


Radio Astronomy

Lecture 5

The Radio Telescope

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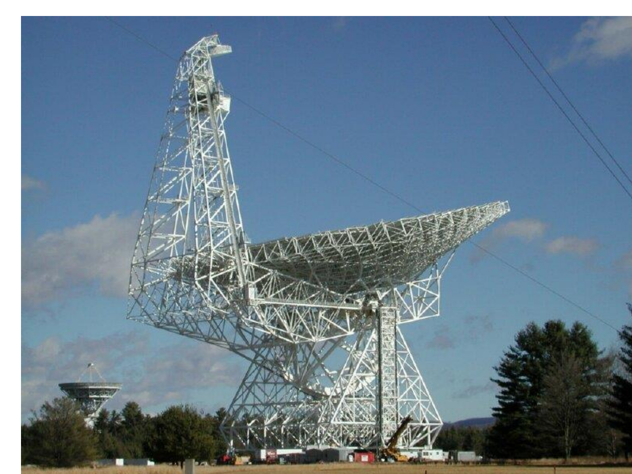


Lecture Outline

- 1st hour:
  Basic terms in radio astronomy - Flux density, intensity, brightness temperature
  Tools - Fourier Transforms, decibels etc.
  Radio telescope systems - antennas, reflectors, performance

- 2nd hour:
  Receivers – performance, types, sensitivity

Basic terms and formulae

- Properties of "black-body radiation" (you should all be familiar with this!)
  - functional form is called the "Planck function":

$$B_\nu(T) = \frac{2h\nu^3/c^2}{e^{h\nu/kT} - 1} \qquad (1)$$

Units of spectral energy density are Watts per Hz per sq. metre per steradian

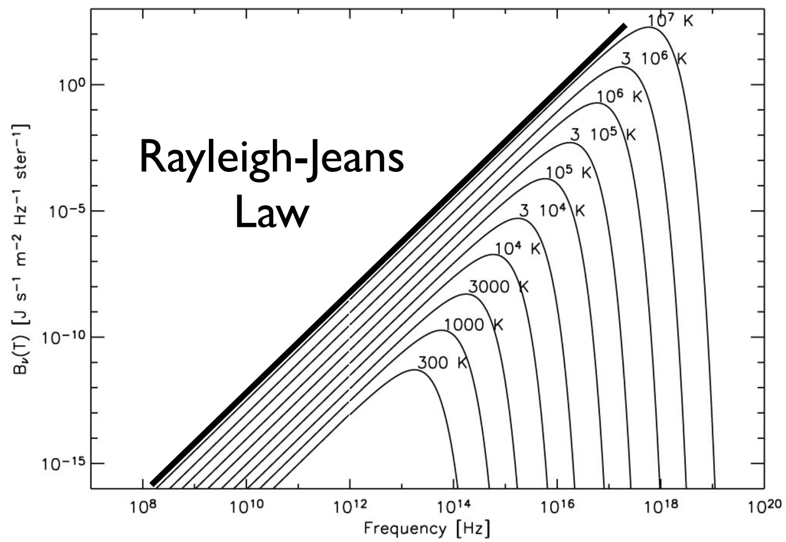
Radio photons are pretty wimpy: $h\nu/kT << 1$ $\quad e^{h\nu/kT} \sim 1 + h\nu/kT + ...$

$$\Longrightarrow \; B_\nu(T) = 2kT\nu^2/c^2 \qquad (2)$$

- Eqn(2) is known as the Raleigh-Jeans law i.e. at low frequencies the intensity increases with the square of the frequency.

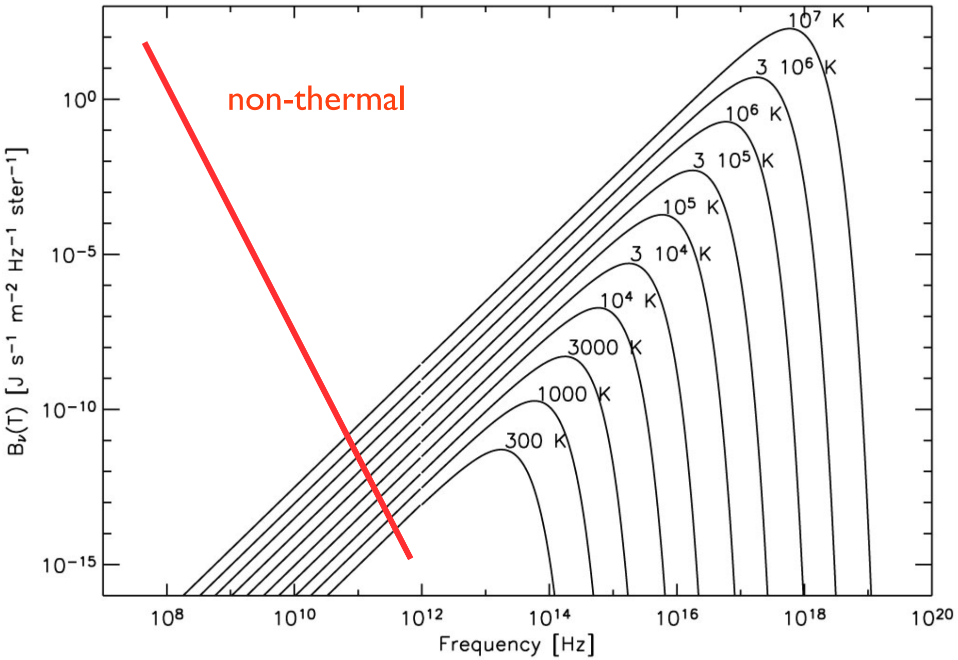
  Note that the R-J law holds all the way through the radio regime for any reasonable temperature

  k is boltzmann's constant = 1.38E-23 $m^2\,kg\,s^{-2}\,K^{-1}$



In the Raleigh-Jeans limit, a black body has a temperature given as (2): $T = B_\nu(T)c^2/2k\nu^2$

Radiation mechanisms in radio astronomy are often non-thermal (Reber's discovery):



But this does not stop radio astronomers also talking about the "brightness temperature" of a source: i.e. the equivalent or effective temperature that a blackbody would need to have in order to be that bright!

$$T_b = B_\nu(T)c^2/2k\nu^2$$

e.g. for Cyg-A, high resolution (small solid angle) VLBI observations measure T_b ~ 100E6 Kelvin - this is not a physical temperature but a measure of the energy density of the electrons and magnetic fields that generate radio emission via non-thermal emission mechanisms (synchrotron).

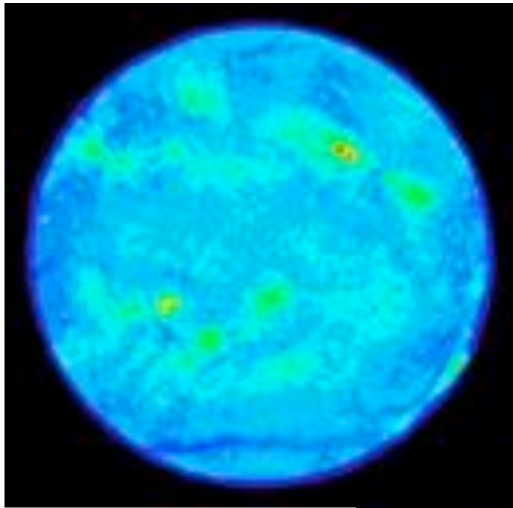
More examples of brightness temperatures:

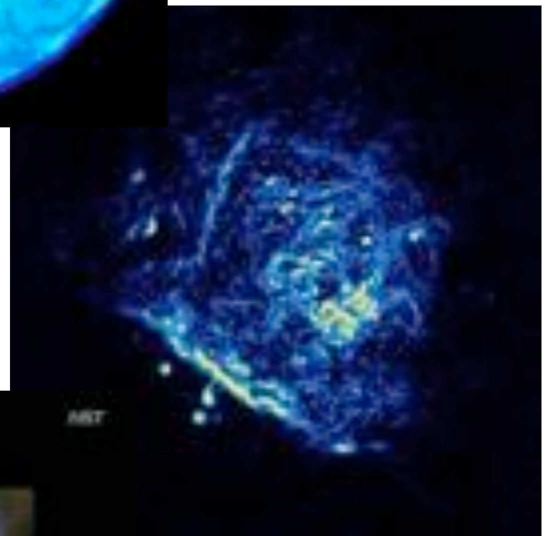
"Blank" sky ~ 2.73 K (big bang BB radiation)

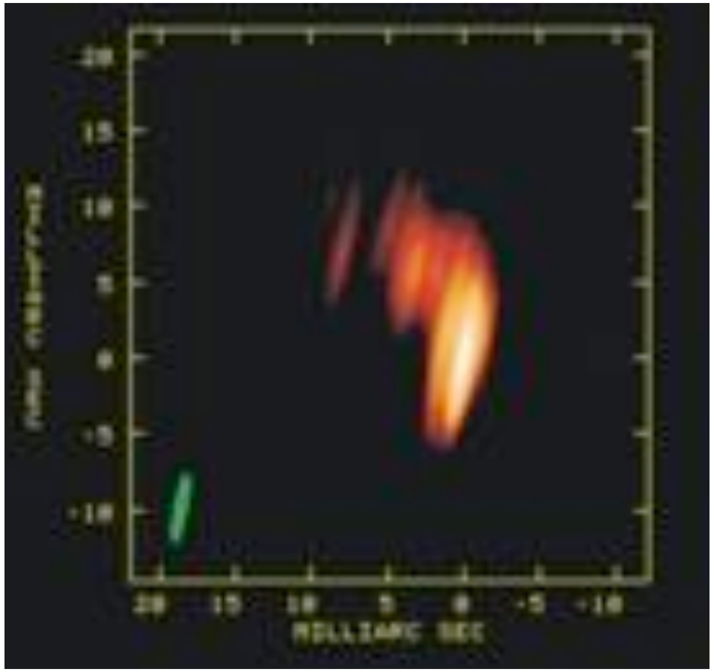
Sun at 300 MHz, 500000 K

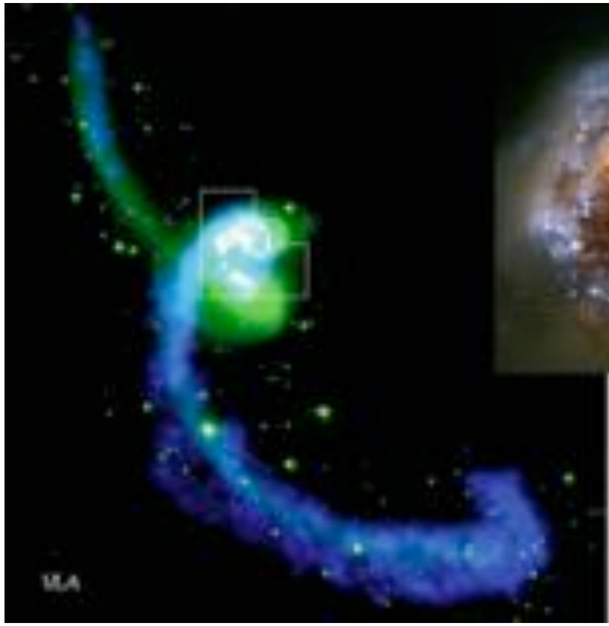
Orion Nebula at 300 GHz ~ 10-100 K ("warm" molecular clouds)

Quasars at 5 GHz ~ 10^{12} K (synchrotron)









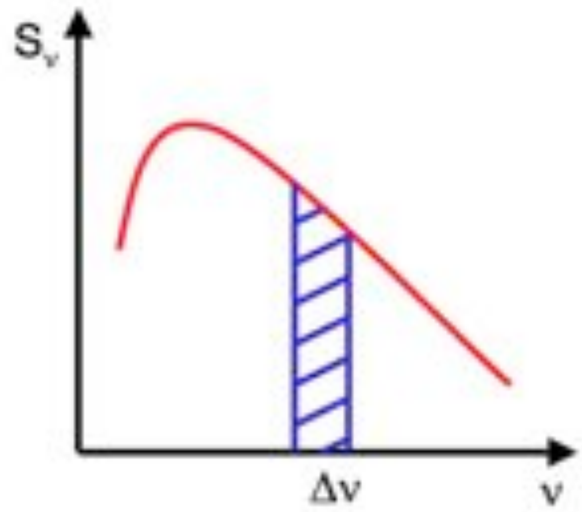
Radio telescopes/astronomers most often measure the (Spectral) "Flux Density" of a source.

