PROCEEDINGS OF SCIENCE

Radio telescope antennas: from single dish to multi-element interferometer

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I give here an overview of the variety of instruments that have been developed in the past seventy-five years in the attempt to understand those phenomena in the Universe, which, invisible to the eye, are, nevertheless, of fundamental importance for Astrophysics and Cosmology. As it always happens, the better the instrumentation has been, the more numerous have been the questions to be answered and the more stringent have been the requests for new instrumentation. This has brought from Radio Telescopes made with one single antenna to space interferometers. This sort of feed-back between science and technology has again reached the point in which existing instruments are no more adequate for the present day Astrophysical problematics and new more complex gigantic designs are required

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1. Introduction

The importance of Radio Astronomy does not need to be illustrated here. I just remember that radio waves can pass through InterStellar medium dusts, which are instead opaque to visible light, that lines emitted in the radio band, first of all HI line at 21 cm, allow to investigate the cold InterStellar media, that many astronomical phenomena, such as, for instance, the Microwave Background Radiation and the Pulsars, would have never been discovered and studied without observations in the radio band.

These are subjects of (under)graduate courses and the most advanced problematics are dealt with also in this book, therefore in my lesson I'll give for granted that at least some of these topics are familiar to you.

This first lesson is mainly aimed at introducing to radio astronomical observational techniques those of you who are not yet familiar with the subject. Therefore I'll be as simple and qualitative as possible, trying, in the meanwhile, to let you appreciate the great importance of excellent instrumentation to make high level *radio* astrophysics. I'll avoid formulas as much as possible and I'll try to explain the concepts instead. In the `ppt` version[1] you will also find many example figures which I'm not going to reproduce here. I'll refer to those slides when appropriate.

Those of you, instead, who already know this topic can just jump to the next more specialized lessons.

I'll begin with a few historical notes and illustrate some of the initial radio telescope designs. Then I'll proceed with discussing the need for high sensitivity and resolution, and present some of the original attempts to reach these goals with huge monolithic instruments. Further, we shall see that the maximum sizes that a single radio telescope can reach (up to ∼one mile) are unsatisfactorily small. The realization of such a limit has brought to the application, in the radio band, of the technique of interferometry, whose fundamentals are those of optical physics. This has been a milestone of modern radio astronomical techniques since, in principle, radio interferometers can be large at will. However, they also have drawbacks, therefore, for the purposes of the modern Astrophysics and Cosmology (see the Scientific Section in this book) more advanced radio telescope configurations are necessary.

2. Historical notes

Radio Astronomy is a quite recent science. Besides the visual band (0.4–0.8 μm) the Earth atmosphere is transparent[2] also to radio waves between $\sim$ 10 m to $\sim$ 1 cm. However, while optical Astronomy is as old as the human specie, since our eyes are the appropriate instruments to visually observe and study the sky (maybe with the help of a telescope), radio astronomy had to wait for the

[1] to be found in http://www.ira.inaf.it/%7Eschool_loc/presentations/PPT/Carla_Fanti.ppt

[2] At longer wavelength it is the ionosphere which stops radio waves. The electron density is high enough that the refraction index becomes imaginary and radio waves are reflected back to the Universe. From the ground, this ionospheric properties have been used in the past for long wave radio broadcasting. At short λ instead the radio waves are stopped by air molecules which have proper frequencies equal to that of the radiation and absorb it. To observe at very small wavelengths it is then necessary to go on high mountains where the atmospheric thickness is small.

comprehension of electromagnetism and of its applications to be mature and the appropriate technology developed. But even when these circumstances began to be realized, still nobody thought to search for emission of signals from astronomical objects in the radio band and their discovery was a typical serendipitous one.

Astronomers were aware that, in principle, black (gray) body thermal radiation from stars had to have a tail of emission also at radio wavelengths, but the expected signal was so low to be uninteresting. In fact today we know that the (much stronger) astronomical radio emissions we detect are non-thermal and are produced by phenomena unknown at that time, like super-relativistic electrons spiraling in weak magnetic fields (*Synchrotron* mechanism)[3] or spectral lines produced by *maser* mechanism.

It occurred to the physicist Karl Jansky, in 1932 (e.g. [2]) to make by chance the first radio observations using a rotating antenna array. To Jansky, working for the Bell Telephone Laboratories, had been given the task of finding the origin of some transmission disturbances, but when he finally detected a "... faint steady hiss ..." at $\lambda \approx 14.5$ m, he discovered that it arrived always from the same direction and always at the same time (actually about 4 minutes earlier) every day. He was lucky, but also smart enough to realize that this signal was unequivocally of extraterrestrial origin, since it raised and set as the stars do. In spite of the greatest skepticism, the most plausible origin was located in the Center of our own Galaxy (the Milky Way). On May, 5, 1933 the news was even published on the first page of *The New York Times*.

This extraordinary discovery stimulated the interest of a small number of *amateurs*, chiefly engineers. One of the most active was Grote Reber who built in the back yard of his house an 8-m radio telescope with which he could map the radio emission from the Milky Way, at the two frequencies of 160 and 480 MHz (e.g. [4]). The Galactic Center, CassA and CygA (the two apparently strongest radio sources after the Galactic Center) and several other structures, well known nowadays, are already visible in his maps.

In spite of this enthusiastic activity, it took a while before radio astronomy began to be a proper field of research: it was the technology developed during the last world war which gave the impulse to the new science.

The pictures of a dozen of historical radio telescope can be found on the `ppt` version (slides from #5 to #32)

3. Main goals for Astrophysics at radio wavelengths

After the first pioneering epochs in which the main interest was to figure out how common the phenomenon "radio source" was, it became clear, as radio astronomy progressed, that the Universe appeared full of weak radio sources. Sometimes they are also intrinsically weak, but more often they are intrinsically very powerful (typically up to 10^{45} erg s^{-1}) but seem weak because they are very distant (see "Science Highlights" this book). Going to weak apparent intensity means, in most cases, to go farther away and hence back in time. Therefore one of the goals is

1) high sensitivity.

[3] This radiation was named after its discovery, in the optical band, in a General Electric synchrotron accelerator in 1946.

in order to observe weaker and weaker objects. In general this implies the use of large collecting areas (but see Sect. 4.3) and low-noise wide band receivers (Sect. 4.2).

It has also been discovered that radio sources have structures more and more complex and details which become more and more numerous and prominent as the instrument resolving power (Sect. 4.1.1) gets better. The study of these details is very important to understand the radio source underlying physics. The other main goal is then

2) high resolution.

One has to consider, however, that radio wavelengths are up to millions of times longer than optical wavelengths, therefore, in order to achieve resolutions ($\propto \lambda/D$, Sect. 4.1.1) comparable to that of optical telescopes one would need radio telescopes with sizes (D) up to millions of times larger than optical telescopes. At a first sight this seems impossible. We shall see in Sect. 5 how this is obtained.

As radio astronomy has progressed it has become evident that in many cases (see "Science Highlights" sections in this book) one needs to image large regions of sky. This is more efficiently obtained with a wide radio telescope *primary beam* or *field of view* – FOV ($\propto \lambda/D_{\mathrm{ant}}$) [4]. This would require small antennas, which is in conflict with the requirements from point *1)*. In addition, in the case of Interferometers (Sect. 5.1) receiver band has to be narrow to avoid image distortions far from the field center. This also would be in conflict with the request of point *1)*, unless receiver band is made of many narrow frequency strips. Therefore one needs

3) large fields of view – wide field mapping capabilities

In the lesson by Muxlow (this book) you shall learn how and how much this can be pushed.

Synchrotron radio emission is mostly wide-band, i.e. continuous, and the observed source properties depend on wavelengths. This, along with source strength and structure, sheds light on the underlying source physics. Therefore a

4) wide frequency coverage

is very important.

Radio lines, like HI (21 cm), OH (18 cm), H_2O (1.3 cm), CO (2.6 mm) and several others are emitted by the cold gas in galaxies. The study of this gas phase is only possible (or better done) in the radio band and allows to understand the physics or the dynamics of these objects. Hence

5) high frequency resolution

is also aimed to (Morganti, this book).

The physics of radio sources may be better understood if one takes advantage of one major property of synchrotron emission: polarization. The possibility to measure

6) polarization

allows to evaluate the distribution and geometry of magnetic fields present in radio sources, of the external magnetic field and the density of the external thermal electrons, hence the properties of *InterStellar* and *InterGalactic* medium.

[4] the suffix "ant" specifies that for interferometers (Sect.5.1) the FOV depends on the size of the largest antenna.

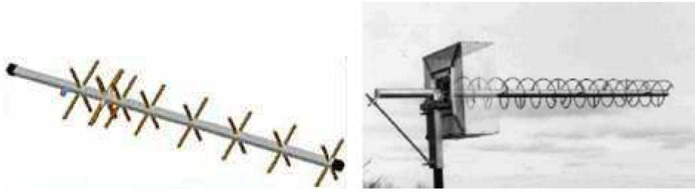


Figure 1: Examples of a Yagi and of a Helical antenna. Note in the first one the pairs of orthogonal dipoles to measure the two components of the polarized radiation.

These are the main goals for doing proper astrophysics and cosmology at radio wavelengths and radio telescopes must be built in a way which satisfies these requests as much as possible with respect to the astrophysical problems of interest.

I'll limit myself to discuss the first two points and mention the third one since the others are approached by other lessons.

4. How is a Radio Telescope made

The Radio Telescope is the instrument which collects and detects radio waves and converts them into a measurable quantity, usually an electric tension.

At long wavelengths (> 1 m), usually "wire antennas" are appropriate for this. They are simple to build, do not require a great structural accuracy and have an effective area (Sect. 4.1.2) roughly proportional to λ^2. We may have one *dipole* or an array of dipoles (Fig. 7), each of which is just a metal bar (usually one half of the radiation wavelength); a *Yagi antenna* (Fig. 1, or arrays of, slide #18) which is a combination of dipoles of different length or an *Helix* (Fig.1) or arrays of.

At shorter wavelengths these models are no more appropriate and radio telescopes have been modeled, originally, after optical telescopes, although the used technology is very different (Fig. 2). The roles of collection and detection are split. There is a – mirror – (in radio language *antenna*) which collects the astronomical signal and reflects and focus it to a device (or ensemble of devices) – receiver – (or *feed* [5]), located at the mirror focus[6], which detects the signal and converts it into an electric tension. The feed can be as simple as a "half-wave" dipole but usually more complex designs are used (see e.g. Fig. 2 right).

Often, a secondary properly shaped mirror is put in between the reflecting surface and the geometric (*primary*) focus, in order to reflect the radiation back towards the mirror. In this case, the feeds to be used at different frequencies are put in a "vertex room" i.e. a little cabin behind the mirror, which has a much easier and comfortable access when some maintenance or frequency change is necessary. This has also the advantage of increasing the radio telescope focal length (Sect. 4.1.3)

[5]Note that some terms, like this one, seem inappropriate for the function they describe, but this is because they derive from the use of antennas in transmission.

