APERTURE SYNTHESIS
(Spatial Interferometry)
WITH THE VERY LARGE TELESCOPE

An Interim Report
Presented by the ESO/VLT Working Group on Interferometry
October 1985
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Giving to the European Very Large Telescope (VLT) an interferometric capability is certainly a new and ambitious objective, but also somewhat controversial. This interim report is therefore rather elaborate, since it has to anticipate many questions and to face problems of an unusual nature among optical astronomers, although they are the daily bread of radioastronomers.

Beyond the scope of assessing the value of the Array concept for the VLT, this report aims to stimulate discussion within and contributions from the European scientific community, for the final report to be as realistic as possible by mid. 1986.

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Chapter 1: OPTICAL INTERFEROMETRY: AN OVERVIEW

1.1 Scientific perspectives

High angular resolution is reached in a very uneven fashion at different wavelengths (Fig. -1-1-1). Radio interferometry at centimetric wavelengths reaches currently a fraction of an arc second with Very Large Array maps while Very Long Baseline interferometry reaches milliarc second resolution and can now provide images at high sensitivity.

Breaking the seeing limit at optical wavelengths was first achieved over 60 years ago. After a long period of eclipse, optical interferometry, through various observational techniques, allows now to reach milliarc second resolution from ground observatories at visible and infrared wavelengths. When a gain of two to four orders of magnitude in spatial resolution is within reach, it is difficult to make detailed prediction about its scientific potential. Yet, preliminary results obtained with modest size telescopes and extrapolation of observation made to date establish at a reasonable degree of confidence, on the conservative side, what a Very Large Telescope could achieve in an interferometric mode.

With a few hundred meters baseline, its angular resolution would span from 0.5 marc sec in the blue to 33 marc sec at 20 µm. Its capability to produce good images with a clean beam would, on the long term, be comparable to the one of the VLA or of the Quasat project (Fig. 1-1-2). Hundreds or possibly thousands of spectral channels could be simultaneously explored.

Its sensitivity is critically dependent on the development of adaptive optics. Relatively easy in the infrared, such a development is less obvious at visible wavelengths, unless new methods, using artificial reference sources, are implemented. Only such developments would fully exploit the capability offered by the interferometric use of large mirrors 8 to 10 meters in size.

At visible wavelengths, observation of objects brighter than $m_V=14$ is straightforward. The surface of stars brighter than $m_V=9$ could be imaged in the continuum and in spectral lines, showing mass losses, activity and surface variability.

Techniques of adaptive optics and fringe tracking could push these limits to $m_V=19$ at least in some favorable cases, and open new fields such as
astrometric measurements on galaxies. Jets, marginally resolved today, and nuclei components in galaxies or quasars could be mapped, hence completing the VLA and VLBI measurements. Superluminal motions could be accessible in the visible, as well as gravitational lensing, where relative motion of the components could be detected.

![Wavelength λ(μm)](image)

Fig. 1-1-1. General view of angular resolution in the infrared (current and planned instruments). Diffraction-limited resolution obtained with speckle techniques on large ground-based telescopes is represented, as is the wavelength dependence of seeing (1 arcsec at 0.5 μm). Atmospheric absorption bands in the near IR are shown (.....). The resolution range for the near-IR ground-based interferometers is shaded (Léna, 1985).

At infrared wavelengths, atmospheric limitations are not as severe and adaptive optics is likely to become a commonplace facility on the VLT. This, combined with reduced atmospheric phase errors and relaxation on mechanical tolerances, allows a rather optimistic view on the sensitivity that the VLT could achieve. At a resolution of 5 to 20 milliarc second the most important contribution would be to obtain a clear picture of accretion discs, bipolar flows, dust ejection and condensation processes on galactic objects embedded in molecular clouds or in thick dust cocoons. Multiple objects, fragmentation and clumps would be
observable in the whole sequence from relatively cold (150 K) early fragments to objects in a proto- or preplanetary stage. Objects similar to IRc2, placed as far as 4 Kpc, T Tau at 3 Kpc or MWC 349 in Messier 31 could be resolved and possibly mapped. Such observations would allow to build a consistent picture of the compact, cold and dense blobs of matter which accompany the star formation process and probably lead to the formation of planets.

Galactic nuclei could be mapped for a whole range of objects such as NGC 1068 and other Seyfert galaxies. Closer to us, asteroids and cometary nuclei could be well resolved and the latter be spectroscopically studied while they approach the Sun.

Since the infrared quality will be a leading criterium in site selection for the VLT, the submillimetric transmission could reach wavelengths as short as 0.5 mm (600 GHz), bridging here the gap with existing millimetric interferometers.
The original model of the source.
Note that the source extends over 8 milliarcseconds, i.e. roughly 80x80 beam areas, and that the source contains compact and extended features as well as some well separated compact features. This model was chosen to provide a stringent test of Quasat.

The VLBA map of the source (declination 45°, Flux density 5.6 Jy). Note that much of the fine detail is lost. Contours: +/-1.1, 2.2, 4.4, etc. mJy/beam.

The Quasat plus VLBA map of the source (declination 45°). This map reproduces the essential features of the source down to a few times the thermal noise level for regions outside the source (see Figure 6). It can be seen that the combination of Quasat plus VLBA does a reasonable job on both extended and compact regions of emission. However, the extended regions show variations significantly above the noise level. This can be substantially improved by adding more ground stations. Contours: +/-0.3, 0.65, 1.3, etc. mJy/beam.

Fig. 1-1-2. Source restitution by the QUASAT array (orbiting VLBI observatory). Scaling these maps to VLT resolution needs to degrade the angular scale by a factor 4 at 0.5 μm, or 40 at 5 μm [Readhead et al., 1984, ESA SP-211].
1.2 Current and planned systems

Since the original observations by Michelson and Pease (1921), optical interferometry was extended to longer baselines by Hanbury Brown et al. (1974) by the method of intensity interferometry. This technique is however limited by low sensitivity. In 1975 Labeyrie succeeded in observing interference fringes produced by two independent small telescopes, showing that long baseline interferometry is feasible at optical wavelengths in a Michelson mode. The instrument now at CERGA consists of two 26 cm telescopes in a steady concrete alt–alt mount. It is now currently operated in the visible with variable baselines up to 67 m long (Koechlin and Rabbia 1985) and in the near infrared (di Benedetto and Conti 1983, di Benedetto 1985). This experiment has brought considerable impetus to optical interferometry all over the world. Similar instruments have been built in several places such as the University of Sydney (Davis 1979, 1981), the University of Maryland (Currie 1977, 1979), and Mount Wilson Observatory (Shao and Staelin 1977, 1980).

Meanwhile the multiple mirror telescope technology was developed in the U.S. as a solution for future giant telescopes. Although the resulting MMT was not intended initially to work in an interferometric mode, its cophasing was successfully achieved by Beckers, Hege and coworkers (Beckers and Hege, 1984), proving that interferometry can be done by coupling large telescopes together. A long baseline interferometer with two 1.5 m telescopes in a concrete mount is now completed at CERGA (Labeyrie et al. 1984). It is intended to work in the infrared as well as in the visible. Heterodyne interferometry at 11 microns has been demonstrated by Townes and coworkers (Townes 1984) and by Gay and coworkers (Assus et al. 1979), but Michelson interferometry is clearly superior at shorter wavelengths. The CERGA heterodyne interferometry project has been cancelled.

Long baseline interferometers are now being planned at Georgia State University, at the University of Arizona, at the University of Hawaii, at the University of California, and at the University of Erlangen, Germany.

Interferometric techniques which are now successfully developed and used for ground base observations will work in space with a much higher efficiency. Although active optics technology may allow to build diffraction-limited lightweight monolithic mirrors larger than the size of the Space Telescope.
coupling telescopes interferometrically appears to be the best approach to high angular resolution in space. This has raised considerable interest during the last few years as shown by the proceedings of two recent colloquia (Reasenberg 1984, Olthof 1984). The most ambitious projects such as TRIO (ESA) or SAMSI (NASA) envisage optical coupling of independent spacecraft over kilometric baselines but rigid structures such as OASIS are also considered (ESA). Ground-based interferometers will not be able to compete with such instruments. However, because of the difficulties introduced by the use of cryogenic liquids in space and because atmospheric distortions are much less severe in the infrared than in the visible, infrared interferometry will likely be developed on the ground well before it is used in space.

The angular resolution provided by the Space Telescope will be less good than the one already obtained at optical wavelengths from the ground, but the absence of atmospheric disturbances will allow a better sensitivity.

1.3 Some achievements

Obtaining images or spectra at the diffraction limit either with a single large pupils or with two pupils from the ground basically suffers from the same limitation: the fluctuation of the index of refraction of the Earth's atmosphere at optical wavelengths. A great deal of knowledge has been accumulated during the last ten years in the acquisition and treatment of images distorted by the atmospheric turbulence (speckle interferometry), both at visible and infrared wavelengths. The results have been thoroughly reviewed (Chelli 1984, Woolf 1982).

At visible wavelengths, spectroscopic binaries are resolved and accurate masses may be deduced, the orientation of asteroids rotation axis is determined (relationship between elliptical and rotation velocity), the orbit of Charon is established, multiple objects images are obtained (Fig. 1–3–1), a faint envelope has been mapped around α Orionis (Fig. 1–3–2), and limiting magnitudes such as given in Table 1–3–1 are now routinely reached by speckle cameras.
Fig. 1-3-1. Holographic speckle interferometry of R136a in the 30 Dor nebula.

a: One of the 8000 reduced speckle interferograms.

b: Reconstructed diffraction-limited image.

(See The Messenger No. 40 for more details.)

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Fig. 1-3-2. Image of αOri at 535 nm ($\Delta \lambda \sim 10$ nm). The image was reconstructed from interferometric observations made at the 3.6 m CFH Telescope in Hawaii, using the Gull and Daniel algorithm. Circumstellar shell is obvious. (Roddier and Roddier 1985.)

5 mm = 10 milli-arcsec.
In the infrared, most large telescopes in the world (4-5 m class) are now equipped with a dedicated specklegraph. Outstanding results have been obtained, such as the discovery (MacCarthy et al. 1984) of a very cold companion to a nearby star (V88 A and B) or the unambiguous resolution (Chelli et al. 1984) of the heating source (IRc2) of the Kleinman-Low nebula in Orion with 0.1 arc sec resolution or slightly better.

Table 1-3-1

<table>
<thead>
<tr>
<th>Table 1-3-1</th>
<th>Speckle sensitivity in the visible</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>$^{n}_v$</td>
</tr>
<tr>
<td>Asteroids</td>
<td>8-11</td>
</tr>
<tr>
<td>Pluto/Charon</td>
<td>15/17</td>
</tr>
<tr>
<td>Double stars</td>
<td>1-13</td>
</tr>
<tr>
<td>Star surfaces</td>
<td>1-5</td>
</tr>
<tr>
<td>Central objects in HII regions</td>
<td>10-16</td>
</tr>
<tr>
<td>Galactic nuclei</td>
<td>12</td>
</tr>
<tr>
<td>Quasars</td>
<td>13-17</td>
</tr>
</tbody>
</table>

Two-telescopes interferometry is only at its beginning stage of development, but its potential is well demonstrated by the measurement of over 15 stellar diameters at 600 nm from 2 to 15 milliarc sec. and the determination of limb darkening coefficients and orbital elements, as shown by Fig. 1-3-4 and Table 1-3-3 (Koechlin et al. 1979. Faucherre et al. 1983).
Table 1-3-3
Giant stars resolved with the small interferometer
(67 m baseline at CERGA)

<table>
<thead>
<tr>
<th>name</th>
<th>measured angular diameter (millisecond of arc)</th>
<th>R/R_g</th>
<th>effective temp. (Kelvin)</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>.55 μm</td>
<td>2.2 μm</td>
<td>.55 μm</td>
</tr>
<tr>
<td>α Cas</td>
<td>5.4±0.6</td>
<td>14.4±0.5</td>
<td>2678</td>
</tr>
<tr>
<td>β And</td>
<td>13.2±1.7</td>
<td>3319</td>
<td>3800±250</td>
</tr>
<tr>
<td>γ And</td>
<td>6.8±0.6</td>
<td>50514</td>
<td>4600±250</td>
</tr>
<tr>
<td>α Per</td>
<td>2.9±0.4</td>
<td>5559</td>
<td>7000±600</td>
</tr>
<tr>
<td>α Cyg</td>
<td>2.7±0.3</td>
<td>145745</td>
<td>8200±600</td>
</tr>
<tr>
<td>α Ari</td>
<td>7.6±1</td>
<td>1535</td>
<td>4300±350</td>
</tr>
<tr>
<td>β Gem</td>
<td>7.8±0.6</td>
<td>8±2</td>
<td>4900±220</td>
</tr>
<tr>
<td>β UMi</td>
<td>8.9±1</td>
<td>30±9</td>
<td>4220±300</td>
</tr>
<tr>
<td>γ Dra</td>
<td>8.7±0.8</td>
<td>10.2±1.4</td>
<td>45±10</td>
</tr>
<tr>
<td>δ Dra</td>
<td>3.8±0.3</td>
<td>15±5</td>
<td>4300±230</td>
</tr>
<tr>
<td>μ Gem</td>
<td>14.6±0.8</td>
<td>20.7±0.4</td>
<td>4530±220</td>
</tr>
<tr>
<td>α Tau</td>
<td>20.7±0.4</td>
<td>15±5</td>
<td>3960±270</td>
</tr>
<tr>
<td>α Boo</td>
<td>21.5±1.2</td>
<td>12±2</td>
<td>4420±120</td>
</tr>
<tr>
<td>α Aur A</td>
<td>8.0±1.2</td>
<td>7±2</td>
<td>5400±200</td>
</tr>
<tr>
<td>α Aur B</td>
<td>4.8±1.5</td>
<td>2.6±0.2</td>
<td>5950±200</td>
</tr>
</tbody>
</table>