The Flux density (S), is the power received (P) within a certain frequency band (dv), via a certain effective collecting area (A) with efficiency η :

$$S = 2\frac{P}{\eta A d\nu}$$

(Watts m^{-2} Hz^{-1}) [3]

The factor of 2 above is because the measurement of power is traditionally only made via one polarisation channel, and it is assumed that the other hand will contribute the same amount of power.



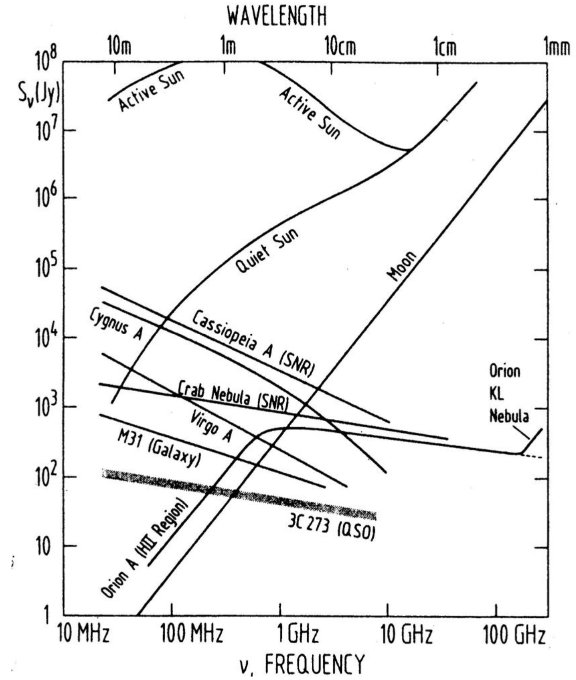
The unit of Flux density is the Jansky (Jy): 1 Jy = 10^{-26} Watts m^{-2} Hz^{-1}. Typical units include milliJansky (mJy) and microJy (uJy). NanoJy requires SKA!

Since the Flux density is proportional to the distance, d^{-2}, between the observer-source, it is not an intrinsic property of the source i.e. it does not reflect the real source luminosity e.g quiet sun has a flux density of 1000 Jy at 6cm.

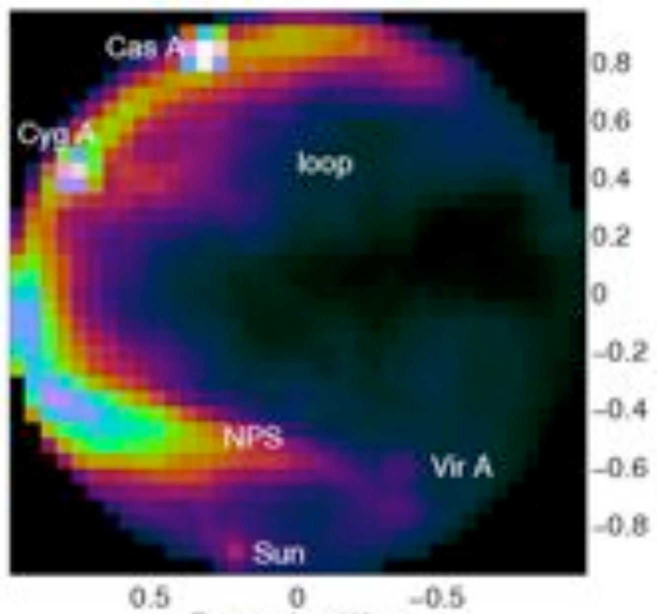
Surface brightness is the Flux Density per unit solid angle (Watts m^{-2} Hz^{-1} Sr^{-1})

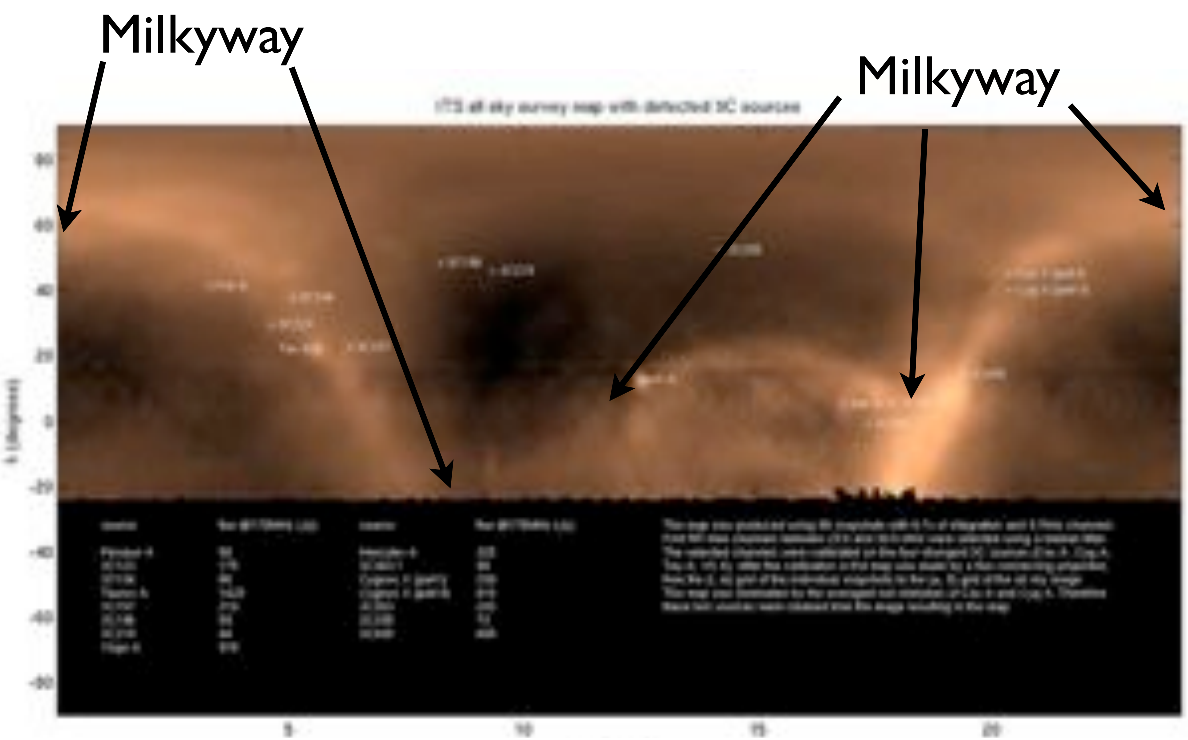
The dominant sources seen in the radio sky are the Sun, supernova remnants, radio galaxies, the Milky Way. The quiet Sun has a typical flux density of 10^5 Jy while the next strongest sources are the radio galaxy Cygnus-A (Cyg-A) and the supernova remnant Cassiopeia-A (Cas-A), both of which have flux densities of 10E4 Jy.

e.g. sources as bright as 1 Jy are relatively rare. For a Westerbork antenna D ~ 25 metres; efficiency ~ 0.7 (==> effective collecting Area ~ 340 sq metres); observing bandwidth 20 MHz. A source of 1 Jy produces a signal of only ~ 7E-17 Watts!



LOFAR images of whole sky (Wijnholds et al.):





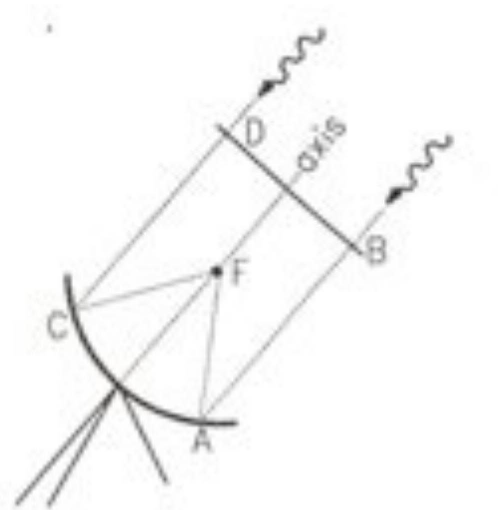
Radio Telescopes (Antennas)

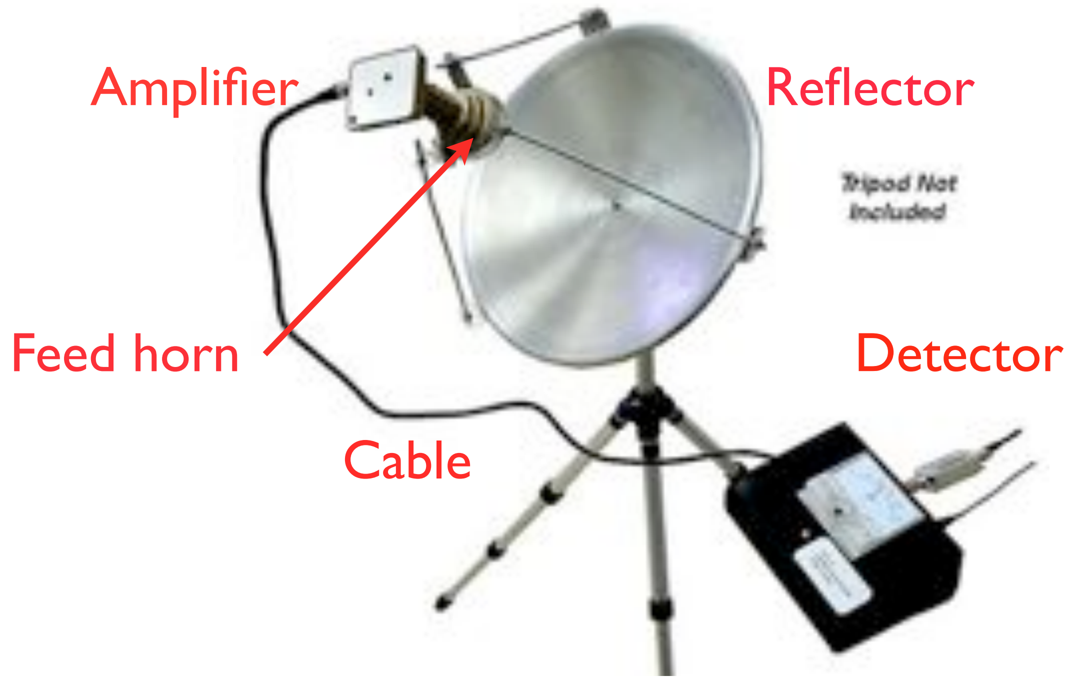
Radio photons are too wimpy to do very much - we cannot usually detect individual photons

- e.g. optical photons of 600 nanometre => 2 eV or 20000 Kelvin (hv/kT)
- e.g. radio photons of 1 metre => 0.000001 eV or 0.012 Kelvin

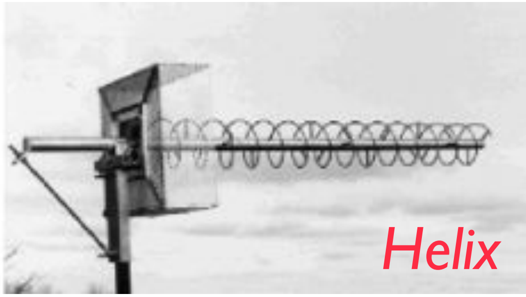
➡ Photon counting in the radio is not usually an option, we must think classically in terms of measuring the source electric field etc.

i.e. the best we can do is for the incoming EM-wave electric field oscillations to induce voltage oscillations in a conductor.