[6]If the mirror is a rotation paraboloid the focus is a point, in case of parabolic cylinders (see Sect. 4.3.1) each transverse section of the cylinder is a parabola, with its own focus and therefore, for the whole antenna, we speak of *focal line*, Fig. 8.

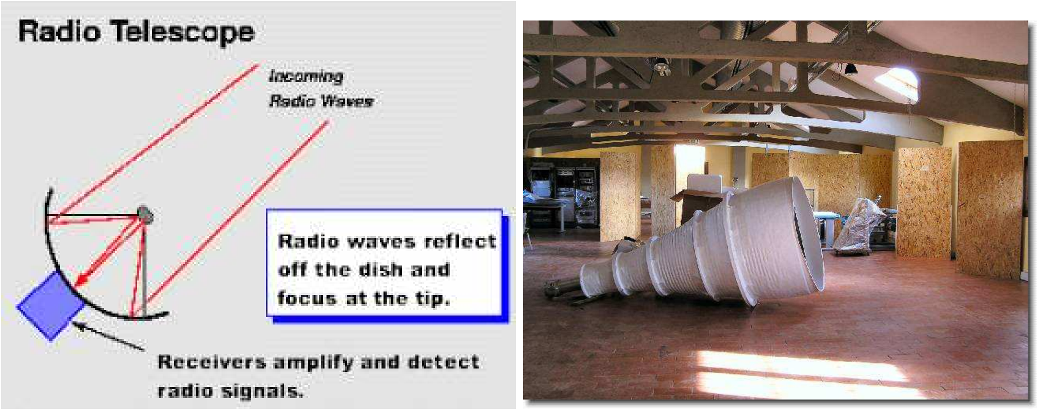


Figure 2: *left)* Scheme of a Radio Telescope (for simplicity a rotational paraboloid): the reflecting surface is the *mirror*. In this figure the reflected radiation is shown to be intercepted by the secondary mirror and reflected back. The *feed*, therefore, is not put at the geometric focus of the paraboloid but at or behind the mirror. *right)* "Horn" feed.

In the years, however, it became clear that such a simple design, usually referred to as *single dish*, had severe limits for many purposes and more complex configurations, involving several antennas, were required. We usually refer to the latter as *multi-element* radio telescopes.

4.1 The antenna

The *antenna* is one of the building blocks of any radio telescope. It may be defined as "the region of transition between a free-space wave and a guided wave" ([3]).

The term "antenna" refers to a variety of designs. Antenna characterizing parameters are its *power pattern* or *beam* (Sect. 4.1.1) and its *aperture efficiency* (Sect. 4.1.2). Other non negligible characteristics are illustrated in Sect. 4.1.3.

Although many different antenna designs have been and are being used, in the following subsections, for simplicity, easiness of understanding and of comparison with optical telescopes, I'll refer mostly just to paraboloids.

4.1.1 Power Pattern – Beam

This is the response of a radio telescope to a point source. It is what in Optics is called "diffraction pattern" from an aperture and optical astronomers call "Point Spread Function" or PSF. In the case of an *ideal paraboloid* the beam is just the Airy pattern, with a central Airy disk containing 83% of the radiation surrounded by fainter and fainter concentric rings. For other radio telescopes (e.g. interferometers, Sect. 5.1) the beam is more complex.

Usually a longitudinal section of the beam is represented, as in Fig. 3 *right*, where the received power is plotted, in polar coordinates, as a function of the angle of arrival of the signal, θ. In this plot we distinguish a *main lobe* (the intensity of the central Airy disk) and *secondary lobes* or *sidelobes* (the intensity of the Airy rings). The first sidelobe, in theory, is usually a few percent of the peak, but aperture obstructions (see Sect. 4.1.2)) or other surface inaccuracies make it higher. This deteriorates the image quality and one tries to keep sidelobes as low as possible. Using a proper feed design one can do in such a way that the sidelobes are reduced ("taper"). This however is obtained at the expenses of also reducing the effective area (Sect. 4.1.2). The beam of an interferometer instead, has high sidelobes (up ~15%) but some numerical corrections can be applied in the imaging step (see Dallacasa, this book).

The instrument resolution, defined as the "capability of separating two point sources", is usually measured by the width of the main lobe at half power (*Half Power Beam Width*, HPBW). It is well known from Optics that the resolution (HPBW) of an instrument of diameter D depends on the ratio λ/D and not on λ or D individually.

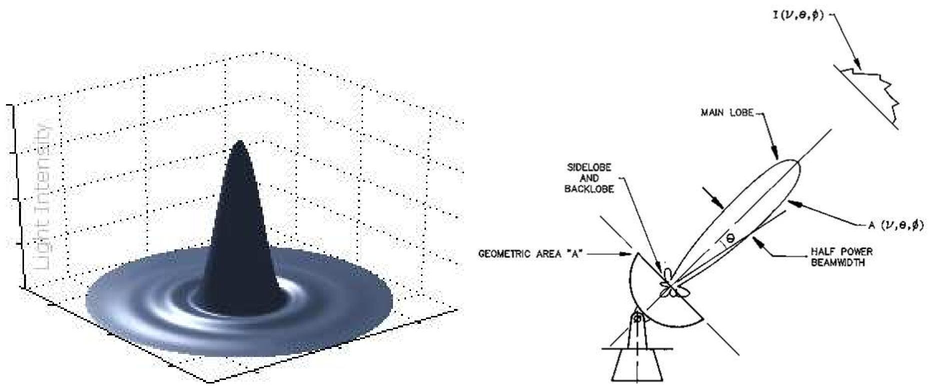


Figure 3: 3D representation of an Airy pattern (*left*) and longitudinal section of the Beam in polar coordinates.

Note that while for ground based optical telescopes the resolution is set by the *seeing*, usually quite worse than the diffraction PSF width, for a radio telescope the resolution is given by its (diffraction) power pattern main lobe. We say that radio telescopes are *diffraction limited*.

A "narrow" beam not only allows studying fine details, but also to have an accurate determination of the radio source position. This is essential, for instance, to find unambiguously the optical counterpart of radio sources, which only can tell us what kind of object we are observing at radio wavelengths (a galaxy, a quasar, a supernova, a star, etc). The positional accuracy is usually taken as 10% of HPBW, and mechanical or software steering of the antenna should not introduce any uncertainty larger than this.

4.1.2 Aperture efficiency

We distinguish between the *geometric area*, A_g, which depends only on the linear sizes of the antenna ($\pi D^2/4$ for a paraboloid of diameter D) and the *effective area*, A_e, which represents the area which actually collects and reflects the radiation. The ratio $\eta = A_e/A_g \leq 1$ is called antenna (or aperture) efficiency. If we indicate with F a radio source *flux density*[7], i.e. the power received on Earth by an *ideal* antenna in 1 m^2 per Hz, then the total power collected by a *real* antenna will be $P = FA_e\Delta\nu = F\eta A_g\Delta\nu$, where $\Delta\nu$ is the receiver bandwidth. Obviously the larger η, A_g and $\Delta\nu$ the better is an antenna performing, and since mechanical problems and costs depend on the antenna size (to the power of 2 or 3) one has to do all efforts to collect as much as possible of the incoming radiation before considering increasing D. Several old radio telescopes, (e.g. Lovell at Jodrell Bank, slide #24, or Parkes, slide #31) have been resurfaced for the purpose of increasing

[7]The measure unit is *Jansky (Jy)*; 1 Jy = 10^{-26} watt m^{-2} Hz^{-1}.

η. In most cases $0.5 < \eta < 0.7$ at best but, since it is wavelength dependent (see below), in can be even less than 0.2 at *mm* wavelengths.

The total surface efficiency is the product of many factors (the major ones listed below), i.e.

$$\eta = \eta_{\rm su}\ \eta_{\rm bl} \eta_{\rm sp}\ \eta_{....}$$

where "su" is for surface, "bl" for blockage, "sp" for spillover.

• – `surface quality`

i) deviations from the theoretical (parabolic) shape,

ii) surface distortions/corrugations,

iii) holes,

iv) . . .

Apart from inaccuracies in the surface construction, gravitational and wind forces, and temperature changes play a major role for point *i)*: when the geometrical shape is modified also the focus in general changes position. For point *ii)* one has to consider that a large paraboloid is not one single piece, but it may be made with wire mesh or be the assembly of numerous shaped panels, which might not be properly positioned. Effects *i)* and *ii)* cause an imperfect convergence of the radiation at the focus, and the feed collects less power than the one reflected by the mirror. i.e. the antenna behaves as if it were smaller. Some radio telescopes are designed in such a way that a paraboloid is changed into another paraboloid and the feed is allowed to move to follow the focus positional changes. In the last fifteen years active/adaptive surfaces have been realized (e.g. GBT, USA, Fig. 5 or the Noto Radio Telescope, Italy) to re-position, more or less in real time, the panels.

The surface quality limits the range of usable wavelengths. If we express the surface quality as standard deviation (σ) from the theoretical shape, a good rule of thumb is that the minimum usable wavelength is $\lambda \sim 20\sigma$: to be able to work at λ =1 cm the surface has to be more accurate than 0.5 mm. On a 100 m telescope this is not too easy to achieve. In the Northern Cross, instead, since it was to be used at 408 MHz ($\sim$ 75 cm) the inaccuracies of its surface were still acceptable if $\lesssim$ 4 cm. Hence a solid surface mirror was unnecessary and wires 4 cm apart could have been used, making the mirror quite lighter. Actually the surface is made by thin steel wires 2 cm apart for redundancy (see Fig. 5 right, where the ice highlights the wires). Looking at the telescope pictures you will also notice that e.g. Effelsberg or Parkes instruments (Fig. 4) have a dish whose interior looks "continuous" while the external ring is made by a wire mesh: this means that at long wavelengths the whole dish is usable while at short λ only the inner part can be used. This technical solution allows a good compromise between resolution, sensitivity and costs.

Finally holes, *iii)*, are usually the spaces between panels, necessary to let water or snow drain. They reduce the geometric area if they do not respect the $\lambda \gtrsim 20\sigma$ condition.

• – `blockage`: the feed itself and the tripod or quadrupod which hold it up (Fig. 4) not only create a "shadow" on the mirror, i.e. regions which cannot be reached by the incoming radiation, but also deviate part of it because of diffraction phenomena. Therefore, from the point of view of the collected power, the mirror performs like a smaller *ideal* mirror. In some telescope designs this problem has been reduced by building only a section of paraboloid (or of parabolic cylinder, Sect.