Recent results with on-line fringe reduction (after Koechlin 1984):

<table>
<thead>
<tr>
<th>name</th>
<th></th>
<th></th>
<th></th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td>α Cyg</td>
<td>2.7±0.3</td>
<td>A21a</td>
<td></td>
<td></td>
</tr>
<tr>
<td>β And</td>
<td>14.0±1.3</td>
<td>MOIII</td>
<td></td>
<td></td>
</tr>
<tr>
<td>α Per</td>
<td>3.8±0.4</td>
<td>FSIb</td>
<td></td>
<td></td>
</tr>
<tr>
<td>ε Cyg</td>
<td>4.8±0.5</td>
<td>KOV</td>
<td></td>
<td></td>
</tr>
</tbody>
</table>

Figure 1.3.4

Orbit of α Aurigae B as determined with the 26 cm interferometer at Cerga (Koechlin et al. 1979).
1.4 Image formation in interferometry

A single telescope provides a spatial frequency coverage which is the two-dimensional autocorrelation function of the pupil measured in units of wavelength. In other words, the diffraction-limited image contains all spatial frequencies from zero to the telescope cut off frequency $\lambda/D$ ($D=$telescope diameter), weighted by a monotonously decreasing function (the Modulation Transfer Function or MTF). The spatial frequency domain is called the $u$-$v$ plane, after its two coordinates axis $u$ and $v$. The atmospheric turbulence introduces random phase errors over the pupil, which produce a speckled image (the "seeing" disc). The atmospheric MTF is now a random function, which highly attenuates the high spatial frequencies. Restoring this information is the goal of speckle interferometry, and this technique has now succeeded in reconstructing with a good accuracy (a few percents) both amplitude and phase of the object complex spatial spectrum up to the frequency $\lambda/D$. This mode of operation is already a powerful prospect for the VLT, and it will be described later in detail (Sec. 4).

Radio aperture synthesis has fully developed the original idea of Fizeau. The wave-front is sampled in two well separated points by two separate telescopes on a base $L$. This samples the $(u,v)$ plane at the frequency $L/\lambda$, and measures one value of the object spectrum, called the complex visibility (amplitude and phase). The $(u,v)$ coverage is obtained by varying the projected baseline $L$ over the sky either by moving the telescopes or by letting the diurnal motion do it, or both. This coverage is usually incomplete, leaving holes or empty stripes in the $(u,v)$ plane. The treatment of radio signals is made easy by the existence of phase-stable amplifiers, that allow to split the broad-band input signal into many channels for separate analysis. The absence of such amplifiers for optical signals is one of the main differences between the two fields.

Radio apertures synthesis arrays of a few kilometer baseline can be made reasonably phase-stable. They can be calibrated by observing objects of known brightness, shape and position from time to time between the observation of unknown objects.

For longer baselines, especially VLBI, phase-stability is out of question. Even for smaller arrays, the phase-stability is not always considered sufficient. In these cases, the observed object itself must be used for calibration, in a process
called SELFCAL (or phase closure, or hybrid mapping).

An absolute requirement for this technique is that most of the measurements errors can be traced back to the individual telescopes.

Fundamentally, trustable information is not obtained on the object itself but on its complex spatial spectrum, where real error bars can be given on amplitude and phase. Therefore, it is in the u.v plane that a confrontation should be achieved between models and observation. Fortunately or not, astronomer's eyes and minds are used to work with images. To reconstruct an image from a partially sampled u-v plane where visibilities are affected with errors introduces ambiguities and artefacts.

**Image reconstruction**

The final image is obtained by Fourier inversion of the measured visibility data. The feasibility of this image reconstruction process depends on the following factors:

- **a)** The u-v coverage: Since the u-v plane is incompletely sampled, the final image will contain the object convolved with some point-spread function (PSF). If the u-v-coverage is very incomplete, this PSF has high sidelobes that may obscure faint objects in the vicinity of strong sources and provide unaesthetic images of single objects.

- **b)** The amount of phase information: Visibility phases contain more information than amplitudes, but are unfortunately much harder to measure. Situation of increasing power are: no phases at all (amplitudes only), phase derivatives with respect to spatial frequency, phase closure and absolute phases. The latter is impossible with non-phase-stable instruments.

- **c)** The noise in visibility data: One must distinguish between random noise and systematic errors. The latter can often be corrected, provided they are correlated over more than one u-v sample.

- **d)** The complexity of the object itself: A small number of well separated point-sources requires less information for its reconstruction than an extended object.

In radio astronomy, telescope phase errors are estimated with the help of a tentative model of the object (a priori assumption). After correcting the data from the estimated errors, a new and hopefully better model is generated, for instance, by means of the CLEAN algorithm or by the Maximum Entropy Mapping.
These processes give a good convergence for arrays with a large number of telescopes (the VLA has 27 antennae). If the array has some redundant spacings, a model is not required to estimate telescope errors. One simply uses the fact that any differences between the output phases of identical interferometers must be caused entirely by measurements errors. This model-independence is very powerful, especially for extended objects that are difficult to model. The requirement that all errors must be telescope-based (which also implies a high S/N) remains the same however. It is important to realise that, as long as the PSF is accurately known and the object is relatively simple, excellent images can be produced with a very poor u-v coverage. The result gets worse only if high sidelobes coincide with other objects with a strength comparable to the sidelobes.

Radio aperture synthesis can now be called mature. It can produce maps with milliarcsecond resolution (VLBI), a dynamic range of up to $10^4$ (Westerbork), and truly breathtaking images of very complex and extended sources like the one of Cygnus A produced by the VLA (Fig. 1-4-1). This latter array was only designed originally to detect point-sources, but developed into a much more powerful instrument.

Current optical interferometers, such as the 0.26 or the 1.5 m CERGA systems, are only measuring amplitudes. The use of phase derivatives with respect to spatial frequency [i.e. the quantity $\nabla \phi(u,v)$] is now well developed in speckle interferometry (the Knox-Thompson method, 1974). The use of other techniques such as phase-derivation with respect to the wave frequency (spectral phase derivatives), or as phase closure has not been developed yet in optical interferometry partially because of the very small number of telescopes used in the interferometers existing to date.

Figure 1-4-1

![VLA image of Cygnus A at λ6cm.](image)

1.5 *From radio to optical interferometry*

The quality of radio interferometry results from its ability to measure complex visibilities over a certain domain of the \( u-v \) spatial frequency plane. Then, various mapping techniques are used to improve the image quality.

At optical wavelengths, the added complexity comes from two facts, due to atmospheric effects:
- a) the phase of the wavefront issued from a point source is, in general, no longer constant over the surface of each pupil;
- b) these phases are randomly varying with time, over short time scales (10 to 100 milliseconds).

Hence getting the amplitude of the complex visibility (or fringes contrast) is straightforward, but atmospheric phase disturbances do prevent a direct phase measurement of the phase of the wave-front.

<table>
<thead>
<tr>
<th>Table 1-5-1</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>Chromatic dependence of coherence area</strong></td>
</tr>
<tr>
<td>(1 arc sec seeing)</td>
</tr>
<tr>
<td>( \lambda (\mu m) ) &amp; 0.5 &amp; 2.2 &amp; 5 &amp; 10 &amp; 20</td>
</tr>
<tr>
<td>( r_0 (m) ) &amp; 0.1 &amp; 0.59 &amp; 1.58 &amp; 3.64 &amp; 8.36</td>
</tr>
</tbody>
</table>

The most favourable situation is encountered at longer wavelengths, where the size of the atmospheric coherence area \( (r_0) \) becomes comparable to the size of the pupil itself (Table 1-5-1).

At shorter wavelengths, or when the condition

*atmospheric coherence area > telescope diameter*

is not fulfilled, adaptive optic, operated in the focal plane is, in principle, able to restore in real time full coherence over the whole pupil. Whenever the full coherence of the wave front can not be achieved over each pupil, one must break the pupil into sub-areas, each one being coherent; the final result is much less favorable in terms of signal-to-noise ratio.

Comparing with radio wavelengths, where phase errors are constant with time and uniform over individual telescopes, it indeed appears that phase retrieval at optical wavelengths shall be more difficult.

Radio interferometry uses a single detector with possibly multiple spectral channels. Optical interferometry needs at least as many pixels, on each spectral...
channel, as there are coherent areas over the pupil. Since u–v coordinates are measured in wavelengths, different spectral frequencies represent different u–v points when they are measured with the same interferometer. As long as the object appearance remains approximately the same at different frequencies, this spectral information may help to determine the phase, or rather the phase derivative $\overline{v}_\phi$ with respect to spatial frequency.

Another approach to determine the phases is instantaneous phase closure, as in radio interferometry: one measures the phase derivative over the u–v plane by dividing the pupil in sub-areas.

Finally optical interferometry requires mechanical accuracies and stabilities of a fraction of the wavelength, as it shall be discussed below in detail.

**Simultaneous measurements**

The techniques that are used with such notable success in radio aperture synthesis have the following requirement: the number of independent measurement errors must be smaller than the number of visibilities measured. This gives the constraints that allow us to generate a model of the observed object iteratively. In the presence of rapidly varying atmospheric phase errors, consecutive measurements will have different errors (unless some method of phase-tracking is possible). The only way to reduce the number of independent phase errors in this case is to measure a number of visibilities simultaneously. There are several ways to do this:

1) With a single interferometer (2 telescopes), one can measure the phase-difference between 2 u–v points that are less than one telescope diameter apart. In practice this provide the phase-derivative.

2) With only 2 telescopes it is also possible to measure the visibility at multiple frequencies by dispersing the light. Since u–v coordinates are measured in wavelengths, each frequency corresponds to a different u–v point.

3) An array of N telescopes can provide N(N–1)/2 different interferometers. If all measurements errors can be assigned to individual telescopes, the number of independent errors is N, which is smaller than N(N–1)/2 if N ≥ 3. The measured phases can be combined to build error-free quantities called "closure phases".

Of these three, the array is the most powerful, but requires simultaneous use of all telescopes. Feasible in principle, these methods remain to be tested in
optical interferometry. Already, speckle work with large telescopes is a considerable help in understanding the wavefront phase perturbations and their remedy through image reconstruction methods. Speckle masking (Lohmann et al. 1983) for instance, a parent method to phase closure, has already succeeded in image reconstruction at optical wavelengths.
Chapter 2: A DETAILED APPROACH TO OPTICAL INTERFEROMETRY

The goal of an interferometer is to measure the complex visibility (amplitude and phase) of the object Fourier transform in the frequency \((u, v)\) plane. The largest extension of the baseline sets the ultimate resolution, while the density of the \((u, v)\) coverage and the signal-to-noise ratio sets the image quality. Once the complex visibility is obtained, conventional techniques well known to radioastronomers may be applied to restore the image.

An optical interferometer must be carefully investigated wherever it differs from a radio interferometer: beam combination, access to the complex visibility, nature of detectors, ultimate sensitivity. Theoretical and experimental approaches to optical interferometry (Labeyrie et al. 1984, Roddier and Léna 1984, Woolf 1984) provide today a firm ground to establish the configuration in which telescopes of 8–10 m size could be used efficiently and to derive conservative performances. In this chapter, we shall discuss in detail what an ideal interferometer would be, in terms of working condition, image production, and ultimate sensitivity. We defer to the next chapter the specific discussion on the VLT, where indeed many compromises would have to be done.

2.1 Beam combination

Each telescope provides an afocal beam, which is fed into a common "beam combination area". The different beams interfere over the detector, which measures the complex visibility.

When only two fixed telescopes are used, the "beam combination area" is mounted on sliding tracks, in order to keep constant the optical path from the source to the detector through either telescopes, correcting for the diurnal motion. When more than two fixed telescopes are used, optical delay lines become necessary.

Optical delay lines are however not as easy to build as radio ones. Trombone-like arrangements of four mirrors are required, implying losses, polarisation problems, extra-maintenance costs, etc.
Delay lines are unnecessary if the telescopes are arrayed along an ellipse, intersection of the ground plane with a paraboloid aimed at the star. The central station where the beams are recombined is at the focus of this paraboloid.

The ellipse however keeps deforming itself during observation, since the axis of the paraboloid has to track the star motion. This implies that the telescopes must move during the observation if delay lines are to be avoided. The motion does not have to be accurate since small delay line elements, with travel in the 1–10 mm range, can easily be incorporated in the central table. Telescope positions (motion inaccuracies, vibrations...) can be monitored by laser beams and the compensation of errors be made by the small delay lines.