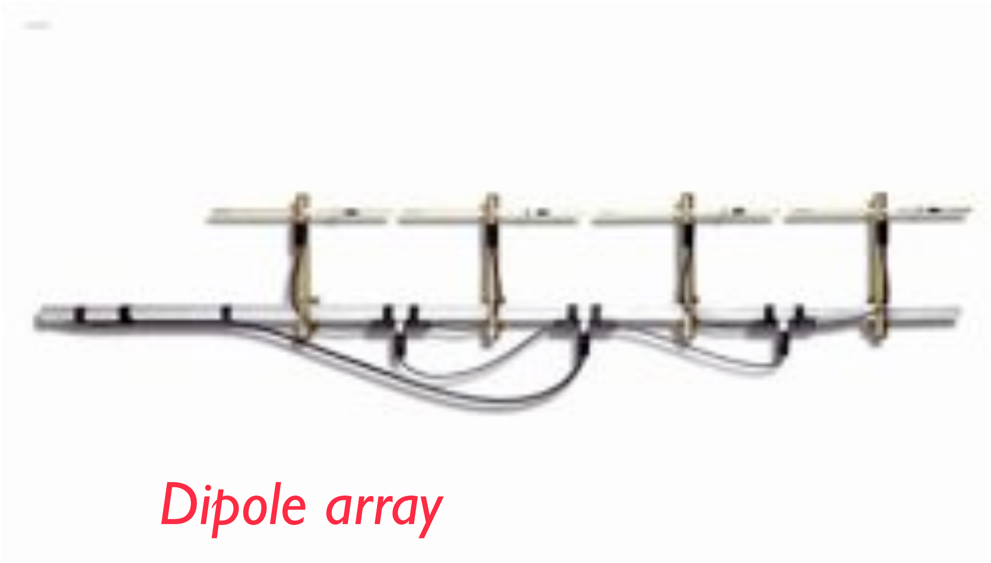




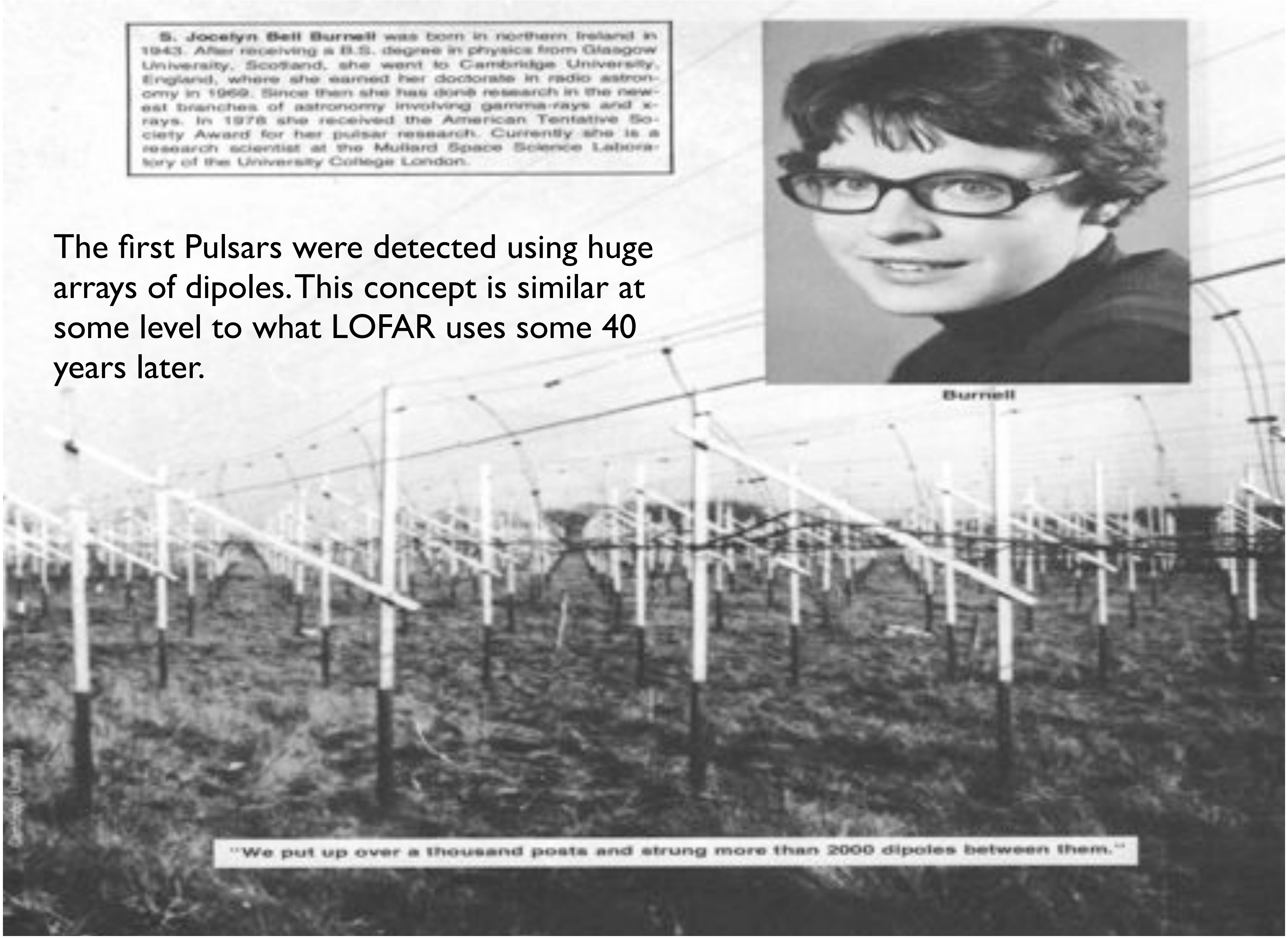
Effective area can also be increased by (for example) using a Yagi antenna. Parasitic antennas direct the wave towards the dipole. Log-periodic antennas are suitable as receptors of broad-band signals.



The gain of a single dipole can be greatly improved by combining together the output of many dipoles arranged in an array (right):



The first Pulsars were detected using huge arrays of dipoles. This concept is similar at some level to what LOFAR uses some 40 years later.



LOFAR Low-Band antenna use droopy (inverted vee) dipoles:

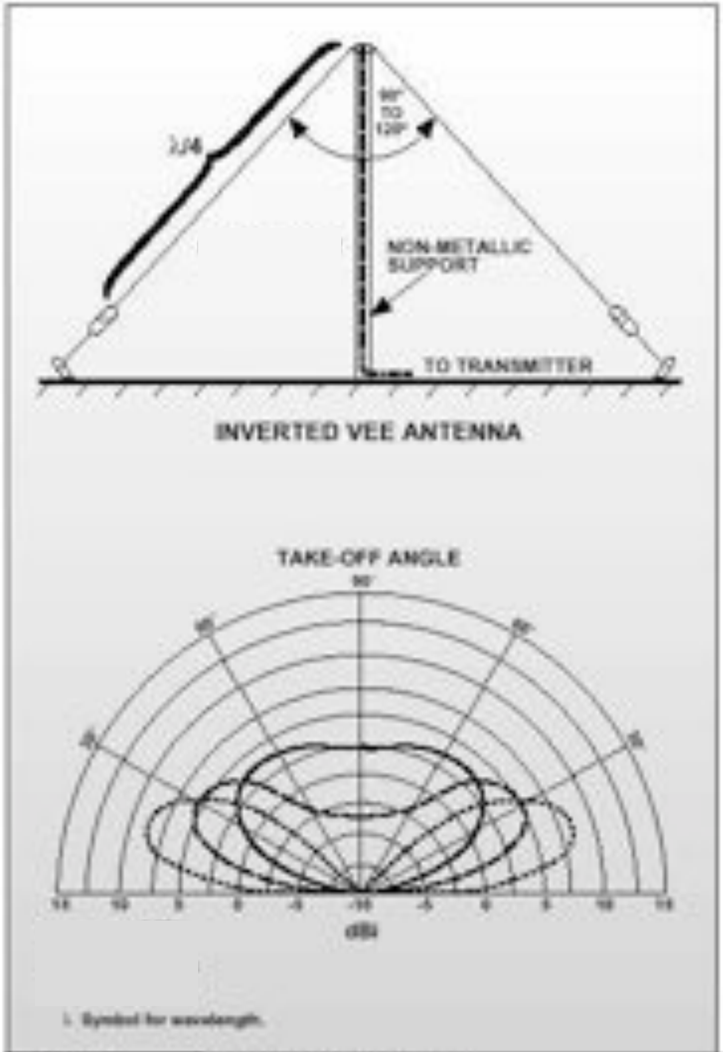
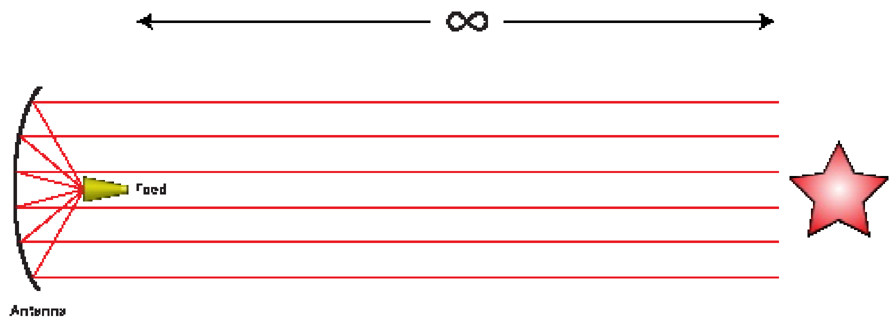


