use a contemporary cosmology with the following parameters: \(H_0 = 71\ \text{km} \text{s}^{-1} \text{Mpc}^{-1}\), \(\Omega_M = 0.27\) and \(\Omega_{\Lambda} = 0.73\). For the spectral index, we use the positive convention: \(S\propto\nu^{\alpha}\).

2. Observations and data reduction

We observed 3C 273 with the National Radio Astronomy Observatory’s Very Long Baseline Array at five epochs in 2003 (February 28th, May 11th, July 2nd, September 7th and November 23rd) for nine hours at each epoch, using six different frequencies (5, 8.4, 15, 22, 43 and 86 GHz). The observations were carried out in the dual-polarisation mode enabling us also to study the linear polarisation of the source. However, we do not discuss the polarisation data here; it will be published in Paper IV (Savolainen et al. in preparation). At the first two epochs, 3C 279 was observed as a polarisation calibrator and as a flux calibrator. At the three later epochs, we added two more EVPA calibrators to the program: OJ 287 and 1611+343.

The data were calibrated at Tuorla Observatory using NRAO’s Astronomical Image Processing System (AIPS; Bridle & Greisen 1994) and the Caltech Difmap package (Shepherd 1997) was used for imaging. Standard procedures of calibration, imaging and self-calibration were employed, and final images were produced with the Perl library FITSPi2.

Accuracy of a priori flux scaling was checked by comparing extrapolated zero baseline flux densities of a compact calibrator source, 3C 279, to near-simultaneous observations from VLA polarisation monitoring program (Taylor & Myers 2000) and from Metsähovi Radio Observatory’s quasar monitoring program (Teräsranta et al. 2004). The comparison shows that the a priori amplitude calibration of our data at 43 GHz is accurate to ~5%, which is better than the often quoted nominal value of 10%. A detailed discussion of the amplitude calibration and setting of the flux scale will be given in Paper II.

The angular resolution of the VLBI data is a function of observing frequency, \(uv\)-coverage and signal-to-noise ratio. Only 7 antennas out of 10 were able to observe at 86 GHz during our campaign (Brewster and St. Croix did not have 3 mm receivers, and the baselines to Hancock did not produce fringes at any observation epoch), and hence, the longest baselines (Mauna Kea – St. Croix, Mauna Kea – Hancock) were lost. This, combined with a significantly lower SNR, results in poorer angular resolution at 86 GHz than at 43 GHz, and thus, unfortunately, 3 mm data does not provide any added value for the kinematical study as compared to 43 GHz. (However, for the spectral analysis 3 mm data is invaluable, as will be shown in Paper II.) At lower frequencies, the \(uv\)-coverage is identical to the one at 43 GHz, and, although the SNR is better, the eventually achieved angular resolution in model fitting is poorer. Currently, there is no model fitting procedure capable of fitting multiple frequencies simultaneously, which makes utilising multifrequency data in the kinematical analysis quite difficult due to different positional uncertainties at different frequencies, and due to a frequency dependent core position. For these reasons, the kinematical analysis, which provides the

\footnote{The VLBA is a facility of the National Radio Astronomy Observatory, operated by Associated Universities, Inc., under cooperative agreement with the US National Science Foundation.}

\footnote{http://personal.denison.edu/~homand/}
needed template for further study of spectra and polarisation of the radio core and the components close to it, is done at the single frequency maximising the angular resolution i.e. 43 GHz. Obviously, the results are preliminary in that the later analysis of the spectra and polarisation will provide a check on the kinematics presented in this paper.

2.1. Modelling and estimation of model errors

The final CLEAN images of 3C 273 (Fig. 1) are not easy to parameterise for quantitative analysis of e.g. proper motions and flux density evolution of different regions in the jet. To make a quantitative analysis feasible, we fitted models consisting of a number of elliptical and circular Gaussian components to fully self-calibrated visibility data (i.e. in the $(u,v)$-plane) using the “modelfit” task in Difmap. We sought to obtain the simplest model that gave a good fit to the visibilities and produced, when convolved with the restoring beam, a brightness distribution similar to that of the CLEAN image. 3C 273 is a complex source with a number of details, and due to its structure that could not easily be represented by well-separated, discrete components, model-fitting proved to be rather difficult. Due to this complexity, we added more constraints to the fitting process by demanding consistency between epochs, i.e. most components should be traceable over time. The final models show a good fit to the visibility data, and they nicely reproduce the structure in the CLEAN images (compare Figs. 1 and 2, the first showing CLEAN images and the latter presenting model components convolved with a Gaussian beam). However, these models (like any result from the model-fitting of the VLBI data) are not unique solutions, but rather show one plausible representation of the data.

Reliable estimation of uncertainties for parameters of the model fitted to VLBI data is a notoriously difficult task. This is due to loss of knowledge about the number of degrees of freedom in the data, because the data are no longer independent after self-calibration is applied. This makes the traditional $\chi^2$-methods useless, unless a well-justified estimate of degrees of freedom can be made. Often there is also strong interdependency between the components – for instance, adjacent components’ sizes and fluxes can be highly dependent on each other. This can be solved by using a method first described by Tzioumis et al. (1989), where the value of the model parameter under scrutiny is adjusted by a small amount and fixed. The model-fitting algorithm is then run and results are inspected. This cycle is repeated until a clear discrepancy between the model and the visibilities is achieved. Difwrap (Lovell 2000) is a program to automatically carry out this iteration, and we have used it to assess uncertainties for parameters of the model components. However, there is a serious problem also with this approach. Since the number of degrees of freedom remains unknown, we cannot use any statistically justified limit to determine what is a significant deviation from the data; we must rely on highly subjective visual inspection of the visibilities and of the residual map. We found that a reasonable criterion for a significant discrepancy between the data and model is the first appearance of such an emission structure, which we would have cleaned, had it appeared during the imaging process. For faint components, the parameters can change substantially with
insignificant increase in the difference between the model and the visibilities, and hence, with low flux components we cannot rely on the results given by Difwrap, but we need to use other means to estimate the uncertainties. The above-described procedure for determination of uncertainties in model parameters does not address errors arising from significant changes in the model (e.g. changes in the number of components) or from imaging artefacts created by badly driven imaging/self-calibration iteration.

Homan et al. (2001) estimated the positional errors of the components by using the variance about the best-fit proper motion model. We applied their method to our data in order to estimate the positional uncertainties of weak components and to check the plausibility of the errors given by Difwrap for stronger components. We have listed the average positional error $\langle \Delta r_{DW} \rangle$ from the Difwrap analysis and the r.m.s. variation about the best-fit proper motion model $\sigma(r)$ for each component that could be reliably followed between the epochs in Table 2. The errors are comparable – for components B1–B4\(^3\) they are almost identical and for components C1–C3 the uncertainties obtained by Difwrap are 0.01–0.02 mas larger than the r.m.s. fluctuations about the proper motion model. Diffuse components A1, A2 and A3 have Difwrap positional errors approximately twice as large as the variance about the best-fit model. We conclude that Difwrap yields reliable estimates for the positional uncertainties of the compact and rather bright components. As was mentioned earlier, it is difficult to obtain an error estimate for weak components with this method, and for diffuse components, the errors seem to be exaggerated. The average errors given in Table 2 agree well with the often quoted “1/5 of the beam size” estimate for positional accuracy

\(^3\) The naming convention of the components in this paper is our own and does not follow any previous practice.
Uncertainty in the flux density of a Gaussian component strongly depends on possible "flux leakage" between adjacent components, which is, in principle, taken into account in the Difwrap analysis. This flux leakage results in large and often asymmetric uncertainties, as can be seen in Figs. 11 and 12, where flux density evolution of the individual components is presented. The error bars in the figures correspond to the total uncertainty in the component flux density, where we have quadratically added the model error given by Difwrap and a 5% amplitude calibration error. The errors are on average 10% for moderately bright and rather isolated components with no significant flux leakage from other components, and ~15% for components with the leakage problem. For some bright and well-isolated components, flux density errors less than 10% were found, but in general uncertainties reported here are larger than those obtained by the $\chi^2$-method, which does not take into account the interdependency between nearby components’ flux densities (e.g. Jorstad et al. 2001). Naturally, the caveat in the Difwrap method is the subjective visual inspection needed to place the error boundaries, but we note that Homan et al. (2002), who estimated flux density errors by examining the correlation of flux variations between two observing frequencies and taking an uncorrelated part as a measure of uncertainty, found flux density errors ranging from 5% to 20% at 22 GHz with 5%–10% being typical values for well-defined components. Our error estimates for component flux densities seem to agree with these values.

Uncertainties in the component sizes were also determined with Difwrap. Plausible values for errors were found only for bright components, and again, asymmetric uncertainties emerged commonly. The uncertainties in the size of the major axis of the component ranged from 0.01 mas to 0.15 mas corresponding to approximately 5%–30% relative uncertainty. Errors in the component axial ratios were difficult to determine with Difwrap, and results were obtained only for a small number of components. The axial ratio uncertainties range from 0.05 to 0.20, corresponding to a relative uncertainty of about 20%–60%. The large errors in the component sizes are again due to “leakage” between adjacent components. For well-isolated components, uncertainties in size would be smaller.

### 3. The structure of the parsec scale jet

In our 43 GHz, high dynamic range images (Fig. 1; see also Table 1) we can track the jet up to a distance of ~11 mas from the core. The “wiggling” structure reported in several other studies (e.g. Mantovani et al. 1999; Krichbaum et al. 1990; Báth et al. 1991; Zensus et al. 1988) is evident also in our images – the bright features are not collinear.

At each epoch, we consider the compact, easternmost component as the stationary radio core, and all proper motions are measured relative to it. As can be seen from Fig. 2, the innermost region “C” with the core and several departing components (C1–C3) extends about 1 mas to south-west. Besides the core, there is also another stationary component, S1, located ~0.15 mas from the core. Identification of the components close to the core over the epochs is problematic due to the number of knots close to each other, and also because new ejections are going on during our observing period. We have plotted the components within 1.5 mas from the core in Fig. 3 with the size of the symbol indicating the relative flux of each Gaussian component. The figure shows four moving components (C1–C3 and B5), and two stationary components (the core and S1), with consistent flux evolution. The figure also confirms that component C3 is ejected from the core during our monitoring period. Components C1–C3 may not necessarily represent independent and distinct features, but they also can be interpreted as a brightening of the inner jet – e.g. due to continuous injection of energetic particles in the base of the jet. VLBI data with better angular resolution could settle this, but, unfortunately, our 86 GHz observations suffer from the loss of the longest VLBA baselines, and thus, they cannot provide the necessary resolution. However, this should not affect the following kinematical analysis.

In the complex emission region at about 1–2.5 mas to southwest from the core, the jet is resolved in the direction transverse to the flow. We have identified five components (B1–B5) that are traceable over the epochs. The components can be followed from epoch to epoch in Fig. 2. B3 is represented by two Gaussian components at the fourth epoch, and B5 is not seen after the third epoch, but otherwise we consider that our identification is robust. B4 seems to catch up with the slower knot B3 and collides with it (see Figs. 6 and 7).

### Table 1. Parameters of the images in Fig. 1.

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<tbody>
<tr>
<td></td>
<td>[yr]</td>
<td>[mas]</td>
<td>[mas]</td>
<td>['']</td>
<td>[mJy beam$^{-1}$]</td>
</tr>
<tr>
<td>2003.16</td>
<td>0.44</td>
<td>0.18</td>
<td>~7.1</td>
<td>2.4</td>
<td>3961</td>
</tr>
<tr>
<td>2003.36</td>
<td>0.40</td>
<td>0.18</td>
<td>~3.5</td>
<td>2.2</td>
<td>2750</td>
</tr>
<tr>
<td>2003.50</td>
<td>0.43</td>
<td>0.19</td>
<td>~4.6</td>
<td>2.0</td>
<td>2600</td>
</tr>
<tr>
<td>2003.68</td>
<td>0.39</td>
<td>0.17</td>
<td>~6.3</td>
<td>1.7</td>
<td>2241</td>
</tr>
<tr>
<td>2003.90</td>
<td>0.44</td>
<td>0.18</td>
<td>~5.5</td>
<td>2.4</td>
<td>1997</td>
</tr>
</tbody>
</table>

### Table 2. Average positional uncertainties for the model components from the Difwrap ($\langle\Delta r_{\text{DW}}\rangle$) and from the variance analysis ($\sigma(r)$).

<table>
<thead>
<tr>
<th>Component</th>
<th>$\langle\Delta r_{\text{DW}}\rangle$</th>
<th>$\sigma(r)$</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>[mas]</td>
<td>[mas]</td>
</tr>
<tr>
<td>C1</td>
<td>0.03</td>
<td>0.01</td>
</tr>
<tr>
<td>C2</td>
<td>0.04</td>
<td>0.02</td>
</tr>
<tr>
<td>C3</td>
<td>0.04</td>
<td>0.02</td>
</tr>
<tr>
<td>B1</td>
<td>0.04</td>
<td>0.04</td>
</tr>
<tr>
<td>B2</td>
<td>0.03</td>
<td>0.03</td>
</tr>
<tr>
<td>B3</td>
<td>0.05</td>
<td>0.05</td>
</tr>
<tr>
<td>B4</td>
<td>0.05</td>
<td>0.03</td>
</tr>
<tr>
<td>B5</td>
<td>–</td>
<td>0.06</td>
</tr>
<tr>
<td>A1</td>
<td>0.10</td>
<td>0.05</td>
</tr>
<tr>
<td>A2</td>
<td>0.11</td>
<td>0.06</td>
</tr>
<tr>
<td>A3</td>
<td>0.06</td>
<td>0.02</td>
</tr>
</tbody>
</table>
The jet broadens significantly after ~1.5 mas and, as shown later in Sect. 4, the trajectory of the component B1 does not extrapolate to the core (Fig. 2). We discuss the possible explanations for this anomalous region “B” in Sect. 6. Farther down the jet, there are two more emission regions visible in the 43 GHz images, both of them showing a curved structure (Fig. 1). The first one is located 3–6 mas southwest of the core and we have identified two components A2 and A3 in it. This feature likely corresponds to knots B1, bs, b1, B2 and b2 of Jorstad et al. (2005), with our A2 being their B1 and A3 being their B2. However, this cross-identification is not certain, since the region is complex and, as described in Jorstad et al. (2005), it contains components with differing speeds and directions as well as secondary components connected to the main disturbances. The second emission region in the outer part of the jet is located at 9–11 mas from the core, and we call this component with a curved shape A1 (Fig. 1).

4. Kinematics of the jet

The data allow us to follow the evolution of the components identified in the previous section, and to measure their proper motion, which is the basis for analysing physical properties of the jet. We follow the practice of Jorstad et al. (2005), and define the average proper motion as a vector \( \langle \mu, \Phi \rangle \), where \( \langle \mu \rangle = \sqrt{\langle \mu_{\text{RA}} \rangle^2 + \langle \mu_{\text{Dec}} \rangle^2} \) is the mean angular speed and \( \Phi = \tan^{-1}(\langle \mu_{\text{Dec}} \rangle / \langle \mu_{\text{RA}} \rangle) \) is the direction of motion. In all but one case, the average coordinate motions \( \langle \mu_{\text{RA}} \rangle \) and \( \langle \mu_{\text{Dec}} \rangle \) are obtained by a linear least-squares fit to the relative (RA, Dec) positions of a component over the observation epochs. For component C1, a second-order polynomial was fitted instead of a first-order one, because it reduced the \( \chi^2 \) for the RA-coordinate fit from 4.7 (with 3 degrees of freedom) to 0.2 (with 2 degrees of freedom). Also, the (RA, Dec)-plot in Fig. 4 indicates a non-ballistic trajectory for component C1. The amount of acceleration measured for C1 is 0.21 ± 0.18 mas yr\(^{-2} \) – a marginal detection (Fig. 5).

In Figs. 4, 6, and 8, we have plotted the trajectories in the (RA, Dec)-plane for components C1–C3, B1–B5 and A1–A3, respectively. As can be seen from these figures, there are three components showing non-radial motion, C1, B1, and A2. As was discussed earlier, C1 seems to have a curved trajectory, while B1 and A2 have straight trajectories, which, however, do not extrapolate back to the core. B1 has an especially interesting path, since its motion seems to be parallel to that of B2, but clearly deviating from those of B3 and B4, which have a direction of ~25° more southern than B1. Also, B1 and B2 have similar speeds, while B3 and B4 are both faster. B1 might have formed from B2 or B3, possibly through an interaction between relativistic shocks connected with these components and the underlying flow, or it might have changed its trajectory.