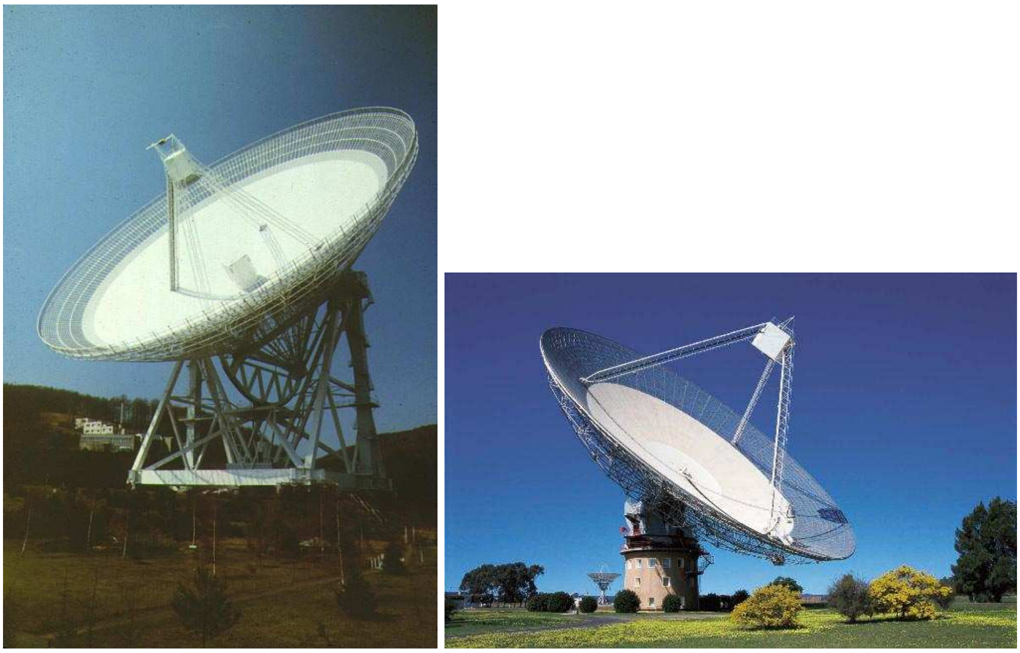


Figure 4: Effelsberg (*left*) and Parkes Radio Telescopes.

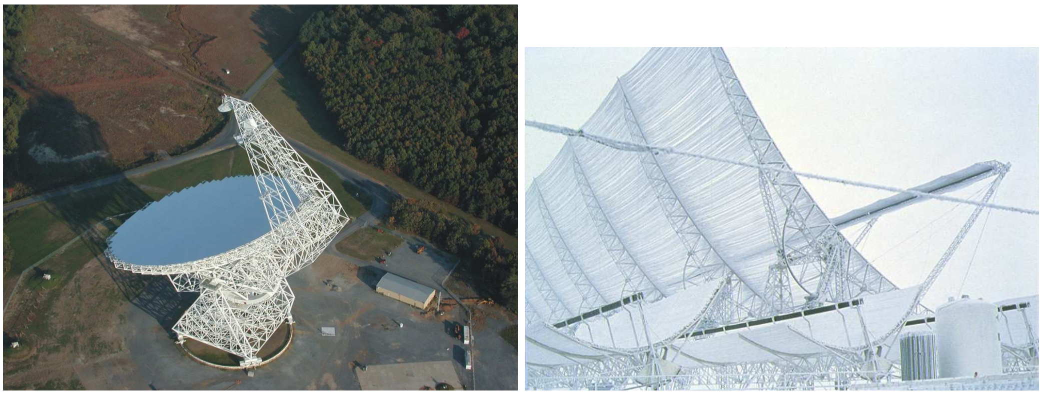


Figure 5: The GBT (Green Bank Telescope) and Northern Cross iced mirror: note the position of the focus in the first radio telescope and of the focal line in the second one.

4.3.1), in such a way that the focus (or focal line) is projected out of the mirror. This of course implies an unbalanced instrument with all the related mechanical problems (see Fig.5).

• – `spillover`[8]: the feed at the focus (Fig. 2) is itself an antenna with its own power pattern (Sect. 4.1.1). Ideally this power pattern should be a "box" of a dimension such that the entire Airy patter is covered, in such a way that the whole focused radiation is collected. Actually, at difference of an optical telescope, the feed does not collect the power of the reflected radiation, but the electric field, whose Airy pattern has positive and negative rings. Therefore, even if the feed beam is wide enough, the gathered electric field is roughly that of the main peak, since positive and negative

[8] see note on page 5.

rings largely cancel out, leaving out a bit less than 16% of the collected energy. Also in this case the mirror performs like a smaller *ideal* mirror since not all the radiation reaching the antenna is delivered to the receiver. One should be able to add “in phase” positive and negative rings. This is more or less obtained (I don’t go into details) with “Horn” feeds (Fig. 2 right).

Another reason for having a “good” feed beam is that if some of the feed sidelobes point out of the mirror they may receive spurious signals from the ground, from the sky etc. increasing the noise (Sect. 4.2).

• – `minor effects` not discussed here.

4.1.3 Other Antenna Properties

The paraboloid is, in some respects, an optimal antenna shape, since it has a single focal point requiring a single feed, has a circularly symmetric beam, it can be fully steered etc, but it is mechanically difficult to realize and to steer, therefore it is very expensive. In principle, the easiest way to point and follow through the sky a radio source, from rise to set, is to have an *equatorial* mount (Fig. 6 left), in which one movement axis (the polar axis) is parallel to the Earth axis, to be able to follow the sky in right ascension, and the other, is perpendicular to it, to point the antenna in declination. This mount allows to point the antenna anywhere, and the steering is simply mechanical. For a big dish, however, the supporting structure is large, heavy and expensive.

An alternative antenna mount is the one called *alt-azimuth* (Fig. 6 right) in which the *azimuth axis* is perpendicular and the *elevation axis* parallel to the ground. In this case much easier and cheaper solutions are possible. With this mounting the gravitational and wind effects (Sect. 4.1.2) are much reduced. In order to accurately steer the antenna and follow a region of sky, however, a computer is required. This explains why until early sixties there have been very few fully steerable paraboloids.

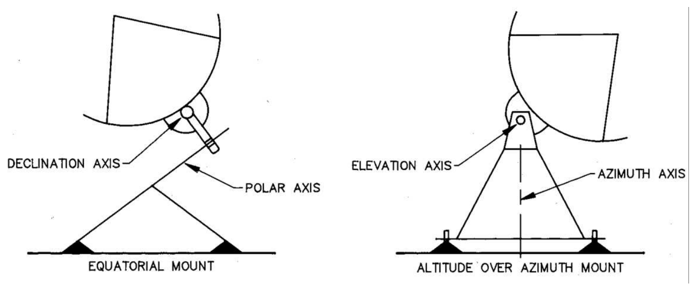


Figure 6: Types of antenna mountings.

A major limit with single parabolic dishes is the modest achievable resolution (Sect. 4.1.1). For instance, the Effelsberg radio telescope (Fig. 4), 100 m in diameter, at $\lambda = 20$ cm has a resolution of ~ 7 arcmin, and we know that details in the radio source structures are much smaller, even much less than 1 arcsec.

Another problem is due to the fact that radio telescope antennas of any shape, at difference of optical telescopes, cannot have a long focal length (F/D): for the Effelsberg telescope, for instance, a focal length of 10 would mean to have the focus, where the receiver is put, 1 km on top of the reflecting surface ... Radio telescopes usually have $F/D \sim 0.4$. This implies that the FOV free from coma is very small ($\approx$ two beam sizes for $F/D \sim 0.4$). The focal length can be increased by using a secondary mirror (Fig. 2 left), but even in this case at the focus of a radio telescope one can observe at most a few independent points or resolution elements ("pixels"), one HPBW apart, at difference of optical telescopes in which one can put a CCD of millions of pixels. This means that to image with a single dish an extended radio source one has to point the antenna at several different sky positions. Nowadays multi-beam capabilities are being developed, but still the imaging situation, with single dishes, is not comparable with that of optical instruments.

4.2 Receiver Noise

Here I'll mention receivers only for the points which are relevant to the observational aspects. Detailed lessons on this subject are to be found elsewhere in this book.

It is well known that any electronic device, working at a temperature different from zero Kelvin, produces some power. This power has a random behavior, characterized by its r.m.s. and is referred to as "noise power". Hence also a receiver produces a "noise power", characterized by a standard deviation $\sigma_{\rm R}$ (or "noise", N) dependent on its own noise power $P_{\rm R}$. The power provided by the astronomical signal ($P \propto F A_{\rm e} \Delta\nu$, Sect. 4.1.2) has itself a stochastic nature and fluctuates at random about its average value $<P>$, with standard deviation $\sigma_{\rm P}$. Unless we are in presence of very strong radio sources ($F >$ a few hundreds Jy) $\sigma_{\rm P} \ll \sigma_{\rm R}$.

For a given instrument both $\sigma_{\rm R}$ and $\sigma_{\rm P}$ decrease with $\sqrt{\tau\Delta\nu}$, where τ is the observing time (Comoretto, this book).

We say we detect a signal (S) when $<P>$ largely overcomes the noise (N), e.g. if $<P> \gtrsim 3-5\ \sigma_{\rm R}$. In this case we say we have a high signal-to-noise ratio (S/N). In terms of radio telescope properties this ratio is expressed as:

$$\frac{S}{N} \propto \left(\frac{F}{P_{\rm R}}\right) A_{\rm e} \sqrt{\tau\Delta\nu}$$

To reach a high S/N one needs:

1) small $P_{\rm R}$ (or $\sigma_{\rm R}$) i.e. "low-noise" receivers

2) large $\Delta\nu$. Actually it cannot be too large in order not to have too large flux density variations within the band due to the source spectrum. For instance for synchrotron emission $F \propto \nu^{-\alpha}$, with $\alpha \sim 0.7$, hence $\Delta\nu/<\nu>= 0.1$ implies a $\sim$10% flux density variation between the two band extremes.

3) large $A_{\rm e}$ (i.e. large $A_{\rm g}$ and η close to unity)

4) long τ .

4.3 How large can a single antenna be?

We have seen in Sect. 4.2 that the possibility of detecting weak signals (point *1)* in Sect. 3) requires low-noise and large bandwidth receivers plus *large collecting areas*. Condition 4) in Sect.4.2 implies the capability to continuously observe a radio source as long as necessary, i.e. it requires *steerable antennas*.

The need to image radio sources with great detail (point *2)* in Sect. 3) also requires large size antennas, since HPBW$\propto \lambda/D$. Actually one might think to solve this last problem by using only small λ, but the need to have observations at as many wavelengths as possible prevents the latter from being a general approach. Moreover receivers are easier to build and work better at longer wavelengths.

How large can we go to achieve a good resolution with single "dish" instruments?