Figure 3-7. *Inverted Vee Antenna*

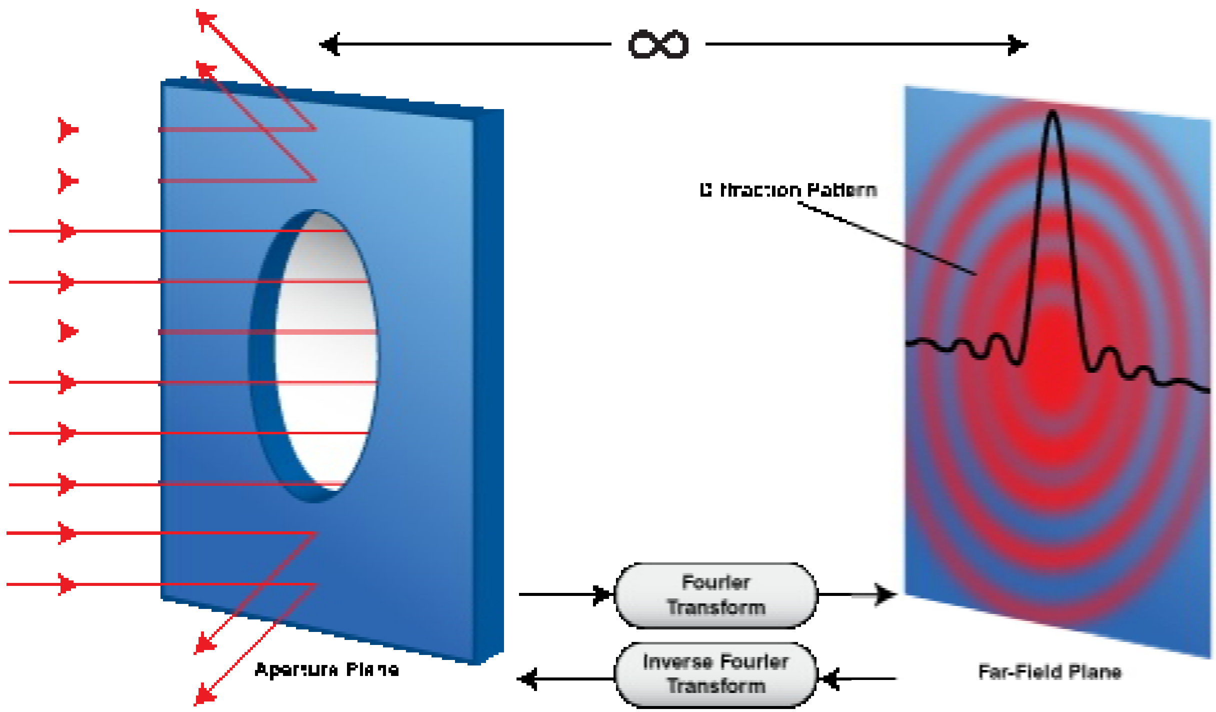
When considering the properties of an radio astronomy antenna its often useful to think in terms of it transmitting (rather than receiving).

Two theorems...

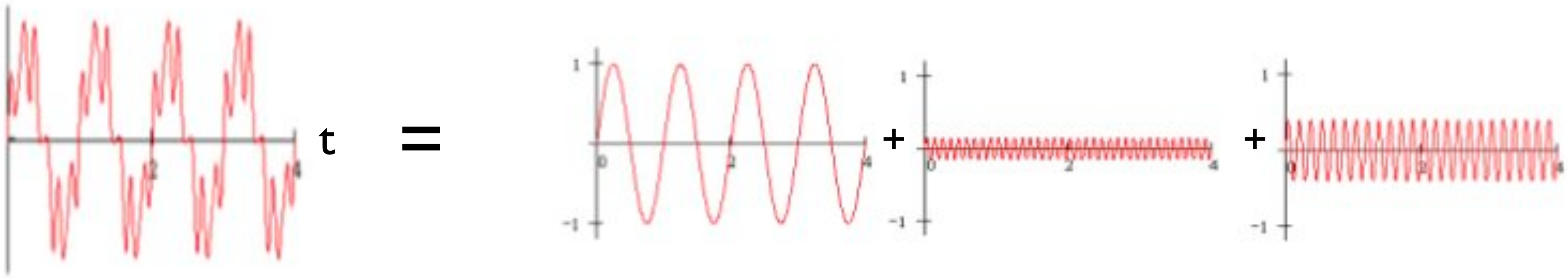
- Reciprocity theorem: Performance of an antenna when collecting radiation from a point source at infinity may be studied by considering its properties as a *transmitter*:



- Far-field pattern (the antenna's "beam") is the Fourier Transform of aperture plane electric field distribution:

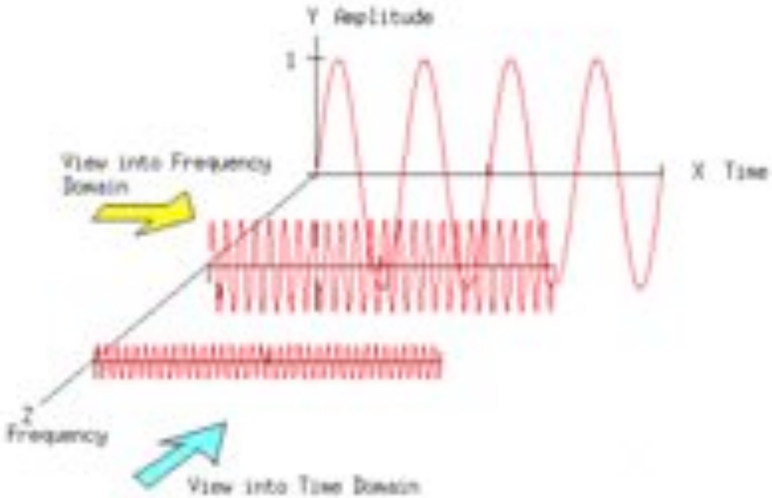


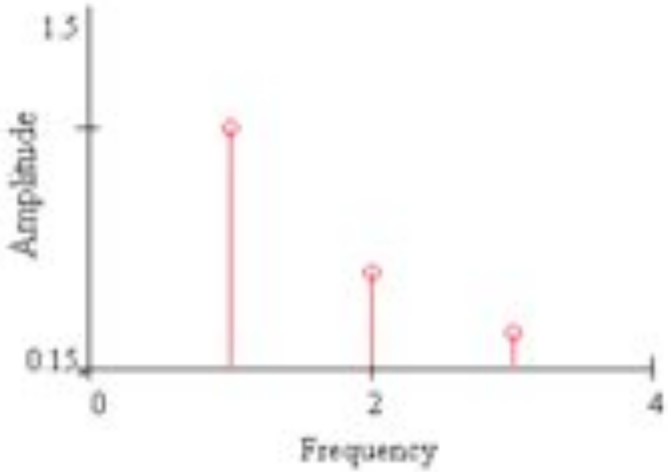
ASIDE: The use of Fourier transforms in radio astronomy is ubiquitous. In the 19th Century, Baron Fourier noticed that its possible to create some very complicated signals by summing up very simple sine and cosine waves of various amplitudes and frequency:



Complex signal *3 simple sine waves added together*

We can also view the same signals in the frequency domain (z-axis):





Most of you will be familiar with FTs from the time domain to the frequency domain:

$$F(\nu) = \int_{-\infty}^{\infty} f(t) e^{-2\pi i \nu t} dt$$ is the Fourier transform of *f(t)*

$$f(t) = \int_{-\infty}^{\infty} F(\nu) e^{2\pi i \nu t} d\nu$$ is the inverse FT of *F(ν)*

Basically, the idea is that any periodic function can be expressed as the sum of a series of sines and cosines of varying amplitude and phase. In other words, f(t) can be built up from the spectral distribution F(v) which is the power at frequency v.

A good and complete reference is “The Fourier Transform and its applications”, Ronald Bracewell.

The fourier transform of a function e.g. f is often denoted as $\mathscr{F}$(f) and the inverse is $\mathscr{F}^{-1}$ (f).

a function which is *wide* in one domain is *narrow* in the other, and vice-versa:

If $\mathscr{F}$ f(x) = F(s) some basic and general properties of FTs include:

1. $\mathscr{F}$ f(ax) = 1/a F(s/a) - *scaling property of FTs*

2. $\mathscr{F}$ $f(x-x_0) = F(s)e^{i2pi\, x_o s}$ - *shifting property of FTs*

3. Suppose that g(x) = f(x) * h(x), then:

G(s) = F(s) H(s) - *convolution theorem of FTs*

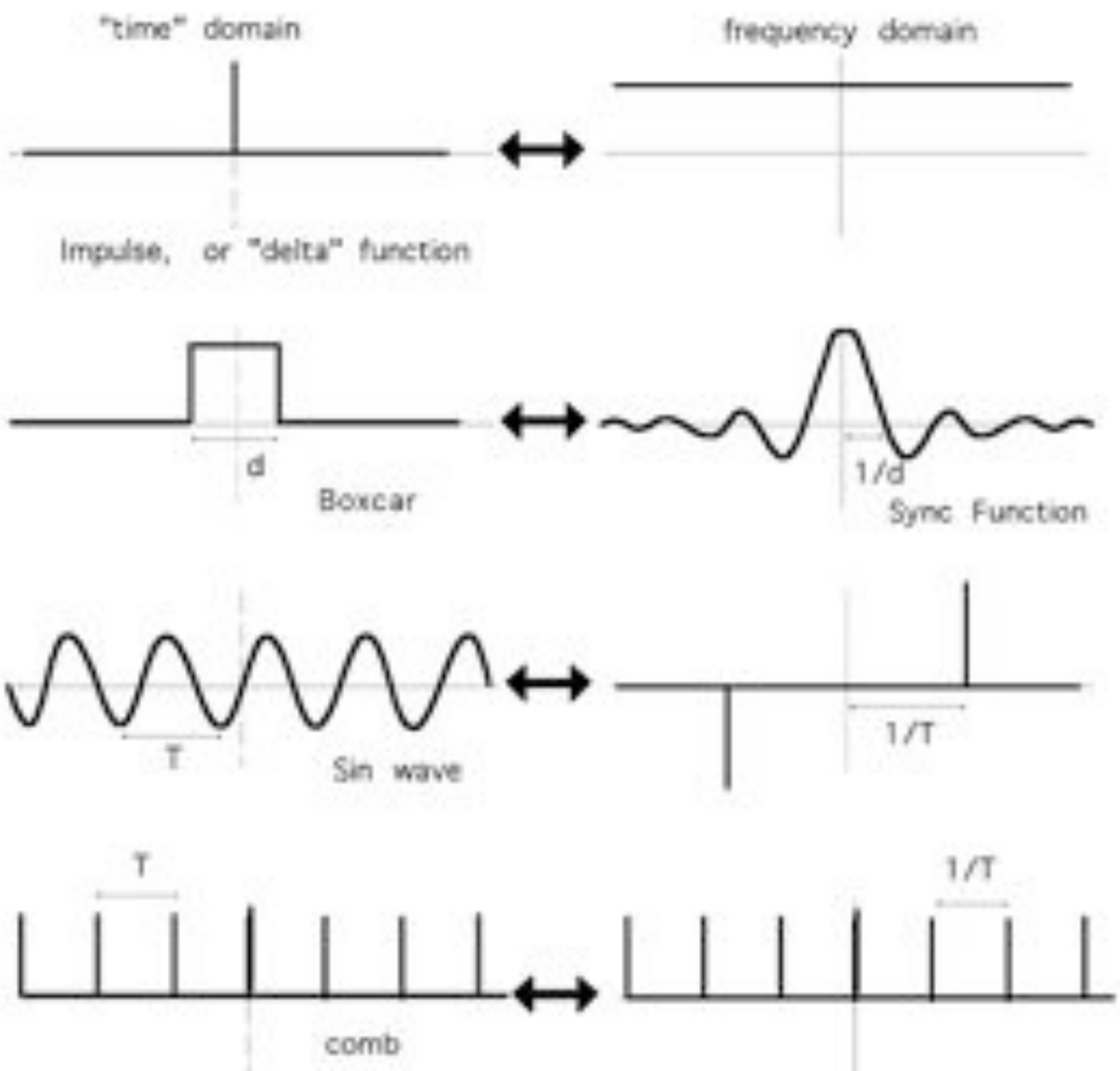
4. Consider the correlation of 2 signals:

$$h(x) = \int_{-\infty}^{\infty} f(u)\, g(x+u)\, du$$

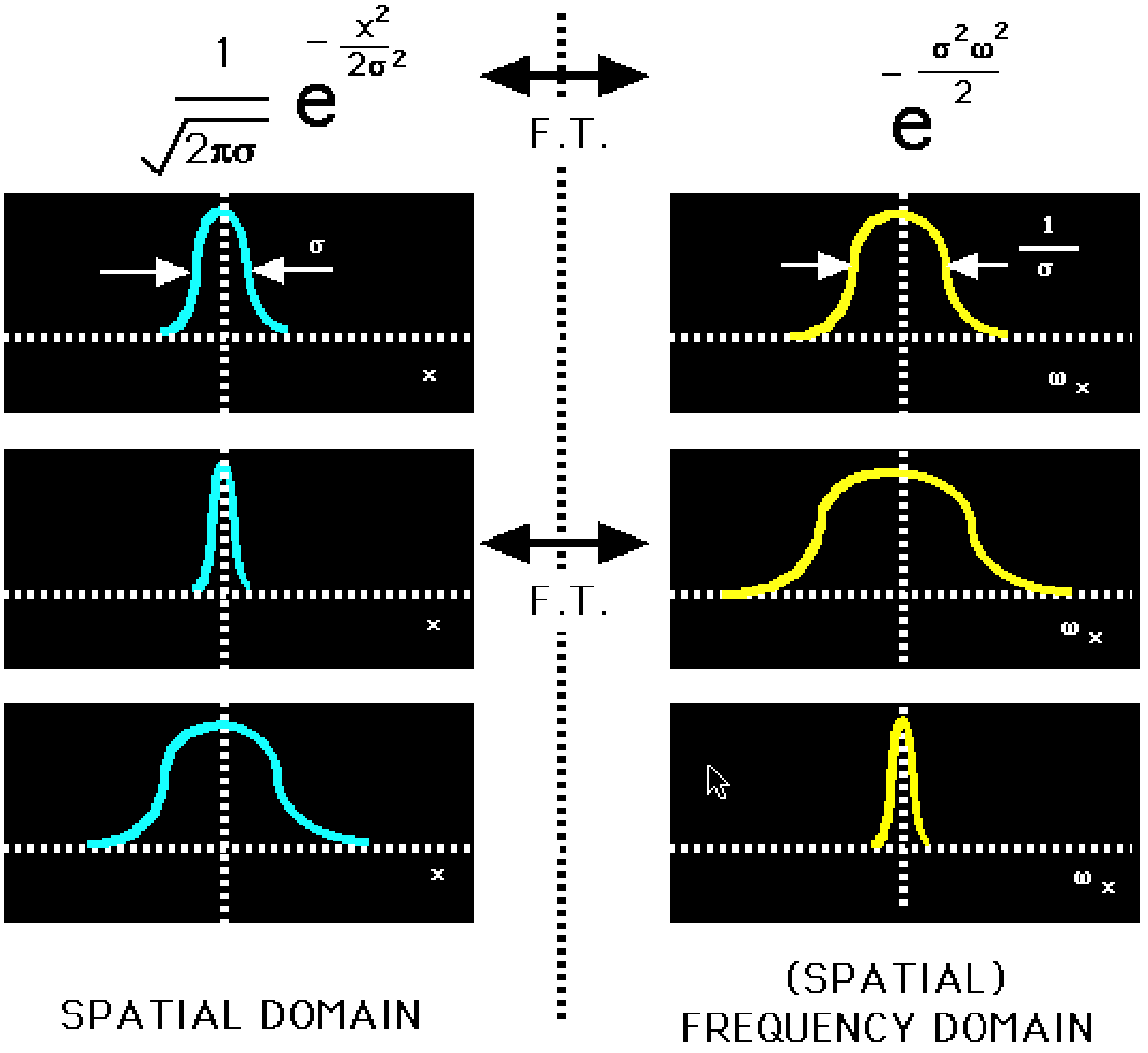
then $\mathscr{F}$ f(x) = F(s) $G^*(s)$ - *Correlation theorem.*

We will make a lot of use of (3) and (4) above. Note that in (4) if the functions f and g are the same, the integral is known as the autocorrelation function.

Some useful FT functions you need to know:



An easy one to remember is that the FT of Gaussian is another Gaussian:



A general thing about FTs to keep in mind, is that a function which is *wide* in one domain is *narrow* in the other, and vice-versa (see above).

In parabolic telescopes (see below) incoming EM-wave electric field oscillations induce voltage oscillations at the *antenna focus*, in a device called a *feed*.

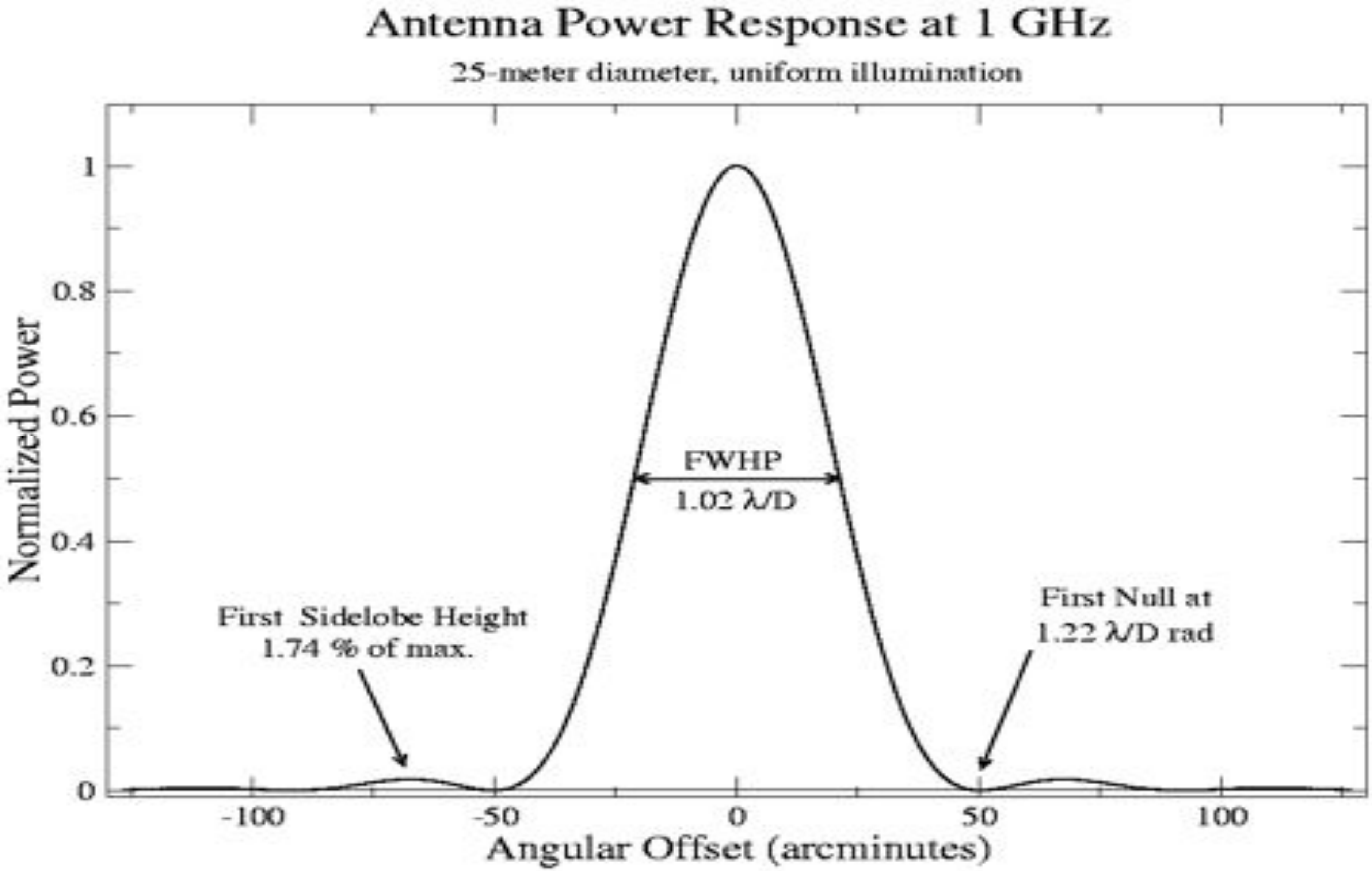
Radio sources are so far away the incoming signals can be assumed to be plane waves.

At cm and mm wavelengths, parabolic collectors are usually optimal for focusing the incoming plane-waves at the focus - where the feed is placed.

alt-az mount

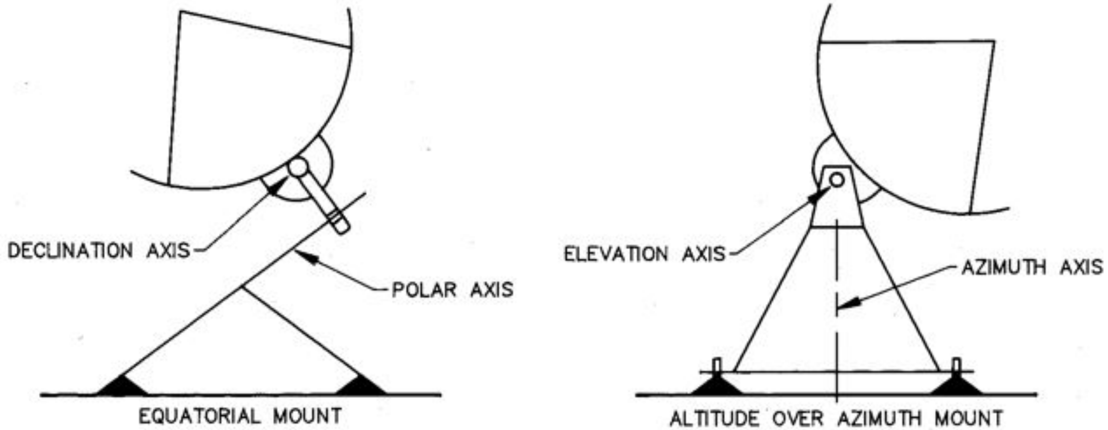
Equitorial (polar) mount

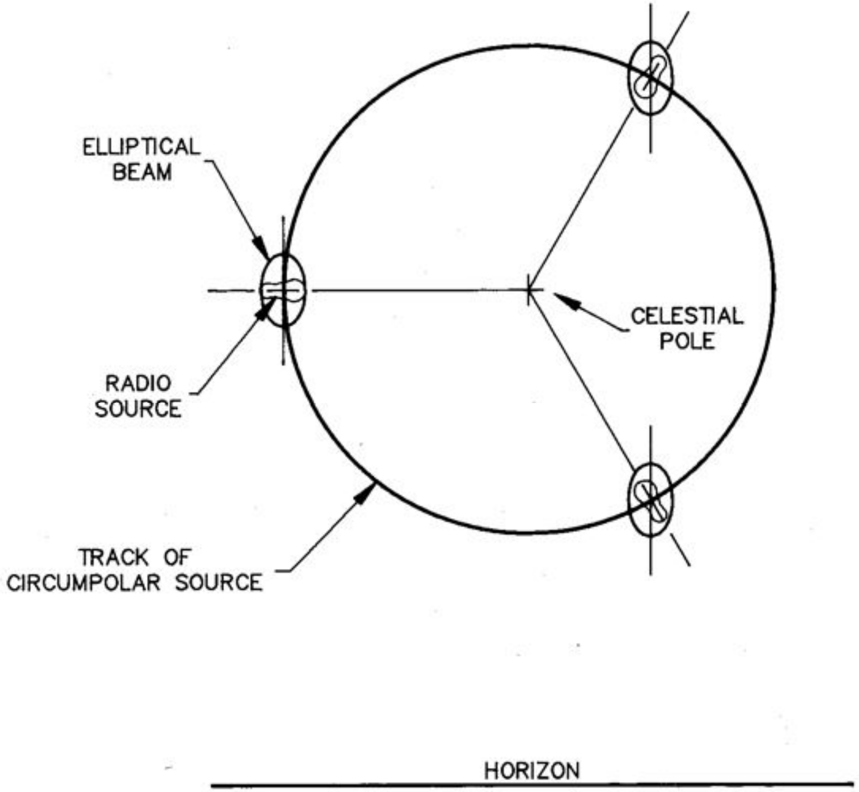
The response of a uniformly illuminated circular parabolic antenna of 25-metre diameter, at a frequency of 1 GHz.