In Figs. 5, 7, and 9, we have plotted the separation from the core as a function of time for the components. Table 4 lists the maximum measured flux density, \( S_{\text{max}} \), the average proper...
motion, \((\mu, \Phi)\), the apparent superluminal speed, \(\beta_{\text{app}}\), and the extrapolated epoch of zero-separation, \(T_0\), for the components. Ejection time is not given for the component S1, which is stationary, nor the components B1 and A2, whose trajectories do not extrapolate back to the core. The uncertainties in the average proper motion are derived from the uncertainties of the fitted polynomial coefficients.

The superluminal velocities for the moving components range from 4.6 \(h^{-1}\) c to 13.0 \(h^{-1}\) c (with \(H_0 = 100\ h\)). This range of velocities is similar to those previously reported by e.g. Zensus et al. (1990), Krichbaum et al. (2000), and Jorstad et al. (2005). The largest observed apparent speed, 13.0 \(h^{-1}\) c for B5, is based only on three epochs of observations, and the component is weak having flux density below 0.25 Jy. This renders the kinematical result somewhat uncertain, and the high proper motion of B5 needs to be confirmed by checking against the lower frequency and polarisation data in the follow-up papers. Components B1 and B2 with \(\Phi \approx -115^\circ\) have velocities that are significantly slower than those of the more southern components, B3–B5, with \(\Phi \approx -145^\circ\). A similar velocity gradient transverse to the jet, with components in the northern side having lower speeds, has been reported by Jorstad et al. (2005).

### 5. Physical parameters of the jet

Doppler boosting has an effect on most of the observable properties of a relativistic jet. Hence, in order to compare observations with predictions from theory, the amount of Doppler boosting needs to be measured first. Unfortunately, the Doppler factor of the jet,

\[
\delta = \left[ \Gamma (1 - \beta \cos \theta) \right]^{-1},
\]

Table 4. Proper motion results.

<table>
<thead>
<tr>
<th>Component</th>
<th>( S_{\text{max}} ) [Jy]</th>
<th>( \Phi ) [\celsius]</th>
<th>( \langle \mu \rangle ) [mas yr(^{-1})]</th>
<th>( \langle \beta_{\text{app}} \rangle ) ([h^{-1} c])</th>
<th>( T_0 ) [yr]</th>
</tr>
</thead>
<tbody>
<tr>
<td>S1</td>
<td>1.28</td>
<td>30.6</td>
<td>0.05 ± 0.05</td>
<td>0.4 ± 0.4</td>
<td>–</td>
</tr>
<tr>
<td>C1</td>
<td>1.34</td>
<td>−139.4 ± 0.3</td>
<td>1.01 ± 0.05</td>
<td>7.3 ± 0.4</td>
<td>2002.76 ± 0.04</td>
</tr>
<tr>
<td>C2</td>
<td>4.38</td>
<td>−150.5 ± 0.7</td>
<td>0.79 ± 0.07</td>
<td>5.7 ± 0.5</td>
<td>2002.97 ± 0.05</td>
</tr>
<tr>
<td>C3</td>
<td>3.00</td>
<td>−147.5 ± 1.0</td>
<td>0.95 ± 0.13</td>
<td>6.9 ± 0.9</td>
<td>2003.36 ± 0.03</td>
</tr>
<tr>
<td>B1</td>
<td>0.80</td>
<td>−115.1 ± 0.7</td>
<td>0.71 ± 0.06</td>
<td>5.1 ± 0.4</td>
<td>–</td>
</tr>
<tr>
<td>B2</td>
<td>1.89</td>
<td>−113.3 ± 0.6</td>
<td>0.63 ± 0.05</td>
<td>4.6 ± 0.4</td>
<td>2001.09 ± 0.19</td>
</tr>
<tr>
<td>B3</td>
<td>1.69</td>
<td>−141.2 ± 0.5</td>
<td>0.99 ± 0.06</td>
<td>7.2 ± 0.4</td>
<td>2001.63 ± 0.11</td>
</tr>
<tr>
<td>B4</td>
<td>1.13</td>
<td>−141.3 ± 0.4</td>
<td>1.34 ± 0.07</td>
<td>9.7 ± 0.5</td>
<td>2002.40 ± 0.06</td>
</tr>
<tr>
<td>B5</td>
<td>0.25</td>
<td>−149.6 ± 1.1</td>
<td>1.80 ± 0.25</td>
<td>13.0 ± 1.8</td>
<td>2002.83 ± 0.10</td>
</tr>
<tr>
<td>A1</td>
<td>0.87</td>
<td>−116.8 ± 1.0</td>
<td>0.95 ± 0.13</td>
<td>6.9 ± 0.9</td>
<td>1992.2 ± 1.5</td>
</tr>
<tr>
<td>A2</td>
<td>0.32</td>
<td>−108.0 ± 1.2</td>
<td>1.17 ± 0.18</td>
<td>8.5 ± 1.3</td>
<td>–</td>
</tr>
<tr>
<td>A3</td>
<td>0.48</td>
<td>−126.2 ± 2.3</td>
<td>1.07 ± 0.32</td>
<td>7.8 ± 2.3</td>
<td>1999.4 ± 1.2</td>
</tr>
</tbody>
</table>

Fig. 9. Separation of components A1–A3 from the core vs. epoch of observation. The solid lines represent the best-fit to the component motion (see Table 4).

where \( \Gamma = 1/\sqrt{1 - \beta^2} \) is the bulk Lorentz factor and \( \theta \) is the angle between the jet flow direction and our line of sight, is difficult to measure accurately (see Lähteenmäki & Valtaoja 1999, for a discussion of problems in determining \( \theta \)). In this section, we narrow down the possible range of values for \( \delta, \Gamma \) and \( \theta \) of 3C 273 in early 2003, so that meaningful comparison with the hard X-ray flux measured by the INTEGRAL satellite can be made (Paper II).

The apparent component velocities \( \beta_{\text{app}} \) given in Table 4 already put limits on \( \Gamma \) and \( \theta \). Namely, for all \( \beta_{\text{app}} > 1 \) there is a lower limit for the Lorentz factor \( \Gamma_{\text{min}} = \sqrt{1 + \beta_{\text{app}}^2} \) and an upper limit for the viewing angle, \( \sin \theta_{\text{max}} = \langle \beta_{\text{app}} \rangle / (1 + \beta_{\text{app}}^2) \). \( \Gamma_{\text{min}} \) and \( \theta_{\text{max}} \) for each component are listed in Table 5 (assuming \( h = 0.71 \)). These values are calculated using lower limits of \( \beta_{\text{app}} \). As can be seen from Table 5, the minimum value for \( \Gamma \) in the region close to the VLBI core in 2003 (components C1–C3) is 7.4 and \( \theta_{\text{max}} = 15.5^\circ \).

Table 5. Limits for \( \Gamma \) and \( \theta \).

<table>
<thead>
<tr>
<th>Component</th>
<th>( \Gamma_{\text{min}} )</th>
<th>( \theta_{\text{max}} ) [\celsius]</th>
</tr>
</thead>
<tbody>
<tr>
<td>C1</td>
<td>9.8</td>
<td>11.7</td>
</tr>
<tr>
<td>C2</td>
<td>7.4</td>
<td>15.5</td>
</tr>
<tr>
<td>C3</td>
<td>8.5</td>
<td>13.5</td>
</tr>
<tr>
<td>B1</td>
<td>6.7</td>
<td>17.2</td>
</tr>
<tr>
<td>B2</td>
<td>6.0</td>
<td>19.2</td>
</tr>
<tr>
<td>B3</td>
<td>9.6</td>
<td>11.9</td>
</tr>
<tr>
<td>B4</td>
<td>13.0</td>
<td>8.8</td>
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<tr>
<td>B5</td>
<td>15.8</td>
<td>7.3</td>
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<tr>
<td>A1</td>
<td>8.5</td>
<td>13.5</td>
</tr>
<tr>
<td>A2</td>
<td>10.2</td>
<td>11.3</td>
</tr>
<tr>
<td>A3</td>
<td>7.8</td>
<td>14.7</td>
</tr>
</tbody>
</table>

5.1. Variability time scales and Doppler factors

37 GHz total flux density data from the Metsähovi Radio Observatory monitoring program for 2002–2004 is shown in Fig. 10. A small flare occurred at the beginning of 2003. It has about four times smaller amplitude than the large outburst in 1991 (Robson et al. 1993), making it only a minor but still useful event to estimate the variability brightness temperature and the Doppler factor of the jet, as we will show. As is clear from the ejection times listed in Table 4, components C1–C3 can be connected to this event. The flux density evolution of C2 (see Figs. 10 and 11) closely matches the flare in the Metsähovi flux density curve. C1 has an a...
a sudden drop at the third epoch to be related to the ejection of C3. Based on the flux curves in Fig. 10, it seems to be a reasonable assumption that the flare seen in the single-dish flux density data in January 2003 is mainly caused by a recently ejected component C2.

Observed flux density variability – both in single-dish data and in the VLBI components – can be used to calculate the size of the emission region. In the shock-in-jet model for radio variability, the nearly symmetric flares observed at high radio frequencies (Teräsranta & Valtaoja 1994) suggest that the flux density variability is controlled by light travel delays across the shocked region (Jorstad et al. 2005; Sokolov et al. 2004) i.e. the flux density variability time scale defined as $\tau = \text{d}t/\text{d}n(S)$ corresponds to the light crossing time across the region. This is true if the light crossing time is longer than the radiative cooling time but shorter than the adiabatic cooling time. Jorstad et al. (2005) verified this by comparing flux density variability time scales calculated for VLBI components with time scales of variability in apparent sizes of the same components. They found that the majority of components had a shorter flux variability time scale than that predicted for adiabatic expansion, implying that at high radio frequencies the decay in flux density is driven by radiative cooling. The above assumption allows us to identify $\tau$ with the light-crossing time across the emission region and hence, to estimate the size of the region, which can be then compared with the size measured from the VLBI data to calculate the Doppler factor.

First, we estimate the variability time scale $\tau_{\text{TFD}}$ from the total flux density data. In order to make a simple parametrisation of the flare in January 2003, we have decomposed the flux curve into a constant baseline and separate flares consisting of exponential rise, sharp peak and exponential decay (see Valtaoja et al. 1999, for details). Model flares are of the form

$$\Delta S(t) = \begin{cases} \Delta S_{\text{max}} e^{(t-t_{\text{peak}})/\tau_{\text{TFD}}}, & t < t_{\text{peak}} \\ \Delta S_{\text{max}} e^{(t-t_{\text{peak}})/(1.3\tau_{\text{TFD}})}, & t > t_{\text{peak}}. \end{cases}$$

Our best fit (dashed line) is presented in Fig. 10, consisting of two small flares and a constant baseline of 11.5 Jy. Parameters of this fit are given in Table 6. The variability time scale for the 2003 flare is $\tau_{\text{TFD}} = 175$ days. The uncertainty in $\tau_{\text{TFD}}$ is rather hard to estimate, but according to Lähteenmäki et al. (1999), a conservative upper limit for $\sigma(\tau_{\text{TFD}}) \leq \tau_{\text{TFD}}/4$. An angular size $a$ corresponding to $\tau_{\text{TFD}} = 175 \pm 44$ d is given by

$$a = \frac{c(1+z)}{D_L} \frac{\delta}{\xi},$$

where $D_L$ is luminosity distance:

$$D_L = \frac{c(1+z)}{H_0} \xi(\Omega_M, \Omega_\Lambda, z),$$

with function $\xi$ given by

$$\xi(\Omega_M, \Omega_\Lambda, z) = \int_0^z \frac{\text{d}z'}{\sqrt{(1+z')^2(1+\Omega_M z')-\Omega_\Lambda(2+z')}},$$

(this applies when $\Omega_M + \Omega_\Lambda = 1$; see Carroll et al. 1992). For January 2003 flare, $a = (0.047 \pm 0.012)\delta$ mas, which can be
The 50%-visibility points coincide for a Gaussian profile and for \( \delta \times 0.5 \pm \tau \pm \) variability time scales and Doppler factors from the X±±T.η.δ.2004± in a region of few standard deviations around is the maximum amplitude of the outburst in Janskys, 2003.6δY±±=±T.η.δ.2003.4=±а=±max=ffmax∆а.[year]τδδ.290.6. Taking the geometric mean of major and minor axes and multiplying the result by 1.8, we get a size equivalent to an optically thin sphere, \( a_{\text{VLBI}} = 0.22 \pm 0.05 \) mas. Comparison with \( a \) from variability gives \( \delta = 4.7 \pm 1.6 \) for C2.

The variability time scale can also be determined solely on the basis of flux density variability of a single VLBI component. As can be seen from Figs. 11 and 12, components B2, B3, B4, C2 and C3 show enough flux density variability that they can be used in estimating the variability time scale. The flux density time scales, \( \tau_{\text{VLBI}} \), the sizes at the time of maximum flux densities, \( a_{\text{VLBI}} \) (with geometry of an optically thin sphere assumed, i.e. the figures are multiplied by 1.8), and the calculated Doppler factors \( \delta \) are presented in Table 7 for these components. In calculating \( \tau_{\text{VLBI}} \) and \( \delta \), we have taken into account the asymmetric uncertainties in flux densities and sizes of several components (see e.g. D’Agostini 2004, for a short review about asymmetric uncertainties and their correct handling). The figures given in Table 7 are probabilistic expected values, not “best values” obtained by direct calculation, and the uncertainties correspond to probabilistic standard deviations.

A non-linear dependence of the output quantity \( Y \) on the input quantity \( X \) in a region of few standard deviations around the expected value of \( X \) is another source of asymmetric uncertainties, which also needs to be properly taken into account when doing calculations with measured quantities. In this paper, when calculating derived quantities like brightness temperature, which has a strongly non-linear dependence on the size of the component, we have used a second-order approximation for the error propagation formula instead of the conventional linear one. Moreover, the non-linearity does not only affect the errors, but it can also shift the expected value of \( Y \) with respect to the “best value” obtained by direct calculation (D’Agostini 2004). In the following, for all the derived quantities, expected values and standard deviations are reported instead of directly calculated values, since they are much more useful for any statistical analysis.

The two variability time scales for C2, estimated from the total flux density and from the VLBI data, differ by 60 days, which affects the calculated Doppler factors – from the total flux density variability we get \( \delta (C2) = 4.7 \pm 1.6 \) as compared to \( \delta (C2) = 7.5 \pm 2.6 \) calculated from the variability in the VLBI data. A weighted average of these, \( \delta (C2) = 5.5 \pm 1.9 \).

\( \) The 50%-visibility points coincide for a Gaussian profile and for the profile of an optically thin sphere with diameter of 1.8 times the FWHM of the Gaussian profile.

### Table 6. Decomposition of the single-dish flux curve into self-similar flares.

<table>
<thead>
<tr>
<th>Flare</th>
<th>( t_{\text{peak}} ) [year]</th>
<th>( \Delta S_{\text{max}} ) [Jy]</th>
<th>( \tau_{\text{FHD}} ) [days]</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>2003.03</td>
<td>9.72</td>
<td>175 ± 44</td>
</tr>
<tr>
<td>2</td>
<td>2002.06</td>
<td>6.76</td>
<td>174 ± 44</td>
</tr>
<tr>
<td>Constant baseline</td>
<td></td>
<td>11.5 Jy</td>
<td></td>
</tr>
</tbody>
</table>

compared with the size of C2 measured from the VLBI data. The full width at half maximum (FWHM) size of the major axis of the Gaussian component C2 is \( 0.23 \pm 0.04 \) mas and the axial ratio is \( \eta = 0.29 \pm 0.06 \). Taking the geometric mean of major and minor axes and multiplying the result by 1.8, we get a size equivalent to an optically thin sphere, \( a_{\text{VLBI}} = 0.22 \pm 0.05 \) mas. Comparison with \( a \) from variability gives \( \delta = 4.7 \pm 1.6 \) for C2.

The variability time scale can also be determined solely on the basis of flux density variability of a single VLBI component. As can be seen from Figs. 11 and 12, components B2, B3, B4, C2 and C3 show enough flux density variability that they can be used in estimating the variability time scale. The flux density time scales, \( \tau_{\text{VLBI}} \), the sizes at the time of maximum flux densities, \( a_{\text{VLBI}} \) (with geometry of an optically thin sphere assumed, i.e. the figures are multiplied by 1.8), and the calculated Doppler factors \( \delta \) are presented in Table 7 for these components. In calculating \( \tau_{\text{VLBI}} \) and \( \delta \), we have taken into account the asymmetric uncertainties in flux densities and sizes of several components (see e.g. D’Agostini 2004, for a short review about asymmetric uncertainties and their correct handling). The figures given in Table 7 are probabilistic expected values, not “best values” obtained by direct calculation, and the uncertainties correspond to probabilistic standard deviations.