Suppose we plan a resolution similar to that of most (old, seeing limited) ground based optical telescopes, i.e 1 arcsec= 5 10^{-6} rad. The antenna diameter should be $D = 2\ 10^5\lambda$. This means:

D =0.2 km at $\lambda = 0.1$ cm

$D = 2$ km at $\lambda = 1$ cm

D =40 km at $\lambda = 20$ cm

. . . I'll let you comment on this!

But how large can a single antenna be?

4.3.1 Transit instruments

Mechanical problems make it impossible to build fully steerable instruments much larger than $\sim$100 m. Think for instance of a hypothetical paraboloid 500 m in diameter. The movement pivot should be put at more than 250 m from the ground and the focus at $\sim$500 m: unfeasible! No crane exists to allow changing feeds quickly etc, not to speak of the foundations necessary to hold such a heavy instrument.

The difficulty of steering coupled with the need for large resolutions has brought, in the early years, to the construction of "steady" instruments, at most only partially steerable. Most existing radio telescopes of this type can be oriented in declination mechanically (like the E-W arm of the Northern cross) or electrically[9]. The sky transits from East to West above the instrument and for this reason they are called *transit* instruments. Once the declination is set all what radiates within a sky strip large as the HPBW in N-S direction is recorded. Each source is observed for a short time (the width of the E-W beam main lobe in seconds of time). A region of sky can be observed only once (eventually twice for circumpolar radio sources) during 24 hours and to increase the sensitivity (long τ) it is necessary to repeat the observation several times in different days [10].

The old 300 ft Green Bank telescope (unfortunately collapsed on November, 15, 1988, slide #82) was a paraboloid steerable only in declination.

Some other radio telescopes were a dipole cross (like the Mills Cross in Australia, Fig. 7) or parabolic-section cylinders (like the elements of the Northern Cross, Medicina, Italy, Fig. 8). A cylinder can be virtually long at will. Actually, since it has to be *tangent* to the Earth curved surface also for it there are construction limitations. However, cylinders up to $\sim$ 1 mile long have been built. The resolution here is "good" only in the direction of the major axis. To have a good

[9] I don't give any detail here, but just pass the information that it is possible, using the appropriate electrical delays, to have a beam which points at different positions on the sky. It is a technique similar to that of the multi-beam.

[10] In the case of the E-W arm of the Northern Cross three independent beams, one beam size apart, (the additional two pointing East and West of the central beam) are electrically realized. In this way three almost simultaneous observations are obtained to be then averaged thus improving the S/N by a factor $\sqrt{3}$. The further advantage is that interferences are immediately recognized since they occur simultaneously on the three beams, at difference of the radio sources.

resolution in both directions one needs a cross type radio telescope (see Figs. 7 and 8), i.e. the combination (actually the "correlation", Sect. 5.1.2) of two perpendicular cylinders (or dipole arrays).

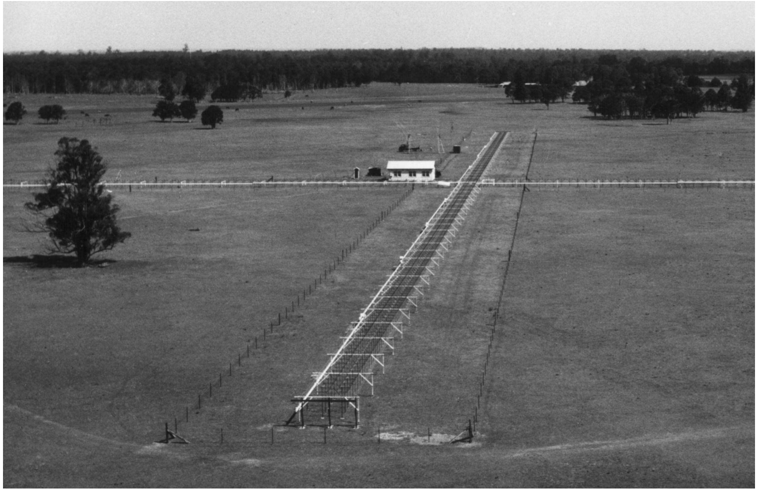


Figure 7: Mills Cross Radio Telescope (electric pointing).

Figure 8: Northern Cross Radio Telescope with the 32 m parabola in the background.

The Arecibo (Puertorico) radio telescope (Fig. 9) is a spherical reflector 300 m in diameter, but it is *not steerable* at all. It is situated in a large natural hole in the ground, a sort of small valley, totally immobile. As part of a sphere, instead of a focus at one point (the instrument would observe at the zenith only for ever !....) it has a focal surface which is a sphere concentric with the mirror and half its radius. The movable part of this instrument is then the feed, which continuously moves on the focal surface to go to pick up the signal from different directions, where it is instantaneously focused during the transit of the source above the radio telescope.

4.4 Need to Balance Sensitivity and Resolution: the Confusion Problem

Even if it were practically possible to build radio telescopes large at will we might, at the end,

Figure 9: Arecibo Radio Telescope – note the movable feed.

be facing the problem of *confusion*. With this term we indicate a situation in which there is more than one radio source within one beam, and, by definition, we are no more able to distinguish if the flux density we measure is due to one or to several objects.

Typically this happens at low flux densities in radio telescopes of large size and therefore of, potentially (it depends also on the receiver...), large sensitivity. Indeed when we increase A_g (and then the sensitivity) we can observe, with the same S/N, weaker radio sources. At the same time, we decrease the telescope beam size. The source number counts, i.e. the number of sources with flux density greater than a given value F, has the form $N(F) \propto F^{-\delta}$ with $\delta > 1$, and the steepness of the source number counts conspires in such a way to increase the number of sources/beam as we go to fainter and fainter flux densities. For instance, if we double A_g, the number of sources per beam will increase by the factor $2^{\delta-1}$. With a very large collecting area we risk to reach the confusion limit at flux densities much higher than the flux density limit set by the receiver noise. In this case we are not able to make full use of the increased *antenna* sensitivity.

A rule of thumb is that one should not have more that one source, on average, every 50-100 beam areas. This guarantees that the probability of having by chance two sources in a beam is negligible (this assumes that radio sources have a random distribution, i.e. don't cluster). The flux density at which this occurs is the lowest one we can observe without running into the confusion problem, no matter if the receiver is so good that we could still have high S/N even at much lower flux density levels. In this situation there is no point in having a larger antenna: we gain nothing!

For instance the E–W arm of the Northern Cross is "confusion limited" already at $\sim$2 Jy, while the receiver noise is so low that it would allow detection of signals of $\sim$ 0.1 Jy !... In the whole Cross, instead, confusion and sensitivity are well matched.

Concluding, confusion is a limit to building huge single instruments. It is therefore necessary to balance sensitivity and resolution.

5. Aperture Synthesis

The need to realize high resolutions coupled with the fact that it is impossible to build individual antennas large as required by the desired HPBW has the natural consequence that *collecting area and resolving power must be disjoined*

This is obtained with sparse arrays, in which many (small) antennas are distributed over a large area and the signals from each of them properly combined.

This is known as *aperture synthesis*, a technique to simulate with small dishes a large aperture. Here the word "synthesis" has the same meaning as in chemistry (synthesis of a compound), electronic music (synthesis of a sound), etc. As far as observations are concerned, we "build" an aperture as large as the ground area over which the (small) antennas are placed.

The basic building block of this technique is the *interferometer* (Sect. 5.1), in which the antennas are used in pairs, and their outputs properly combined.

With N antennas one gets, instantaneously, $N(N-1)/2$, usually different, antenna pairs. The synthesized aperture will have roughly the same resolution of a physical antenna of linear size equal to the maximum distance between any two antennas, $D_{\rm Max}$, as we learned from Optics for the diffraction grating, but a geometric area usually smaller than the sum of all antennas collecting areas[11].

Note that in aperture synthesis instruments the HPBW, $\propto \lambda/D_{\rm Max}$, is much smaller than the FOV, which is $\propto \lambda/D_{\rm ant}$. This allows to image with high resolution a wide sky area, easily placing in it some millions of independent resolution elements ("pixels")

Also cross-type radio telescopes work as aperture synthesis instruments. For instance in the Northern Cross each arm is divided into electrically independent "blocks" (six in E-W and eight in N-S) and each E-W block is combined with each N-S block, thus providing 48 pairs and simulating, as far as resolution is concerned, a square aperture of $\sim 600 \times 600$ m^2.

I let you note that this technique has been "exported" to other frequency domains like Infrared (VLTI, Very Large Telescope Interferometer, Cerro Paranal - Atacama desert, 2635 m above sea level) or mm/submm (ALMA, Atacama Large Mm/submm Array, 5000 m above sea level, under construction) just to mention two of the recentmost projects.

5.1 How does an interferometer work?

5.1.1 Simple Adding Interferometer

A *simple two-element adding* interferometer is nothing else, in principle, than two holes in a mask. In Optics to combine the light from the two holes one converging lens is used and fringes are seen on a screen. Here electric cables convey the signal from the two antennas to an electronic device, but the result is the same.

In Fig. 10 you find the scheme of a simple two-element interferometer whose elements are separated by a distance D (m). The wavefront of the radiation, coming from a direction which forms an angle θ from the interferometer axis, hits first the antenna on the right in the figure (that I indicate with A) then the antenna on the left (B) after a time

$$\tau_{\rm g} = \frac{D \sin\theta(t)}{c}$$

This delay is called *geometrical delay* since it depends only on the interferometer geometry and changes with time while the source moves above the radio telescope.

[11] In a many element interferometer (Sect. 5.3), it can be ∼30 % smaller that the total available collecting surface

We shall see in the following that many interferometer observational parameters depend on a *vector* **b** oriented from A to B with length $|\mathbf{b}| = (D/\lambda)\cos\theta$. This is called *projected baseline* and represents the apparent separation and orientation of the two antennas on the plane of the sky (i.e. as seen from the source). Paradoxically an interferometer could have a zero (projected) baseline even if the two antennas are kilometers apart. Note that the baseline does not depend on D only, but on D/λ. So, for instance, $|\mathbf{b}| = 1000$ rad^{-1} corresponds to $D\cos\theta = 200$ m for $\lambda = 2$ cm or 2 km for $\lambda = 20$ cm. More often it is preferred to use for the baseline, instead of the vector notation, the baseline two components, (u,v), on the plane perpendicular to the source line of sight, with v axis pointing to North and the u axis to East. This plane is universally known as u-v *plane*. Note that here the sky is approximated to a plane. This is acceptable for small FOVs, but when this is no more possible a third baseline component w has to be considered (see footnote on page 18 for the implications).