A system of radially arranged tracks can meet these requirements, but ideally superior flexibility would be obtained with a smooth platform of 300x300 meters carrying the four telescopes on air bearings.

Among the existing types of telescopes mounts, those best adapted to such attractive designs are the alt-alt and the spherical. They indeed provide a horizontal coudé output with only 3 mirrors (primary, Cass and Central flat). Instead, the alt-az mount requires seven mirrors and the vertical coudé beam cannot be reflected horizontally in all directions. Alt-az mounts would be compatible with radial tracks, in which case the coudé beam could propagate in a tunnel below the tracks. They do not seem adequate for the platform concept, where the coudé beams would propagate in free air or tubes above the platform. The free-air propagation of coudé beams is affected by seeing, but considerably less than the vertical propagation of the wide beam coming into the telescope.

The performances of Michelson interferometry with large ground-based telescopes operating at visible wavelengths have been discussed by Greenaway (1979a) and more recently by Roddier and Léna (1984). Expressions are given for the signal-to-noise ratio on the fringe visibilities. An important result is that it is more efficient to use the telescopes pairwise rather than to superimpose all the images unless the configuration is highly redundant (i.e. many telescope pairs contribute to the same spatial frequency) (Greenaway 1979b).

In a non-redundant configuration, the use of telescope by pairs has two disadvantages: first, it does not allow to apply phase-closure methods, classical at radiowavelengths and eliminating atmospheric phase errors. Second, it provides, at a given time, a u-v coverage limited to a single point or to a small
neighbourhood of a single point.

In the pair operation, two basic configurations are possible: either use a single pair at a time, or use all possible pairs.

For simplicity, we restrict the discussion to single pair combination.

**Beam superposition**

Images from the several telescopes can be recombined in a variety of ways:

1. The beams can have different orientations, with no angular overlap at the point where they cross each other, either in the pupil or in the image plane. This is achievable with any number of telescope. The pattern of angles adopted between the beams may or not reproduce the pattern of apertures. In the former case, referred to as the Fizeau case, optical path differences are invariant in the field, at least to the first order. In the latter case, known as the Michelson case, a first-order dependance of path difference with respect to field position occurs and fringes are observed in the field. Fizeau arrangements have advantages for compact arrays, but require many pixels in the case of diluted arrays. Michelson
arrangements are of interest for diluted apertures used in narrow fields. They also provide such possibilities as having a redundant entrance aperture with a non-redundant exit aperture (Olthof 1984).

2- Coaxial recombination with beam splitters. This is an extreme case of Michelson-type recombination. It can provide a flat interference field if the waves have uniform phase, which does not occur in the visible when dealing with large apertures, unless adaptive optical corrections are made.

Image plane and pupil plane interferometry have their advantages, but it is likely that the latter will ultimately be the most favorable for large telescopes interferometry (Chelli and Maratti 1985).

Adaptive optics

Interferometric operation is optimized when a uniform phase is achieved over the individual pupils to be combined. This can be achieved with adaptive optics. A closed-loop control system, that senses the actual phase errors, and an active optical element that reshapes the optical wavefront by adding a set of controllable optical path differences allow a real-time compensation of the wavefront. Basically two different correction strategies are possible: zonal or modal correction. If the modal approach is used rather than a zonal decomposition, significant correction is possible even with a limited number of modes, allowing partial or full pupil correction.

An adaptive system contains four basic elements: an optical train and receiver, a phase-error sensor, a servo-control system, and an optical phase shifting element. The distortion of the received wavefront is usually compensated by reflecting the light beam at a deformable mirror. The surface is adjusted in real-time to compensate the path length aberrations. The information required to deform the mirror is obtained by analyzing the reflected beam with a wavefront sensor. Two methods in wavefront sensing have been developed further for this type of application: the Shack–Hartmann test and the shearing interferometer. With these wavefront measurements a map of wavefront errors determines the signal required to drive the deformable mirror by expanding the spatial $f_n(r)$ and temporal $a_n(t)$ dependent correction function $C(r, t)$

$$C(r, t) = \mathcal{L} a_n(t) \cdot f_n(r)$$
A typical set of spatial functions are the Zernicke polynomials, which are usually used to describe the common optical aberrations. For atmospheric turbulence compensation they are only optimal for the modes with lower numbers. High order spatial and temporal frequency fluctuations can be more efficiently handled by the Karhunen–Loève functions.

Typical values for the spatial and temporal constraints in the visible wavelength range are:

- subaperture size: 10 cm
- time constant: 10 milliseconds

For 8–10 m telescopes this would lead to approximately 6000 controlled subapertures working at frequencies up to 1000 Hz. More realistic and technically feasible is the correction at infrared wavelengths. Systems with 150 to 200 subapertures have been developed and tested for military applications (e.g. Itel Corp., AOA, CGE, MBB, etc...).

In order to avoid wavefront sensing at infrared wavelengths, experiments are in progress to use the low frequencies in the visible transfer function for the correction in the infrared. So far this approach looks very promising.

Existing technology and developments expected during the next few years make it absolutely likely that the single VLT telescopes will be equipped with adaptive correction devices allowing a full pupil correction at wavelength longer than 4 \( \mu \)m. These systems will give an interferometer the required phase uniformity over the individual pupils. At shorter wavelengths, partial MTF correction will still be obtained. Furthermore, it has to be investigated whether the deformable mirrors can additionally be used for the stabilisation of the interferometric beam combination train.

**Multiple spectral channels**

Radio interferometers have multispectral capabilities: Westerbork has 32 000 channels. Nobeyama has 20 000. VLA has more than 10 000. The radio arrays choose to distribute these channels over different polarization (to improve signal-to-noise ratio) or over sine and cosine terms (to measure phase) or to measure many baselines simultaneously. They could also put all the channels on one or two baselines and immediately reach \( 10^4 \) spectral channels.

At optical wavelengths, it may or not be of interest to display separately the interferometric information collected in different colors. Indeed, the
color-dependant morphology of stars, galaxies, etc... has interesting variations associated to the various spectral lines. Like radio arrays, optical ones can observe 10 000 or more spectral channels simultaneously. A variety of spectrographic techniques can achieve the required separation of wavelengths, from prismatic dispersion to more elaborate Fabry-Perot and possibly Fourier transform spectrometer techniques using directly the fringe modulation (J. M. Mariotti 1985). Also, field grating techniques can provide a larger number of monochromatic images in different colors.

Differential interferometry may be conceived as an extrapolation of differential speckle interferometry.

Phase and amplitudes may be measured at a number of adjacent wavelengths. If the assumption is made that the object shape is wavelength independant over a certain spectral interval, it becomes possible to determine the amplitude and the phase gradient $\phi(u,v)$ of the source (spectral phase-gradient estimate).

Detection requirements

Severe requirements on detectors in term of number of pixels are present, as soon as adaptive optics can not provide an uniform phase over the pupils. These requirements are evaluated in Table 2-1-1.

<table>
<thead>
<tr>
<th>Wavelength ($\mu$m)</th>
<th>0.4</th>
<th>0.8</th>
<th>2</th>
<th>5</th>
<th>10</th>
<th>20</th>
</tr>
</thead>
<tbody>
<tr>
<td>Complete adaptive optics</td>
<td>1</td>
<td>1</td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>No adaptive optics (7m telescopes)</td>
<td>$\frac{420}{\lambda_{\mu m}}$</td>
<td>$\frac{140}{\lambda_{\mu m}}$</td>
<td>$\frac{840x280}{\lambda_{\mu m}}$</td>
<td>$\frac{210x70}{\lambda_{\mu m}}$</td>
<td>$\frac{42x14}{\lambda_{\mu m}}$</td>
<td>$\frac{21x7}{\lambda_{\mu m}}$</td>
</tr>
</tbody>
</table>

(*) under typical (1 arc sec) seeing condition
The need to freeze atmospheric phase drifts implies fast detection, faster in the visible (1 ms) than in the infrared (10 to 100 ms). Mosaics of image detectors are needed to record several spectral channels simultaneously.

**Measuring \( \dot{\varphi} \)**

Atmospheric phase fluctuations between pupils prevent the direct determination of the phase of the complex visibility. But interferences between sub-areas of individual pupils provide again the phase derivative of the object in a given domain (spatial phase gradient estimate) as shown in Fig. 2-1-2.

![Figure 2-1-2](image_url)

Forming pairs between sub-areas (\( r_o \) in size, or smaller) of the two interfering pupils, in order to obtain phase information.

\[
\delta \phi_1 = \phi_1' - \phi_1 = \Phi_{\text{source}}(\vec{W}) - \Delta \phi \\
\delta \phi_2 = \phi_2' - \phi_1 = \Phi_{\text{source}}(\vec{W}') - \Delta \phi \\
\Delta \phi = \text{atmospheric phase error} \\
\delta \phi_2 - \delta \phi_1 = \Phi_{\text{source}} \cdot (\vec{W}' - \vec{W})
\]

There is no direct equivalent of the technique in radio interferometry, since the beam throughput always use the full pupil area.

It is interesting to consider the wavelength dependence of random atmospheric phase excursions between two pupils placed at a distance \( L \) apart. For stationary condition, the rms value \( \sigma_\phi \) is given by

\[
\sigma_\phi(x_d) = 6.4 \left[ \frac{L}{x_o(\lambda)} \right]^{5/6}
\]

as long as \( L \) is smaller than the outer scale of turbulence which is probably of the
order of a few tenths of meter at an astronomical site (Table 2-1-2)

<table>
<thead>
<tr>
<th>Table 2-1-2</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>Rms phase excursion between two pupils, 50 m apart</strong></td>
</tr>
<tr>
<td><strong>[homogeneous turbulence, L(outer scale) &gt; 50m]</strong></td>
</tr>
<tr>
<td>( \lambda ) (( \mu )m)</td>
</tr>
<tr>
<td>( \sigma )</td>
</tr>
</tbody>
</table>

(*) under typical (1 arc sec) seeing condition

in practice, scarce data are currently available (Mariotti and Benedetto 1984), but they show values significantly smaller (an order of magnitude) than the "worst case" given by the above formula. In any case, the gain in phase stability is considerable at infrared wavelengths.

**Field of view**

An MMT offers a PSF which has very limited side lobes (typically those of an Airy function). Wide-field (arc sec) speckle interferometry is possible since the exit pupil is always exactly homothetic to the entrance pupil. The situation is different in a thin array, where the PSF will present high side lobes. Objects which have a typical extension of the order of the resolution will not suffer from confusion or reduced photometry accuracy, as will larger objects, unless cleaning procedures are used a posteriori, as in radioastronomy.

The condition of path equalization for two telescope interferometry, when ever the object is on-axis or off-axis is fulfilled with an MMT (similar in this respect to a standard telescope), while it varies with a multitelescope interferometer. The field-of-view of the interferometer will be limited to the size of the seeing disc (\( \sim 1-2^\circ \)) (Waller 1980). Corrective action may be taken at the cost of additional optics (Enard 1984), which may, on the other hand, introduce higher thermal loads for IR use.

**Phase closure**

This technique is of common use in radio arrays. In an array of \( N \) antennae, up to \( N(N-1)/2 \) simultaneous interferometers baselines can be used.
If all the antenna pairs are combined, this factorization reduces the number of unknown complex gains that must be determined by calibration from \( N(N-1)/2 \) to \( N \). When \( N < N(N-1)/2 \), it is thus possible, in principle, to determine both the unknown instrumental (and atmospheric) complex gains and some parameters of the source from the observation (self-calibration methods).

In optical interferometry, \( N \) is no longer the number of telescopes, but the total number of coherent areas \( \sigma_0(\lambda) \) in size over all the pupils. With 4 telescopes, \( N \) may vary between several thousands (at 0.5 \( \mu m \)) to a few units (at 20 \( \mu m \)).

Bispectral analysis or triple-correlation (also called speckle masking) is a generalization of phase-closure methods. Indeed the bi-spectrum transfer function is reached because of phase-closure

\[
\tilde{S}(\tilde{u}) \tilde{S}(\tilde{v}) \tilde{S}^*(\tilde{u}+\tilde{v})
\]

real \( \tilde{u} \tilde{v} \)

point source bi-spectrum \( \tilde{u}+\tilde{v} \)

phase closure

**Compact pupils** (non monolithic)

Beckers (1984) makes a distinction between monolithic arrays in which the internal pathlengths remain invariant while tracking the source, and non-monolithic arrays in which the internal pathlengths are continuously adjusted to compensate for external pathlengths variations. He also distinguishes between compact arrays providing full coverage of the frequency plane within an approximately circular outer bound, and thin arrays providing only partial coverage. He proposes a classification in three types of arrays. Type 0 are compact monolithic such as the MMT. Type 1 are thin monolithic such as the Michelson and Pease experiment or the University of Arizona proposed versatile array. Type 2 are thin non-monolithic. This distinction between thin and compact arrays is similar to the distinction made in early radioastronomy between "grating" synthesis and "aperture" synthesis.

Within Type 2 arrays, the **diluted** type gives a coverage which presents holes, while the **compact** type gives a \((u,v)\) coverage which presents a complete continuity (Fig. 2–1–3)
Compact array: the pupil autocorrelation function has no hole.