A non-linear dependence of the output quantity \( Y \) on the input quantity \( X \) in a region of few standard deviations around the expected value of \( X \) is another source of asymmetric uncertainties, which also needs to be properly taken into account when doing calculations with measured quantities. In this paper, when calculating derived quantities like brightness temperature, which has a strongly non-linear dependence on the size of the component, we have used a second-order approximation for the error propagation formula instead of the conventional linear one. Moreover, the non-linearity does not only affect the errors, but it can also shift the expected value of \( Y \) with respect to the “best value” obtained by direct calculation (D’Agostini 2004). In the following, for all the derived quantities, expected values and standard deviations are reported instead of directly calculated values, since they are much more useful for any statistical analysis.

The two variability time scales for C2, estimated from the total flux density and from the VLBI data, differ by 60 days, which affects the calculated Doppler factors – from the total flux density variability we get \( \delta (C2) = 4.7 \pm 1.6 \) as compared to \( \delta (C2) = 7.5 \pm 2.6 \) calculated from the variability in the VLBI data. A weighted average of these, \( \delta (C2) = 5.5 \pm 1.9 \).

\( \) The 50%-visibility points coincide for a Gaussian profile and for the profile of an optically thin sphere with diameter of 1.8 times the FWHM of the Gaussian profile.

### Table 7. Variability time scales and Doppler factors from the VLBI data.

<table>
<thead>
<tr>
<th>Component</th>
<th>( \tau_{\text{VLBI}} ) [days]</th>
<th>( a_{\text{VLBI}} ) [mas]</th>
<th>( \delta )</th>
</tr>
</thead>
<tbody>
<tr>
<td>C2</td>
<td>110 ± 30</td>
<td>0.22 ± 0.05</td>
<td>7.5 ± 2.6</td>
</tr>
<tr>
<td>C3</td>
<td>120 ± 10</td>
<td>0.12 ± 0.04</td>
<td>4.1 ± 1.4</td>
</tr>
<tr>
<td>B2</td>
<td>110 ± 20</td>
<td>0.36 ± 0.05</td>
<td>12.2 ± 2.8</td>
</tr>
<tr>
<td>B3</td>
<td>160 ± 30</td>
<td>0.18 ± 0.04</td>
<td>4.2 ± 1.7</td>
</tr>
<tr>
<td>B4</td>
<td>100 ± 30</td>
<td>0.23 ± 0.05</td>
<td>8.6 ± 3.2</td>
</tr>
</tbody>
</table>

agrees well with other Doppler factor values for 3C 273 derived elsewhere: \( \delta_{\text{var}} = 5.7 \) (Lähteenmäki & Valtaoja 1999), \( \delta_{\text{eq}} = 8.3 \), \( \delta_{\text{SSC}} = 4.6 \) (Güijosa & Daly 1996), and \( \delta_{\text{SSC}} = 6.0 \) (Ghisellini et al. 1993). The variability Doppler factors of other components listed in Table 7 vary significantly, with C3 and B3 having the smallest values, \( \delta \approx 4 \), and B2 having the highest, \( \delta (B2) = 12.2 \pm 2.8 \). This is not very surprising, since also the apparent velocities of these components differ considerably.

### 5.2. Brightness temperatures and equipartition

Brightness temperature is another quantity of interest in sources like 3C 273. From the total flux density variability, we can estimate the variability brightness temperature (Lähteenmäki et al. 1999)

\[
T_{b,\text{var}} = 5.28 \times 10^{20} \cdot h^{-2} \frac{\Delta S_{\text{max}}}{v^2 T_{\text{FHD}}} (1+z)^2 \xi(\Omega_M, \Omega_\Lambda, z)^2, 
\]

where \( v \) is the observation frequency in GHz, \( z \) is redshift, \( \Delta S_{\text{max}} \) is the maximum amplitude of the outburst in Janskys, and \( T_{\text{FHD}} \) is the observed variability timescale in days. The numerical factor in Eq. (6) comes from assuming that the component is an optically thin, homogeneous sphere. Lähteenmäki et al. (1999) estimate a conservative upper limit to the error in \( T_{b,\text{var}} \) to be \( \Delta T/T < 0.50 \). For the flare in January 2003, \( T_{b,\text{var}} = (6.6 \pm 3.3) \times 10^{12} \) K.
Table 8. Component properties.

<table>
<thead>
<tr>
<th>Component</th>
<th>$T_{b,\text{VLBI}}$ [K]</th>
<th>$\delta$</th>
<th>$T_{b,\text{int}}$ [K]</th>
<th>$\Gamma$</th>
<th>$\theta$ [$\text{'}$]</th>
</tr>
</thead>
<tbody>
<tr>
<td>C1</td>
<td>$(7.9 \pm 5.5) \times 10^{10}$</td>
<td>–</td>
<td>$(3 \pm 2) \times 10^{10}$</td>
<td>9.8 ± 3.2</td>
<td>9.8 ± 2.1</td>
</tr>
<tr>
<td>C2</td>
<td>$(1.7 \pm 0.7) \times 10^{11}$</td>
<td>5.5 ± 1.9</td>
<td>$(8 \pm 3) \times 10^{10}$</td>
<td>14 ± 5</td>
<td>10.0 ± 1.5</td>
</tr>
<tr>
<td>C3</td>
<td>$(3.4 \pm 1.4) \times 10^{11}$</td>
<td>4.1 ± 1.4</td>
<td>$(2 \pm 1) \times 10^{9}$</td>
<td>8.0 ± 1.0</td>
<td>4.3 ± 1.6</td>
</tr>
<tr>
<td>B1</td>
<td>$(4.0 \pm 2.3) \times 10^{10}$</td>
<td>–</td>
<td>$(3 \pm 2) \times 10^{10}$</td>
<td>17 ± 7</td>
<td>9.7 ± 0.8</td>
</tr>
<tr>
<td>B2</td>
<td>$(2.5 \pm 0.7) \times 10^{10}$</td>
<td>12.2 ± 2.8</td>
<td>$(5 \pm 2) \times 10^{9}$</td>
<td>18 ± 8</td>
<td>6.0 ± 1.2</td>
</tr>
<tr>
<td>B3</td>
<td>$(1.2 \pm 0.7) \times 10^{11}$</td>
<td>4.2 ± 1.7</td>
<td>$(5 \pm 2) \times 10^{9}$</td>
<td>18 ± 8</td>
<td>6.0 ± 1.2</td>
</tr>
<tr>
<td>B4</td>
<td>$(4.0 \pm 1.8) \times 10^{10}$</td>
<td>8.6 ± 3.2</td>
<td>$(5 \pm 2) \times 10^{9}$</td>
<td>18 ± 8</td>
<td>6.0 ± 1.2</td>
</tr>
</tbody>
</table>

The observed variability brightness temperature is proportional to $\delta^3$, i.e. $T_{b,\text{var}} = T_{b,\text{int}} \delta^3$. Using $\delta = 5.5 \pm 1.9$, obtained earlier for component C2, we get a value $T_{b,\text{int}} \approx (8 \pm 6) \times 10^{10}$ K of intrinsic brightness temperature (the strong non-linear dependence of $T_{b,\text{int}}$ on $\delta$ shifts the expected value of $T_{b,\text{int}}$ significantly from the directly calculated “best value” of $(4 \pm 1) \times 10^{10}$ K; D’Agostini 2004), which is consistent with the source being near equipartition of energy between the radiating particles and the magnetic field. According to Readhead (1994) the equipartition brightness temperature, $T_{b,\text{eq}}$, depends mainly on the redshift of the source and weakly on the observed optically thin spectral index, the peak flux density and the observed frequency at the peak. For the redshift of 3C 273, $T_{b,\text{eq}} \approx 8 \times 10^{10}$ K for any reasonable values of $\alpha$, $S_m$ and $\nu_m$. This exactly matches the $T_{b,\text{int}}$ calculated for the January 2003 flare, indicating that near the core the source could have been in equipartition. Our result is consistent with the findings of Lähteenmäki et al. (1999), who have shown that during high radio frequency flares all sources in their sample have intrinsic brightness temperature close to the equipartition value.

Interferometric brightness temperatures of the components are also of interest. The maximum brightness temperature for an elliptical component with a Gaussian surface brightness profile is

$$T_{b,\text{VLBI}} = 1.22 \times 10^{12} \frac{S_{\text{VLBI}}(1 + z)}{\eta \sigma_{\text{maj}}^2}$$

(7)

where $S_{\text{VLBI}}$ is the flux density in Janskys, $\nu$ is the observed frequency in GHz, $z$ is redshift, $\sigma_{\text{maj}}$ is the FWHM size of the major axis in mas, and $\eta$ is the axial ratio of the component. In order to consider a surface brightness distribution, which is more physical than a Gaussian, we convert the Gaussian brightness temperature to an optically thin sphere brightness temperature by multiplying the former by a factor of 0.67. Table 8 lists the maximum interferometric brightness temperatures for components within 2.5 mas from the core and with non-zero areas. Again, we have taken into account the effect of asymmetric errors in $S_{\text{VLBI}}$, $\sigma_{\text{maj}}$, and $\eta$, as well as the non-linearity of the Eq. (7), for error propagation. The table also contains the average Doppler factor and $T_{b,\text{int}}$ calculated by dividing $T_{b,\text{VLBI}}$ by $\delta$. $T_{b,\text{int}}$ for C2 given in Table 8 differs from $(8 \pm 6) \times 10^{10}$ K that was quoted earlier, because it is calculated using the interferometric brightness temperature instead of variability brightness temperature. However, the difference is not significant, considering the overall uncertainties in measuring $T_{b,\text{int}}$.

5.3. Lorentz factors and viewing angles

Since we have measured $\delta$ and $\beta_{\text{app}}$ for a number of components, it is possible to calculate the values of $\Gamma$ and $\theta$. The Lorentz factor is given by equation

$$\Gamma = \frac{\beta_{\text{app}}^2 + \delta^2 + 1}{2\delta}$$

(8)

and the angle between the jet and our line of sight by equation

$$\theta = \arctan \frac{2\beta_{\text{app}}}{\beta_{\text{app}}^2 + \delta^2 - 1}.$$  

(9)

The most reliable value of the Doppler factor was obtained for component C2. With $\beta_{\text{app}}(C2) = 8.0 \pm 0.7$ and $\delta(C2) = 5.5 \pm 1.9$, we get $\Gamma = 9.8 \pm 3.2$ and $\theta = 9.8 \pm 2.1^\circ$.

In Table 8 we have gathered Lorentz factors and viewing angles also for other components besides C2. The results indicate that a transverse velocity gradient across the jet does exist, even when the viewing angle difference is taken into account. B2 has $\Gamma = 8$, while B3 and B4 show $\Gamma \sim 17$, and even the lower limit of Lorentz factor for B4 is larger than $\Gamma(B2)$ (see Table 5).

6. Discussion

3C 273 is the closest quasar, and due to its proximity it provides one of the best chances to study the physics of a strong relativistic outflow in detail. The jet in 3C 273 is complicated, with a number of details which would be smoothed away if the source was located at higher redshift. This complexity also makes it hard to interpret. For instance, several interferometric studies on the parsec scale jet of 3C 273 in the 1980s and 1990s have reported a “wiggling” structure with time-variable ejection angles and component velocities (e.g. Zensus et al. 1988; Krichbaum et al. 1990; Bååth et al. 1991; Abraham et al. 1996; Mantovani et al. 1999). In the following, we will explore some possible scenarios that could lie behind the observed curving structure, and discuss the observational evidence for and against them.
6.1. Precession of the jet nozzle

Abraham & Romero (1999) proposed a simple model of a ballistic jet with a precessing nozzle to explain the variable ejection angles and component velocities. Their model fits well the early VLBI data describing components ejected between 1960–1990. To compare our more recent high frequency data with their model, we have plotted the apparent speeds and directions of motion for the components as a function of their ejection epochs in Fig. 13 (excluding the faint B5). Overlaid in the figure are the predicted curves from the Abraham & Romero model. It is clear that the component velocities and directions show much faster variations than the precession model with a period of $\sim 15$ years predicts. No periodicity can be claimed on the basis of our data, although the apparent velocities may show pseudo-sinusoidal variation. However, the directions of the components’ proper motions show no clear pattern, and the apparent direction changes as fast as $\sim 30^\circ$ in half a year.

There is one important difference between our data and that used by Abraham & Romero (1999) to construct their model. Most observations cited in their paper were carried out at lower frequencies (mainly 10.7 GHz) with significantly poorer resolution. Many of the components in our 43 GHz maps are small and often part of a larger emission structure, which is likely to show up as a single component on lower resolution maps. Such a component could show drastically different ejection angles than the individual features in it due to averaging. We have tested whether our data could fit the precession model of Abraham & Romero, were it observed with poorer resolution. The 43 GHz uv-data were tapered into a resolution comparable to VLBA observations at 10.7 GHz and then model-fitted (an example is shown in Fig. 14). In Fig. 15 we have plotted the apparent velocities and directions of motion for three prominent components in the tapered data as a function of their ejection epochs. We have also included the data from Abraham & Romero (1999), scaled to the cosmology used in this paper, and superimposed their precession model to the figure. As can be seen, we cannot exclude the possibility that the average jet direction changes periodically, and that it can be approximately described by the Abraham & Romero model.

6.2. Ballistic or not?

The simplest possible model describing proper motions in a relativistic jet is the one where knots ejected from the core move at constant velocity following a ballistic trajectory. On the contrary, non-ballistic motion in the jet implies a more complex physical system with dynamics governed by e.g. magnetic fields, fluid instabilities or interaction with the ambient medium.

Most of the components observed in our data follow, at least on a timescale of nine months, a straight path starting from the core with a constant velocity. However, there are a few important exceptions. Component C1 shows a curved trajectory with slight bending towards the north (Fig. 4). Its position at the last epoch deviates $\sim 0.11$ mas from the straight line path extrapolated from the first three epochs and it does not follow the path taken earlier by components B3–B5. Follow-up observations tracking the motion of C1 will show if the bending continues and C1 becomes a “northern component” with a
Trajectory similar to B2 or B3. The component B1 of Jorstad et al. (2005) has a curved trajectory with these characteristics, indicating that the behaviour of C1, if it actually continues to bend, is not unique. The other two exceptions to radial motion are components B1 and A2, which do not extrapolate back to the core. B1 lies ~0.6 mas south of B2 and moves parallel to it. This component may have been split from B3 or B2, or it may have changed its direction earlier – similarly to C1.

Several studies report non-ballistic motion in 3C 273. Kellermann et al. (2004) found a non-radial trajectory for one of the components in their 2 cm VLBA Survey monitoring, Homan et al. (2001) also reported a curved path for one component, and Jorstad et al. (2005) identified several knots following a bent trajectory. Figure 5 in Krichbaum et al. (2000) is a very interesting plot, where the long term motion of 12 components is presented in the (RA, Dec)-plane. The components show slightly curved trajectories with both convex and concave shapes. It is apparent from this figure that had the components been followed only for a year, one would have described most of them as ballistic. This is in agreement with our data, and we consider it likely that several components have non-ballistic trajectories, although with large radii of curvature. One more piece of evidence supporting this view comes from the apparent opening angle of the jet. At 1–2 mas from the core, we can estimate the apparent opening angle of the jet from the motion directions of B2 and B3, yielding 28° in the plane of the sky. In a 5 GHz image taken at the first epoch, the jet is well-collimated up to 50 mas and its opening angle on the scale of 10–50 mas is significantly smaller than 28° (Paper II; preliminary 5 GHz data is published in Savolainen et al. 2004, see their Fig. 1). In the kpc scale jet, the half-power width remains constant (1.0") from 13" to 20" along the jet implying highly collimated outflow and, again, a small opening angle (Conway et al. 1993).

If the large observed opening angle of 28° at 1–2 mas from the core is not due to a rare event of changing jet direction (see Sect. 6.4), the components B2 and B3 are not likely to continue on ballistic trajectories for long. Instead, either collimation or increase in the viewing angle is expected to occur, because of the smaller apparent opening angle farther downstream. In either case, the component trajectory is not ballistic.