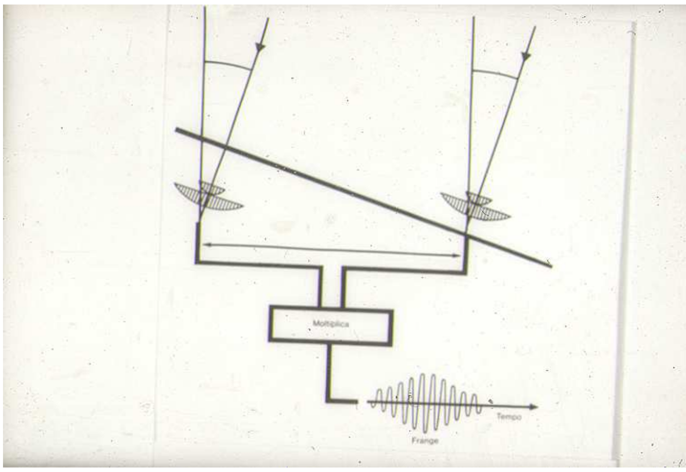


Figure 10: Interferometer scheme. The radiation is shown to arrive from a direction forming an angle θ with the interferometer axis. The wavefront reaches first the antenna on the right (antenna A). The oscillating line represents the electric field product (*fringes*, Eq. 5.2). The fringe amplitude modulation in this scheme is due to the radio source "entering" and "exiting" the individual antenna beam pattern. Usually interferometers track the radio source and this effect disappears.

If we indicate with:

$$\mathscr{E}_{\mathrm{A}}(t) = \mathscr{E}_0 \sin(2\pi\nu t) \;\text{ and }\; \mathscr{E}_{\mathrm{B}}(t) = \mathscr{E}_0 \sin[2\pi\nu(t-\tau_{\mathrm{g}})] \tag{5.1}$$

the electric fields from antennas A and B respectively, the obtained power is given by:

$$(\mathscr{E}_{\mathrm{A}} + \mathscr{E}_{\mathrm{B}})^2 = {\mathscr{E}_{\mathrm{A}}}^2 + {\mathscr{E}_{\mathrm{B}}}^2 + 2\mathscr{E}_{\mathrm{A}}\mathscr{E}_{\mathrm{B}} \tag{5.2}$$

The frequency ν here is not the astronomical frequency but the so called *Intermediate Frequency, IF*. This is a much lower frequency (mostly 30–100 MHz) obtained by mixing the astronomical signal with an artificial signal produced by a *Local Oscillator, L.O.* This has numerous advantages including that the telescope *back-end* does not need to be changed when changing observing frequency (I'm not going to discuss this; see Comoretto this book, instead)

The two terms ${\mathscr{E}_{\rm A}}^2$ and ${\mathscr{E}_{\rm B}}^2$ are "total power" terms, and include all signals detected by the individual antennas, including spurious signals, like thermal emission from the ground or from the sky, human made disturbances, etc. The product instead is the interesting term. In fact, only *coherent signals*, i.e. those which maintain a phase difference constant with time, like two points on the same wavefront, remain after a short integration time. Therefore it is the product of the electric fields the term containing only the astronomical information. Expanding this product and approximating to zero terms of the type $\cos(at+b)$, which vanish in one integration time[12], we obtain:

$$P \propto \mathscr{E}_{\rm A}\mathscr{E}_{\rm B} \approx \mathscr{E}_0^2 \cos(2\pi\tau_{\rm g}) = \mathscr{E}_0^2 \cos[2\pi(D/\lambda)\sin\theta(t)]$$

This is a *quasi-sinusoidal* function of t which represents what in Optics is called *interference fringe* pattern (Fig. 10). In the radio band the same word "fringes" is kept.

It is easy to show that the fringe period (i.e. the distance in time between two contiguous maxima) depends on $(D/\lambda)\cos\theta(t)$, the *projected baseline*.

5.1.2 Correlation Interferometer

Clearly an interferometer does not produce anything similar to a radio source image. However fringe *Amplitude* and *Phase* carry the information on radio source *structure and position.* Therefore an interferometer which provides the *product only* of the two electric fields is what needed.

In the early years, in the adding interferometers phase switching receivers were built in order to cancel out the unwanted squared fields, but since over thirty years electronics has allowed to directly produce the desired multiplication. Considering that the receiver output is an average over the instrument integration time, from Eq. 5.1 we can re-write the field product as:

$$P \propto \int_{\rm int.time} \mathscr{E}_{\rm A}(t)\mathscr{E}_{\rm B}(t-\tau_{\rm g})dt \tag{5.3}$$

Eq. 5.3 is the mathematical definition of "correlation" between $\mathscr{E}_{\rm A}$ and $\mathscr{E}_{\rm B}$ therefore the "multiplying" interferometer is more commonly called "correlation" interferometer. Eq. 5.3 is the fundamental equation of interferometry.

5.2 Visibility Function and its relation with Sky Brightness

By observing a radio source with several interferometers, characterized by different (projected on the sky) baselines, one obtains a set of complex values which represent a discrete complex function called *Visibility Function*. I'll indicate it with $V(\mathtt{u},\mathtt{v}) = A(\mathtt{u},\mathtt{v})e^{i\psi(\mathtt{u},\mathtt{v})}$, where $A(\mathtt{u},\mathtt{v})$ and $\psi(\mathtt{u},\mathtt{v})$ are respectively the fringe amplitude and phase for *each baseline* (or u,v pair).

It can be shown that each interferometer output, $V(\mathtt{u},\mathtt{v})$, is just *one value of the* bidimensional *Fourier Transform* (FT) of the sky brightness, $B(\theta,\phi)$, where θ and ϕ are the sky coordinates:

$$V(\mathtt{u},\mathtt{v}) \propto \int B(\theta,\phi)e^{-2\pi i(\mathtt{u}\theta+\mathtt{v}\phi)}\mathrm{d}\theta\mathrm{d}\phi \tag{5.4}$$

[12] Recall that the signal detection always implies non-zero integration time. Even a fraction of a second is always a very long time with respect to the IF periods (e.g. 100 MHz correspond to 10^{-8} sec) and can be approximate to ∞. This allows one to approximate to zero sin or cos expressions, when appropriate.

Equation 5.4 tells that it does not matter if interferometry does not provide, almost immediately, an image of the sky like single dishes, because we can obtain it with the *discrete* inverse-FT of the visibility function (Eq. 5.5) computed for different baseline values (but see Sect. 6.2).

$$B(\theta_m, \phi_n) \propto \sum_{k,j} V(\mathtt{u}_k, \mathtt{v}_j) e^{2\pi i(\mathtt{u}_k \theta_m + \mathtt{v}_j \phi_n)} \tag{5.5}$$

provided we have enough samples of $V(\mathtt{u},\mathtt{v})$[13].

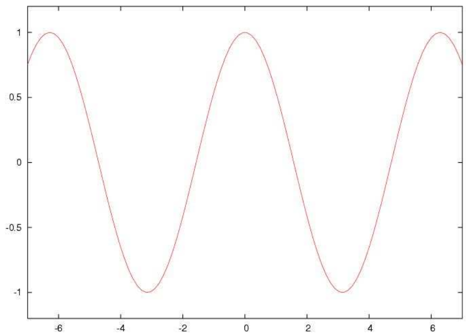
5.2.1 One-dimensional examples

Let's see how we proceed (one dimension only for simplicity).

The brightness distribution of a point source of flux density F at position θ_0 can be described by a δ-function as $B(\theta) = F\delta(\theta-\theta_0)$. Its FT is the continuous complex function $V(\mathtt{u}) = Fe^{-2\pi i \mathtt{u}\theta_0} = F[\cos(2\pi\mathtt{u}\theta_0) + i\sin(2\pi\mathtt{u}\theta_0)]$, a function with $A(\mathtt{u}) = |V(\mathtt{u})| = F$ (constant in u) and phase $\psi(\mathtt{u}) = 2\pi\mathtt{u}\theta_0$, i.e. linear in u with a slope $\propto \theta_0$, the source position. For a slightly extended (Gaussian) source $|V(\mathtt{u})|$ would be a Gaussian, etc.

At this point, to simplify further the discussion, I assume $\theta_0 = 0$, hence $\psi(\mathtt{u}) = 0$, $V(\mathtt{u})$ real and only cos-functions in the summation of Eq. 5.5.

A variable length interferometer will sample one or more discrete values of the visibility functions from which to attempt to reconstruct the source brightness.



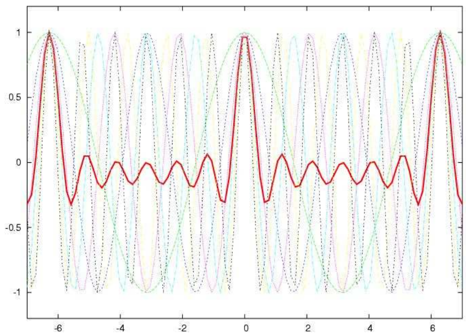


Figure 11: Approach to the reconstruction via FT of a point source at $\theta_0 = 0$ (i.e. $\psi = 0$ and cosines only in the visibility function) with *one* (left) and with *a few* (right) visibility points. The thin lines represent the used visibilities.The abscissa is the sky angular coordinate while the other axis gives $B(\theta)$ (both in arbitrary units).

Clearly with one baseline only we would have in hand just *one* cosine of period 1/u , (Fig. 11 left) of amplitude $A(\mathtt{u}) = F$ but from this value only it would be impossible to understand what type of a source it is. One could guess to be dealing with a point source of flux density F. If the source total flux, F_0, is known from other data, the ratio F/F_0 could give an estimate of the source extension, assumed to be a Gaussian. Instead, with many baselines, we can add together all the cosines, of different periods and equal amplitudes $A(\mathtt{u})$, and approach the real source brightness distribution (Fig. 11 right).

[13]Should component w be required (Sect. 5.1.1) a 3D Fourier Transform would be necessary in place of Eq. 5.5. This would mean an enormous computational problem !... but see Muxlow (this book) for the adopted approaches.