Diluted (thin) array: there are holes in the u-v coverage.

An interference pattern produced by a compact type of array is basically similar to a speckle image given by the large pupil of a single telescope. In such a case, phase continuity is ensured over the whole covered area of the (u, v) plane, and the complex visibility of the source can be fully reconstructed. Various methods achieve this goal; speckle masking is one of them (Fig. 2-1-4) [Lohmann et al., 1983].

Speckle masking observation of the close spectroscopic double star Psi SGR (One of 100 evaluated speckle interferograms and reconstructed image, see Weigelt and Wirnitzer, Optics Lett. 8, 389, 1983, for more details).

The advantage of speckle masking is the fact that no point source is required near the object for image reconstruction.
Combining unequal pupils

It can be shown (Roddier and Léna 1984) that the combination of unequal pupils does not in general bring a significant advantage in term of signal-to-noise ratio. The signal-to-noise ratio given by a pair of telescopes of unequal size is, in the best case, only \( \sqrt{2} \) times larger than the one given by a pair of telescopes of the smaller size. In the worst case (numerous speckles in the image), the largest telescope does not even bring this improvement. It is only at short infrared wavelengths (2–3 \( \mu \)m) that the classical situation, encountered in radioastronomy, may be obtained: the signal-to-noise ratio varies as the geometric mean \( \sqrt{S} \) of the two areas.

Combining a smaller telescope with one of the dishes of the VLT may therefore be useful in terms of improved u–v coverage, but it will not take benefit of the largest pupil in the pair.

2.2 Image reconstruction

From the analysis of the combined beams, one may safely assume that either one of the following measurements may be achieved.

a) Amplitude \( |V_\lambda(u,v)| \) of the complex visibility \( V_\lambda \). No phase information is available. This is the most straight case. Noise will cause an rms error \( \sigma_\lambda, |V| \).

b) Amplitude \( |V_\lambda(u,v)| \) and phase gradient \( \Phi_\lambda(u,v) \). The phase gradient may be determined either by the spatial or by the spectral method. Measurement errors are \( \sigma_\lambda, |V| \) and \( \sigma_\lambda, |\Phi| \).

c) Amplitude \( |V_\lambda(u,v)| \) and phase \( \Phi_\lambda(u,v) \). This optimum case could be obtained directly only in a favorable phase closure configuration, or in the speckle masking configuration discussed above.

(u, v) coverage

The autocorrelation of the pupil, projected on the sky, gives the instantaneous (u, v) coverage of the instrument. With a type 0 array (MMT), this coverage is independent of tracking and time. The Fourier transform of this pupil function is the point-spread function (PSF) of the interferometer.
Any array of individual telescopes (non-monolithic arrays) provides a coverage of the ($u, v$) plane which is time- and declination-dependent. For a given source declination, fixed telescopes provide a fixed coverage. Movable telescopes obviously extend this coverage.

A goal of interferometer observations will be to make 2-dimensional images, with a roughly circular beam (point spread function). At centimeter wavelength, this can be done, for high-declination sources only, by east-west arrays, which synthetize the image as the earth rotates. At visible and IR wavelengths, however, the atmospheric absorption is large, preventing optical/IR arrays from tracking sources down to the horizon, as centimeter radio arrays do. A typical cutoff is 2 air masses, or elevation 30 deg. This means that a one-dimensional visible/IR array will always have large sectors missing from its $u-v$ coverage, even for high-declination sources.

This major differences to radio synthesis telescopes, combined with the many important sources located near the celestial equator (Orion, 2C273, 3C279, the galactic center, etc...), where a linear array is anyway limited, makes it imperative for a useful optical interferometer to have two-dimensional coverage.

To simulate image reconstruction, various configurations were studied, all with 4 telescopes:

- **Thin arrays.** The ($u, v$) coverage is not continuous. Three different interferometer configurations were considered (Fig. 2-2-1):
  - The first (QEW100) has four fixed telescopes on a line at azimuth 100 deg. The departure from a strictly east-west line leaves some north-south resolution even at the equator. The four telescopes are placed so as to provide the six space harmonics of 25 m up to 150 m.
  - The second test configuration (FIXY) is a wye with four fixed telescopes at the crossing and at the tips of three arms placed at 120 deg from one another. One of the three arms is north-south.
  - The third configuration (VARY) has one fixed telescope at the centre of the wye and three movable telescope, one on each arm of the wye. Vary has been tested in two different configurations thus providing twelve space harmonics against six for the first two interferometers under test.
The three configurations used to test the VLT-PSF in the interferometric mode. The four telescopes (10m) are on scale with the selected baselines.

The u-v coverage of these configurations, placed at a latitude of -25°, has been computed for sources at various declination and the point spread function is directly deduced (Fig. 2-2-2).

A model source, having four components of different intensities and sizes is then used for simulation. Observations of this source have been simulated with the three interferometers described above. Point spread functions have been used to restore the images through the CLEAN algorithm which is one of the methods to recover the missing space harmonics by interpolation in visibility space. This process does not extrapolate in visibility space that is to say it does not provide improved resolution relative to the actual observations. The source was assumed at the field center and the effect of attenuation by the diffraction pattern of a single 8 m aperture was not included because it does not affect the conclusions. Fig. 2-2-3 shows the restorations at declinations -65 deg and -15 deg for the three test arrays and three different amounts or random phase errors: 0°, 30° and 60° rms. Fig. 2-2-4 gives similar data, using the CLEAN algorithm.

(Figs. 2-2-2, 2-2-3 and 2-2-4 are located at the end of this report.)
Because it is likely that only gradients will be measured, phase offsets (in addition to random phase errors) should have been assumed between the various (u, v) tracks in a case like QEW100 where (u, v) tracks never cross one another. However, in the case of all the wye configurations, (u, v) tracks are crossing one another. It can therefore be assumed that this kind of partial redundancy will allow the construction of a consistent set of phases from the observations all over the (u, v) plane within two arbitrary phases which do not prevent image reconstruction. The two phases just prevent absolute positioning of the source. Random phase offsets between the (u, v) tracks make the restorations worse but do not affect the conclusions stated below.

The conclusions are the following: the restoration with QEW100 is quite unacceptable at all declinations. The restoration is acceptable with FIXY at high declination and moderate phase errors only. The restoration becomes acceptable for 30 and 60 deg rms phase errors and all declinations with VARY.

It again appears that some two-dimensional coverage is essential for image quality. Moreover, if many (u, v) tracks cross one another, a common relative phase reference is easier to obtain and phase gradient information is sufficient to reconstruct the images.

b- Compact arrays: In this configuration, the continuity in the (u, v) plane allows the use of speckle masking techniques. If all telescope are simultaneously used, the compact array provides a better u–v coverage, hence a better PSF. Fig. 2–2–5 shows a digital simulation of image reconstruction from interferograms by speckle masking. In this experiment photon noise corresponding to an average of 10 photons per pixel has been simulated. The phase and amplitude may then be completely reconstructed.
Figure 2-2-5

Image reconstruction from interferograms by speckle masking. The figures show (1) the pupil function, (2) the laboratory object, (3) one of the 30 evaluated interferograms (photon noise corresponding to an average of 10 photons per pixel), (4) theoretical and reconstructed Fourier phase, and (5) the reconstructed image. (See Hofmann and Weigelt, 1984 for more details)
c- Hybrid concept: One problem with the thin array is the gap in interferometric coverage caused by the minimum possible separation of telescopes. This is another important difference to radio arrays, due to the fact that visible/IR telescopes have greater f/D ratios than radio dishes, and hence longer tubes. Therefore, relative to their sizes, independent visible/IR telescopes cannot be spaced as closely as radio dishes. This sets fundamental limits to the imaging capabilities of the interferometer.

One solution to this problem is two mirrors on the same mount. This would give a rigid structure ensuring optical path stability for interferometry over baselines of 0 to 16 meters. This arrangement can also easily provide the adjacent spacings at longer baselines; for example, interferometry with another telescope 40 m away automatically gives baselines of 40 m and 48 m. For complicated astronomical sources, a complete coverage of all spacings, with no gaps, could thus be built up by the use of additional, movable telescopes. For simple sources, the dual-mirror telescope gives built-in quasi-redundancy for each spacing, which helps to sort out ambiguities and possibly, phases. This concept is an hybrid combination of MMT (2 dishes on the same mount) and thin array (additional dishes).

Table 2-2-1 provides an overview of the various possibilities.

<table>
<thead>
<tr>
<th>Type of array</th>
<th>u-v coverage</th>
<th>PSF</th>
<th>Maximum baseline</th>
<th>Available information</th>
<th>Image restoration</th>
</tr>
</thead>
<tbody>
<tr>
<td>MMT</td>
<td>continuous</td>
<td>excellent</td>
<td>25 m</td>
<td>yes yes yes</td>
<td>speckle masking</td>
</tr>
<tr>
<td>Compact linear array</td>
<td>continuous</td>
<td>anisotropic</td>
<td>50 m</td>
<td>yes yes yes</td>
<td>speckle masking</td>
</tr>
<tr>
<td>Compact 2D-array</td>
<td>continuous</td>
<td>isotropic</td>
<td>excellent</td>
<td>40 m</td>
<td>yes yes yes</td>
</tr>
<tr>
<td>Hybrid array</td>
<td>continuous</td>
<td>isotropic</td>
<td>good</td>
<td>150 m</td>
<td>yes yes yes</td>
</tr>
<tr>
<td>Thin linear array QEW (Fixed telescopes)</td>
<td>anisotropic</td>
<td>isotropic</td>
<td>150 m</td>
<td>yes yes no</td>
<td>difficult</td>
</tr>
<tr>
<td>Thin linear array QEW one N-S telescope</td>
<td>more isotropic</td>
<td>isotropic</td>
<td>good</td>
<td>150 m</td>
<td>yes yes some</td>
</tr>
<tr>
<td>Fixed Y-array</td>
<td>holes</td>
<td>isotropic</td>
<td>good</td>
<td>150 m</td>
<td>yes yes some</td>
</tr>
<tr>
<td>Movable Y-array</td>
<td>no holes</td>
<td>isotropic</td>
<td>excellent</td>
<td>150 m</td>
<td>yes yes yes</td>
</tr>
</tbody>
</table>

(*) Inverse of cut-off frequency D/λ.
Higher resolution is obtained if some a priori information is available.
2.3 Sensitivity performance and limiting magnitudes

The limited amount of experience available in interferometry would lead here to a conservative approach. Yet, it may be noted that speckle interferometry achievements in terms of sensitivity are extremely good and actually close to theoretical limits. Since the limiting factors are basically the same, there is no reason to doubt the ultimate performances of an interferometer.

It is necessary to treat separately visible and infrared wavelengths, because the limiting factors are different.

**Infrared case**

From a theoretical point of view, interferometry in the infrared is characterized by the decreasing number of atmospheric coherence areas, the longer atmospheric coherence times and the thermal photon noise which limits the performance of the detectors. With detectors arrays of limited pixel numbers, areas of coherence can be resolved, and image stabilisation can be applied with rather of low spectral coherence requirements, permitting relatively wide spectral filters. In particular below 10 \( \mu \text{m} \) detectors with high performance under low background conditions are required, as used in cryogenic infrared instrumentation in space telescopes. Such detectors are fast enough to match the motions of the areas.

Though discrete detectors are useful, in particular at long wavelengths, where only small numbers are required, monolithic detector arrays become desirable as the wavelength decreases. Advantages are in the larger filling factor, the more uniform performance of the elements and the multiplexed readout, which limits the cryogenic load. Drawback is at present their larger intrinsic noise compared to single detectors (about 10 times).

Existing data show that discrete elements in connection with cooled J-Fet amplifiers exhibit noise levels down to \( 10^{-17} \) WHz\(^{-0.5} \) and even lower values have been reported. For arrays, values of \( 10^{-16} \) have been reached but lower values can be expected, depending on reduction of the readout noise. At present CID and CCD readouts attain noise values of 1500 to 3000 electrons at quantum efficiencies of \( \eta = 0.2 - 0.4 \); recent new developments using TMOS integrated direct voltage readout (DVR) techniques promise even noise levels down to 100 electrons (giving NEP = \( 10^{-17} \) at 10 \( \mu \text{m} \) if \( \eta = 0.2 \)). Such numbers depend on
temperature and type of readout and have lowest values at low frequencies (1-10 Hz). More detailed information is presented in Table 2-3-1 for detectors with very low intrinsic noise which form the basis of recent array developments.