6.3. Plasma instabilities

The “wiggling” structure in the jet of 3C 273 is present from the subparsec (Bååth et al. 1991) to kiloparsec scale (Conway et al. 1993), strongly indicating that there is a wavelike “normal mode” configuration in the jet. Lobanov & Zensus (2001) reported the discovery of two threadlike patterns in the jet resembling a double helix in their space-VLBI image. They successfully fitted the structure with a Kelvin-Helmholtz instability model consisting of five different instability modes, which they identify as helical and elliptical surface and body modes. The helical surface mode in their model has significantly shorter wavelength than its anticipated characteristic wavelength, implying that the mode is driven externally. Lobanov & Zensus associate this driving mechanism with the 15-year period of Abraham & Romero (1999). This is in accordance with our observations: the average jet direction agrees with the precession model, but individual components show much more rapid and complex variations in speed and direction. Lobanov & Zensus predicted that a substantial velocity gradient should be present in the jet. Such a gradient was observed in this work as was shown in the previous section.

If the magnetic energy flux in the jet is comparable to or larger than the kinetic energy flux, the flow will, in addition to Kelvin-Helmholtz modes, be susceptible to current-driven instabilities. Nakamura & Meier (2004) carried out three-dimensional magnetohydrodynamic simulations that indicate that the growth of the current-driven asymmetric ($m = 1$) mode in a Poynting-flux dominated jet produces a three-dimensional helical structure resembling that of “wiggling” VLBI jets. Although their simulation was non-relativistic, and thus needs confirmation of the growth of current-driven instability modes in the relativistic regime, the idea of current-driven instabilities in a Poynting-flux dominated jet being behind the observed structure in 3C 273 is tempting, since the current-carrying outflow would also explain the rotation measure gradient across the jet observed by several groups. Asada et al. (2002) were the...
first to find a gradient of about 80 rad m$^{-2}$ mas$^{-1}$ at about 7 mas from the core, and recently Zavala & Taylor (2005) reported a gradient of 500 rad m$^{-2}$ mas$^{-1}$ at approximately the same location in the jet by using data from a higher resolution study. Attridge et al. (2005) performed polarisation VLBI observations of 3C 273 at 43 and 86 GHz and their data reveal a rotation measure gradient of $\sim$10$^5$ rad m$^{-2}$ mas$^{-1}$ at 0.9 mas from the core. These observations can be explained by a helical magnetic field wrapping around the jet, as is expected in Poynting-flux dominated jet models. Both Zavala & Taylor (2005) and Attridge et al. (2005) conclude that the high fractional polarisation observed in the jet rules out internal Faraday rotation, and thus the synchrotron-emitting particles need to be segregated from the region of the helical magnetic field. In a Poynting-flux dominated jet, a strong magnetic sheath is formed between the axial current flowing in the jet and the return current flowing outside (Nakamura & Meier 2004). In this magnetic sheath the field lines are highly twisted and it would be a natural place to produce the observed rotation measure gradients outside the actual synchrotron emission regions.

6.4. Nature of region “B”

The most interesting structure in our images (Figs. 1 and 2) is the region at 1–2 mas from the core, where there are several bright components at equal distances from the core with vastly different directions of motion. This structure is already present in 43 and 86 GHz images of Attridge et al. (2005) observed at epoch 2002.35, and the authors suggest that they have caught the jet of 3C 273 in the act of changing its direction. They also conjecture that the southern component Q7/W7 in their data is younger and faster than the northern component Q6/W6. We identify components Q6/W6 and Q7/W7 in Attridge et al. (2005) with our components B2 and B3, respectively, and confirm their conjecture about the relative age and velocity of the components. Their first hypothesis concerning the change of jet direction is more problematic. Supporting evidence for this change is that the components B3–B5 and C1–C3, ejected after B2, all have directions of motion between $\text{PA} = -140^\circ$ and $\text{PA} = -150^\circ$ as compared to $\text{PA} = -113^\circ$ of B2. However, the change would have had to happen in a short time, in about six months (Fig. 13). It is hard to come up with a mechanism able to change the direction of the entire jet so abruptly. We have compared the positions of the components Q6/W6 and Q7/W7 of Attridge et al. (2005) with the positions of B2 and B3, and Q7/W7 nicely falls into the line extrapolated from the proper motion of B3, but for Q6/W6 the situation is more complicated. Q6/W6 can be joined with B2 by a straight line, but this path does not extrapolate back to the core implying that Q6/W6/B2 may not follow a ballistic trajectory. Also, the trajectory of component C1 (ejected after B2) shows some hints of being deviated from the path taken earlier by B3–B5, as was discussed before. If C1 continues to turn towards the “northern track”, it may follow a trajectory similar to B1 and B2.

What possible explanations are there for the observed vastly different directions of component motion in region “B” other than the changing jet direction? One possibility is that the double helical structure of Lobanov & Zensus (2001) is already prominent at 1 mas from the core. In their analysis there are two maxima in the brightness distributions across the jet, and a wavelike structure would fit the curved paths of our C1 and the component B1 of Jorstad et al. (2005). Further observations following the motions of B2 and B3 are needed to confirm or disprove this explanation, but we consider it as a simple and viable alternative to an abruptly changing jet direction.

Another potential alternative explanation is the “lighthouse model” by Camenzind & Krockenberger (1992), who propose that the fast optical flux variations in the quasars may be due to a rapid rotation of the plasma knots within the jet. The rotation changes the viewing angle, and consequently induces a periodic variability of the Doppler factor near the beginning of the jet. If the jet has a non-negligible opening angle, the local rotation period increases and the trajectories of the components asymptotically approach straight lines as the knots move downstream. However, within a few milliarcseconds from the core, there could still be detectable helical motion, and the knots in the region “B” in 3C 273 could be components having non-negligible angular momentum and following different paths. Again, a longer monitoring of the component trajectories is needed to test this idea.

The jet could also broaden due to a drop in confining pressure (either external gas pressure or magnetic hoop stress) at 1 mas from the core. However, as the jet remains collimated in the downstream region, we would expect the pressure drop to be followed by a recollimation, and consequently, a jet of oscillating cross-section to form (Daly & Marscher 1988). No oscillation of the jet width can be clearly claimed on the basis of either our data or that of previous VLBI studies. Particularly, in Jorstad et al. (2005), where a broadening of the jet – similar to the one seen in our images – was observed in 1998, there is no evidence of such an oscillation in the following two years. Still, the observations are not easy to interpret in this case, and we cannot conclusively reject a pressure drop scenario either.

7. Conclusions

We have presented five 43 GHz total intensity images of 3C 273 covering a time period of nine months in 2003 and belonging to a larger set of VLBA multifrequency monitoring observations aimed to complement simultaneous high energy campaign with the INTEGRAL satellite. We fitted the data at each epoch with a model consisting of a number of Gaussian components in order to analyse the kinematics of the jet. Particular attention was paid to estimating the uncertainties in the model parameters, and the Difwrap program was found to provide reliable estimates for positional errors of the components.

The images in Figs. 1 and 2 show an intriguing feature at 1–2 mas from the core, where the jet is resolved in direction transverse to the flow. In this broad part of the jet, there are bright knots with vastly different directions of motion. The jet may have changed its direction here, as proposed by Attridge et al. (2005), although we consider it more likely that we are seeing an upstream part of the double helical structure observed by Lobanov & Zensus (2001).
We have analysed the component kinematics in the parsec scale jet and found velocities in the range of $4.6-13.0 \, h^{-1} \, c$. There is an apparent velocity gradient across the jet with northern components moving slower than southern ones. Thus, we can confirm the earlier report of this gradient by Jorstad et al. (2005). We have also found curved and non-radial motions in the jet, although most of the components show radial motion on a time scale of nine months. Taking into account other studies reporting non-ballistic motion in 3C 273 (Krichbaum et al. 2000; Homan et al. 2001; Kellermann et al. 2004; and Jorstad et al. 2005), we consider it likely that several components in the jet have non-ballistic trajectories, although with large radii of curvature. This also is in accordance with the fluid dynamical interpretation of motion in 3C 273.

By using flux density variability and light travel time arguments, we have estimated the Doppler factors for the prominent jet components and combined them with the apparent velocities to calculate the Lorentz factors and viewing angles. For instance, for the newly ejected component C2, we get an accurate and reliable value of the Doppler factor, $\delta(C2) = 5.5 \pm 1.9$. The Doppler factors will be used in Paper II together with component spectra to calculate the anticipated amount of hard X-rays due to the synchrotron self-Compton mechanism. The Lorentz factors obtained in this paper range from 8 to 18, and show that due to the synchrotron self-Compton mechanism. The Lorentz factors obtained in this paper range from 8 to 18, and show that due to the synchrotron self-Compton mechanism.

We can, with confidence, identify the ejection of component C2 with the 2003 flare, and components C1 and C3 are perhaps associated with the large outbursts in 1983, 1988, 1991 and 1997. The source underwent a mild flare in early 2003, at the time when INTEGRAL observed it. Unfortunately, the flare was weak compared to the large outbursts in 1983, 1988, 1991 and 1997. We can, with confidence, identify the ejection of component C2 with the 2003 flare, and components C1 and C3 are perhaps connected to this event.

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References

Bridle, A. H., & Greisen, E. W. 1994, AIPS Memo 87, NRAO
Courvoisier, T. J.-L. 1998, A&ARv, 9, 1
Savolainen, T., Wiik, K., & Valtaoja, E. 2004, in Proc. of the 5th INTEGRAL Workshop, ed. V. Schönhöfer, G. Lichti, & C. Winkler, ESA SP-552, 559
Schmidt, M. 1963, Nature, 197, 1040
Taylor, G. B., & Myers, S. T. 2000, VLBA Scientific Memo 26
Multifrequency VLBA monitoring of 3C 273 during the INTEGRAL campaign in 2003

II. Extraction of the VLBI component spectra

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ABSTRACT

Aims. The synchrotron spectra of the emission features in the VLBI jet of 3C 273 are studied.
Methods. We observed 3C 273 with the Very Long Baseline Array simultaneously at 5.0, 8.4, 15.3, 22.2, 43.2 and 86.2 GHz on February 28, 2003 as a part of polarimetric multifrequency VLBI monitoring carried out during a simultaneous campaign with the INTEGRAL γ-ray satellite in 2003. This multifrequency data set is used to carefully construct spectra of different emission features in the parsec scale jet by using a model-transfer method. The method is presented and its uncertainties are discussed.
Results. Spectra of 16 emission features in the VLBI jet are presented. Four of these show frequency-dependent angular sizes, probably due to radiative losses. Eight features show spectral turnover in our frequency range, and we have calculated their magnetic field density, relativistic electron energy density and the anticipated amount of synchrotron self-Compton flux density. Errors in these parameters were determined by Monte Carlo methods. The core has a magnetic field of the order of 1 Gauss and the field density decreases roughly inversely proportional to the distance from the core. A strong gradient in the magnetic field density, coincident with a transverse velocity structure at ∼1.5 mas from the core, was found: the slower superluminal component B2 on the northern side of the jet has a magnetic field density two orders of magnitude lower than faster southern components B3 and B4. A comparison of our results with the INTEGRAL observations indicates that SSC-emission from the radio components in the parsec scale jet is generally weak compared to the observed hard X-ray flux density. An exception is the peculiar feature B2, which seems to be a strong SSC-emitter, and which could, in principle, produce the observed hard X-ray flux density. An anomalous nature of this component, however, leaves a doubt about its true X-ray strength.

Key words. Galaxies: active – Galaxies: jets – quasars: individual: 3C 273 – Techniques: interferometric

1. Introduction

Traditionally, Very Long Baseline Interferometry (VLBI) observations of compact extragalactic jets have mostly concentrated on imaging the sources in total intensity at a given frequency band in order to determine the source morphology and to monitor possible motions in the jet. These studies suffice to measure the kinematics of the emission features in the jet (usually called “components”), which has led to the discovery of a source population showing apparent superluminal motion – a clear indication of relativistic velocities being present in these sources (see Zensus 1997 for a comprehensive review). However, kinematics alone gives only a limited view and more information on the physical state of the matter and fields in the jets is needed to understand their workings. VLBI observations can provide two further means of exploring the jets: polarisation and radio spectra. Polarisation VLBI observations are nowadays a routine practice (see e.g. Taylor et al. 2005, Lister & Homan 2005, Jorstad et al. 2005, Homan et al. 2002), and, due to frequency agility of the Very Long Baseline Array (VLBA), also spectral studies of the parsec scale jets with a wide frequency coverage have become feasible after commissioning of this facility.

Pioneering observations of the AGN spectrum on the parsec scale were done already before the advent of the VLBA (Cotton et al. 1980, Bartel et al. 1984, Marscher & Broderick 1985, Marscher 1988), but they often had a rather limited frequency
coverage, and observations at different frequencies were not simultaneous. VLBA’s capability to make practically simultaneous multifrequency observations over a wide range of frequencies – and yet with high dynamic range – has changed this. However, although the potential of the spectral analysis with the VLBA has been successfully demonstrated by e.g. Walker et al. (2000), Marr et al. (2001) and Vermeulen et al. (2003), only a very small number of sources have been studied so far. This is probably partly due to technical difficulties, such as uneven \( \nu \)-coverages at different observing bands, and partly due to laborious data reduction needed. A detailed discussion of the problems involved in the extraction of spectral information from multifrequency VLBI campaigns can be found in Lobanov (1998).

The radio spectra of the emission features in the parsec scale jet are important observables. In the framework of incoherent, random pitch-angle synchrotron radiation model (e.g. Marscher 1987), the observed spectra together with measurements of the emission region’s angular size provide means to derive the physical conditions in the jet – such as magnetic field density and energy distribution of the radiating electron population. Ultimately, the total kinetic luminosity of the jet – a fundamental parameter constraining possible mechanisms powering the jets – can be estimated, provided that additional information of the jet plasma composition is available.

In addition to broad band synchrotron radiation ranging from radio to UV, flat spectrum radio quasars and BL Lac objects (commonly grouped together as blazars) are also observed to emit copious amounts of high energy X-rays and \( \gamma \)-rays (Mukherjee et al. 1997, Hartman et al. 1999). Although inverse Compton (IC) scattering of soft photons from relativistic electrons moving in the jet is currently a favoured model for the high energy emission, within this model there is a substantial controversy about the origin of seed photons and location of the emission site. For instance, the seed photons could be synchrotron photons emitted by the same electrons which scatter them later, (synchrotron self-Compton model, SSC; Maraschi et al. 1992, Bloom & Marscher 1996), or they could be photons external to the jet originating in the accretion disk or in the broad line region clouds (external Compton model; Dermer et al. 1992, Ghisellini & Madau 1996, Sikora et al. 1994). Beside these leptonic models there are also a number of models where \( \gamma \)-rays are due to hadronic processes initiated by relativistic protons co-accelerated with electrons (Mannheim 1993).

The emission processes in blazars have been studied by fitting their spectral energy distributions (SED) with one-zone synchrotron-IC models, where it is usually assumed that the high energy emission comes from the part of the jet that is very close to the central engine (\( \leq 0.1 \) pc from the black hole) and within the broad line region (e.g. Mastichiadis & Kirk 1997, Hartman et al. 2001, Böttcher & Reimer et al. 2004). These models fit the inverse Compton bump and IR-optical part of the synchrotron spectrum rather well but do not adequately describe the radio-mm emission, which is thought to arise over a parsec downstream of the \( \gamma \)-ray emission site. However, there are a number of studies reporting correlations between millimetre wavelength and \( \gamma \)-ray emission in blazars (e.g. Reich et al. 1993, Valtaoja & Teräsranta 1995, 1996, Mücke et al. 1996).

Lähteenväki & Valtaoja (2003) conclude that \( \gamma \)-ray flares occur during the rising phase or the peak of a radio flare, and that on average the \( \gamma \)-ray emitting region is at the distance of 4.9 pc downstream from the core. Jorstad et al. (2001b) find a correlation between \( \gamma \)-ray flares and ejections of superluminal components, and based on the delay between the time when a superluminal radio component is coincident with the radio core and the epoch of the maximum observed \( \gamma \)-ray flux, they conclude that the radio and high energy events occur within the same shocked region located near but probably slightly downstream of the core. This contradicts with the assumptions usually made in the SED modelling, and the source of the target photons and the actual site of the high energy emission thus are currently open issues.

Measuring energy densities of the relativistic particle population and the magnetic field in the parsec scale jet by multifrequency VLBI observations provides one method for testing the SSC model. If the relativistic Doppler factor of the jet is known, it is possible to predict how much high energy radiation each jet feature should produce through SSC mechanism by measuring its size, the synchrotron turnover frequency, and the turnover flux density. The prediction can then be compared with near-simultaneous satellite observations of the X/\( \gamma \)-ray flux density, allowing us to distinguish which, if any, of the components in the parsec scale jet is a major source of the X/\( \gamma \)-ray emission.