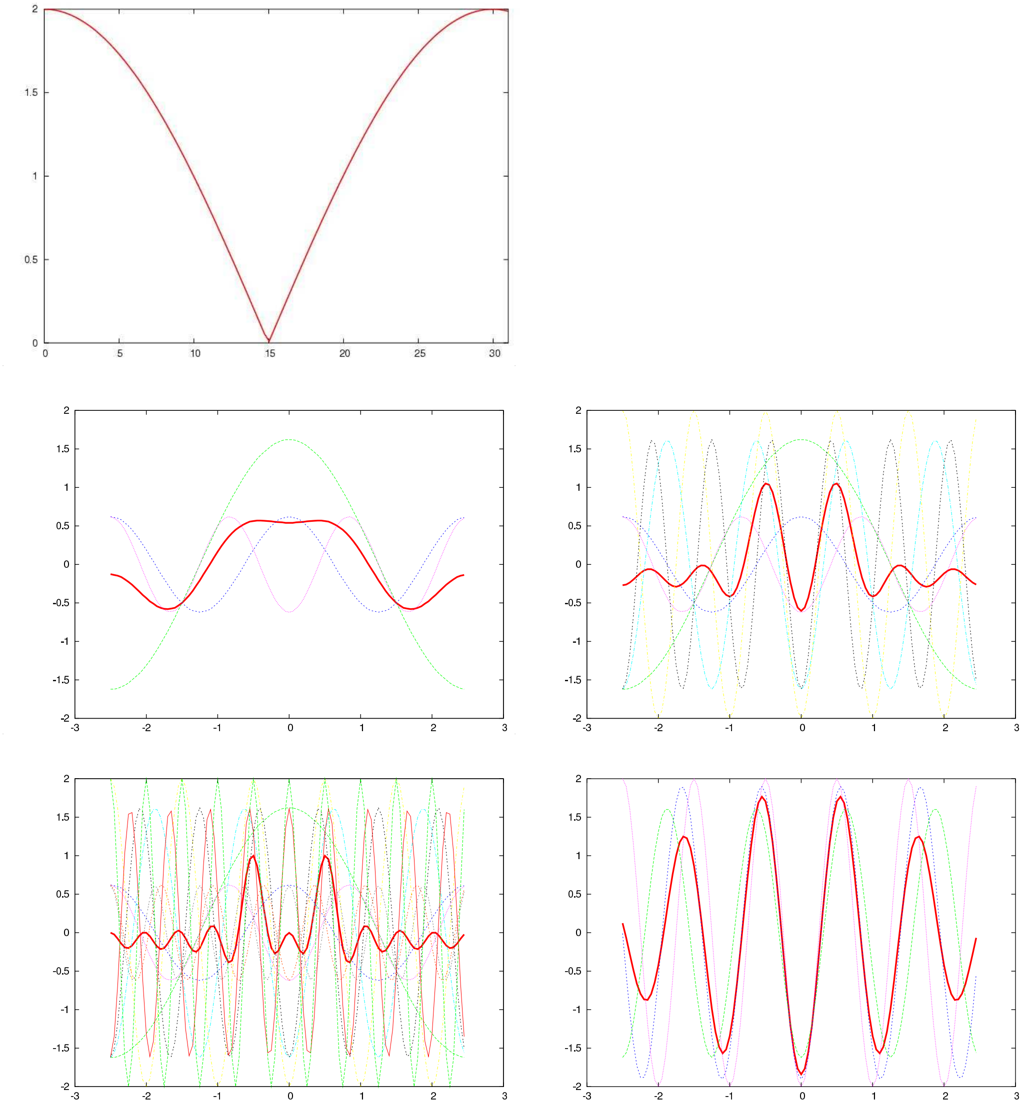


Figure 12: Topmost panel: amplitude of the Visibility function of a symmetric double source with equal components symmetric with respect to $\theta_0 = 0$ and 1 arcmin in separation. The phase (not plotted) is *zero* up to the visibility amplitude minimum, where it *jumps* to 1π. In abscissa the baseline is given, in units of kλ (0 to 30). The following four panels are reconstruction (via FT) obtained using different samplings of the Visibility function. The thin sinusoids represent the sampled Fourier component, the thick line the computed FT. The Visibility samplings are, in the order: (3, 6, 9 kλ, *top left)*, (3, 6, 9, 12, 15, 18 kλ, *top right)*, (3, 6, 9, 12, 15, 18, 21, 24, 27, 30 kλ, *bottom left)*, and (24, 27, 30 kλ, *bottom right)*.The θ coordinate is in arcmin.

Note that the figure is not exactly a δ-function: the central peak does not have zero width and there are small oscillations visible on both of its sides. These are artifacts due to the limited extension of the used baselines. To obtain exactly a δ-function you may think that it is necessary to add measurements up to $|\mathbf{b}| = \infty$! Actually not even this is correct, since the interferometer does not provide all the values of the visibility function, but just samples of it. The fact that discrete values of $V(\mathtt{u})$ are used, creates the secondary maxima, which, if the step $\Delta\mathsf{u}$ is constant, are as high as the main peak and $1/\Delta\mathsf{u}$ radians apart. This is a well known result from the diffraction grating theory.

In Fig. 12 is an example of a symmetric double source: two pointlike equal components in positions $\pm$ 0.5 arcmin with respect to the origin. The figure shows the amplitude of the visibility function (first panel), three brightness distributions obtained using an increasing number of baselines (see caption to the figure) and one with the three longest baselines only. Note that with the shortest baselines only we barely recognize the double structure, while this becomes more and more clear, with the two peaks at $\pm$ 0.5 arcmin from the center, as we add larger $|\mathbf{b}|$ values. The small oscillations are artifacts due to the limited baseline extension ($<$30 kλ). In the last panel, instead, we have no way to tell the source structure.

It is evident from the above discussion and examples that one single baseline, giving one visibility point only (or, if you prefer, one Fourier component only), or long baselines only are not enough to provide information on a source brightness distribution.

Multi-baseline observations are therefore necessary.

On the mathematical point of view (Eq. 5.5) there is no restriction on obtaining several different baselines simultaneously or at different times. The only limitations are connected to the eventual radio source flux density variability, different instrumental or weather conditions from day to day, but mainly to the time consuming observations, in case one obtains only a few baselines at a time.

In addition radio sources are bidimensional brightness structures, therefore bidimensional u,v distributions are also required.

5.3 Multi-element Interferometer

To illustrate which was the situation in the very early times I'll describe the way of operating of the two-element variable spacing Caltech interferometer ($\sim$1960). Two 90 ft antennas could be positioned at fixed stations on two perpendicular railroad tracks oriented in East-West (200, 400, 800 and 1600 ft) and in North-South (200, 400 and 800 ft) in such a way to realize four and three spacings respectively in each direction. For some time, depending on the astronomer requests, the interferometer worked in a given configuration, producing *one baseline* data for each observed source. Then the configuration was changed (baseline length variation or swap from East-West to North-South or vice-versa) and for some more time the interferometer produced data from the *second* baseline, and so forth (see slides from #133 to #141 in the `ppt` version). At the end, up to seven baselines with the corresponding visibility data points were made available to the astronomers, who, more or less by hand (FT was not of great utility because of the small number of visibilities), with the help of source models, could study their source structures.

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I like to recall you, that in spite of the paucity of data and inadequacy of resources, A. Moffet and P. Maltby, in their PhD theses, observed at 31 cm a large sample of $\sim$ 130 radio sources from various low resolution catalogs. Using a model-fitting approach they could prove that the vast majority of powerful radio sources have a double structure. This was the start of real astrophysics.

Nowadays interferometers count many elements. In these interferometers, the electric fields from the N antennas are sent to a multi-entry correlator which multiplies them in pairs, producing $N(N-1)/2$ sets of data. The individual pair outputs, $V(\mathtt{u},\mathtt{v})$, will then be properly combined according to Eq. 5.5 to synthesize an aperture.

Pictures of two famous "patriarch" large interferometers are reported in Fig. 13 and Fig. 14. They are the Westerbork Synthesis Radio Telescope (WSRT, in the Netherlands, $\sim$ 1970, 14 antennas arranged in East-West, 1D) and the Very Large Array (VLA, Socorro, New Mexico, $\sim$ 1980, 27 antennas, in a "Y' shaped configuration, 2D). Do not forget, however, ATCA – Australian Telescope Compact Array (1D), GMRT – Giant Meter Radio Telescope, India (2D), and many others. Some or all the telescopes in these interferometers can be moved on tracks to realize arrays of various compactness.

Figure 13: Aerial view of the Earth rotation synthesis WSRT ($\sim$ 2.8 km). Here four movable antennas are combined with ten fixed antennas and give 40 different baselines. It is also possible to combine all the 14 antennas and obtain 91 baselines with different degrees of redundancy.

6. Earth rotation Aperture Synthesis

The real enormous improvement was the *Earth Rotation Aperture Synthesis* introduced in 1965 by Sir Martin Ryle (for which he also got the Nobel Prize) that he himself, with the Cambridge (GB) group, extensively applied with the three-element One-mile Cambridge interferometer. The term "Earth Rotation Aperture Synthesis" derives from the fact that the Earth, during its diurnal rotation, carries the interferometer around, thus continuously changing the position of the antennas

Figure 14: The Earth rotation synthesis VLA, 27 antennas (351 baselines) distributed over up to 36 km.

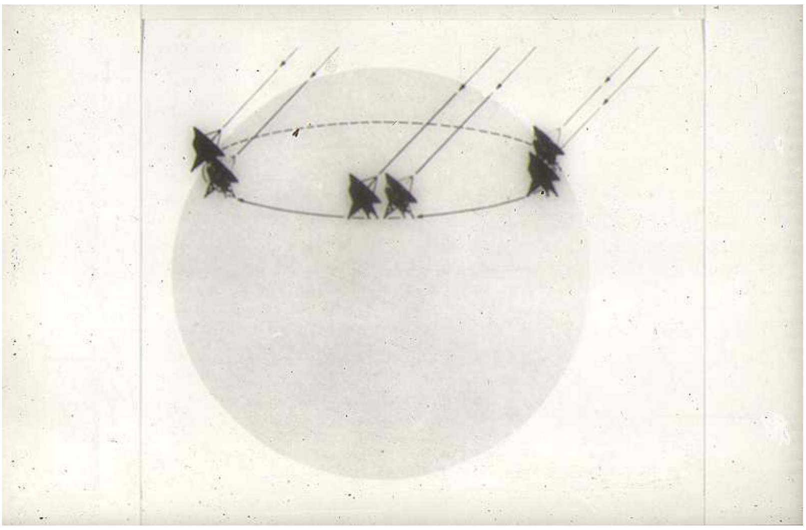

Figure 15: Principle of Earth Rotation Aperture Synthesis.

with respect to the pointed radio source (Fig. 15). So the VLA, for instance, instead of 351 "instantaneous" baselines, can produce many times more, depending on the length of the observation (observing time / integration time, see note on page 17).

During an observation the tip of each baseline **b** describes in the u-v plane an (arc of an) ellipse depending on the observation scheme. Namely the projected baseline length changes with the time and this has the consequence of giving different resolutions in different sky directions (HPBW$\propto 1/|\mathbf{b}|_\lambda$).

One can observe a source from rise to set (*full synthesis*) but more often just for a few time intervals, (*snap-shot mode*), if necessary evenly spaced,[14] are enough to derive a reliable source

[14]Since the VLA has a bidimensional (2D) antenna configuration just one time interval may be appropriate. For E-W

image. Examples of u-v *coverages* obtained with the VLA for a radio source at $\delta = 45^\circ$ are shown in Fig. 16.