Different types of mount:

Modern antennas are mostly alt-az because they are cheaper to build.



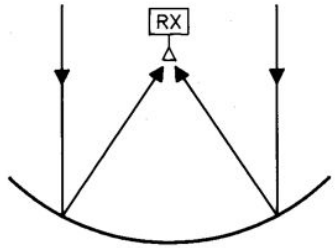


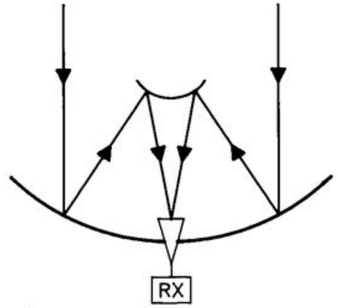
Disadvantage of alt-az telescopes is that the orientation of the telescope beam changes as the source moves across the sky.

For polarisation measurements, this must first be corrected for (usually in software).

Reflector types

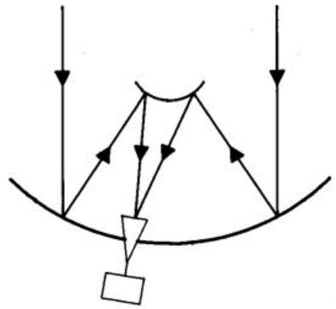
Prime Focus

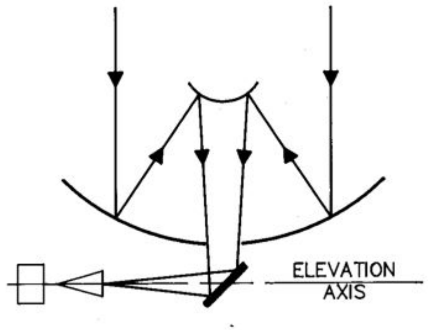




Cassegrain Focus

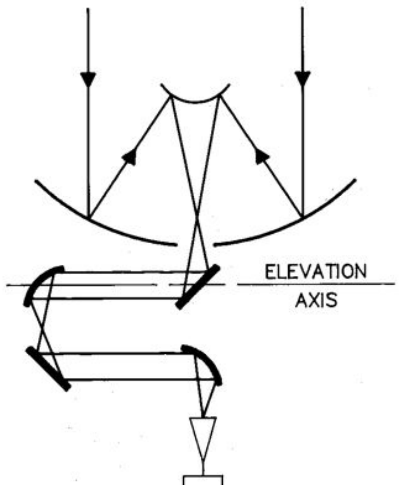
Offset Cassegrain

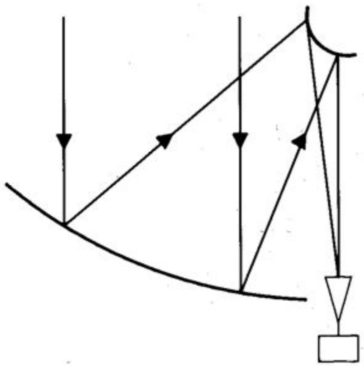




Naysmith

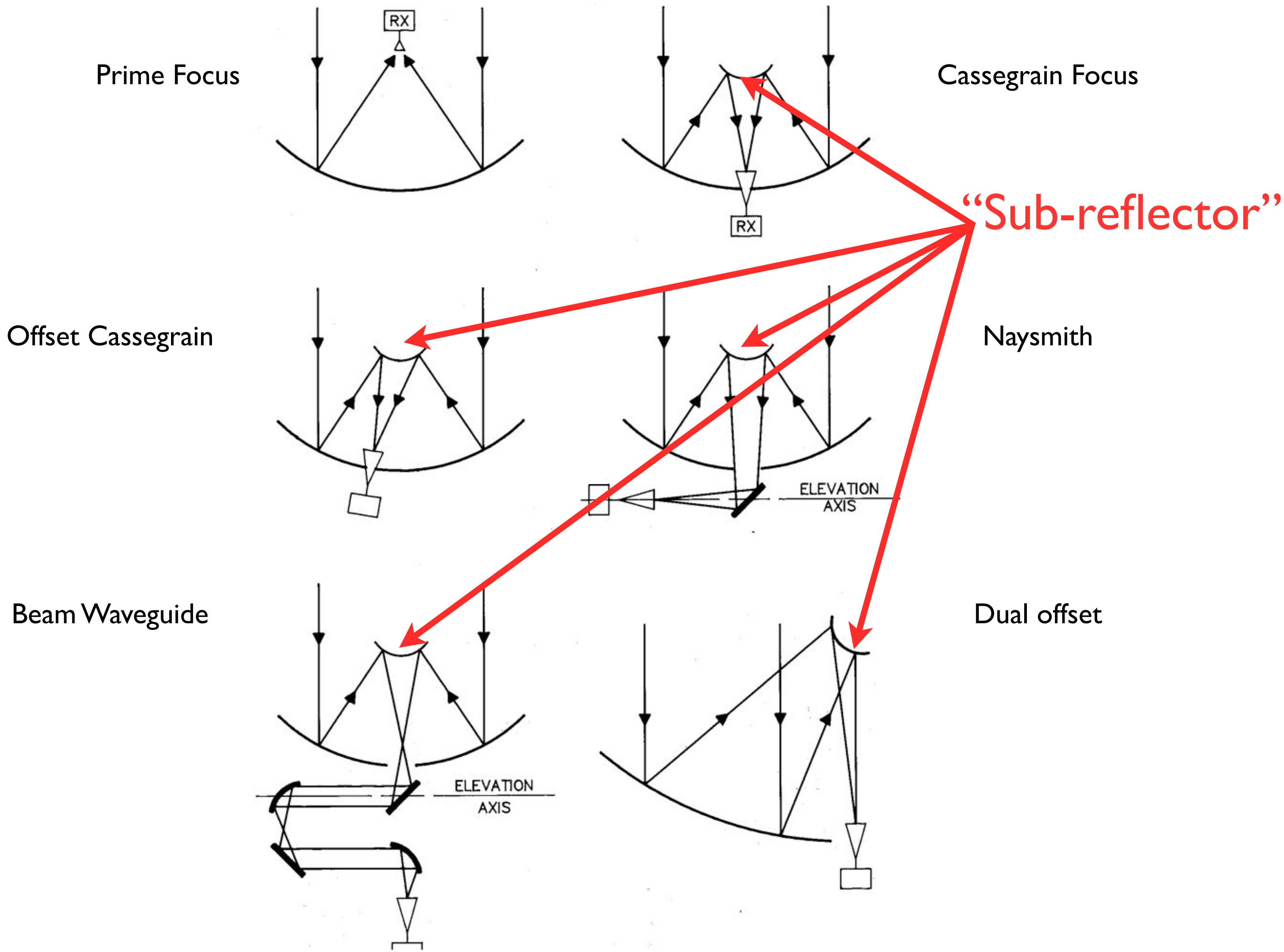
Beam Waveguide





Dual offset

Reflector types



Reflector types

Prime Focus
e.g. GMRT

Cassegrain Focus
e.g. Mopra (AT)

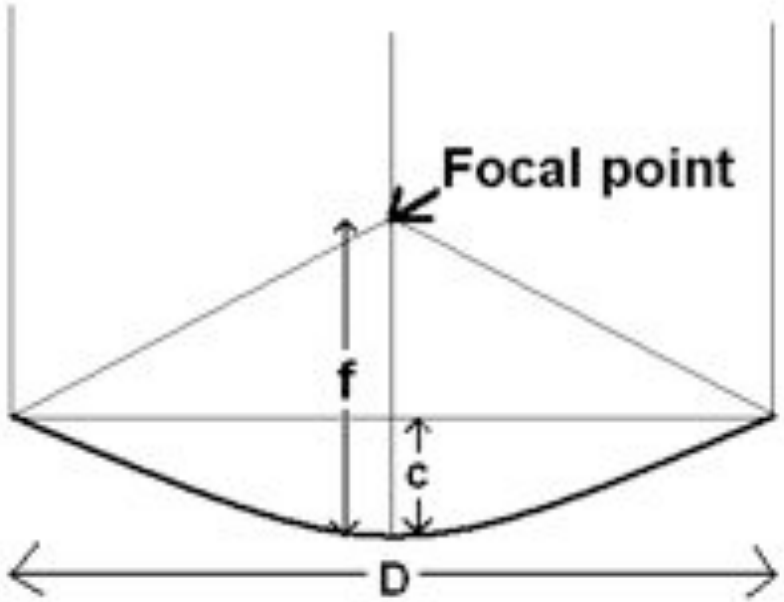
Offset Cassegrain
e.g. VLA and
ALMA

Naysmith
e.g. OVRO

Beam Waveguide
e.g. NRO

Dual gregorian
offset
e.g. ATA

Parabolic Reflector



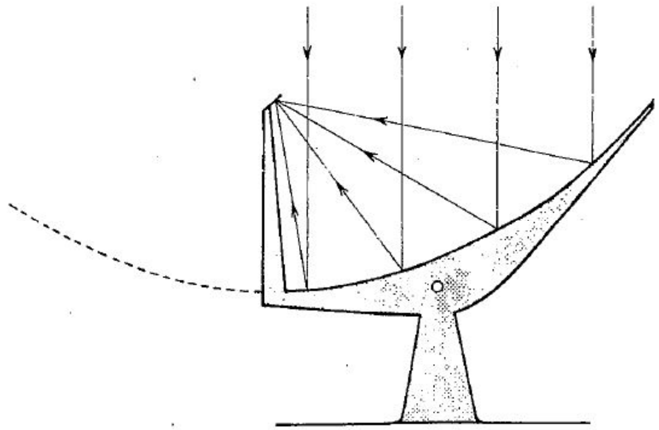
The focal length, f is given by:

$$f = D^2/16c$$

Typically the f/D ("f over D") ratio of a radio telescope is about 0.5. e.g. the WSRT antennas.

This typical value also ensures a rigid structure.

- Prime focus:
  - can be used across full frequency range of antenna but access to receiver is restricted
- Cassegrain (and other non-prime focus e.g. Naysmith and waveguide) provide good access to receivers but low-frequency receivers become impractically large and must be placed at prime focus.
- Off-axis Cassegrain (e.g. VLA antennas) enables frequency flexibility; receivers located in a circle can be quickly rotated to the focus. However, the assymetry of the offset optics introduces nasty polarisation characteristics that can limit imaging results.