In the year 2003, the quasar 3C 273 (\( z = 0.158 \); see Courvoisier 1998 for a comprehensive review of the source properties) was a target of a multiwavelength campaign organised to support the observations done with the INTEGRAL \( \gamma \)-ray satellite (Courvoisier et al. 2003, 2004). At high energies, 3C 273 was found to be very faint during January 2003, the \( 3 – 500 \) keV flux being at the lowest level ever observed up to that date (Courvoisier et al. 2003). Interestingly, simultaneous multiwavelength observations showed that the source was bright in the optical but faint in the radio (Courvoisier et al. 2004). To complement this campaign with imaging data of the parsec scale jet, we carried out a polarimetric monitoring of 3C 273 at six frequencies using the VLBA. These observations were described in Savolainen et al. (2006; hereafter Paper I) where a kinematical analysis of the motions of 11 components in the inner 10 mas was presented. We determined the relativistic Doppler and Lorentz factors of five individual emission features and found a substantial velocity gradient across the jet at about 1.5 mas from the core.

In the current paper, we present 5.0, 8.4, 15.3, 22.2, 43.2 and 86.2 GHz Stokes I images of 3C 273 observed on the 28th of February 2003, and describe in detail our method for extracting spectra of individual emission features from this VLBI data set. Spectra of 16 emission features in the jet are presented, and their reliability is discussed. The measured synchrotron turnover frequencies and flux densities together with emission region sizes are used to estimate the physical conditions in the jet, i.e. the magnetic field density and the relativistic electron energy density. We assess the uncertainties in these quantities by using a Monte-Carlo approach. We also calculate the anticipated amount of synchrotron self-Compton emission from the individual emission features and compare the prediction with the near-simultaneous X/\( \gamma \)-ray observations made by
INTEGRAL. In order to keep the number of pages in the current paper reasonable, only the spectra obtained at the first observation epoch are discussed here. The spectral evolution of the components will be studied in Paper III of the series.

Throughout the paper, we use a contemporary cosmology with the following parameters: $H_0 = 71$ km s$^{-1}$ Mpc$^{-1}$, $\Omega_M = 0.27$ and $\Omega_{\Lambda} = 0.73$. This corresponds to a luminosity distance of 747 Mpc at redshift of 3C 273 giving a linear scale of 2.7 pc / 1 mas. For the spectral index $\alpha$, we use the positive convention: $S_{\nu} \propto \nu^{\alpha}$.

2. Observations and data reduction

3C 273 was observed with the VLBA five times in 2003 (February 28th, May 11th, July 2nd, September 7th and November 23rd) for nine hours at each epoch (see Paper I for a more detailed description of the campaign). Observations were made at six different frequencies (5.0, 8.4, 15.3, 22.2, 43.2 and 86.2 GHz) recording in dual-polarisation mode. Individual scans at different frequencies were interleaved in order to obtain practically simultaneous multifrequency data. The scan length was 5 minutes at each frequency and the total integration time of 3C 273 was 65 minutes per frequency. Even though there might be some loss in the dynamic range due to the adopted “snapshot” type observing strategy (when compared to a full-track observation), we were able to achieve a good peak-to-rms ratio of $\sim 9000 : 1$ at 8 GHz. The reduced $uv$-coverage due to interleaved frequencies does not seem to be a serious problem, since the obtained dynamic ranges of $\sim 1000 – 9000 : 1$ at 5 – 43 GHz are more than adequate for the purposes of this study.

At the first epoch, we obtained 86 GHz data only from six telescopes, since Brewster and St. Croix did not have 3 mm receivers and baselines to Hancock and Pie Town did not produce any fringes. As we also had to flag a large fraction of the data from Mauna Kea, a rather poor $uv$-coverage at 86 GHz for the February observation was obtained. At all the other frequencies, the performance of the VLBA and, consequently, the correlated data were excellent.

The data were calibrated using NRAO’s Astronomical Image Processing System (AIPS; Bridle & Greisen 1994) and the Caltech Difmap package (Shepherd 1997) was used for imaging, self-calibration and model-fitting. Standard procedures were used in a priori amplitude calibration, single band delay determination, fringe fitting, self-calibration, and imaging. For 15 – 86 GHz data, a correction for atmospheric opacity was applied along with a priori amplitude calibration. As discussed in the following section, the calibration of the flux density scale was done with extra care since it is a crucial step in extracting correct component spectra from a multi-frequency VLBI data set.

2.1. Calibration of the flux density scale

Since the data will be used to construct spectra of different emission regions in the jet, we need an accurate flux density scale calibration at all frequencies. At low frequencies, the a priori amplitude calibration of the VLBA telescopes is known to be better than about 10%. The gain correction factors after applying the amplitude self-calibration at frequencies 5 – 22 GHz remain below 10%, which is consistent with the above assumption. At 43 GHz the individual corrections are slightly larger, but no systematic effect, which could mean an error in the flux density scale, is found.

In order to accurately calibrate the flux density scale, we compared the extrapolated zero baseline flux density of our calibrator source 3C 279 at 5, 8, 22 and 43 GHz, to the flux densities from VLA polarisation monitoring program (Taylor & Myers 2000) interpolated to the correct epoch. The flux densities from the automatic reduction process of the VLA monitoring program are accurate to $\sim 5\%$ at 5 and 8 GHz according to Taylor & Myers (2000). However, they warn about significant deviations in the flux densities at 22 and 43 GHz in some occasions due to problems with pointing, gain curves, or poor weather for a given session. As can be seen from Table 1 and Fig. 1, the VLA flux densities are systematically 4 – 7% larger than the VLBA flux densities at all frequencies. This is most likely due to a small amount of large-scale emission, which is resolved out in the VLBA images (see de Pater & Perley 1983 for VLA images of 3C 279). Since the variation of $S_{\text{VLA}}/S_{\text{VLBA}}$ between the frequencies is negligible, we conclude that the flux density scale calibration of our VLBA data at 5, 8, 22 and 43 GHz is better than expected from the nominal accuracy of the a priori amplitude calibration, being accurate to $\sim 5\%$ at 5, 8, 22 GHz and 43 GHz. At 15 GHz, similar accuracy is expected.

At 86 GHz, accurate amplitude calibration is more problematic and the gain correction factors after amplitude self-calibration are several tens of percents. Fortunately, we have single-dish flux density measurements of the calibrator 3C 279 at 90 GHz made with the Swedish-ESO Submillimetre Telescope (SEST) on the 15th and 17th of February and on the 2nd of April 2003. The averaged single-dish flux density for 3C 279 from February and April observations are $S_{\text{SEST}} = 13.3 \pm 0.7$ Jy and $S_{\text{SEST}} = 13.6 \pm 0.3$ Jy, respectively. According to the data from the Metsähovi Radio Observatory, the source

\begin{table}[h]
\centering
\caption{Flux densities of the calibrator source 3C 279 with the VLBA and the VLA}
\begin{tabular}{|c|c|c|c|}
\hline
Freq. & $S_{\text{VLBA}}$ & $S_{\text{VLA}}$ & $S_{\text{VLA}}/S_{\text{VLBA}}$ \\
(GHz) & (Jy) & (Jy) & \\
\hline
5 & 13.5 & 14.5 & 1.07 \\
8 & 13.6 & 14.3 & 1.05 \\
15 & 14.3 & – & – \\
22 & 14.3 & 15.2 & 1.06 \\
43 & 13.4 & 14.0 & 1.04 \\
\hline
\end{tabular}
\end{table}

\footnotetext[1]{At 86 GHz, the dynamic range is lower, $\sim 200 : 1$, due to high thermal noise, poor antenna efficiency and smaller number of antennas.}

\footnotetext[2]{The 3 mm system at Hancock suffered from technical problems. At Pie Town, there was a pointing problem as well as snowing during the observation.}

\footnotetext[3]{Final images were produced using Dan Homan’s Perl scripts (http://personal.denison.edu/~homand/).}
unresolved source or a slightly resolved source with a good estimate of the resolved emission missing from the VLBI image. Although being core-dominated, 3C 279 has kpc-scale structure which has to be taken into account. From Fig. 1 we can readily obtain an estimate of the amount of the missing flux in the large scale jet structure: it is ~ 6% between the VLBA and VLA, and ~ 10% between the VLBA and single-dish antennas. The single-dish observations have the same spectral index, \( \alpha = -0.1 \), between 37 and 90 GHz, as the VLA and VLBA observations have between 22 and 43 GHz. It is thus likely that the large discrepancy in the flux densities between the 86 GHz VLBA data and the 90 GHz SEST data is due to a gain calibration error in the VLBA data and not due to an excessive amount of emission that is resolved out in the 86 GHz VLBA image. The 86 GHz CLEAN image of 3C 279 (Fig. 2) shows structure up to \( \approx 1.5 \) mas from the core, while at the lower frequencies (8 – 43 GHz) emission is detected up to 5 mas. Fortunately, we can estimate the spectral index from 8 – 43 GHz data, \( \alpha = -0.8 \), for the extended structure between 0.5 – 5 mas from the core, and the expected amount of emission in this structure at 86 GHz is only \( \approx 0.9 \) Jy. This can be compared with the measured 0.6 Jy at 86 GHz, which implies that the observed large discrepancy in the flux density between the VLBA and SEST measurements is most likely not due to a resolved structure. Instead, we attribute only 10% of the 90 GHz single-dish flux density to the structure that is resolved out in our VLBA data. With this estimate, we obtain a gain correction factor of 1.4 for the 86 GHz VLBA measurement. Due to several uncertainties (exact amount of resolved flux, used interpolation) in this calibration, we conservatively estimate the corrected VLBA flux densities at 86 GHz to be accurate only to \( \approx 20\% \).

2.2. Modelling the source structure

The final CLEAN images of the first epoch observations are presented in Figs. 3 and 4. These images contain a wealth of information and they can be used – after a proper alignment, \( \nu \)-tapering and convolution with a common restoring beam – to construct maps of the jet spectral index (Lobanov 1998). However, we have not chosen this method for the spectral analysis, because describing kinematic and emission properties of the parsec scale jet is difficult with CLEAN components, and because it is not a very practical approach for comparing data with theory. Instead, a manageable number of spectra characterising the different emission regions in the jet is desirable at this point. Secondly, our model-fitting-based method, which uses \textit{a priori} information of the source structure from higher frequencies, makes it possible to measure spectra also for those features that have a size scale smaller than the resolving beam at low frequencies. This approach is similar to the “super resolution” technique (Bertero & De Mol 1996), and it allows us to extract spectral information in finer angular scales than is possible with the spectral index mapping.

Our approach is to derive the spectra of the prominent emission features in the jet using the flux densities measured by fitting the emission structure with two-dimensional Gaussian model components. This method has been earlier used for stud-

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Fig. 1. Flux densities of the calibrator source 3C 279 with the VLBA (diamonds), VLA (circles) and single-dish telescopes (Metsähovi and SEST; squares). We have interpolated the VLA and single-dish observations to match the VLBA observation epoch. As can be seen, the absolute flux density calibration of the VLBA data is very good at 5, 8, 22 and 43 GHz. The apparent shift between the VLA and VLBA data at these frequencies is probably due to large scale emission resolved out in the VLBA observations. However, the absolute flux density scale of the 86 GHz VLBA measurement clearly needs to be corrected.

Fig. 2. Naturally weighted CLEAN image of the calibrator source 3C 279 at 86 GHz observed on February 28 2003. The map peak flux density is 6.44 Jy beam\(^{-1}\), the off-source rms noise is 10.2 mJy beam\(^{-1}\), and the lowest contour level corresponds to 3 times the rms noise. The restoring beam size is 0.52 \( \times \) 0.11 mas with P.A. of \(-18.1^\circ\).
Naturally weighted 5, 8 and 15 GHz CLEAN images of 3C 273 observed on February 28th 2003. The contours start at $5\times$ the off-source rms noise and increase in steps of $\times2$. The restoring beam size, the peak intensity, and the off-source rms noise for each image are given in Table 2. The components are labeled with letters a, A, B, and C and a sequence number, e.g. “A1”. The features “a1–a5” refer to those components, for which no kinematical model was presented in Paper I. The regions B and C contain several components which are not plotted individually in the low frequency maps. See for Fig. 5 for a more detailed image of these emission regions. The radio core belongs to region C. Note that the images are not aligned.
Fig. 4. Naturally weighted 22, 43 and 86 GHz CLEAN images of 3C 273 observed on Feb 28 2003. See Fig. 3 for explanation.

ever, done (see also Paper I). Our B2 and B3 correspond to Q6/W6 and Q7/W7 of Attridge et al. (2005), respectively. The region containing components A2+A3+a4+a5 in our data corresponds to knots B1, bs, b1, B2 and b2 of Jorstad et al. (2005), with our A2 likely being their B1/bs. Our A3 could be their B2, but for a4 and a5 we cannot confidently point counterparts. A prominent component A1 most likely corresponds to U4/K4 of Homan et al. (2001), and B1/G1 of Jorstad et al. (2001a). Figure 4 in Krichbaum et al. (2001) suggests that this component has accelerated along the jet and corresponds to C10 of Abraham et al. (1996). The component a3 (which actually has two brightness peaks and thus may represent two knots) probably corresponds to C9 of Krichbaum et al. (1990, 2001). One cannot say much about components a1 and a2, but one of them could be related to either C7 or C8 of Zensus et al. (1990).

In the 5 GHz image, the jet is visible up to \( \sim 50 \) mas from the core, and the overall morphology in 5–22 GHz images is tubular, indicating a very small jet opening angle beyond region B, i.e. \( \sim 2 \) mas from the core. Another prominent feature apparent in the multifrequency images is the “wiggling” structure of the jet; the regions of bright emission are not collinear, and the ridge-line deviates both left and right from the average jet direction. The “wiggling” jet at 43 GHz was already discussed in Paper I, where it was suggested that plasma instabil-
Reliable extraction of spectral information from VLBI data is a difficult task, which requires understanding of several influencing factors and careful planning and reduction of the observations. The problematic aspects include limited and uneven sampling of the \(uv\)-plane, image alignment, accuracy of the flux density calibration, source variability, and a narrow range of observing frequencies (see Lobanov 1998 for a detailed discussion). Fortunately, many of these problems can be avoided when VLBA is used. As is evident from section 2.1, the calibration of the VLBA flux density scale is very accurate at all the observing frequencies except 86 GHz. Source variability is not a problem either, since VLBA allows fast switching between the observing frequencies resulting in a practically simultaneous multifrequency data set. We took the full advantage of this feature and chose to observe 3C 273 at six different frequencies ranging from 5 to 86 GHz. This is an unprecedented simultaneous frequency coverage for spectral observations with the VLBI technique, and it provides us with a sufficient number of data points for accurately measuring the spectral turnover of those VLBI components that have synchrotron peak within our frequency range. The remaining problems are uneven sampling at different frequencies, image alignment and uncertainties in the model-fitting procedure.

### 3. Method for extracting the component spectra

As can be seen from Figs. 3 and 4, the jet of 3C 273 is full of details at all the observed frequencies. We do not even try to extract spectra for all the details present in the images, but rather seek to obtain a simple parameterisation for the most prominent emission features. As shown in Paper I, the structure within 2 mas from the core is very interesting, and because it likely contains components having synchrotron peaks within our observed frequency range, we have tried to obtain spectra for all the bright components visible in the 43 GHz image of this region. This means that at lower frequencies we need to fit the visibilities a model with components having sizes and mutual separations significantly smaller than the resolving beam. Such an out-of-band extrapolation of spatial frequencies is possible, because we use the 43 GHz source model as \textit{a priori} information of the emission structure at lower frequencies, and because of the high signal-to-noise ratio of the observations (see Kovalev et al., 2005 and references therein). The underlying assumption of this technique is that the brightness centroids of the optically thin features do not show positional changes with frequency (i.e. there are no spectral index gradients across the components). Our method can be outlined as follows:

1) First, we have slightly simplified the 43 GHz source model presented in Paper I (see Fig. 5). In order to minimise the number of free parameters in the model, most components are substituted by circular Gaussians and elliptical components are used only when absolutely necessary (for components A1 and A3). Also, one weak component, which represents diffuse emission between B2 and B4 in Paper I, is omitted, and A1 is represented only by one elliptical component instead of three. This simplification increases the noise level in the residual map by a factor of four, and the excluding of the diffuse component from the model leaves \(\sim 0.1\) Jy of emission between components B2 and B4, which corresponds to \(\sim 5\%\) of the flux density of these components. However, we consider this to have only a very minor effect on the spectra if compared to the increased robustness of the model with fewer free parameters. Components a1–a3, which are not visible at 43 GHz, are naturally modelled at lower frequencies only.