### Table 2-3-1

Summary of Detector Developments

<table>
<thead>
<tr>
<th>Material</th>
<th>Range (µm)</th>
<th>Details</th>
</tr>
</thead>
</table>
| InSb     | 2-5       | Discrete: NEP $10^{-17}$ WHz-0.5  
Array technology at  
- SAT (France) with 8*8 pixels (1); CID readout;  
  NEP 1.4 $10^{-16}$, Design goal 3 $10^{-17}$ WHz-0.5  
- SAT with 32*32 pixels (1); CID  
- SBRC (US) with 32*32 pixels (1); CCD; NEP 5 $10^{-17}$ WHz-0.5  
- AEG-Telefunken (Germany) with 32*32 pixels (2) |
| Si:In    | 3-7       | Array technology at  
- LIR (France), 8*8 pixels (1); Multiplexed J-Fet readout  
  LIR 32*32 development;  
  TMOS or CCD (hybrid), expected noise 100 electrons  
- Rockwell (US) with 64*64 pixels, DVR.  
- AEG (Germany) with 64*2 pixels (3), other formats possible |
| Si:Ga    | 5-17      | Discrete: NEP $10^{-16} - 3 \times 10^{-17}$ WHz-0.5 (4)  
Array technology at  
- AEG (Germany) with 64*2 pixels (3), other format possible  
- LIR (France) with 8*8 pixels (1), multiplexed J-Fet  
- LIR (France) 32*32 development (1), TMOS-DVR;  
  expected noise 100 electrons. |
| Si:Bi    | 6-18      | Discrete: NEP $3 \times 10^{-17}$ WHz-0.5 (6)  
Array development reported from Aerojet (US) with 32*32 and 64*2 formats (5), NEP 3 $10^{-17}$ WHz-0.5 |
| Si:As    | 10-24     | Discrete: NEP $4 \times 10^{-17}$ WHz-0.5 (6).  
Array developments unknown |
| Si:P     | 12-28     | Discrete: NEP $3 \times 10^{-16} - 2 \times 10^{-17}$ WHz-0.5 (4).  
Array technology at  
- AEG (Germany) (3), 64*2 pixels, evt. other formats. |

References:
1 ISO-camera proposal to ESA, Jan 1985.  
2 3rd Int Conf. Infrared Physics, Zurich 1984, Kohlbacher et al.  
3 ISO-short Wavelength Spectrometer Proposal, Jan 1985  
5 ibid, Werner & Mc Creight.  
With image stabilisation a stable interference pattern over long base lines is obtained within the coherence time $\tau$ under rather reduced spectral coherence requirements (Roddier & Léna, 1984), and the limiting magnitude mainly depends on the value of $\tau$ and the detector noise. At short wavelengths this is the pure intrinsic noise. At longer wavelengths, where thermal noise limits the sensitivity, arrays become very profitable: The coherence areas within the telescope pupils are not averaged out by one detector. This allows to track the phase changes of the coherence areas which interfere. As a result integration of the signals over long times becomes feasible. In the short wavelength region if pure detector noise dominates, the advantage is much less because the signal is diluted over pixels without a decrease of the noise, though integration is still possible.

![Image of NGC 1068](image)

**Fig. 2-3-1.** Interferometric sensitivity in the infrared, with a pair of either $D = 1.5\,\text{m}$ or $D = 8\,\text{m}$ telescopes. Seeing $1\,\text{arc\,sec}$, atmospheric coherence time $100\,\text{msec}$ at $5\,\mu\text{m}$, visibility $S/N$ ratio $= 5$, $8\,\text{m}$ telescope pupil phased by adaptive optics. Integration times are respectively $1\,\text{s}(1.5\,\text{m})$, $1\,\text{s}$ and $1\,\text{hr}(8\,\text{m})$. Spectral bandwidth $\Delta \lambda / \lambda = r_0(\lambda) / D$ (Roddier and Léna, 1984).

Long time integration (hours) with phase drifts control requires much development, but would lead indeed to a much higher sensitivity.
Visible case

At visible wavelengths, detector noise becomes completely negligible, but the atmospheric coherence area is small ($r_0 \sim 10$ ms). Fast detectors, with a large number of pixels are therefore necessary.

Limiting magnitudes have been estimated for typical seeing conditions (1 arcsec seeing) assuming that fringes can be frozen on 0.02 second exposures. Several cases have been considered. The limiting magnitude is found to be of the order of $m=4$ for visual fringe detection independently of the telescopes size. Photoelectric fringe tracking should allow to observe objects up to at least $m=9$ (Tango and Twiss 1980; Shao and Staelin 1980), possibly $m=14$ (Roddier and Léna 1984), with very little dependence on the telescope diameter. The improvement given by a large telescope becomes significant when adaptive optics reduces the number of speckles in the image. "Classical" adaptive optics in the visible requires a bright object ($m_V < 9$) to be implemented: as in the radio methods of self-calibration or phase-closure, the adaptive seeing corrections require a signal-to-noise ratio greater than unity on the short (0.02s) exposure. Therefore no improvement can be, in principle, expected from the use of a large telescope, as far as limiting magnitude is concerned in adaptive optics.

On the other hand, a range in magnitude from 13 to 20 and possibly more is accessible if adaptive seeing correction proves feasible with artificial reference sources. The feasibility of using laser pulse techniques to provide such references sources should be further investigated, since it could influence the philosophy of the VLT. (Foy and Labeyrie 1985)

It may therefore be considered that "bright" objects ($m_V < 9-14$) would be of easy access with a VLT but "faint" objects definitely require significant developpements before they can be measured.

For "bright" objects, the use of a large telescope as compared with a smaller one brings a favorable gain in observing time: the time required to obtain a given signal-to-noise ratio on visibility varies as the reciprocal of telescope area.

Comparing with space platforms

The interferometric mode of the VLT has the great advantage that it can yield higher angular resolution than the Hubble Space Telescope and other space telescopes and ground-based telescopes of the near future. Only a long-baseline
interferometer in space, for example TRIO (Labeyrie et al. 1982), could yield higher resolution. Table 2–3–2 compares a VLT interferometer with 100 m baseline with the Hubble Space Telescope.

<table>
<thead>
<tr>
<th>telescope/method</th>
<th>resolution</th>
<th>limiting magnitude $m_v$</th>
</tr>
</thead>
<tbody>
<tr>
<td>100m VLT interferometer</td>
<td>0.001&quot; (at 500nm)</td>
<td>14-18 (very seeing-dependent)</td>
</tr>
<tr>
<td>Hubble Space Telescope: direct imaging</td>
<td>0.05&quot;</td>
<td>28-30</td>
</tr>
<tr>
<td>Hubble Space Telescope: roll deconvolution (Walter and Weigelt, 1985)</td>
<td>0.015&quot;</td>
<td>22</td>
</tr>
</tbody>
</table>

At visible wavelengths, an interferometer placed in space is only suffering from phase errors due to relative drifts of its individual telescopes. The sensitivity is greatly enhanced (Fig. 2–3–2; Roddier 1983).

In the infrared, the factor which limits the sensitivity is not so much the atmosphere than the background noise due to the thermal emission of the telescope. A space instrument brings significant gain only when it is cooled. This is illustrated on Fig. 2–3–3. Because cooling the whole optics is a costly and difficult task in space, it appears that ground based interferometry with large telescopes is an interesting step.
Expected signal-to-noise ratio for the object energy spatial spectrum, in a Michelson-type interferometric experiment, in the visible, as a function of the stellar magnitude. We assume a detector quantum efficiency of 0.2, an optical bandwidth of 60 nm, a fringe visibility of 0.5. We further assume that the telescope optics give diffraction-limited images when the aperture is reduced to an area of 1 m² and that a total observing time of 20 min is divided into subexposures of duration to allow for pointing errors (Roddier, 1983).

Comparison of ground and space interferometers sensitivities at infrared wavelengths. The ground-based 8m uses adaptive optics, the 1m space telescope is cooled at 100 K. Integration times are indicated (Léna, 1984)
Chapitre 3 : THE VLT CONCEPT AND ITS INTERFEROMETRY MODE

The initial VLT concept of a fixed linear array provides indeed an interferometric capability, which has triggered the present study. In order to keep flexibility, this study is placed in a broader frame: given 4 pupils, 8–10 m in diameter, what are the different configurations which provide high angular resolution? It is fairly obvious that an MMT configuration is straightforward: most of the difficulties and challenges lie in the coupling of separate telescopes. Whether these telescopes should be close-packed or far apart, in line or in a Y-shape, fixed or movable is part of the flexibility which has to be kept at this stage in order to search for the optimum configuration.

3.1 Single telescope constraints

3.1.1 Image formation

Interferometry first requires a high quality of imaging given by the individual telescopes.

Although the optical specification of the VLT unit telescopes has not yet been formally fixed, it is expected it will be something like:

80% of the geometrical energy within 0.1 arcsec at visible wavelengths (500 nm)

This specification is based on the New Technology Telescope specification which expects to reach diffraction limited performance for low spatial frequencies and 80% within 0.15 arcsec for high spatial frequencies (ripple).

Active optics

The active optics system envisaged is (following the NTT) a closed loop system for correction of low band pass errors with time frequencies lower than the time corresponding to integration of atmospheric seeing (Wilson 1982; Franz and Wilson 1982; Wilson 1983). The aim is to correct all low spatial frequency modes (those influencable by flexure of optics or tube) to the diffraction limit all the time. The limitation will then be non-correctable high frequency (ripple) effects left by the optician, on which a severe specification will be set.
The NTT optics test set-up (1 m thin mirror) has now proven the practical functioning of this system, so that the development for the VLT of a system based on similar principles is seen as straightforward.

The possible extension of the same system to the correction of the effect of wind buffets is under study.

**Guiding**

In the NTT the active optics system works with an image analyser which uses the offset-guide star **from time to time** by mirror switching between the image and the auto-guider. In the larger sense, the auto-guider is simply part of the active optics, correcting tracking errors in real time and also any small wavefront tilts arising from the active optics corrections. The NTT auto-guider will be a well-tested TV system under development (Tarenghi and Ziebell 1982) using a window about 4 x 4 arcsec and an integration time (currently) of about 1-2 s. Higher frequency tracking errors must therefore be < 0.1 arcsec.

**Adaptive optics**

A considerable gain in sensitivity of the interferometric mode is expected if atmospheric distortion of the images may be partially or totally corrected.

"Adaptive" optics is used here (following Woolf 1984) as a term referring to the correction of high temporal band errors caused by atmospheric turbulence.

Because of the limitations of the isoplanatic angle, a general solution for visible wavelengths will still seem quite utopic.

It has been recently proposed (Foy and Labeyrie 1985) that the fundamental problem of the availability of a reference star within the isoplanatic angle might be solved by shooting laser pulses through the observing telescope or an auxiliary telescope to a layer in the atmosphere about 80 km high which reemits by scattering or excitation of sodium atoms sufficient light for the telescope to pick it up as a point source, providing a reference point source with seeing information. If this could work in practice (and such laser ranging has been done for other purposes), it would represent a breakthrough in the problem of the isoplanatic angle.

Adaptive correction of image motion can be done by projecting the pupil onto a tiltable plane mirror by a Fabry lens of size equal to or larger than the
isoplanatic field. Higher spatial frequencies require real-time deformation of the mirror or of the optical thickness of a deformable plate. At high time frequencies, the problems are formidable because of the rate of image analysis (information rate) and correction even for a single isoplanatic field. However, Hardy's results (Hardy 1981), show it can be done if the reference source-isoplanatic angle problem could be solved.

The infrared case is much more favorable longward of 3 μm given the properties of atmospheric turbulence. The large IR isoplanatic field (up to 2–3 arc min at 20 μm) would allow to use an offset bright visible star to correct the central faint IR image (Léna et al. 1986). The number of elements to correct on a pupil is also manageable (a few tenths).

In any event, progress in adaptive optics, at least for IR, will be essential for the VLT since, assuming the active optics achieves its aims, the atmospheric seeing will be the major limitation. The gain for the S/N interferometry goes at least with the inverse cube of the PSF (Roddier 1975; Dainty 1975; Roddier and Léna 1984) so no other parameter can bring such an advantage.

3.1.2 Mechanical stability

It is assumed that each telescope provides an afocal beam to a central "beam combination-area", where appropriate motion compensates for Earth rotation. Optical path differences are introduced by the atmosphere on each beam. The temporal power spectrum of these fluctuations is seeing-dependant, and can be reasonably estimated on the basis of turbulence models, now well confirmed by speckle work. Let \( \tau(\nu) \) be the rms path fluctuation (in micrometer) at the frequency \( \nu \). At telescope exit, the afocal beam suffers from other perturbations: the wave front has additional path fluctuation with respect to the ground frame reference, induced by mirrors (primary, secondary or tertiary) vibration or by the mount motion, which have their own spectrum \( \tau^\text{M}(\nu) \). The system is atmosphere limited if the condition

\[
\tau^\text{M}(\nu) < \sup \left[ \frac{\lambda}{20}, \tau(\nu) \right]
\]

is fulfilled.
Figure 3-1-2

Expected standard deviation $\sigma$ of the optical path fluctuations during an exposure time $T$. Full line: lucky observer model. Broken line: average model.

Figure 3-1-2 (Roddier 1985) is a graph of $\tau(\nu)$, from which one may derive the mechanical stability requirement.

The mechanical vibration constraints have been discussed in detail by Roddier (Roddier 1985), relative to the optical path fluctuations induced by the atmosphere. In general, the time spectrum of mechanically induced fluctuations must lie below the time spectrum of atmospherically induced fluctuations. This means that the rms values of the mechanical fluctuations of the optical path
should lie below the lines of Fig. 3-1-2 [continuous line: "lucky observer" conditions according to Barletti (Barletti et al. 1976); dotted line: average conditions]. At high frequencies, the limit is a horizontal line corresponding to fluctuations of \( \leq 0.025 \mu m \) for visible wavelengths (\( \lambda/20 \)). For IR observations at \( \lambda = 2.2 \mu m \) corresponding to a frequency of about 1 Hz, amplitudes of the order of 0.1 \( \mu m \) would be acceptable.