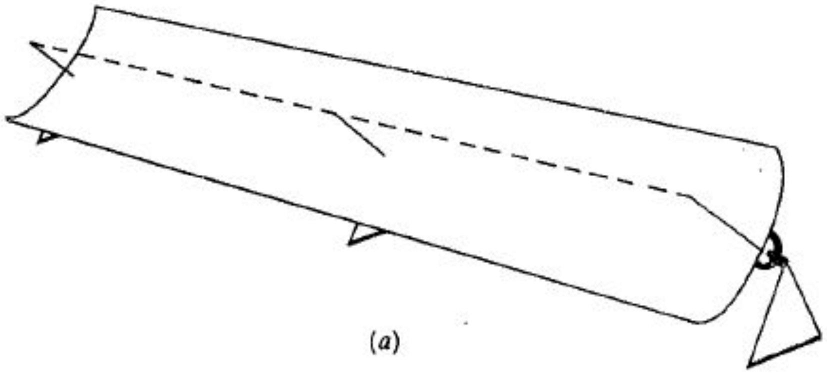
- Offset gregorian (left & above) has no blockage of the aperture.

A good example of this system includes the largest telescope in the world: the Greenbank Telescope (GBT)

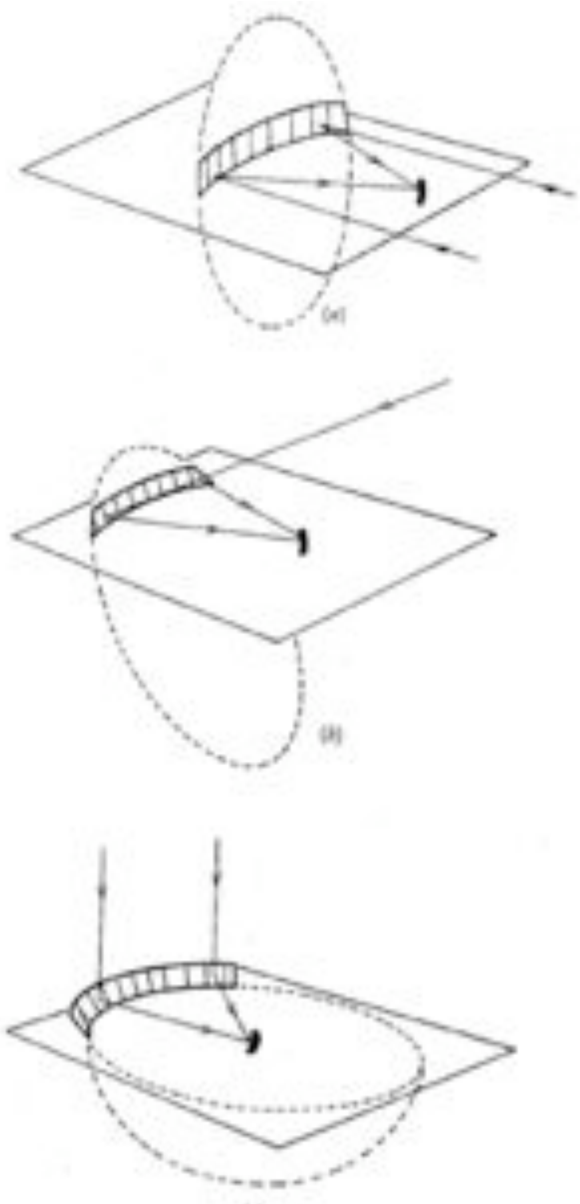
Aside: Some less conventional (weird!)reflector types

The Jodrell Bank Mark 2 telescope (1964). Was considered to be a prototype of the then planned giant 300-metre MkIV. The aperture is elliptical - the idea was that a 300-metre would require an elliptical surface in order to reduce the height of the structure off the ground.

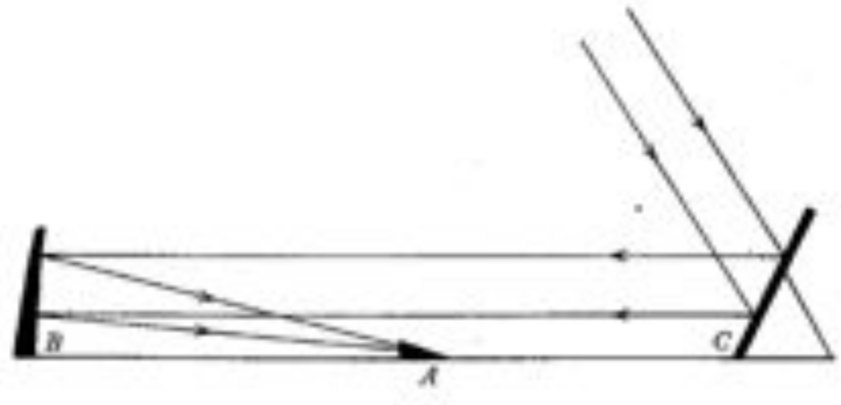
The off-axis cylinder radio telescope at Ooty. India (1970)



Instead of building a large paraboloid the Kraus antenna employs (sub-) sections of a parabolic surface. To steer the beam the reflector needs to be tilted:



Fixed paraboloid:



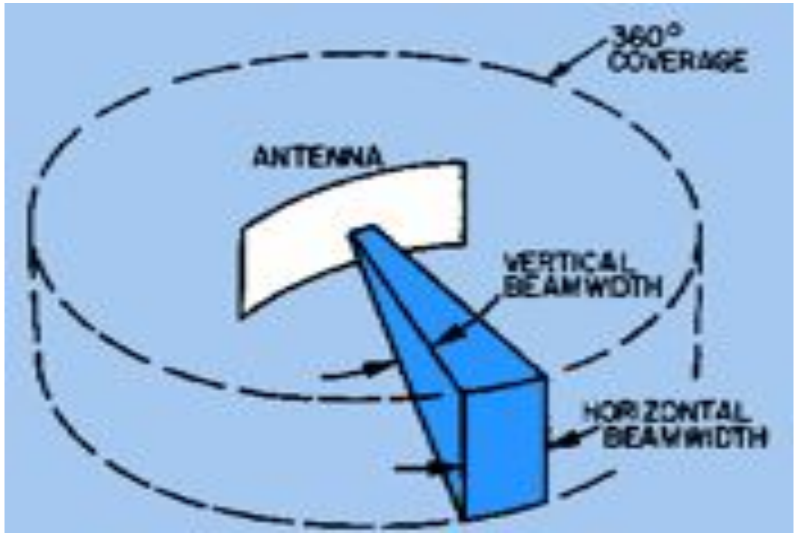
The big-ear antenna built by Kraus

Other similar examples: Ratan 600 - Russia

Nancay (France):

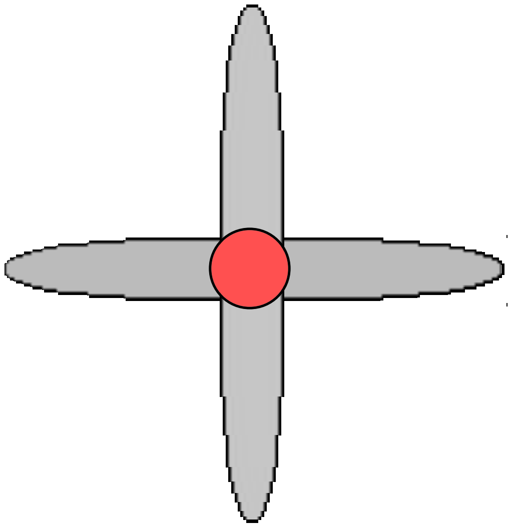
Cross antennas

Instead of building an entire parabola, Cross Antennas employ only a narrow section of a parabola, we get a beam narrow in the antenna's wide direction, and broad in the other direction: .



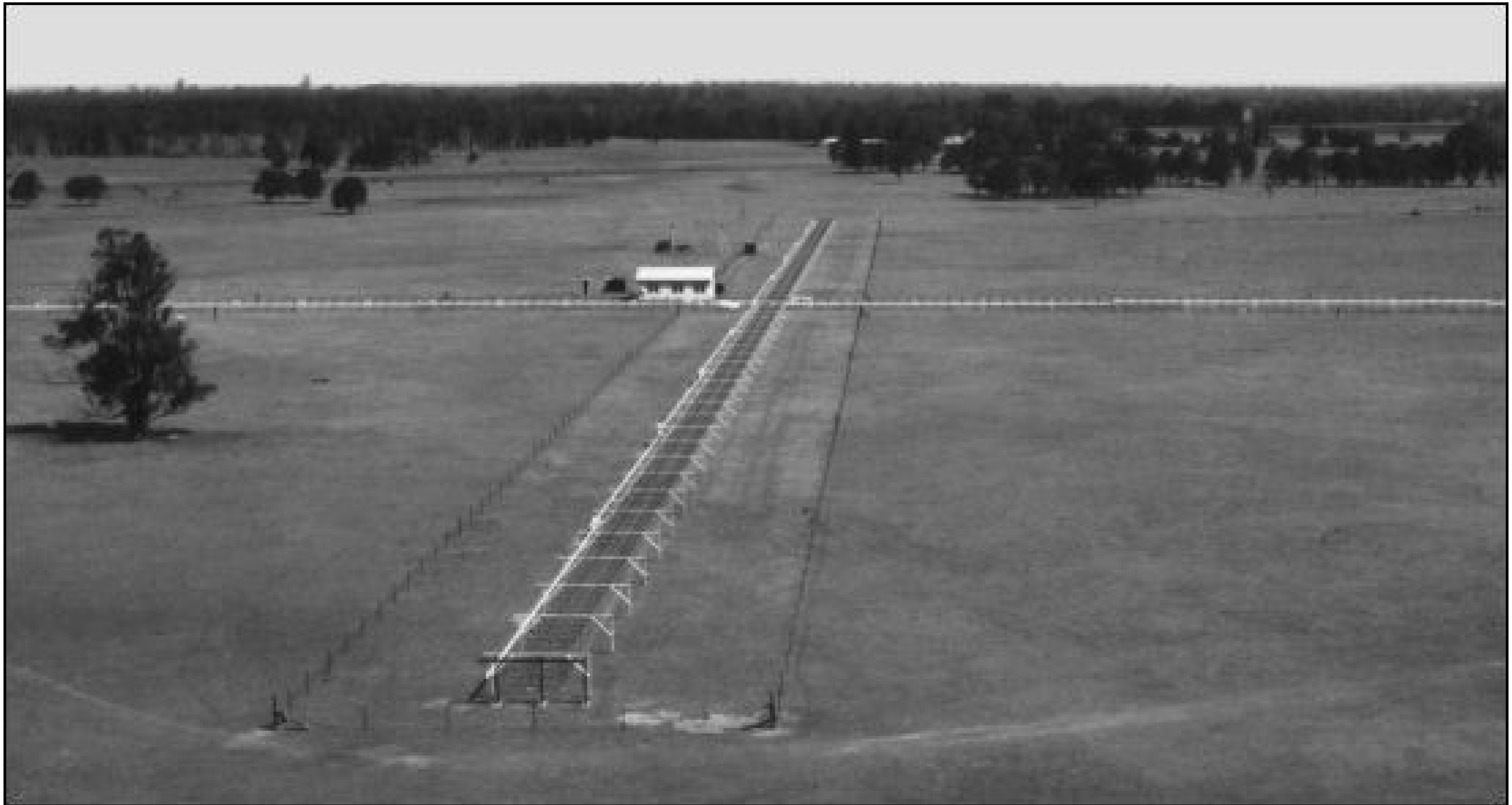
By observing a source with two orthogonal beams, we can get a 2-d image of the sky.

The cross antenna response is similar to the crossing point of the two beams (but with very high side-lobes):



The first Cross telescope - was built by Bernie Mills in Australia.