2) The 43 GHz model is transferred to other frequencies. Here we assume that optically thin components have frequency independent positions, i.e. their brightness centroids are co-spatial at all frequencies. The relative positions of the components are fixed to the values obtained from model-fitting at 43 GHz, and the models are aligned by using the above assumption and determining the position offsets with bright and well-defined components B2 and A1.

3) Model-fitting is run at each frequency letting only component sizes and flux densities in the transferred model to vary. If any two components have a separation smaller than \(\sim 1/5\) of the beam size at the given frequency, these are replaced by single component (see Paper I). That is, we do not try to extrapolate beyond \(uv\)-radius corresponding to \(\sim 1/5\) of the beam size. Fig. 5 shows the transferred model of the innermost jet region at different frequencies.

4) We assume that the angular size of a component is either constant or varies smoothly over the frequencies. The sizes of the components are inspected as a function of frequency and
fitted with a constant value or with a power-law after removing clear outliers. Here we try to remove any bad size measurements which may have occurred in the model-fitting. We also try to estimate a possible frequency dependence of the component angular size, which may occur if e.g. the electron energy distribution is inhomogeneous within the component due to radiative losses (Marscher 1987, Sokolov et al. 2004). It is important to know if there are any frequency-dependent angular sizes, since the component size to be used in the subsequent analysis of synchrotron emission should be measured at the turnover frequency.

5) The angular sizes of the components are fixed to the values derived from the power-law fit (or to a constant value), and the model-fit is run again with the component flux densities as the only free parameters. This results in the final spectra for the components. The assumption of a smoothly varying component angular size allows us to use only a small number of parameters in the final model, resulting in robust estimates of the component spectra even in the size scales significantly smaller than the resolving beam.

The results of this spectral extraction method are presented in sections 4 and 5, but before discussing them, we look at possible error sources of the method.

3.2. The effect of uv-sampling

Unlike reconfigurable arrays such as VLA, VLBI networks cannot be scaled. This results in significantly different uv-coverages at different observing frequencies and poses a potentially hazardous problem to the attempts to infer spectral properties of astronomical sources from multifrequency VLBI data. Swain (1999) and Lobanov (1998) have studied the effect of uneven uv-sampling in the case of spectral imaging. Swain (1999) points out that simply convolving images at different frequencies with a common beam does not guarantee that significant errors in the spectral images will not be made. To study quantitatively the effect of limited and uneven uv-sampling, Lobanov (1998) created visibility datasets simulating a typical multifrequency VLBA observation and found that the signal-to-noise ratio of the pixel is the main factor determining the size of the flux density errors due to uneven uv-sampling: for pixels with SNR \(\gtrsim 7\), the fractional errors do not exceed 10%, while below that the errors increase rapidly (Lobanov measures SNR with respect to the self-calibration noise, which is 5–10 times larger than the formal rms noise of the image).

In order to study how much the uneven uv-coverage at different frequencies affects the flux densities and angular sizes of the model-fit components obtained with the method described in the previous section, we produced fake multifrequency uv-dataset with AIPS task UVCON. Simulated VLBA data at each of six frequencies (5, 8, 15, 22, 43 and 86 GHz) were generated from the same CLEAN component model of the structure of 3C 273 at 43 GHz. The simulation included several different types of measurement errors: Gaussian additive (thermal) noise, frequency- and station-dependent phase fluctuations (describing tropospheric phase noise and instrumental instabil-
majority of the components $\Delta S < 0.2$ Jy. Their average absolute error $\langle \Delta S / S \rangle = 0.09$ Jy. Consequently, the fractional errors in flux density are mostly below 10% for components with $S > 1$ Jy, increase to roughly 15% for components with $0.5 > S > 1$ Jy and are up to 60% for the weakest modelled features. The average fractional error in the component flux density for the whole data set, $\langle \Delta S / S \rangle = 13%$.

As can be seen from Fig. 7, the component angular size is a more uncertain quantity, and quite large deviations are possible. If the components having a separation smaller than $\sim 1/5$ of the beam size at the given frequency, as well as those having zero radius, are omitted, the average fractional error in the component angular size $\langle \Delta \alpha / \alpha \rangle = 23%$.

We do not find a correlation between the observing frequency and the measured flux density or the angular size, i.e. the uneven $uv$-coverages at different frequencies do not introduce any systematic error to either of them.

### 3.3. Uncertainties in the final spectra

Estimation of true $1\sigma$ uncertainties of the component flux densities in the model transfer method is difficult. The total uncertainty in the component flux density depends on the accuracy of the flux density scale calibration, errors in the model fitting procedure and errors due to the uneven $uv$-coverage. The errors in the flux density scale calibration $\sigma_{\text{cal}}$ were determined in section 2.1. In the previous section we estimated roughly the effect of uneven $uv$-coverage at different frequencies, $\sigma_{uv}$, from simulated visibility data and found that it is $\lesssim 0.2$ Jy regardless of the component flux density. This estimate, however, also includes uncertainties induced by model-fitting procedure itself, and hence, it partly overlaps with the uncertainties obtained for the individual model components with Difwrap program, $\sigma_{\text{Difwrap}}$ (Lovell 2000; see also Paper I). We used Difwrap to obtain error estimates for each component of the transferred model at each frequency by varying the flux density only while keeping other component parameters fixed. The average uncertainty from Difwrap analysis is $(\sigma_{\text{Difwrap}}) = 0.11$ Jy with a standard deviation of 0.06 Jy. This is of the same order as the errors due to uneven $uv$-coverage obtained for the simulated data sets in the previous section.

We estimate the total uncertainty in the component flux density to be $\sigma_{\text{tot}} = \sqrt{\sigma_{\text{cal}}^2 + \sigma_{\text{Difwrap}}^2 + \sigma_{uv}^2}$, where $\sigma_{\text{cal}}$ is $5\%$ of component flux density at 5–43 GHz and $20\%$ at 86 GHz, $\sigma_{\text{Difwrap}}$ is the uncertainty from the Difwrap analysis, and $\sigma_{uv}$ is the average error found in simulations, 0.09 Jy, for all the components. The final spectra indicate that $\sigma_{\text{tot}}$ indeed is a quite good estimate of the flux density errors of the bright components, while for the weak features it may be too conservative.

### 4. Frequency-dependence of the component angular size

The usual problem in the analysis of a synchrotron source is the highly non-linear dependence of the magnetic field, the relativistic electron energy density and the anticipated amount of synchrotron self-Compton flux density on the synchrotron...
Fig. 8. Component size as a function of frequency in a logarithmic scale. Non-zero sizes are shown with filled circles and arrows mark components with zero radius. Solid lines represent power-law fits to the non-zero component sizes, while dashed lines represent constant size. We note that we do not get reliable size estimates for the core, C1 and C2 at 15 GHz. Therefore, these data points are left out from the plots and constant size is assumed.
Table 3. Fitted spectral and size parameters for inner jet components

<table>
<thead>
<tr>
<th>Component</th>
<th>$r$ (mas)</th>
<th>$S_m$ (Jy)</th>
<th>$v_m$ (GHz)</th>
<th>$\alpha$</th>
<th>$\zeta$</th>
<th>$a(v_m)$ (mas)</th>
<th>$\delta$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Core</td>
<td>0.00</td>
<td>$1.9 \pm 0.2$</td>
<td>$50 \pm 10$</td>
<td>$-0.6 \pm 0.4$</td>
<td>$-0.07 \pm 0.01$</td>
<td>$-0.2 \pm 0.2$</td>
<td>$0.1 \pm 0.03$</td>
</tr>
<tr>
<td>C2</td>
<td>0.15</td>
<td>$3.2 \pm 0.1$</td>
<td>$38 \pm 1$</td>
<td>$-0.4 \pm 0.1$</td>
<td>$-0.09 \pm 0.02$</td>
<td>$0.0 \pm 0.1$</td>
<td>$0.0 \pm 0.03$</td>
</tr>
<tr>
<td>C1</td>
<td>0.30</td>
<td>$2.4 \pm 0.2$</td>
<td>$21 \pm 1$</td>
<td>$-0.8 \pm 0.2$</td>
<td>$-0.3 \pm 0.2$</td>
<td>$0.2 \pm 0.03$</td>
<td>$0.2 \pm 0.03$</td>
</tr>
<tr>
<td>B5</td>
<td>0.54</td>
<td>$1.4 \pm 0.2$</td>
<td>$8 \pm 1$</td>
<td>$-0.0 \pm 0.1$</td>
<td>$-0.0 \pm 0.1$</td>
<td>$0.2 \pm 0.03$</td>
<td>$0.2 \pm 0.03$</td>
</tr>
<tr>
<td>B4</td>
<td>1.07</td>
<td>$2.5 \pm 0.1$</td>
<td>$16.5 \pm 6$</td>
<td>$-0.0 \pm 0.1$</td>
<td>$+0.1 \pm 0.1$</td>
<td>$0.2 \pm 0.03$</td>
<td>$0.2 \pm 0.03$</td>
</tr>
<tr>
<td>B2</td>
<td>1.31</td>
<td>$6.1 \pm 0.3$</td>
<td>$5.2 \pm 0.7$</td>
<td>$-0.6 \pm 0.1$</td>
<td>$-0.2 \pm 0.0$</td>
<td>$0.1 \pm 0.03$</td>
<td>$0.1 \pm 0.03$</td>
</tr>
<tr>
<td>B3</td>
<td>1.54</td>
<td>$3.2 \pm 0.1$</td>
<td>$15.8 \pm 7$</td>
<td>$-0.7 \pm 0.1$</td>
<td>$-0.3 \pm 0.2$</td>
<td>$0.2 \pm 0.03$</td>
<td>$0.2 \pm 0.03$</td>
</tr>
<tr>
<td>B1</td>
<td>1.66</td>
<td>$0.6 \pm 0.2$</td>
<td>$11 \pm 5$</td>
<td>$-0.2 \pm 0.2$</td>
<td>$-0.0 \pm 0.1$</td>
<td>$0.2 \pm 0.03$</td>
<td>$0.2 \pm 0.03$</td>
</tr>
</tbody>
</table>

The Doppler factor of component B2 has a wrong value in Paper I. The correct value is here. See text.

A5 3.42 $-0.9 \pm 0.1$ 0.0 $0.1$ 1.3
A3 4.04 $-0.3 \pm 0.1$ 0.0 $0.2$ 0.3
A4 4.54 $-0.5 \pm 0.2$ $-0.2 \pm 0.2$ 0.5
A2 5.27 $-0.8 \pm 0.1$ $-0.2 \pm 0.1$ 0.9
A1 10.4 $-0.7 \pm 0.1$ $-0.2 \pm 0.1$ 1.3
A3 14.3 $-1.1 \pm 0.1$ $-0.1 \pm 0.1$ 2.5
A2 17.3 $-1.4 \pm 0.3$ $-0.2 \pm 0.1$ 3.0
A1 22.8 $-1.3 \pm 0.2$ $-0.2 \pm 0.1$ 3.5

Component A1 has a broken power-law with a break frequency $v_{break} = 13 \pm 3$ GHz. The spectral index above $v_{break}$ is $-1.0 \pm 0.1$.

At frequencies above the turnover, angular sizes are frequency-dependent if there is a gradient in the magnetic field or in the upper cutoff energy of the electrons within the emission feature. Such a gradient is expected, for example, in the shocked jet models, where emitting electrons are accelerated in a thin shock front. As soon as the electrons leave the acceleration region their energy decreases due to radiative cooling, establishing a gradient in the electron energy distribution across the source (Sokolov et al. 2004). Now the angular size of the component will be smaller at higher frequencies. Marscher (1987) points out that this type of a frequency-dependent angular size will steepen the optically thin spectrum of the component. He considers the case where the depth along the line of sight is frequency-independent and the angular size varies as $a \propto \nu^\zeta$. The optically thin flux density would then change as $S \propto \nu^{\alpha+2\zeta}$. This is a simplification, but shows that the underlying optically thin spectral index, $\alpha$, may be flatter than the measured one.

In non-uniform sources, frequency dependent angular sizes are expected also when the source is partially opaque, i.e. at frequencies below the turnover. Marscher (1977) calculates this effect for several combinations of different magnetic field and electron density gradients, but, unfortunately, reliable measurement of the angular size at frequencies below the turnover is a very challenging task.

In Fig. 8 we have plotted the angular sizes of the model components as a function of frequency in a logarithmic scale for the February 2003 observation of 3C 273. Components A1 and A3 are elliptical, and for those the presented angular size is the geometrical mean of the major and minor axes. We have fitted the angular sizes with a constant value or with a power-law function, $a \propto \nu^\zeta$, and these fits are shown in the figure by solid and dashed lines, respectively. Clear outliers were left out from the fits.

Fig. 8 and fitted values of $\zeta$ in Tables 3 and 4 show that in most cases the component angular sizes are either close to constant across frequencies or slightly decreasing with increasing frequency. Fig. 8 does not contain error bars, since it is rather difficult to obtain reliable $1\sigma$ errors for the component sizes. In calculating errors for the fitted values of $\zeta$, we have assumed that the uncertainties in the angular sizes of components in the crowded regions of A2/A3/a4/a5, B and C are ~ 20%, and that larger and more isolated components have sizes accurate to 10%. With these assumptions, there are four components (out of 16) having $\zeta$ clearly deviating from zero (A1, A2, A1, and B3). The average value of the power-law index for these components is $\langle \zeta \rangle = -0.2$. As will be seen in the following section, all of the frequency-dependent angular sizes are measured for the optically thin part of the spectrum and we consider it likely that the observed frequency-stratification is due to radiative losses within the components.

5. Component spectra

The final component spectra obtained by our model transfer method are shown in Figs. 9, 10 and 11. We have fitted the spectra of the components in regions B and C with a function describing self-absorbed synchrotron radiation emitted by electrons having a power-law energy distribution ($N(E) = N_0 E^{-\delta}$) in a homogeneous magnetic field (e.g. Pacholczyk 1970):

$$S(\nu) = S_m \left( \frac{\nu}{v_m} \right)^{\frac{\alpha}{2}} \frac{1 - \exp(-\tau_m(\nu/v_m)^{\alpha-5/2})}{1 - \exp(-\tau_m)},$$

(1)
where $S_m$ is the maximum flux density reached at the turnover frequency $\nu_m$ corresponding to an optical depth $\tau_m$, and $\alpha = -(x-1)/2$ is the optically thin spectral index. Spherical geometry of the source is assumed. The optical depth at the turnover can be approximated as

$$\tau_m = \frac{3}{2} \left( \sqrt{1 - \frac{8\alpha}{15/2}} - 1 \right)$$

(Törrler et al. 1999). The spectrum describing a homogeneous source (i.e. with an optically thick spectral index of 2.5) can fit all the components in regions B and C, and the data does not require flatter self-absorbed spectra for any of them (although flatter spectral indices are possible for most components). Therefore, we do not consider inhomogeneous component spectra here. We also assume that the optical depth for free-free absorption in the jet is significantly below unity at our observing frequencies; this assumption is supported by the lack of optically thick spectral indices steeper than 2.5 in our spectra.

If the components are closer to each other than $\sim 1/5$ of the beam size at low frequencies (5 and 8 GHz), we have measured the flux density of the whole region in question at those frequencies and tried to find a spectral dissection which best describes the data. In Figs. 9 and 10 we show these dissections for the region C and for the component pair B1 and B3, respectively. In both cases homogeneous synchrotron spectra are assumed for all the components. One might anticipate that the core had an optically thick spectral index less than 2.5, but our data does not require that. Thus, we have used the homogeneous spectrum also for the core, although a flatter self-absorbed spectrum is possible.

The parameters describing the synchrotron spectra ($S_m, \nu_m$ and $\alpha$) of the core and components C1–C2 and B1–B5 are collected in Table 3. The reported uncertainties are formal 1$\sigma$ errors from the fitting program. In the same table, we have also listed the component’s distance from the core, $r$, the index describing the frequency-dependence of the component angular size, $\zeta$ (see the previous section), the component angular size at the turnover frequency $\alpha(\nu_m)$ and the Doppler factor, $\delta$, reported in Paper I.

For the components located at distances larger than 2 mas from the core, $\nu_m$ is below 5 GHz, and we can measure only the optically thin part of the spectrum. The spectra of these components are fitted with a simple power-law, except in the case of A1 where such a simple power-law does not adequately describe the data. The spectrum of this component is fitted with a broken power-law indicating a change in the spectral slope of $\Delta\alpha \approx 0.3$ at $\nu_{\text{break}} \approx 13 \pm 3$ GHz. The spectral indices of the components A1–A3 and a1–a5 are listed in Table 4 together with $r$, $\zeta$, and the maximum size of the component, $a_{\text{max}}$.