Labeyrie (1985) has pointed out the increasing problem of mechanical vibrations with size of telescope. While the problem is simple with 25 cm telescopes, with 1.5 m apertures fluctuations of optical path of about 3 \( \mu m \) rms, with lifetimes over a few seconds, may result from the tracking and be a major problem in preventing fringe detection. Improvement of more than an order of magnitude would be necessary to meet the above tolerances in a passive system.

Labeyrie has also pointed out the specific danger of spider vibration in Cassegrain telescopes. This can be well illustrated by the case of the ESO NTT where the lowest eigenfrequency of the top unit is about 70 Hz. It is far more difficult to calculate the amplitude since the existing signal and the damping conditions must both be defined. However, it is very unlikely to be below 1 \( \mu m \) and may well be several \( \mu m \). Such amplitudes are completely insignificant for normal telescope use but would be fatal for interferometer in a passive system.

Clearly, therefore, the mechanical requirements, above all for vibration at high frequencies, are going to be far harder for telescopes intended for interferometry, even in the IR, than in the normal case. Indeed, they are so hard that it seems more reasonable to consider active compensation rather than the imposition of absolute tolerances. Labeyrie (1985) has considered possible systems for achieving this. Although difficult, this active control should anyway be easier than adaptive optics control of the atmosphere.

Labeyrie's "Boule" mounting may well offer advantages of mechanical stability if very smooth tracking can be achieved and the spider problem solved.

3.1.3 Thermal Background

The reduction of local emissivities to improve the sensitivity is not a specific requirement of interferometry. Therefore requirements established by the VLT infrared Working Group are fully applicable here.

Because of thermal radiation, detectors experience a photon noise in the
infrared that degrades their performance if its value surpasses the intrinsic noise of the detector. For this reason detectors are operated in a low temperature environment that generates a number of photons that can be neglected compared to the number generated by the atmosphere. The noise can be calculated by determining the mean square deviation in the rate of arrival of the photons at the detector. Using Planck's equation for the number of photons per second in a band of $d\nu$ Hz within an etendue of $S \Omega$:

$$N(\nu) = \frac{2\nu^2}{c^2} \cdot \frac{1}{(e^{\frac{h\nu}{kT}} - 1)} S \Omega d\nu$$

We find:

$$\sqrt{\langle N(\nu)^2 \rangle} = N(\nu)^{0.5} \left(1 + \frac{1}{(e^{\frac{h\nu}{kT}} - 1)}\right)^{0.5}$$

Assuming that the photons emitted by a medium of $T=300$ K and an "average" emission coefficient $\epsilon = 0.5$ arrive within $S \Omega = \lambda^2$ through an instrument with transmission $t = 0.1$ the noise can be calculated if $d\nu$ is known.

To guarantee spectral coherence over the interfering areas, $d\nu$ has to be chosen according the seeing conditions in the atmosphere:

$$d\nu = 2.7 \times 10^{14} \frac{r_0(\lambda)}{D^{5/6}}$$

with diameter of the telescope and $r_0(\lambda)$ Frieds parameter at 0.5 $\mu$ which is connected to the seeing at the wavelength. The formula holds in case motions of the image in the FOV are controlled, allowing fringe detection over even long baselines. The noise is given in Fig. 8.1-1-3-1 for two cases:

- $D = 1.5$ m with $r_0(\nu) = 20$ cm (seeing 0.5")
- $D = 8$ m with $r_0(\nu) = 5$ cm (seeing 2").

For reference purposes detector NEP values are drawn, converted to the same noise scale. Because the best detectors have intrinsic noise levels of ca $10^{-17}$ WHz$^{-0.5}$ it appears that photon noise limitations occur for wavelengths $\lambda > 3.5$ $\mu$. The situation changes if square arrays are applied such that one pixel fits the mean area of coherence in the pupil of the telescope. advantageous because
the visibilities are not averaged out on one detector. The thermal background is now diluted over the number of detector elements, which leads to a lower photon noise per element and a comparable lower NEP value. The number of coherence area's in the pupil is given Table 1-5-1.

\[ N_s = 1.09 \times 10^{-10} \left( \frac{D}{\lambda_{0}} \right)^{12/5} (\sigma \text{ in cm}^{-1}) \]

Assuming \( N_s > 1 \) one calculates the NEP values given in Fig. 3-1-3-2. If the intrinsic detector noise is about \( 10^{-17} \text{ WHz}^{-0.5} \) photon noise limitation occurs for \( \lambda > 5 \mu \text{m} \) with 8 m dishes and 2° seeing \((D/r(\lambda) = 160)\) and above \( \lambda > 3 \text{ m} \) with a 1.5 m dish and 0.5° seeing \((D/r_0(\nu) = 7.5)\). As a consequence, dilution of the radiation over N pixels allow to image the interference patterns without any loss above 3–5 μ. In the detector noise limited region, at shorter wavelengths, the S/N for fringe measurement drops, depending on \( D/r(\lambda) \) and the array format. A small number of pixels however will not resolve the coherence areas in case of less good seeing and the fringe contrast decreases.

For \( \lambda > 5 \mu \text{m} \) the number of pixels can be chosen to accomodate also rather bad seeing conditions. If an 8 m telescope is used in the K–window, 32x32 arrays are quite adequate for seeing up to 1.5° and in L more than 2° can be accepted, but the format should be weighted against the number of nights with a good seeing. At 5, 10 and 20 μ arrays of 16x16, 7x7 and 3x3 or larger can be applied for 8 m telescopes and a seeing better than 2° (Roddier and Léna 1984).
3. 2. **Movable telescopes**

The discussion on image reconstruction (Sec. 2) shows the unescapable conclusion that the u–v coverage has to be as dense as possible, and certainly two-dimensional. It is therefore worth investigating the possibility to move the telescopes.

Movable telescopes can have two significations:
- the telescopes could be brought to separate stations in order to vary the baselines. Once they are set, optical delays are compensated as discussed above, the telescopes remaining fixed during the time of the observation.
- the telescopes could be moved all the time during the observation, in order to avoid optical delay lines.

At first, the concept of moving 8–10 m telescopes may seem odd to optical astronomers, while radio astronomers, having gained a large amount of experience in such matters, would consider the concept worth of investigation (Citterio 1985).

For example, the VLA telescope weights 250 tons each, have a pointing accuracy of \(<15^\circ\) and a positioning accuracy of \(<1\) cm. The IRAM–Bures telescopes weights 100 tons, are pointed within \(\pm 1.2^\circ\) and have a positioning accuracy of \(1\) mm.

Concerning the movement of the telescopes two possible configurations are considered:
1–Point to point translation of the telescopes
2–Continuous movement of the telescopes following a predetermined law accuracy.

Conceptual solutions to these problems have been investigated following the approach to take hints from systems already working and proved on other field of applications and to extrapolate them to the interferometry case.

3. 2. 1 **Point to point translation of the telescopes**

In this configuration the telescopes are supposed to be translated on a certain number of hard standing points situated on a flat site. The number and the layout of these points must be defined in order to achieve the best u–v plane coverage. The optical beams from the telescopes are supposed to be combined
in a central station and optical delay lines must be provided for the continuous optical path equalisation.

The solution proposed for this approach is derived from a system which is used for the transport and precise positioning of large mechanical pieces to be machined by large CNC (Computer Numerical Control) machine tools. The system is composed of a platform floating on an air film which is towed in place by means of an external power drive: when in place the platform is precisely positioned by means of pins and precision pads that define respectively the lateral and the vertical position. The position accuracy is of the order of 10 micron.

For the translation of a telescope its base would be connected to a platform of sufficient dimension and stiffness, depending on the weight of the telescope a certain number of air cushions must be mounted under the platform. As an example, for a telescope of 300 tons one can estimate a platform of 50 tons: for a total weight of 350 tons it would be necessary, for instance, 16 aero caster load modules K48NHD working at 3/4 full load. The air flow would be ~ 264 L/sec at 2.2 bar. The total lifting area would be 16.5 square meters.

For every point where the telescope is supposed to be moved, an hard stading area with reference points and pads must be provided with the addition of a clamping mechanism to increase the stiffness of the fixation of the telescope to the ground. The hard standing area must be interconnected by cemented roads with suitable load capability and surface smoothers.

3.2.2 Continuous movement of the telescope

In this configuration the telescopes are supposed to be moved along straight radial with respect to the central station. The movement of the telescope is made following a predetermined law in such a way to have a continuous optical path equalisation. Due to the stringent requirements on the phase stability of the optical wavefront it is expected that an active phase control will be necessary to take care of small fluctuations of the movement.

The solution proposed for this configuration is also derived with reference to large CNC machine tools. For machining large mechanical pieces [e.g. 500 tons] these machines have a rototraversing table which moves in linear hydrostatic bearings. The movement is provided by backlash-free, high efficiency and stiffness hydrostatic worm and rack which in principle do not have
limitations on the length of the movement. The movement is very smooth and the positional accuracy is \(\approx 0.01\) mm. The maximum speed of translation is \(\approx 6 \text{ meter per minute}\). Large horizontal milling and boring machines that have a linear movement of 30 to 40 m have been already built.

In the case of VLT project the base of the telescope would be fixed to the hydrostatic saddle which will move on a linear hydrostatic bearing of a bed having the desired length. The point of major concern related to the hydrostatic solution is the need for a suitable tight protective cover of the bed top against dust, rain, snow, etc. An investigation is in progress to find a suitable solution to this problem. It is obvious that, as in the case of the point to point configuration, the solution with hydrostatic bed required a flat site. Despite some complexity with the realisation of the hydrostatic rail, this solution presents the best combination of rigidity–frictionless required for the interferometry with movable telescopes.

The investigation made shows that, from the pure mechanical point of view, these could be possible solutions to the problem of moving telescopes that should be analysed in detail by a feasibility study.

The cost-effectiveness of these features remains to be evaluated in the context of the overall VLT project.

**The platform concept**

To allow all of the desirable telescope-array configurations of interest for interferometry, but also for non-interferometric observing, a smooth platform on which telescopes can be displaced randomly would be ideal. It would also make delay lines unnecessary for interferometry since telescope motions during observation could be programmed to follow the elliptical tracks ensuring equal optical paths. No such possibility was exploited for large radio arrays since smooth platforms spanning many kilometers raise unsolved problems.

The scale of 100–300 m considered for the VLT is compatible with a square or circular platform. Air pads such as described previously can carry the telescopes on smooth concrete. Among the unsolved problems however are:

1. the moderate accuracy of motion achieved on air pads of this type;
2. traction mechanisms for bidirectional motion.

Concerning the traction problem, bidirectional configurations of linear motors could be investigated. As shown in Figure 3.2.1, an array of magnets immersed in the concrete platform can provide a checker-board pattern of
magnetic fields. Driving forces can be achieved by electric coils located on the telescope support.

Fig. 3-2-1.
Immersed electric coils for a continuous 2-D motion over a platform.

A more conventional approach would use crossed railway tracks (Labeyrie et al. 1984). It however appears less flexible for continuous motion in two dimensions.

3.3 Array configuration

Apart from operational aspect to be discussed below (Sec. 3.5), the optimum configuration of an interferometer has the following requirements, as imposed by image reconstruction procedures:

a) dense coverage of u-v plane, from zero frequency up;

b) two dimensional coverage

c) simultaneous use of telescopes for adequate phase reconstruction

d) maximum redundancy.

It is obvious that a 4-telescopes interferometer such as the VLT can not fulfill all these conditions in the optimal sense. Yet, some configuration are more adequate than others toward the goal of high angular imaging; it is these compromises which are envisaged in this section.

The original linear configuration

The QEW (Quasi East West) baseline is the less unfavorable orientation of
fixed telescopes. A minimum requirement for this configuration is to have at least one movable telescope on a N–S base, and the capability of beam combination with the three other telescopes.

**The fixy configuration**

It has been shown above that a wye configuration can provide a reasonable u–v coverage with fixed telescopes. The main drawback of this thin array solution is a poor coverage of the low spatial frequencies, and a poor phase reconstruction.

**The VARY configuration**

As soon as movable telescopes are available, the capabilities of the interferometric mode are greatly enhanced. With mobility along radial or Y–tracks, good aperture synthesis can be achieved. A full two-dimensional motion over a large platform is even better. A critical point is the coverage of low spatial frequencies. At first, it was estimated that each 8–10 m telescopes could not be brought at a distance closer than 25 m axis-to-axis in order to avoid mechanical interference. This distance criterion is somewhat dependant of the type of mount to be selected, as well as of the f/ratio of the primary: spherical mounts may allow to provide compact coverage of the u–v plane at low frequencies. Another attractive alternative is to place two mirrors on the same mount, the two other telescopes being movable independently and further apart.

These various configuration are summarized on Fig. 3–3–1.