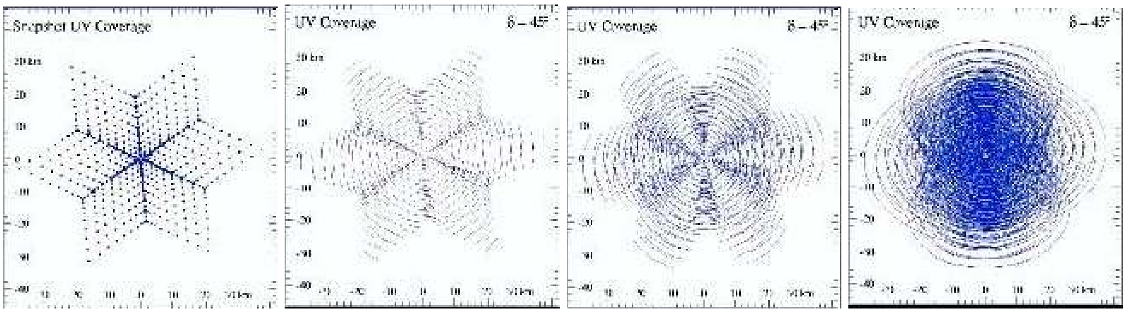

Figure 16: Examples of u-v coverage from the VLA for different observation lengths. From left to right: 10 seconds, 1, 3, 12 hours.

6.1 Need for good u-v coverage

We have seen that a multi-element interferometer is a *Fourier transformer* of the sky brightness distribution and that the inverse-FT of the visibility function provides a first approximation of it. We said that with N antennas one gets, instantaneously, $N(N-1)/2$, usually different, baselines. The point is how to distribute these antennas on the ground in such a way to have a fair distribution of baseline lengths and orientations (given a certain amount of money and human power ...). Naively one could think that the longer the baselines the better. This is only partially true, as proved by the last panel in Fig.12 as compared with the previous one. It is correct to say that the longer the baseline the better is the resolving power, but in order to *image* a radio source with high details both *short* and *long baselines* are usually required. In fact each baseline $|\mathbf{b}|$ radian^{-1} long samples only structures of the order of $1/|\mathbf{b}|$ radians, therefore in order to have information on also the extended structures one needs also small $|\mathbf{b}|$.

When one goes to higher and higher resolution instruments (see Sect. 6.3) for practical and financial problems one has to limit the number of antennas to a few tens at most and these are usually put in such a way to form long baselines, in order to favor high resolutions. The poor sampling of short baselines then explains why we miss more and more extended structures (e.g. Fig. 17) when we image radio sources with increasing resolutions.

Note that one interferometer always lacks baselines shorter than about the size of the largest antenna involved. In fact, the closest one can put two antennas is when they are physically in contact (or begin to shadow one another for certain pointings). Since the baseline is defined by the distance between the two feeds, which, in this last case, is twice the sum of the half sizes of each antenna (or the individual antenna size in the case of identical antennas) this is the minimum baseline one can get. A corollary of this is that the so called *zero baseline* is always lacking. This prevents from measuring with an interferometer the source flux density since, from Eq. 5.4

$$V(0,0) = \int B(\theta,\phi)\mathrm{d}\theta\mathrm{d}\phi$$

instruments, like the WSRT, several snap-shots are required for an acceptable bidimensional baseline distribution.

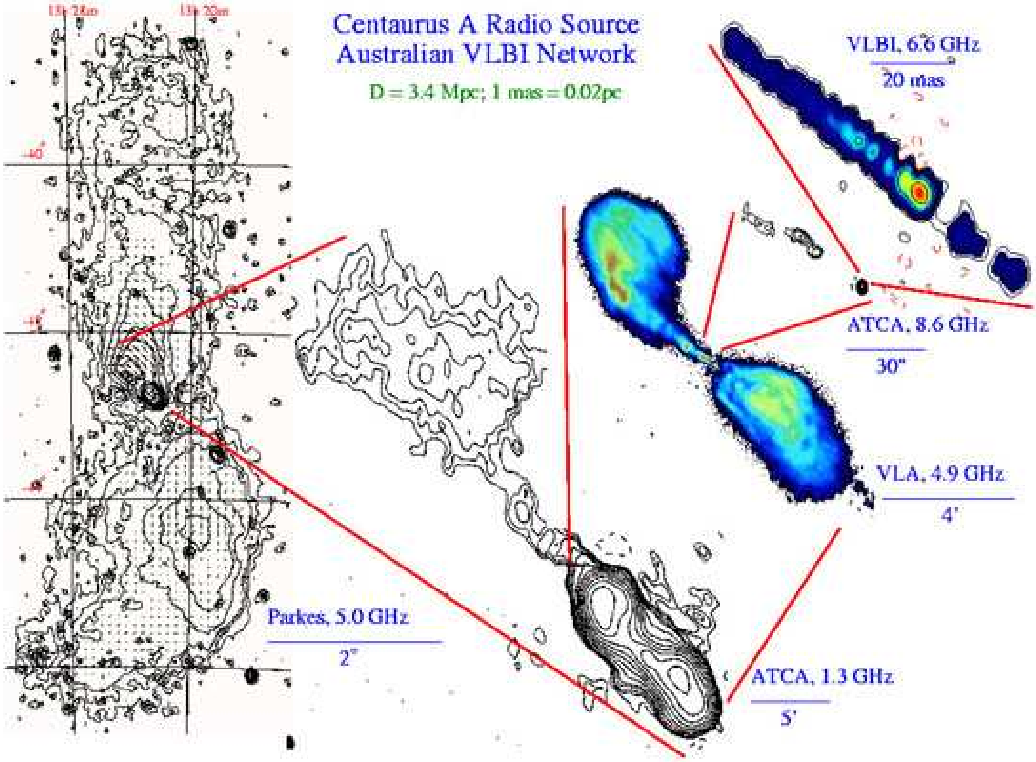

Figure 17: Multi-frequency multi-resolution image of CenA. Frequencies and resolutions (increasing from left to right) are marked on the figure itself. The first image on the left is obtained with a single dish radio telescope (Parkes, Fig. 4) which contains all spacings from zero to $|\mathbf{b}|_{\rm Max}$. The others are obtained with longer and longer (in terms of $\mathbf{b}_\lambda$) interferometers and miss more and more extended features.

is the source total flux density. Image reconstruction algorithms (Sect. 6.2), e.g. CLEAN, however, attempts to recover some of the missing flux density (see Dallacasa, this book)

6.2 Need for Image Restoration

In Fig. 18, I show the synthesized beam and two images as obtained by the inverse-FT of full synthesis (12 hours) observations with the WSRT. The beam shows a narrow high peak at the center (in the image you hardly appreciate it, unfortunately) surrounded by fainter rings, (first ring being $\sim$15 % of the peak) which are the sidelobes, similar to, although higher than, those of a single dish (Sect. 4.1.1). The strong rings, whose amplitude decreases outward, are called *grating lobes* (in the figure are visible the first one and arcs of the second one in the figure corners). These are the major peaks present in the diffraction pattern of an optical grating (see comments to Fig. 11) turned around by the Earth rotation. In radio language we say that they are caused by gaps in the u-v coverage. As you can see in Fig. 18 the images are badly affected by sidelobes and grating lobes which alter the intensity and the shape of radio sources in their proximity and can even hide weak sources. For this reason, in the radio jargon, these are usually referred to as *dirty* beam and *dirty* images. With the VLA, which has a more complex bidimensional antenna distribution, the

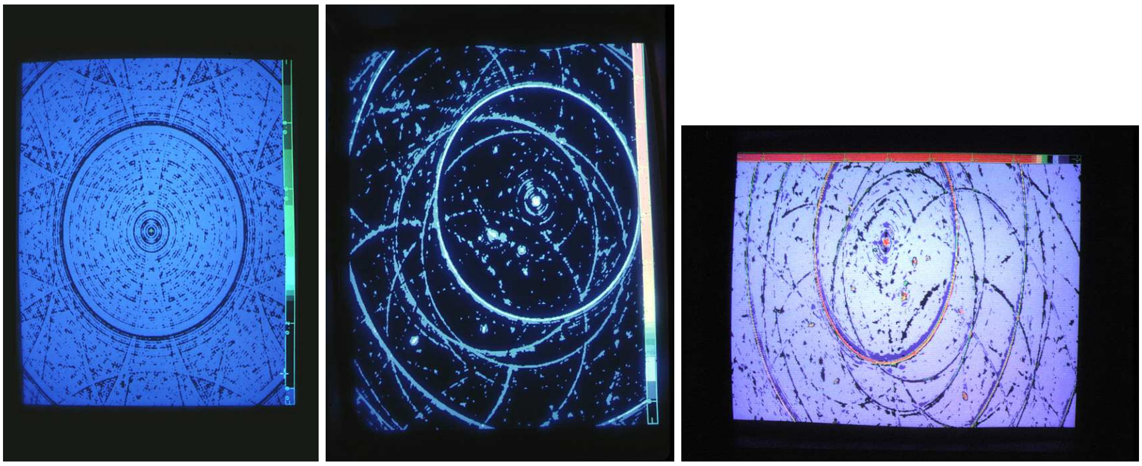

Figure 18: The Dirty beam and two dirty images from 12^{h} WSRT observations.

grating lobes are lower (slide #51 in Dallacasa presentation) but they still deteriorate the image. We have seen in Sect. 5.2.1 that, if Δb_λ radian^{-1} is the average gap length (or the distance in units of λ between two antennas) the distance between two contiguous grating lobes is $\approx 1/\Delta b_\lambda$ radians.

Nowadays that computers are very powerful and relatively cheap it is possible improve images using techniques of *image reconstruction* or *restoration*, i.e. numerical algorithms which allow removal of grating (and side) lobes (e.g. AIPS CLEAN or similar, Dallacasa, this book). This is not exactly the same as not having grating lobes at all, but it is the only solution adopted at present with a good degree of satisfaction and reliability.

6.3 How large can an interferometer be?

In principle there is no limit to the size that an interferometer can have. In practice, however, the technical solutions, correlation and the post-correlation analysis depend very much on the instrument length.

We have (had):

• *Short* – local scale arrays (up to ~ 30 km) like the WSRT (Fig. 13), the VLA (Fig. 14), ATCA, GMRT, ... Here the signals from each antenna pair is brought to the many entry correlator by coaxial cables. The calibration and self-calibration (see Dallacasa, this book) are relatively simple since the atmosphere is reasonably quiet on such scales, at least at $\lambda \gtrsim 6$cm.

• *Long* – regional scale arrays ($\sim 30 - \sim 200$ km), e.g. the MERLIN (Multi Element Radio-Linked Interferometer, six antennas, 15 baselines) in Great Britain (Fig.19). In this case cables are no more appropriate since electric losses would be too high. Radio-Links are used instead. The correlator is located at Jodrell Bank. Also the signal from one single L.O. (Sect. 5.1.1) is sent to each antenna via radio link. The limit of this method is again the distance between antennas, since to deliver the signal from any station to Jodrell Bank and vice-versa several repeaters are necessary for both keeping the radio-signal collimated and to account for the Earth curvature, which has to be approximated by a polygonal light path. At a distance of 100 km from its origin a straight light beam would pass at an altitude of ~ 750 m above the ground. Post correlation analysis begins to be difficult and self-calibration is required because atmosphere may change a lot over such distances.