It is interesting to look for possible trends in the parameters describing the synchrotron spectrum with respect to the distance from the core. In Fig. 12 we have plotted $S_m$ and $\nu_m$ as a function of $r$ in the logarithmic scale. The peak flux density does not show any clear trend, while the peak frequency seems to be steadily decreasing as a function of distance from the core (Törrler et al. 1999).

The Doppler factor reported in Paper I for component B2 ($\delta = 12.2 \pm 2.8$), was incorrect. It was determined using the variability timescale, $\Delta t_v = dt/d \ln(S)$, which was measured from the flux density curve in Fig. 12 of Paper I. However, it turned out that the minimum in the curve at the epoch 2003.68 was likely due to B2 overlapping with the nearby model component (see Fig. 2 in Paper I). Thus, epoch 2003.68 probably does not correspond to the true flux density curve minimum, and it should not be used in the variability timescale calculation. The new value for $\Delta t_v$, ignoring the 2003.68 data, is $190 \pm 40$ days and corresponding $\delta = 7.1 \pm 2.5$. This value of Doppler factor yields a Lorentz factor of $6.6 \pm 1.4$ and a viewing angle of $8.1 \pm 3.0$° for B2. It is notable that the new viewing angle value agrees much better with the other components than $\theta = 4.3 \pm 1.6$° reported in Paper I. The new value of $\delta$ does not affect the conclusion in Paper I stating that there is a significant gradient in component velocity across the jet – on the contrary, with the new value of $\Gamma$, the gradient is even larger.
Fig. 11. Spectra of the VLBI components in the parsec scale jet of 3C 273 observed with the VLBA on February 28th 2003. The solid lines represent spectral fits – either a self-absorbed synchrotron spectrum or a simple power-law (except for A1 where a broken power-law is used). The downward pointing arrow in a3 marks a 3σ upper limit. It is noted that for the core and components C1, C2, B1, B3 and B5, there are more data constraining the spectral fits than what is shown here (see spectral dissections in Figs. 9 and 10).
6. Physical conditions in the parsec scale jet

Problems in understanding the relativistic jets are, for a large part, due to our inability to determine the physical conditions in the jet. While we would like to have accurate measurements of the quantities like bulk velocity, magnetic field density, particle number density, and Alfvén speed in the jet and in the ambient medium, for most of them we have only rather vague estimates. Usually we have even less knowledge about how these quantities vary within the jet.

In the following, we adopt the standard synchrotron theory and assume that all emission sources are uniform and spherical. With this assumption, we can use the observed component spectra in Fig. 11 – together with measured component sizes – to estimate the magnetic field density, \( B \), the electron energy distribution normalisation factor, \( N_0 \), relativistic electron energy density, \( U_e \), and the expected synchrotron-self-Compton flux density at a given frequency, \( S_{\text{SSC}}(\nu) \). The highly non-linear dependence of these quantities on the parameters \( v_m \) and \( a(v_m) \) leads to large uncertainties in the final values, but due to the good accuracy of the spectral parameters measured here, the results are still useful and even the variation between the components can be studied to some extent.

6.1. Magnetic field and electron energy density

Let us assume that each VLBI component in Fig. 5 corresponds to a spherical and uniform synchrotron source with a relativistic electron energy distribution of the form

\[
N(E) = \begin{cases} 
N_0 E^{-\alpha}, & E \leq \gamma_1 m_e c^2 \\
0, & \gamma_1 m_e c^2 < E \leq \gamma_2 m_e c^2 \\
0, & E > \gamma_2 m_e c^2.
\end{cases}
\]

The electrons emit incoherent synchrotron radiation, for which the spectral shape is given by equation 1, and which can be described by three parameters: self-absorption turnover frequency \( v_m \), flux density at turnover \( S_m \) and optically thin spectral index \( \alpha = -(s-1)/2 \). Adopting expressions from Marscher (1987), the magnetic field density within the source is

\[
B = 10^{-5} b(\alpha) a_S(v_m)^4 v_m^3 S_m \left( \frac{\delta}{1 + \frac{1}{\delta}} \right) \ [\text{G}],
\]

and the normalisation factor of the electron energy distribution is

\[
N_0 = n(a) D_{\text{Gpc}}^{-1} a_S(v_m)^{-7(4\alpha)} v_m^{-5(4\alpha)} S_m^{3-2\alpha} \times (1 + \frac{1}{\delta})^{2(3-\alpha)} -2(2-\alpha) \ [\text{erg}^{-2\alpha} \text{cm}^{-3}],
\]

where the diameter of the spherical component at the turnover frequency, \( a_S(v_m) \), is given in milliarcseconds and \( v_m \) in GHz. \( D_{\text{Gpc}} \) is the luminosity distance to the source in Gpc (0.747 Gpc in our case), \( \delta \) is the Doppler factor and \( S_m' \) is the flux density at \( v_m \) extrapolated from the straight-line optically thin slope and given in Janskys. \( S_m' \) can be calculated from \( S_m \) and \( \tau_m \) (given in section 5) by using the following expression:

\[
S_m' = \frac{S_m \tau_m}{1 - e^{-\tau_m}}.
\]

Since we have assumed that emission features have a spherical geometry, the \( FWHM \) values of Gaussian components listed in Table 3 have to be corrected before inserting them into equations 4 and 5. The 50%-visibility points coincide for a Gaussian profile and for the profile of an optically thin sphere with diameter of 1.8 times the \( FWHM \) of the Gaussian profile. Hence, we use \( a_S(v_m) = 1.8a(v_m) \). Marscher (1987) gives terms \( b(\alpha) \) and \( n(\alpha) \) for different values of \( \alpha \) in his Table 1. Here, logarithmic interpolation is used for values of \( \alpha \) between the tabulated ones.
Once $N_0$ is determined, it is possible to integrate $N(E)$ and $N(E)/E$ over $E$ to get the number density and the energy density of the relativistic electrons, respectively. Since the number density is highly dependent on the value of the low energy cut-off, which is currently an unknown quantity, we will calculate only the energy density (for which our ignorance of $\gamma_1$ makes little difference):

$$U_{\text{re}} \approx f(\alpha, \nu_2/\nu_1) D_{\text{Gpc}}^{-1} a_3(\nu_m)^{-9} \times \nu_m^{-2} S_m^{-4} (1+z)^7 \delta^{-5} \text{[erg cm}^{-3}] \tag{7}$$

Again, the function $f(\alpha, \nu_2/\nu_1)$, where $\nu_1$ and $\nu_2$ are low and high frequency cut-offs of the synchrotron spectrum (corresponding to cut-offs $\gamma_1$ and $\gamma_2$ in the electron energy distribution; $\nu_{1,2} = 2.8 \times 10^6 B_{\gamma_{1,2}}^2$, can be found in Marscher (1987). We have adopted $\nu_2/\nu_1 = 10^6$ (corresponding to $\gamma_2/\gamma_1 = 10^3$) for our calculations. However, the exact value of this ratio has very little influence on $U_{\text{re}}$ for $\alpha \lesssim -0.5$, and B1 is the only component for which $\nu_2/\nu_1$ has a notable effect ($U_{\text{re}}(\text{B1})$ changes by a factor of two for an order of magnitude change in $\nu_2/\nu_1$). Besides $U_{\text{re}}$, we are also interested in the energy density of the magnetic field, which can be directly calculated from $B$:

$$U_B = \frac{B^2}{8\pi} \text{[erg cm}^{-3}] \tag{8}$$

Since equations 4, 5 and 7 have a very strong non-linear dependence on the observed quantities, the linear approximation of the error propagation formula cannot be used. Moreover, the strong non-linearity shifts the expected value of the output quantity with respect to the directly calculated “best value” if measurement errors in the input quantities are taken into account (see Paper I). For these reasons, we have employed a Monte Carlo approach to calculate the expected values and standard deviations of $B$, $N_0$, $U_{\text{re}}$ and $U_B$. We drew a million random values for each observed quantity ($\nu_m$, $S_m$, $a_3(\nu_m)$, $\alpha$ and $\delta$) from a truncated normal distribution having an expected value and a standard deviation corresponding to observed values given in Table 3. The lower and upper limits of the truncated normal distribution were chosen so that any non-physical values are avoided, e.g. no negative values were allowed for size, frequency or flux density. Since the Doppler factor was determined only for components C2, B2, B3 and B4 in Paper I, we have used an average value ($\delta = 5.6$ for rest of the components. An uncertainty $\sigma(\delta) = 2.0$ was assumed for these components. For the stationary core, we have considered two cases: $\delta = 5.6$ and $\delta = 1$ (i.e. non-Doppler-enhanced core).

From the random input values, a million output values for $B$, $N_0$, $U_{\text{re}}$ and $U_B$ were calculated. Due to the very long tails of the formed probability density functions, we have taken a logarithm of the distributions before calculating the expected values and standard deviations. Table 5 lists the results. The errors calculated using the Monte Carlo method are admittedly large, but nevertheless, we now have a firm estimate of our level of uncertainty regarding the physical parameters directly calculated from the observed quantities and standard synchrotron theory without an assumption of e.g. minimum energy condition. Also, the uncertainties in the magnetic field density are actually rather modest for the brightest components having well-measured sizes and peak frequencies – only a factor of two for B3 and B4.

Fig. 13 shows $B$ as a function of distance from the core, $r$, in a logarithmic scale. The core has the magnetic field density of an order of 1 Gauss, a value compatible with earlier measurements made using infrared and optical variability (Courvoisier et al. 1988). Generally, the magnetic field becomes weaker as we move farther from the core. However, there is a large difference in $B$ between components B2 and B3, both having roughly the same distance from the core. Component B2, which is located $\approx 0.6$ mas north of B3 and which has both velocity and direction of motion clearly different from B3 and B4 (Paper I), has over two orders of magnitude smaller $B$ than component B3. This gradient in the magnetic field across the jet may be related to the velocity gradient in the same location reported in Paper I. We will further discuss this in section 6.4. In Fig. 13, there are plotted two examples of a power-law dependence $B \propto r^b$ with $b = -1$ and $b = -2$. If component B2, with its anomalously small value of $B$, is excluded, the magnetic field density seems to drop inversely proportional to the distance from the core – or possibly with even a flatter slope. The average magnetic field density at $r = 1$ mas is $\approx 0.1$ G, while in the kpc-scale jet, at $r \approx 10^4$ mas, the equipartition magnetic field is $\approx 10^{-4}$ G (Jester et al. 2005). This gives $b = -0.8$ between milliarcsecond and arcsecond scales, further suggesting that, on average, magnetic field decays with power-law index $b \geq -1$. However, due to the small number of data points and large errors in $B$, this result needs to be confirmed using more data. It would be especially important to determine the magnetic field density at distances $r > 10$ mas, which would require observations at frequencies below 5 GHz in order to measure $\nu_m$ in this part of the jet.
Table 5. Physical parameters of the emission regions

<table>
<thead>
<tr>
<th>Component</th>
<th>(r) ((\text{mas}))</th>
<th>(\log_{10}(B)) ((\text{Gauss}))</th>
<th>(\log_{10}(N_e)) ((\text{erg}^{-26} \text{cm}^{-3}))</th>
<th>(\log_{10}(U_{re})) ((\text{erg cm}^{-3}))</th>
<th>(\log_{10}(U_B)) ((\text{erg cm}^{-3}))</th>
<th>(\log_{10}(U_B/U_{re}))</th>
<th>(\log_{10}(S_{SSC}(100\text{keV})))</th>
</tr>
</thead>
<tbody>
<tr>
<td>Core</td>
<td>0.00</td>
<td>+0.3 ± 0.6</td>
<td>-6.9 ± 1.8</td>
<td>-5.0 ± 1.2</td>
<td>-0.8 ± 1.1</td>
<td>+4.2 ± 1.6</td>
<td>-4.4 ± 1.9</td>
</tr>
<tr>
<td>Core²</td>
<td>0.00</td>
<td>-0.4 ± 0.5</td>
<td>-3.4 ± 1.2</td>
<td>-1.5 ± 0.9</td>
<td>-2.2 ± 1.1</td>
<td>-0.7 ± 1.3</td>
<td>-0.5 ± 1.2</td>
</tr>
<tr>
<td>C2</td>
<td>0.15</td>
<td>-0.4 ± 0.5</td>
<td>-4.2 ± 1.2</td>
<td>-3.8 ± 1.3</td>
<td>-2.1 ± 0.9</td>
<td>+1.7 ± 1.6</td>
<td>-2.1 ± 1.2</td>
</tr>
<tr>
<td>C1</td>
<td>0.30</td>
<td>-1.2 ± 0.7</td>
<td>-6.4 ± 1.9</td>
<td>-3.7 ± 1.6</td>
<td>-3.8 ± 1.3</td>
<td>-0.1 ± 2.1</td>
<td>-3.1 ± 1.9</td>
</tr>
<tr>
<td>B5</td>
<td>0.54</td>
<td>-2.2 ± 1.2</td>
<td>-7.6 ± 3.6</td>
<td>-3.7 ± 2.7</td>
<td>-5.8 ± 2.4</td>
<td>-2.1 ± 3.6</td>
<td>-4.0 ± 3.4</td>
</tr>
<tr>
<td>B4</td>
<td>1.07</td>
<td>-1.0 ± 0.3</td>
<td>-6.1 ± 1.1</td>
<td>-5.2 ± 1.1</td>
<td>-3.3 ± 0.7</td>
<td>+1.9 ± 1.3</td>
<td>-3.7 ± 1.1</td>
</tr>
<tr>
<td>B2</td>
<td>1.31</td>
<td>-3.1 ± 0.5</td>
<td>-3.5 ± 1.4</td>
<td>-2.1 ± 1.3</td>
<td>-7.6 ± 1.0</td>
<td>-5.5 ± 1.6</td>
<td>-0.4 ± 1.3</td>
</tr>
<tr>
<td>B3</td>
<td>1.54</td>
<td>-0.7 ± 0.3</td>
<td>-7.2 ± 1.2</td>
<td>-5.0 ± 1.0</td>
<td>-2.7 ± 0.6</td>
<td>+2.3 ± 1.2</td>
<td>-3.9 ± 1.2</td>
</tr>
<tr>
<td>B1</td>
<td>1.66</td>
<td>-1.9 ± 1.5</td>
<td>-2.2 ± 2.3</td>
<td>-2.2 ± 2.6</td>
<td>-5.2 ± 3.0</td>
<td>-3.0 ± 4.0</td>
<td>-1.6 ± 2.3</td>
</tr>
</tbody>
</table>

² No Doppler boost i.e. \(\delta = 1\).

Fig. 14. Logarithm of the relativistic electron energy density of the components as a function of logarithmic distance from the core. The open circle corresponds to the core with \(\delta = 5.6 ± 2.0\) assumed, while the open square corresponds to the core showing no Doppler boost (\(\delta = 1\)). The core is placed at \(r = a_{core}/2 = 0.035\) mas.

Fig. 15. Logarithm of the ratio of the energy densities in the magnetic field and in the relativistic electrons as a function of logarithmic distance from the core. The open circle corresponds to the core with \(\delta = 5.6 ± 2.0\) assumed, while the open square corresponds to the core showing no Doppler boost (\(\delta = 1\)). The core is placed at \(r = a_{core}/2 = 0.035\) mas. The dashed line marks the equipartition value.

In Figs. 14 and 15 we have plotted \(U_{re}\) and the ratio \(U_B/U_{re}\) as a function of \(r\), respectively. The very large errors in \(U_{re}\) limit analysis, but two things seem evident. Firstly, component B2, which has direction of motion \(\Phi = -113^\circ\), has 2–3 orders of magnitude larger \(U_{re}\) than southern components B3 and B4, which have \(\Phi = -141^\circ\) (Paper I). Secondly, an average value of relativistic electron energy density within 2 mas from the core is about \(10^4\) erg cm\(^{-3}\).

The errors in ratio \(U_B/U_{re}\) are even greater than in \(U_{re}\), since it depends on the seventeenth power of the product \(a(y_m)v_m\). From Fig. 15 it is, however, clear that if the plasma flowing through the core has relativistic bulk velocity, the core has significantly more energy in the magnetic field than in the radiating particle population. B3, B4 and C2 also seem to show a slight magnetic field dominance (although not as clear as the core), while the northern component B2, on the other hand, has the relativistic electron energy density at least \(10^4\) times larger than the magnetic field energy density.