**Combining unequal pupils**

As shown above (Sec. 2–1), the use of a smaller pupil in a pair gives practically, at almost every optical or infrared wavelength, a signal-to-noise ratio which is equivalent to diaphragm the large VLT dish to the size of the smaller pupil. At first glance, it therefore appears that the use of a smaller telescope in the VLT interferometric configuration would be very useful to improve the u–v coverage at the cost of using the VLT large dishes in a non-optimal way, while the combination of VLT dishes between themselves takes full advantage of their large pupil.
Baseline extension

This is a compromise between maximum resolution and good u–v coverage. In other words, the array must not be too thin. Moreover, the site capability limits the array extension. It can be shown that the signal-to-noise ratio drops very rapidly when the object is fully resolved. The resolution of the array should therefore be adapted to the size of the objects to be studied.

At infrared wavelength (2–20 \( \mu m \)), the size of typical objects of interest is listed in Table 3.3.1.

A maximum baseline of 150 m seems therefore appropriate (6 marc sec at 5 \( \mu m \)), site compatible and providing not too thin an array. Whatever the bases, positionnal accuracy and good metrology are essential.

It should be noted that the telescope do not need to be exactly within an horizontal plane.
TABLE 3-3-1

Resolutions required at infrared wavelengths

<table>
<thead>
<tr>
<th>Phenomena</th>
<th>Resolution (milli arc sec) for an object at 1 Kpc</th>
</tr>
</thead>
<tbody>
<tr>
<td>Accretions flows</td>
<td>10,000 - 100</td>
</tr>
<tr>
<td>Bipolar flows</td>
<td></td>
</tr>
<tr>
<td>H$_2$O masers clumps</td>
<td>10</td>
</tr>
<tr>
<td>Fragmentation, Discs, Multiple cold stars</td>
<td></td>
</tr>
<tr>
<td>Stellar winds</td>
<td></td>
</tr>
<tr>
<td>Magnetic structure</td>
<td>1000 - 1</td>
</tr>
<tr>
<td>Protoplanetary evolution, discs, haloes</td>
<td>1000 - 1</td>
</tr>
<tr>
<td>Ultra compact HII regions</td>
<td>100</td>
</tr>
<tr>
<td>Dust condensation processes</td>
<td></td>
</tr>
<tr>
<td>Maser pumping processes (OH, NH$_3$, SiO...)</td>
<td>10,000 - 10</td>
</tr>
<tr>
<td>Evolved stars: dust, chemistry, cycles (Miras)</td>
<td>1000 - 10</td>
</tr>
<tr>
<td>Galactic nuclei: dust, energy production, time evolution, jets</td>
<td>1000 - 10</td>
</tr>
<tr>
<td>Gravitational lenses</td>
<td>1000 - ?</td>
</tr>
</tbody>
</table>

**MMT type**

The relative merits of the "MMT" approach compared with the "Array" approach (limited number of individual telescopes in each case, about 4) have been much debated. The array has, in its very nature, more flexibility for which there are certain prices to be paid. For example, the field that can be transmitted between two telescopes will reduce with increasing separation and will tend to be less with an array than in an MMT. However, the use of the individual telescopes is free and their field unrestricted by the requirements of a single mount. Thus, even without interferometry (phasing of the telescopes), the choice of an array of a small number of telescopes can well be justified, either in the mode of incoherent optical combination or in that of electronic (post-detection) combination.

However, the coherent use of the array telescopes in interferometry
undoubtedly adds substantially to the attraction of the array solution. Even with fixed telescopes in one line, the total baseline of the order of 150 m is about 8 x more than that available in the fixed, compact scheme of an MMT. Although the aperture is inevitably more dilute, adequate u-v plane coverage is possible with 4 telescopes. Step-wise movement over a line, or better a plane, seems technically feasible and adds greatly to the possibilities. Continuous movement would be technically much more ambitious but would produce the ultimate as an interferometer facility.

Beckers (1984) stresses about the same conclusion about type O array (monolithic MMT): easy cophasing of telescopes at submicron precision by an internal metering system for sources of any brightness, compact PSF but more limited angular resolution.

3. 4 Beam combination

The present VLT concept allows the use of pairs of telescopes as long baseline interferometers. A non-redundant distribution will provide up to 6 baselines. With independently mounted telescopes the entrance pupils are not coplanar as the telescopes points off the zenith angle. In order to guarantee phasing of the separate beams, the Lagrange invariants for the individual telescopes in the array must equal and additionally the overall Lagrange invariant of the array must be conserved. This requires optical path length and pupil corrections. To achieve the necessary stability the coherent beam combination path has to be actively stabilized by adaptive optical methods. Techniques analog to laser beam stabilization systems could be applied.

3. 4.1 Delay lines and pupil correctors

By setting the VLT to the interferometric mode the beams from two telescopes can be combined with an equal path length in the interferometric laboratory. There are four options for the path length correction:
(1) use of a trombone as an optical delay line in the combining beams of the individual telescopes and a stationary combining systems or,
(2) use a moving combining system to ensure equal path length.
(3) move the telescopes themselves.
(4) use monomode optical fibers (could be very efficient in the infrared).

Pointing off the zenith angle with two telescopes results in a pupil foreshortening which decreases the synthetic pupil diameter. To maintain the geometrical scaling of the lateral pupil geometry the pupil separation at the combining optics has to be compensated. Also the longitudinal pupil position must be compensated due to the change of the relative location of the optical elements in the system. With variable field optics the images of the pupils can be continuously transformed to the same plane in the combining optics.

3.4.2 Active control of the path correction

The operation of the VLT in the interferometric mode requires a continuous path correction. The required precision and stability is only achievable with closed loop active optical control systems.

In case of continuously movable telescopes the complexity of beam combination and the control of the path corrections would be increased drastically.

3.4.3 Pupil rotation

The linear array concept with telescopes in alt–az mount has an image and pupil rotation as a consequence. This drawback can be overcome by the use of an image de–rotator or an interferometric instrumentation mounted on synchronously rotating stage. The first solution requires transmissive elements or a large number of reflecting surfaces. Therefore the second solution seems to be a preferable one.

3.4.4 Optical efficiency

The optical efficiency of the coherent operation mode depends strongly on the availability of high reflectivity coatings. Recent developments indicate progress for these high efficiency coatings for IR wavelengths.
3.4.5. Polarization

Elliptical vibrations are produced upon reflection of light beams from tilted mirrors. This can reduce the visibility of fringes. The effect is quite predictable on aluminium mirrors, and its wavelength dependence has no sharp variations. High-efficiency mirrors using metal and dielectric coatings, as well as those made of fully dielectric coatings, tend to have more complex variations, including some more sensitive to wavelength and angle of incidence. This can in principle be computed, and coating types with acceptable neutrality in some spectral range could possibly be developed.

However, there are good reasons to observe separately both linear polarisations. This can provide additional astrophysical information, particularly on magnetic stars and those with all types of mirrors. It is nevertheless likely that, due to the large number of mirrors which are necessary for the coherent beam combination and the various reflection angles, polarization sensitive measurements will be difficult.

3.4.6 Field of view

A preliminary design of the coherent beam combination system shows that a field of 15 arcsec seems to be possible.

3.5 Interferometer operation

To form an acceptable image of an astronomical source, there should be at least 12 to 15 tracks in the u-v plane, each track lasting a full night per source, including calibrations. For comparison, the VLA, with 27 dishes, gets 351 tracks in 8 h; Westerbork, with 14 dishes, gets 91 tracks (some redundant) in 12 h; the Cambridge 5-km telescope, with 8 dishes, made many maps with 16 tracks, obtained in 12 h. The final image has roughly as many pixels as the square of the number of tracks. A map with only 3 or 4 tracks will be poor, firstly because there are not enough points to do a Fourier transform, with proper weighting to suppress sidelobes, and secondly because the final image will have
dimensions of only 3x3 or 4x4 pixels.

This simple fact means that an array of 4 fixed telescopes has inherent limitation right from the start, so all efforts should be made to design array telescopes which can be displaced on rails, or to include more telescopes in the array.

However, the need to collect at least 12 to 15 baselines for interesting images, also illustrates another possible difficulty with the VLT interferometric mode. If the interferometry uses only two telescopes per night, while the other telescopes are scheduled for non-interferometric projects, then 12 to 15 nights will be needed to make a map of just one object. As the good season for interferometry may be only three months per year, in which excellent seeing, clear sky, low turbulence and low winds occur 50% of the time, then such a mode of operation allows complete imaging only 3 sources per year, or possibly up to 6-8 if an outstanding site is selected. Moreover, since telescopes are used by pairs, phase closure, or equivalent speckle masking, is not available for phase determination.

Greater productivity is obtained if information is collected on all baselines simultaneously (6 tracks per source per night, for a 4-telescope array). If the array is dedicated to interferometry, none of the telescopes being used for single telescope projects, about 3 months per year, 50% of the time, then 20 sources per year could be mapped.

Another operational consequence of the interferometry is that to collect the 12 to 15 tracks required, the telescopes would probably not make major moves every day. As with radio arrays with a small number of elements, the array would be left in the same configuration for a period or two or three weeks, during which time it would observe many sources, proposed by many different observing groups. After several weeks, the configuration would be changed, and the same sources would be observed again. After several configuration changes, typically lasting a few months, the observers could start to put their maps together. Since it would take a few months to do an observing program, it is most unlikely that observers would come to Chile or even to ESO headquarters during their programs. Instead, the use of an array will force ESO to provide standard instrumentation for interferometry, and to have observatory personnel carry out the programs, as with most radio interferometers.

The only way to avoid this, and to give astronomers a "hand-on" feeling of
participation in the observing, is to have a dedicated array with enough
telescopes to synthesize a map in one night—six to eight telescopes for example.
In this case, it would be possible for astronomers to travel to the observatory, as
for normal observing.

3.6 Site constraints

In addition to the qualities expected from a good astronomical site, interferometry has specific requirements.

Seeing quality

High angular resolution interferometric or adaptive techniques put more stringent conditions on seeing quality than other astronomical observations. These conditions are discussed in an ESO workshop (Roddier 1983). We summarize them here.

First of all the S/N ratio for interferometric observations varies as the inverse second to fourth power of the seeing disk size giving high weight to sites which, at least occasionally, provide exceptional seeing conditions. Stability of seeing conditions is important for accurate calibration of the atmospheric transfer function.

The S/N ratio depends also upon the so-called wavefront boiling time which is related to wind velocities in the atmosphere. Sites with low speed should be preferred. Turbulence near the tropopause can be detrimental because it is often associated with high wind velocities. High altitude turbulence also reduces the size of the isoplanatic patch within which high angular information can be restored. A good site for interferometry should have little high altitude turbulence.
Stellar scintillation is directly related to high altitude turbulence (Loos and Hogge 1979, Roddier 1983) and its measurements is highly recommended in the site testing campaign. Sites showing little stellar scintillation should be preferred.

The outer scale of turbulence specifies the rms phase excursion between telescopes, and should be as small as possible. Very few measurements are yet available.

Seismicity

The origin of the seismic background noise is to be related to different kinds of sources (Citterio 1984).

At long periods (T>20 sec) the earth noise is mostly related to atmospheric and oceanic sources, with the addition of the noise caused by the free oscillation of the earth excited by deep energetic earthquakes.

At short periods more sources must be added to the atmospheric and oceanic one: mostly anthropic sources (such as highways, railways, factories and human activities in general) and, in some areas, endogeneous sources (permanent volcanic noise). The intensity spectrum of the seismic background noise is strongly dependent on the specific sites.

As an example the high frequency activity recorded at some Italian sites shows the following values:

- Periods: 1 to 2 sec; 0.06 to 0.08 sec.
- Amplitude: 0.02 to 0 micron
- Max tilt: 0.006 arcsec.

From these data one can say that in particular conditions the microseismic activity could affect the stability of the phase of the fringes in the visible and in the near infrared. Specific considerations can only be made after direct measurements of the microseismic activity at the candidate sites. The measurements should be distributed through the year in order to monitor the most significant seasonal conditions.

Available area

A surface of 150 x 150 m seems the minimum requirement. If a LINEAR array is selected, with the addition of a single, movable N-S telescope, the N-S baseline could be compromised on a moderate slope (< 20°).
3.7 Submillimetric interferometry with the VLT

Spatial interferometry with the VLT can be done with coherent- and incoherent detectors. For coherent detectors the state of the art is represented by SIS-mixers and Schottky-Diode mixers in quasi-optical mounts. The system noise temperatures that are achieved at present range from 400 K (DSB) at 350 GHz to 2300 K (DSB) at 810 GHz. These heterodyne techniques are improving fast and better performance can be expected for the near future. Interferometry with heterodyne detectors is analogous to radio interferometry as the phase is preserved in the intermediate frequency. At the University of Berkeley, Townes and coworkers are progressing in the construction of an heterodyne interferometer working at 10.6 μm. The most sensitive incoherent detectors in the submm range are $^3$He-cooled bolometers. They are approaching the photon background noise limit at bandwidths matched to the atmospheric windows. It is likely that interferometry with incoherent detectors will be most useful for continuum work, whereas coherent detectors will be more sensitive for line work. The detailed tradeoffs still have to be calculated.
Chapitre 4 : HIGH RESOLUTION IMAGING WITH A SINGLE TELESCOPE

Diffraction limited imaging with a single 8–10 m telescope provides 10 to 20 times less angular resolution, but is nevertheless a very appealing goal, which has little consequence on cost and of management of a VLT.