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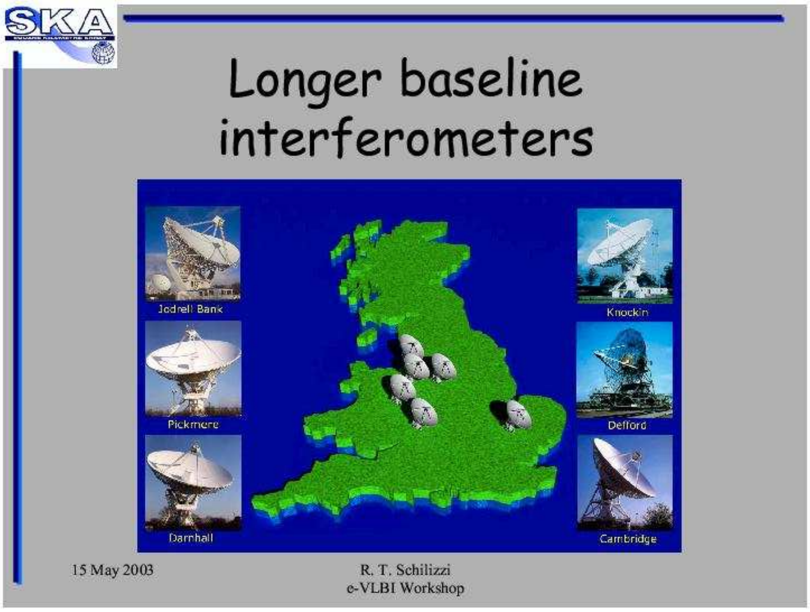

Figure 19: The Multi Element Radio Linked Interferometer (MERLIN).

• *Very Long* – global scale arrays ($>$ 1000 km), i.e. all VLBI (Very Long Baseline Interferometry) arrays. Here antennas are spread over the whole Earth surface. They are either pre-existing antennas properly equipped for doing VLBI observations, or built expressly for this purpose. Radio Links are no more adequate for these distances. Here, at each station, the radio wave electric fields are digitized and "physically" delivered (on a magnetic supports like a tape or a disk) to the site where are to be correlated (e.g. to JIVE, Dwingeloo, The Netherlands; NRAO, Socorro, New Mexico). The correlator in this case is a digital correlator, i.e. a powerful computer which calculates Eq. 5.3. Without entering into details I just mention that the electric fields are to be sampled at least twice in one of their period ($\propto 1/\nu$), so, although the observing frequency (e.g. 5 GHz) is usually down-converted to 2 MHz and the data are recorded with the very minimum of information (one or two bits) still millions of data have to be stored every second. This means that VLBI could not be a routine observing technique until mid '80s when industrial magnetic supports and tape/disk players became available to handle this amount of data. Nowadays VLBI interferometers work routinely.

There presently are three major networks in the world (Fig.20) each including 10-20 antennas: EVN (European VLBI Network) involving stations in Europe, China, South Africa; VLBA (Very Long Baseline Array) in the US (built on purpose); APT (Australia Pacific Telescope), including Australia, Japan, China and South Africa. Occasionally individual telescopes join these networks or part of the networks work together. Since a VLBI array is a huge single instrument, with this technique during an observing run all stations must, by definition, work in synchronism, the rhythm of all operations being marked by the Universal Time (UT) regardless of the local time.

Consider that, to be able to operate, all antennas in a network must receive the same wavefront (although with the appropriate geometrical delay), hence each antenna must be "on source" when

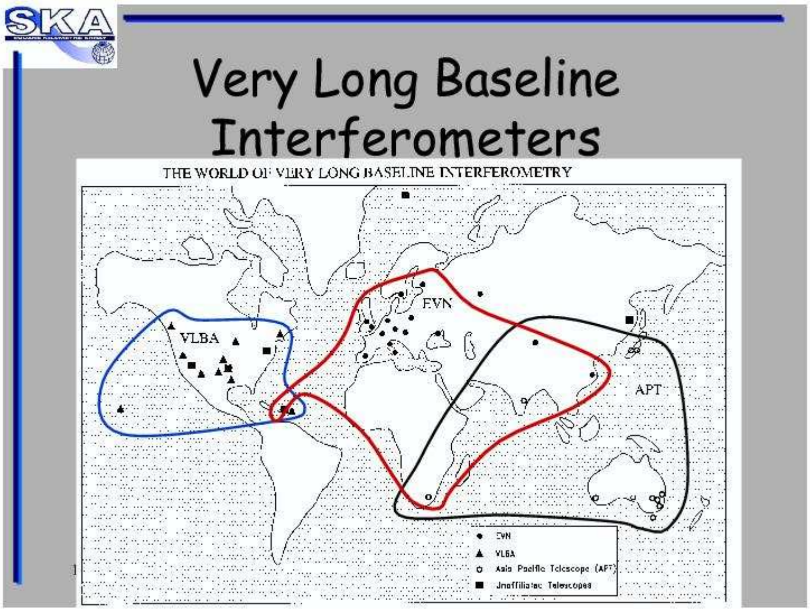

Figure 20: The three major VLBI networks.

also the others are, at least for some time. For the longest baselines this may be happening for short times (e.g. 1 hour only).

Until a couple of years ago tens and tens of tapes had to be brought by truck/plane to the correlator. At present wide band optical fibers allow to immediately transmit the sampled electric fields to the correlator giving the possibility of real time correlation. Not all telescopes involved are equipped at present for this type of operation, but quite soon even Earth-scale interferometers will be again linked instruments, like conventional interferometers, but with order of magnitude of improvements in terms of science and money.

The post-correlation analysis is very complicated, not only because of the atmospheric turbulence, which is not negligible on such large scales (in the radio the so-called isoplanatic patches have sizes of a few tens of km or less, depending on the observing frequency), but also because of Local Oscillators, one at each station. A Local Oscillator always introduces unknown random errors which however cancel out if one single L.O. is used (as in conventional interferometers or MERLIN). With VLBI, instead, these errors can only be removed after many cycles of accurate self-calibration (Dallacasa, this book).

• *space VLBI* – In this case some antennas are on board of satellites orbiting the Earth and operate together with the ground based antennas. The idea of space VLBI goes back to the seventies but the first pioneering experiment, aimed at testing the idea, was made in 1986-88. The 4.9 m antenna on board of the telecommunication satellite TDRSS (Transfer and Data Relay Satellite System) was used at the two frequencies of 2.3 and 15 GHz, for short periods, in combination with the 70 m Tidbinbilla antenna (Australia) and the two Japanese antennas of Usuda (64 m) and Kashima (26 m). The maximum baseline was 2.16 Earth diameters. The experiment was successful and proved that this technique was possible.

The only experiment realized so far for astronomical observations (VSOP, VLBI Space Observatory Program, Fig. 21) had an 8-m deployable wire mesh antenna, working at 18 and 6 cm. Also 1.3 cm had been planned but once in orbit it could not be used properly. VSOP was launched on February, 12, 1997 and operated until 2005 on an orbit of 25 000 km in diameter. Another experiment, VSOP-2 is planned for launch in 2008-2010.

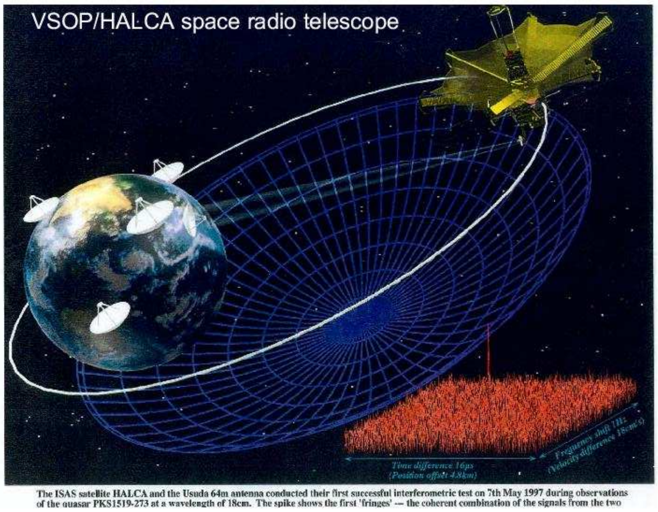

Figure 21: Artist view of VSOP space VLBI: the orbiting antenna is seen to operate together with the ground based antennas thus synthesizing a huge dish.

Space VLBI is the best example of "bad" u-v coverage because of the large gap between ground baselines and orbit-to-Earth baselines. But VSOP was not aimed at producing high quality images (although many excellent results have been obtained) but mainly at searching for high brightness temperature weak radio sources[15] for which extremely high resolutions are required.

7. Summary and conclusions: towards SKA[16]

At the end of this lesson we may go back to the points discussed here and ask ourselves if the observing facilities built in the past years are still satisfactory for the present day astrophysical problematics.

high resolution: present days Multi-element interferometers already seem to solve this problem. Actually we cannot do any better from Earth. What lacks is the simultaneous presence of both long and short baselines to be able to map extended structures with high resolutions. Actually the VLA can do quite well in this respect since its 27 antennas can be distributed over four different configurations (*A* to *D*) each of which is three times more compact than the previous one. The

[15]It can be shown that the minimum Brightness Temperature, $T_{\rm B}$, one can measure with an interferometer of size D is $T_{\rm B} \approx \frac{1}{2k} F\ D^2$,where k is the Boltzmann constant and F a source flux density.

[16]A companion complementary project with characteristics similar to SKA is LOFAR, being expressly built to work at very low frequencies ($<$ 240 MHz)

combination of all four configurations allows an excellent u-v coverage, but, since configuration changes occur every three months, one has to wait for the appropriate epoch to perform the desired observation. One can also combine data from different VLBI networks or from the MERLIN and the VLA, or the MERLIN and a VLBI network, but the procedure is never straightforward and results never optimal.

SKA is planned to have a "dense core" and "dense outstations" in such a way to always make long and short baselines simultaneously present.

What is really required is *high sensitivity*, a factor 50–100 better than what obtained at present. Besides excellent receivers, one needs, therefore, a very large collecting area, which, we have seen, must be realized with a sparse array. In addition the u-v coverage must be such that very low grating lobes are present, to guarantee high image reliability and fidelity.

Therefore very many antennas have to be built and accurately located. Furthermore, in order to realize a wide FOV, which speeds up the imaging of very extended regions of sky, small antennas are a better choice (the so called "large$N$/small$D$" SKA concept)

It is evidently a gigantic project which will involve money, human power and brains from all over the world.

Good luck !!!

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Useful books:
Principles of design of radio telescopes: [1], [3], [5], [6]
Radio Astronomy for beginners: [7]
More specialized Radio Astronomy: [8]

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I wish to thanks all the colleagues whose pictures (from on-line presentations) I've used, with or without their permission

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