In addition to magnetic field and electrons, there should be energy stored in cold protons, if the jet is composed of electron-proton plasma. If one cold proton per a relativistic electron is assumed, the energy density of protons, \(U_p = m_p c^2 \int N(E)dE\), where \(N(E)\) is given in equation 3. The result depends heavily on the low energy cut-off of the relativistic electron distribution, \(\gamma_1\), and therefore cannot be calculated with the current data. However, if the jet is required to be rest mass dominated beyond the core instead of being Poynting flux dominated (i.e. \(U_p > U_B\)), components C2, B3 and B4 must have \(\gamma_1 \leq 10\).
6.2. Synchrotron self-Compton flux density

Using Gould’s (1979) formulae for synchrotron and inverse Compton emission from an optically thin spherical source, Marscher (1987) presents the following formula for calculating the first-order synchrotron self-Compton (SSC) flux density:

\[
S_{SSC}(E_{\text{keV}}) = d(\alpha)\delta(1-2\alpha)\nu_m^{-5-3\alpha}\nu_m^{2(2-\alpha)}
\times (h\nu_{\text{keV}})^{\alpha} \ln\left(\frac{\nu_m}{\nu}\right)^{2(2-\alpha)} [\mu\text{Jy}],
\]

where \( (h\nu)_{\text{keV}} \) is the energy of the scattered photon in keV, and \( d(\alpha) \) is tabulated in Marscher (1987). Equation 9 gives a robust estimate of the expected SSC flux density unless the source is rapidly varying, in which case light-travel time effects have an impact on the result. We have used the same Monte Carlo method which was described in the previous section to calculate the expected amount of SSC flux density at 100 keV from components located within 2 mas from the core. Photon energy of 100 keV was chosen, because we want to compare the predicted SSC flux density with the observed X-ray flux density from the jet. Grandi & Palumbo (2004) studied the X-ray spectral variability of 3C 273 over a 5-year period using BeppoSAX data, and they concluded that, at energies higher than 40 keV, the non-thermal jet component dominates over the thermal Seyfert-like component at all epochs. Therefore, in order to minimize the contribution from the Seyfert-like component (i.e. accretion disk and a hot plasma corona), the comparison is made at a photon energy of 100 keV, at which 3C 273 is still well detected by INTEGRAL (Courvoisier et al. 2003). The ratio \( v_2/v_m \) inside the logarithm in equation 9 is assumed to be 10^4 for all components.

The anticipated SSC flux density at 100 keV for the core and for components C1–C2 and B1–B5 is given in Table 5. Fig. 16 shows \( S_{SSC}(100 \text{ keV}) \) as a function of \( r \) together with the observed 100 keV flux density from the January 2003 INTEGRAL pointings. From the figure, it is clear that the mm-wave core can produce the observed hard X-ray flux density through SSC-mechanism only if it is not relativistically beamed (the maximum \( \delta \) for the core to emit the observed amount of 100 keV X-rays within the errors is ~ 1.5) or if we have overestimated its size (if the core had \( a(v_m) \approx 0.02 \text{ mas} \) instead of 0.07 mas, it would produce enough X-rays through SSC-mechanism to match the INTEGRAL observation). If \( \delta = 5.6 \) is assumed, the upper limit of \( S_{SSC}(100 \text{ keV}) \) for the core is over two orders of magnitude below \( S_{\text{INTEGRAL}}(100 \text{ keV}) \). A weak self-Compton X-ray flux from the core is compatible with the results obtained by Mantovani et al. (2000) in a similar study. It is also compatible with the findings of Unwin et al. (1985) who applied the inhomogeneous-jet model of Königl (1981) for the radio core of 3C 273. Similarly, most of the other components also show weak SSC flux density. At best, if both C1 and C2 had their maximum values, they would make up about 0.14 \( \mu \text{Jy} \) or 25% of the observed 100 keV flux density components.

---

5 For January 2003 (2003-01-05 – 2003-01-18) observations, a joint 2 power-law fit (soft-excess + hard X-ray) of INTEGRAL and XMM-Newton data with OSA 5.0 (INTEGRAL) and SAS 6.5 (XMM-Newton) gives \( S_{\text{INTEGRAL}}(100 \text{ keV}) = (5.74 \pm 0.51) \times 10^{-4} \text{ mJy} \) (Chernyakova et al. 2005; Marc Türl er, private communication).
the timescale of variability in size for components is defined as 
\[ \Delta t = \frac{dt}{\ln(a(S_{\text{max}})/a(S_{\text{min}}))}, \] where \( a(S_{\text{max}}) \) and \( a(S_{\text{min}}) \) are the observed sizes of the component measured at the epoch of its maximum and minimum flux density, respectively, and \( dt \) is the time between \( S_{\text{max}} \) and \( S_{\text{min}} \). For B2 at 43 GHz (see Paper I), \( \Delta t = 2.1 \) yr, which is four times longer than the observed flux variability timescale. The difference between \( \Delta t \) and \( \Delta t_0 \) is larger than is expected for adiabatic losses in optically thin shocked gas (expected \( \Delta t_0 \approx \Delta t_s/1.4 \) for B2 with \( \alpha = -0.6 \); see Marscher & Gear 1985). Since neither synchrotron nor adiabatic losses seem to be strong enough to explain the observed \( \Delta t_0 \), inverse Compton losses are probably dominant in B2. This further suggests that a significant fraction of X-ray emission could originate in component B2. Adiabatic losses are also probably significant, though not dominating.

In January 2003, 3C 273 was in a relatively low state at radio and hard X-ray frequencies, meaning that the jet was in a quiescent stage. According to our results, we identify three sources that in principle could have emitted a significant fraction of the observed 100 keV flux density through SSC-mechanism during this quiet state:

1) The hard X-rays could come from the mm-wave core if its size is significantly overestimated or if it is unbeamed. The first option does not seem likely because the measured size is constant across the frequencies (see Fig. 8). The latter possibility also seems improbable since VLBI observations of radio cores generally suggest beaming. Therefore, we find it unlikely that SSC emission from the mm-wave core would appreciably contribute to the hard X-ray radiation.

2) Since component C2 corresponds to a small flare in the 37 GHz flux density curve peaking in January, it could have had higher brightness temperature and, consequently, higher SSC flux density during the INTEGRAL observation than what is observed with the VLBA on February 28. Hence, C2 possibly was a significant SSC emitter in January when it was closer to the core, but, unfortunately, this is hard to quantify.

3) From Fig. 16, B2 is the prime candidate for the source of a significant SSC-emission. How reliable is this result? Can we consider B2 – one peculiar component located ~ 1.3 mas downstream of the core – to emit practically all the observed 100 keV X-rays, while the SSC flux densities of the other components in the VLBI jet are weaker by several orders of magnitude? It is possible, but the current data does not allow strong conclusions to be drawn about B2’s contribution to the observed X-ray flux density. Firstly, the reported uncertainty in \( S_{\text{SSC}}(100 \text{ keV}) \) alone is well over an order of magnitude. Secondly, if there are errors in the model alignment at low frequencies, these may further increase the uncertainty in \( S_{\text{SSC}}(100 \text{ keV}) \). Large errors in the alignment are not likely, but unfortunately, even rather modest deviations in the flux density of B2 at 5 and 8 GHz due to an alignment error could shift \( \nu_m \) by 1–2 GHz and \( S_m \) by 0.5–1 Jy, resulting in an order of magnitude error in \( S_{\text{SSC}}(100 \text{ keV}) \). This should be kept in mind when looking at the Fig. 16.

6.3. On the assumption of homogeneity

In the previous sections, we have determined physical parameters describing synchrotron emitting plasma – such as \( B, N_0, \) and \( U_B/U_{\text{re}} \) – from the observed parameters by assuming a homogeneous, spherical source. How reliable is the assumption of homogeneity? Gradients in \( B \) and \( N_0 \) within the emission feature should manifest themselves as an optically thick spectral index flatter than +2.5. However, our observations do not require this for any of the components, suggesting that emission features can be considered more or less homogeneous. Even if there were inhomogeneities present, the turnover frequency is determined mainly by the part that has the highest magnetic field and electron density.

In the case of component B3, the observed angular sizes in Fig. 8 suggest that there is frequency stratification above \( \nu_m \), which is likely due to radiative losses (see section 4). This implies that component B3 is actually not homogeneous. However, as long as we are able to measure the angular size of the component at the turnover frequency, there should not be any significant effect on \( B \) and \( N_0 \). Also, the measured brightness temperature should provide a robust estimate of the expected SSC flux density, unless \( T_B \) keeps to increase below \( \nu_m \) due to inhomogeneities. As can be seen from Fig. 10, the spectrum below \( \nu_m \) is not likely to be flatter than \( S \propto \nu^2 \), and thus \( T_B \) has its maximum at \( \nu_m \) and is suited for the SSC-calculation.

The effect of frequency stratification on the optically thin spectral index may be of some concern in the case of B3 since due to the frequency dependent angular size, the observed spectral index may be steeper than the underlying value. In order to get a rough estimate of how much the observed spectrum steepens, we have run again the model-fitting algorithm for component B3 while assuming it to have a constant size over frequency, \( \alpha = 0.15 \) mas. The resulting spectrum has \( S_m = 2.9 \) Jy, \( \nu_m = 14 \) GHz, and \( \alpha = -0.5 \). The turnover frequency and flux density changed very little from values given in Table 3, which shows the robustness of the spectra derived in this paper. With the optically thin spectral index of \( \alpha = -0.5 \), the physical parameters calculated from this spectrum are: \( \log_{10}(B) = -0.8 \), \( \log_{10}(N_0) = -5.4 \), \( \log_{10}(U_B) = -4.4 \), \( \log_{10}(U_B/U_{\text{re}}) = 3.1 \), and \( \log_{10}(S_{\text{SSC}}(100 \text{ keV})) = 2.8 \). These results are consistent with those given Table 5, i.e. the frequency stratification does not seem to add significantly more error into the derived physical parameters than is already contained in them due to measurement uncertainties.

Since the derived physical parameters depend on the line-of-sight depth of the emitting feature, the assumption of component geometry also affects the results. Because we do not know the true geometry, spherical components were assumed in this study. For a non-spherical geometry, the situation needs to be reconsidered, but that is beyond the scope of the current paper.

6.4. Transverse structure in the jet

The kinematic analysis presented in Paper I already showed that the jet has an interesting transverse structure at 1–2 mas from the core. A significant velocity gradient across the jet was
derived from the component proper motions and the variability Doppler factors. The northern component B2 has a bulk Lorentz factor $\Gamma = 6.6 \pm 1.4$ (see footnote 4) while the southern components B3 and B4 have $\Gamma = 17 \pm 7$ and $\Gamma = 18 \pm 8$, respectively. Component B2 also has a direction of motion different from that of components B3 and B4 by $\pm 28^\circ$ (see Fig. 6 in Paper I). Table 5 and Fig. 13 clearly show that there is a significant magnetic field gradient across the jet coincident with the velocity gradient: B2 has two orders of magnitude smaller $B$ than B3 and B4. On the other hand, B2 has larger relativistic electron energy density than B3 or B4, and is clearly far from equipartition state. Small alignment errors at 5 and 8 GHz, mentioned in the context of the SSC emission of B2, cannot remove the gradient in $B$ – although they could make it slightly smaller. Since large, systematic alignment errors are highly unlikely – we would not expect to see such consistent spectra as those shown in Fig. 11 if there were serious alignment problems – we consider the slope in the magnetic field density across the jet to be real.

The difference in the magnetic field density between B2 and the southern components is probably related to the bulk velocity difference and can be interpreted in at least two ways: 1) the VLBI components correspond to shocks, and the larger $\Gamma$ of B3 and B4 means larger compression ratio and therefore larger enhancement of the magnetic field density for these components\(^7\), or 2) we are seeing a spine/sheath structure (with B3 and B4 corresponding to a fast spine and B2 to a slower layer), where the magnetic field density decreases from the high jet-axis value to the lower value in the sheath. In the latter interpretation it is not, however, clear why we would see emission from the layer only on the northern side of the spine. The velocity difference could also arise e.g. in a scenario where B2 moves on the edge of the jet and is slowed down as it entrains matter of the ambient medium. The analysis of the polarisation data should show if the (Faraday rotation corrected) magnetic field orientation changes between B2 and B3, and if B2 indeed can be attributed to a component in a shearing layer between the jet and the ambient medium (Savolainen et al. in preparation).

7. Conclusions

As is evident from Fig. 11, simultaneous multi-frequency VLBA observations can provide good quality spectra of the compact extragalactic jets. The model-transfer method, which uses a priori information of the source structure from the higher frequencies, is proved to be effective and reliable.

The presented component spectra demonstrate how the flat radio spectrum in 3C 273 between 1 and 100 GHz, as observed by single-dish radio telescopes, is composed of a number of synchrotron emitting features, each becoming self-absorbed at progressively lower frequency as they move out along the jet. A simple power-law fit to the peak frequencies gives $\nu_{\text{m}} \propto r^{-0.6 \pm 0.1}$. The jet cannot stay self-similar much beyond the core observed at mm-wavelengths. The core has a turnover frequency of 50 GHz, while the total spectrum is flat up to 60–100 GHz. Hence, we are within about a factor of two from seeing the actual core site, which is not affected by opacity shift – or we are already seeing it.

Four out of sixteen components show frequency-dependent angular sizes with $a \propto \nu^{-0.2}$ above the turnover frequency. This is most likely due to particles suffering radiative losses when they travel away from an excitation front such as a shock.

Using the obtained spectra together with the measured values of the component size and the Doppler factor, we calculated the magnetic field and electron energy density for the emission features located within 2 mas from the core. Monte Carlo approach was used in determining the errors of these parameters. Although the errors are large, ranging from a factor of two to four orders of magnitude, some interesting analysis of the results can be made. The magnetic field density has rather modest uncertainties as can be seen from Fig. 13. The core has a magnetic field density of an order of 1 G, a value reported also by other studies. In general, the magnetic field seems to decrease roughly as $B \propto r^{-1}$. At $\sim 1.5$ mas from the core, there is a large variation in the magnetic field density between the southern components B3/B4 and the northern component B2. The strong gradient in the magnetic field density is coincident with the transverse velocity structure found in Paper I. The larger bulk velocity corresponds to larger $B$ and vice versa. This can be understood if emission features are shocks and the larger $\Gamma$ causes larger compression ratio which enhances the magnetic field more effectively. Another possible explanation is that we are seeing a spine/sheath structure, where $B$ decreases from the high on-axis value in the fast spine to the lower value in the slow sheath.

Although errors in ratio $U_B/U_{\text{ke}}$ are very large, we find that unless the mm-wave core is unbeamed, its magnetic field energy density clearly dominates over that of relativistic electrons. If the jet beyond the core is assumed to be (proton) rest mass dominated instead of Poynting flux dominated, we get the following limit to the low energy cut-off of the electron distribution: $\gamma_1 \lesssim 10$.

In general, the radio components in the parsec scale jet of 3C 273 were weak SSC emitters at the time of our VLBA observation. Specifically, the core has low SSC flux density unless it is unbeamed. In Fig. 16, the only strong SSC source is the northern component B2, which has very low magnetic field density compared to other features within 2 mas from the core. In principle, B2 is able to produce all the hard X-ray emission observed in the 2003 quiescent state of 3C 273. However, due to the large uncertainties involved in determining $S_{\text{ssc}}(100 \text{ keV})$, one cannot draw strong conclusions about B2’s contribution to the X-ray flux density from a single observation. On the other hand, the recently ejected component C2

\(^7\) Here (and throughout the paper) we assume that for all the components, the pattern speed corresponds to the flow speed. However, according to recent results from 3-D relativistic hydrodynamic simulations, this is probably an oversimplification, and the observed component motions are related to a complicated mixture of bulk and phase motions, viewing angle selection effects, and non-linear interactions between perturbations and the underlying jet and/or the ambient medium (Aloy et al. 2003). Therefore, sophisticated numerical modeling of the jet of 3C 273 may be needed in order to understand the complicated structures that we observe.
– which corresponds to a small flare peaking in January – also could have had significant SSC flux density at the time of the INTEGRAL observations in January when C2 was closer to the core. Unfortunately, there is not enough data to quantify this.

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References

Bridle, A. H., & Greisen, E. W. 1994, AIPS Memo 87, NRAO
Lovell, J., 2000, in Astrophysical Phenomena Revealed by Space VLBI, eds. H. Hirabayashi, P.G. Edwards, & D.W. Murphy, Institute of Space and Astronautical Science, 301
Taylor, G. B., & Myers, S. T. 2000, VLBA Scientific Memo 26