4.1 Visible wavelengths

The 8m-telescopes of the ESO VLT will be extremely attractive for high-resolution speckle work since they can yield an angular resolution of about 0.01" in the visible. The limiting magnitude will be very seeing dependent. In nights of very good seeing the limiting magnitude will be about 20^{m}. Objects of 17th magnitude were already observed at 1" seeing, with about 30 min. observing time and with relatively simple speckle cameras.

Various speckle methods are available. The list below describes some of these methods and the information which can be reconstructed by these methods:

Labeyrie’s speckle interferometry (Labeyrie 1970):
  - diffraction-limited autocorrelations.
Knox–Thompson method (Knox and Thompson 1974):
  - diffraction-limited images.
Speckle masking (Weigelt and Wirnitzer 1983):
  - diffraction-limited images
Speckle spectroscopy (Stork and Weigelt 1984):
  - diffraction-limited objective prism spectra
Differential speckle interferometry (Beckers 1982).

Speckle observations with high signal-to-noise ratio will be possible at 20th magnitude if
- a site is selected which has 0.5" seeing or better in many nights (flexible scheduling)
- dome seeing is avoided completely
- the mirror quality is very good (better than 0.2°)
- observations are performed in many spectral channels simultaneously (e.g. by dichroic beam splitters)
- one telescope is used for the object and a second telescope is used to measure the speckle transfer function.

4.2 Pupil-plane interferometry

Pupil-plane interferometry has been suggested as a substitute for speckle interferometry as early as 1972 (Breckinridge 1972; Kenknight 1972). Since there is no attenuation of the amplitude of the object Fourier components in the pupil plane, this technique avoids calibration errors inherent to speckle methods. Moreover averages can be taken over a smaller number of frames since turbulence does not contribute to the variance of the amplitude estimate. For faint objects both methods are photon-noise limited and the S/N ratios have been shown to be ultimately the same (Dainty and Greenaway 1979). For bright objects the S/N ratio per frame saturates to unity in a speckle experiment whereas it keeps increasing as the square root of the number of photons in a pupil plane experiment (Goodman and Belsher 1977). In this case pupil-plane interferometry can be far superior to speckle interferometry.

Stellar diameter measurements have been demonstrated by Currie et al. (1974) using a wavefront folding interferometer (Koster prism). Roddier and Roddier (1983) obtain two-dimensional visibility maps using a rotation shearing interferometer. By processing only 71 frames they have been able to map the visibility function of Alpha Orionis with a S/N ratio superior to that of speckle measurements using several thousand frames, showing evidence for a dust envelope close to the stellar disk (Roddier and Roddier 1983). An image of this envelope has been reconstructed from these data (Roddier and Roddier 1985; Roddier et al. 1984). New data obtained with the same interferometer show evidence for two faint stellar companions to Alpha Orionis (Karovska 1984; Roddier et al. 1985). These results clearly demonstrate the efficiency of pupil-plane interferometry.

Whereas the transfer function for speckle interferometry is the autocorrelation of the pupil transmission function, the transfer function for pupil
plane interferometry depends upon the wavefront shear used. Because of the circular symmetry of telescope pupils, rotational shears are the most convenient since they make use of all available photons. In this case the transfer function has the same shape as the pupil transmission function. Using a 180 degree rotational shear with a full aperture allows to observe simultaneously every spatial frequency up to the telescope frequency cut-off. Because most telescopes have a centrally obscured pupil, additional observations are necessary to recover the low spatial frequency information. An important requirement for this type of observations is that the telescope should have the smallest possible central obstruction.

As for speckle interferometry, this technique requires that every effort should be made to minimize refractive index inhomogeneities along the optical path inside the dome. Experience indicated that any source of heat close to the beam (generally electronics at prime focus and warm wires along spider arms) have a detrimental effect.

4.3 *Infrared wavelengths*

Until now high resolution IR imaging has been restricted to one-dimensional (1-D) speckle interferometry because IR arrays have not been implemented in adequate systems. However this is to be done in the very near future and should be a qualitative improvement (access to regularly sampled 2-D visibilities and therefore easy-to-reconstruct images) but not a significative one in terms of S/N per single resolution element.

Practice has led to the following technical requirements in addition to usual IR constraints:

- excellent tracking to not truncate the image wings despite a field of a few arcsec; good and fast blind pointing for source switching;
- telescope mean MTF i) very stationary ii) very good as aberrations severely degrade the visibility S/N (Dyck et al. 1984); this point makes active optics quite promising if it can keep the mean MTF still stationary;
- temperature gradients cancelled or inversed: apart from bad seeing, the possibility of image degradation by these gradients is the major because it is not stationary, not linear and usually not under control.
The resolution (in the Rayleigh’s sense) of an 8m dish being 54 marcsec at K, 93 marcsec at L and 118 marcsec at M. the effective resolution, based on simple model fitting, can reach 20, 31 and 39 marcsec respectively at K, L and M.

For a 1-D system, the sensitivity is not improved by sizing the collector dish up. on the contrary 0.4 mag is lost when the detector is background–limited. However comparing 4 m and 8 m dishes at same resolution shows a gain of 0.2 mag (at M) to 1.0 mag (at M). For a 2-D system (CID receptor) a gain of 0.75 mag (prop. to D) is obtained. Moreover, quasi diffraction limited operation, reached with excellent seeing at M, is greatly favored by large apertures. Theoretical magnitudes, attained within 1 mag by present systems, are 7.7 at K, 5.8 at L depending on limitation type. 4.5 at M with a seeing of 1.3 arcsec at V (and 1 h integration, 200 ms exposure time, S/N of 10, emissivity of 0.15).

Their \( r_0 \)-dependence (proportionnal to \( r_0^2 \) at 1-D, to \( r_0 \) at 2-D) recalls that the seeing quality has also a huge quantitative impact.

Present techniques (with a computer of \( \approx 0.1 \) Mflops) would already allow to implement:

- autoguiding with the source signal itself;
- on-line best image sorting (Mariotti et al. 1983) and phase computation by Knox and Thompson algorithms (Cheilli et al. 1983);
- on-line preliminary image restoration (ibid.).

A speckle IR system should benefit from the technical effort put on large telescope project and make use of these potential facilities at only price of software work.

The technique itself is still evolving towards more efficient although more complicated configurations which will be fully operational at the time of VLT’s first light, like IR shearing interferometer and differential speckle interferometer (Beckers 1982, Dyck et al. 1983). These “second generation” instruments bring improvements of limiting magnitudes in their own working domain.
Chapitre 5: THE INTERFEROMETRY MODE: ITS IMPLICATIONS ON THE VLT PROJECT.

A design study

The current analysis is by no means a design or a feasibility study. Based on numerous theoretical papers and on limited experience from actual observation, it gives a firm base to undertake such a detailed design study. The present analysis has identified a few areas which would be the major items of such a design study.

(*) - overall telescope stability
(*) - movable telescopes
(*) - adaptive optics (visible and especially infrared)
(*) - active control of path length differences
- "beam combining area": optics and detection
- software for data acquisition
- image reconstruction by interpolation-techniques.
- minimum mechanical/optical complexity.

Although they are all important, only a few items (*) can be considered to lay on a critical path.

It is also obvious that essential knowledge is expected from the existing interferometers or from the interferometric operation of the MMT in Arizona.

About cost

It is not possible at this stage to detail what an interferometric mode would lead in terms of costs for the VLT project. This analysis was carried with the assumption that a 10% cost increase was the maximum acceptable ceiling for this. Only a design study can substantiate this number.

Obvious costs are the ones directly tied to interferometry, such as the provision for beam recombination areas, movable telescopes (if any). Other costs are due to exceptional stability requirements (thermal or mechanical), to active control loops for path length stability, to combining optics, to software development for data analysis.
About focal plane instrumentation

It may appear desirable that the beam-recombining "focal package" be interchangeable, as are spectrographs and photometers on conventional telescopes. Following the construction of a "number-one" focal package by institutes, more sophisticated instruments will probably be developed and made available to the community.

On the other hand, the complexity of this package and the relatively small number of European experts for such instrumentation should probably lead to the formation of a consortium of laboratories, well integrated and in close connection with the VLT project group itself.

About operation

The design study would lead to an estimate of the technical staff requested to run the interferometric mode of the VLT. Beside the normal telescope staff, experience with radio arrays indicates that 10 to 15 full-time people could be necessary to maintain the specific adaptive optics and active loop systems, the metrology, the beam-combining instrumentation and the associated software. This technical group, in close connection with the interferometry consortium, could provide a centralized service to the European community of astronomers.

The supporting consortium of European scientists may take the shape of a centralized laboratory, at least for the initial phase of development, a number of over 30-40 scientists and engineers would be necessary.
A number of developments are necessary to ensure a proper analysis of the VLT interferometry option. It is therefore recommended that appropriate action be taken either within the VLT project group, or at European level, or at national levels, for these subjects to be investigated. 
- In order to test the full range of problems affecting interferometric use of the VLT (stringent limits on mechanical fluctuations due to tracking and secondary mirror support), techniques for displacing optical telescopes on rails, wind and thermal effects, best methods of beam combining, delay lines, operational aspects, etc., it would be desirable to operate a prototype two-telescope interferometer carrying mirrors in the 1.5 to 2 m class, on an acceptable mountain site. This interferometer could be made fully operational on a short (2–3 years) time scale and capitalize all experience gained with the various existing interferometers (I2T and GI2T at CERGA, interferometer at Erlangen, US instruments and current studies).
- Adaptive optics developments should be vigorously initiated and pursued. Its feasibility on faint visible and infrared objects must be evaluated.
- Image reconstruction algorithm should be developed, with emphasis on the specific condition encountered in optical interferometry (phase gradient measurement). This work could be jointly carried with the European Space Agency, since many aspects of image reconstruction will be similar in ground-based and space-borne interferometry.
- Mono mode optical filter.
- Continuation and development of diffraction limited work, at visible and infrared wavelength, on 4 m-class telescopes (for example, high resolution testing station at the 4.2 m Herschel telescope Nasmyth focus).
Optical interferometry offers to explore completely a new field, imaging objects at 1 to 10 mas resolution at visible and infrared wavelengths. It will for the first time bring optical imaging close to VLBI resolution (which reaches 0.3 mas in H$_2$O lines and will attain 0.1 mas at millimetric wavelengths). Space mission shall not be able to compete with such resolution before the year 2005-2010, and the VLT interferometer is probably the best method to prepare these missions on a sound scientific basis.

The preliminary analysis reported here has indeed shown that an array of large telescopes (> 3 m), fully dedicated to interferometry and optimized for it, is undoubtedly the optimum solution. It also shows that, at the cost of some compromises, a VLT interferometric mode appears feasible and productive.

Although a MMT configuration has interesting high resolution capabilities, an array brings a factor of 10 in resolution, with other advantages such as flexibility of use, or wind loading reduction. The interferometric mode of a VLT array would be an obvious complement to other large telescopes projects such as the NNTT in the United States. Joint programs between the NNTT and the VLT would allow to continuously cover the $u$-$v$ spatial frequency plane from 0 up to 1000 arcsec$^{-1}$.

It is quite certain that the longer the wavelength, the easier the use of an interferometric VLT. Deciding the short end cut-off will therefore not be an easy task. It may be that the ultimate compromise, involving cost and feasibility, will place this boundary around 2–3 $\mu$m, but such a decision should be delayed as late as possible.

Many astronomers will inquire whether this is at all realistic. After all, only a few interferometers have now reached the scientific production stage. On the other hand, radio astronomers have already explored and solved many of the problems which may seem unsolvable to optical astronomers. Although limited in scope this report has probably covered all the really critical points of the interferometry mode: basically the mechanical stability and the adaptive optics could well be evaluated by a thorough design study. The operation of the VLT interferometric mode may also be considered as a threat to the normal and conventional use of optical telescopes. Since it appears that the VLT will, in any
case, be a telescope of unusual operationnal configuration. the innovation there may remain within acceptable limits, while leading to exceptionnal discoveries.
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Fig. 2-2-2. u-v coverage (top) and point spread function (bottom) of different array configurations. Angular units are in milliarcsecond per micrometer of wavelength or, as equivalent, in arcsecond per millimeter of wavelength. u-v maps ordinates are in meters (actual telescope positions). Two declinations are studied (-65° and -15°) for three configurations detailed in Fig. 2-1-1 (QEW 100, FIXY and VARY).
a) Model source, having four components of different intensities.

b) Uncleaned maps. The source is placed at $-65^\circ$ or $-15^\circ$ and mapped with the three array configurations (QEW 100, FIXY or VARY). Two rms errors have been postulated ($\sigma_p = 0^\circ$, $60^\circ$).
Fig. 2-2-4. Cleaned maps of the same four-component source, for declination -65° and -15°, with three array configurations (QEW 100, FIXY, VARY). Two rms phase errors have been postulated ($\sigma_\phi = 0^\circ, 60^\circ$).