The Mills cross:

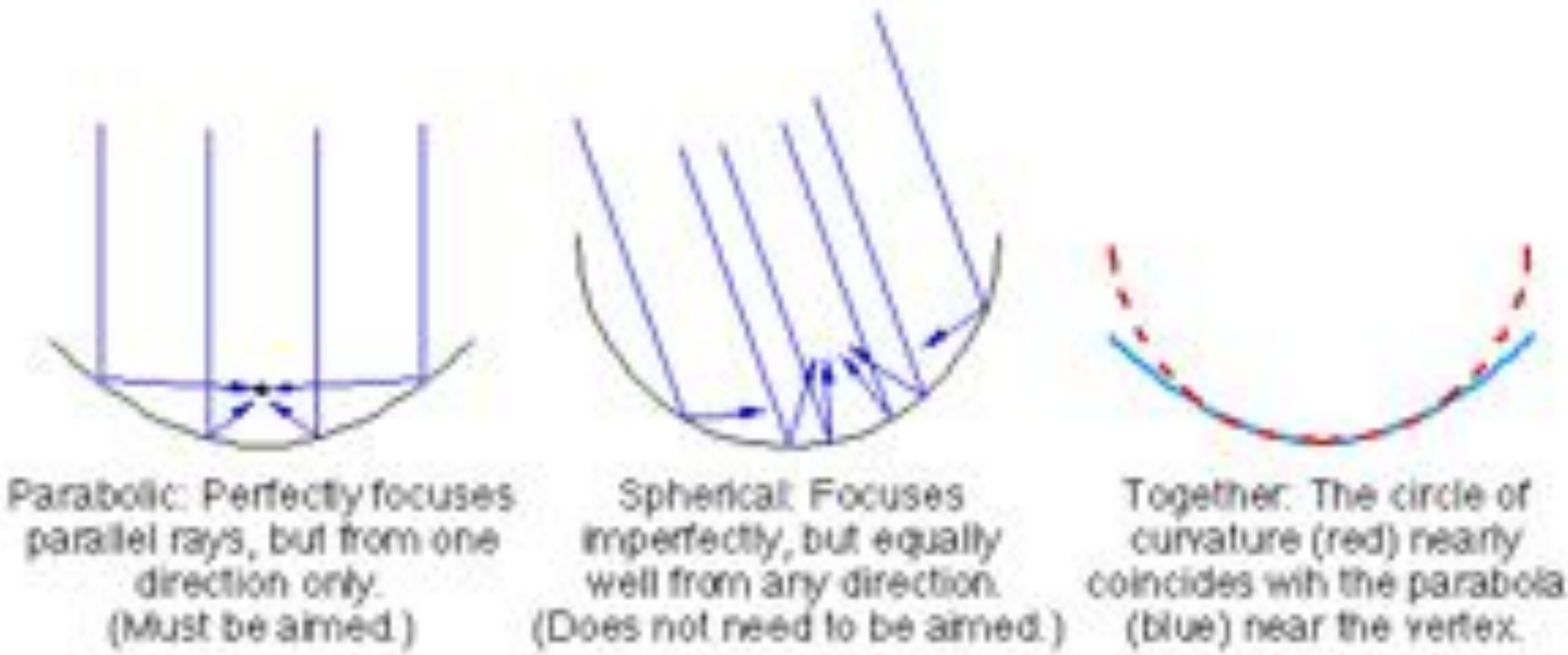


MILLS CROSS (CSIRO)

Many other examples:

The 305-m Arecibo telescope is fixed in the ground. A spherical reflector is therefore employed:

While a parabola has a single focus point, a spherical reflector focus the incoming radio waves on a line:



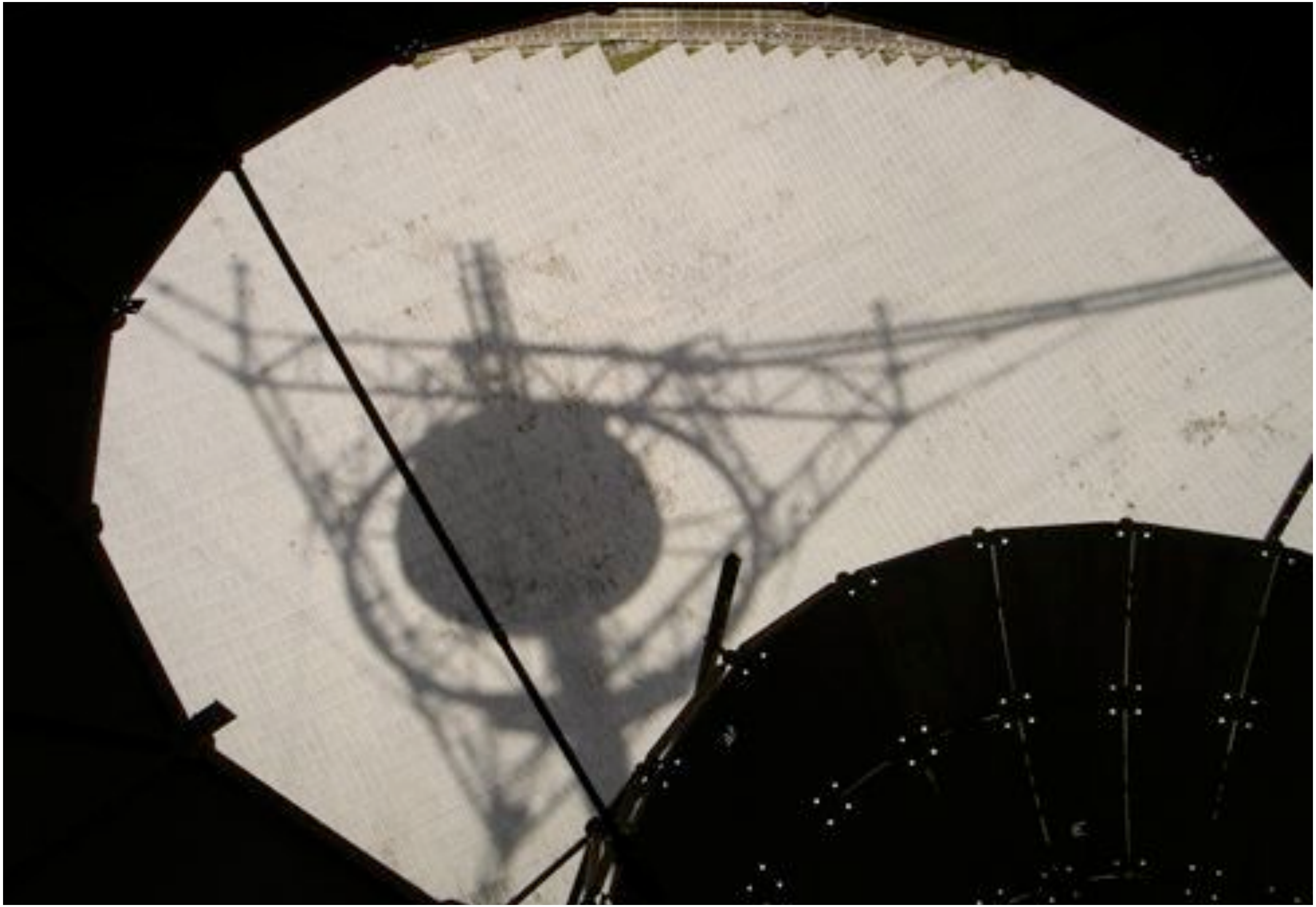
By having a moving secondary a spherical reflector can be pointed in different (but still somewhat limited) directions on the sky.

Note that only part of the total surface area is useable for any given direction.

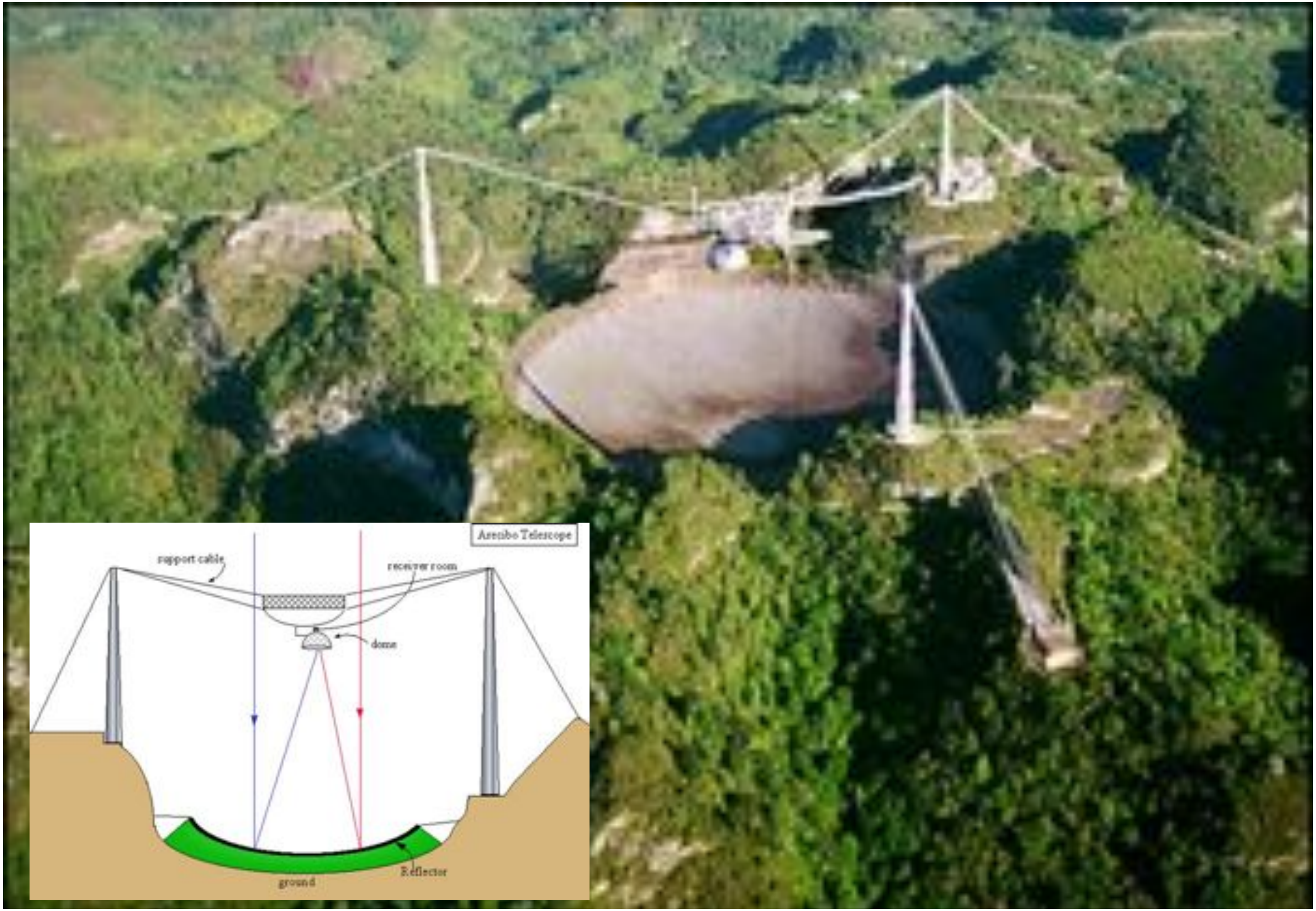
Part of the primary:

Looking on to the surface: quite a lot of litter!

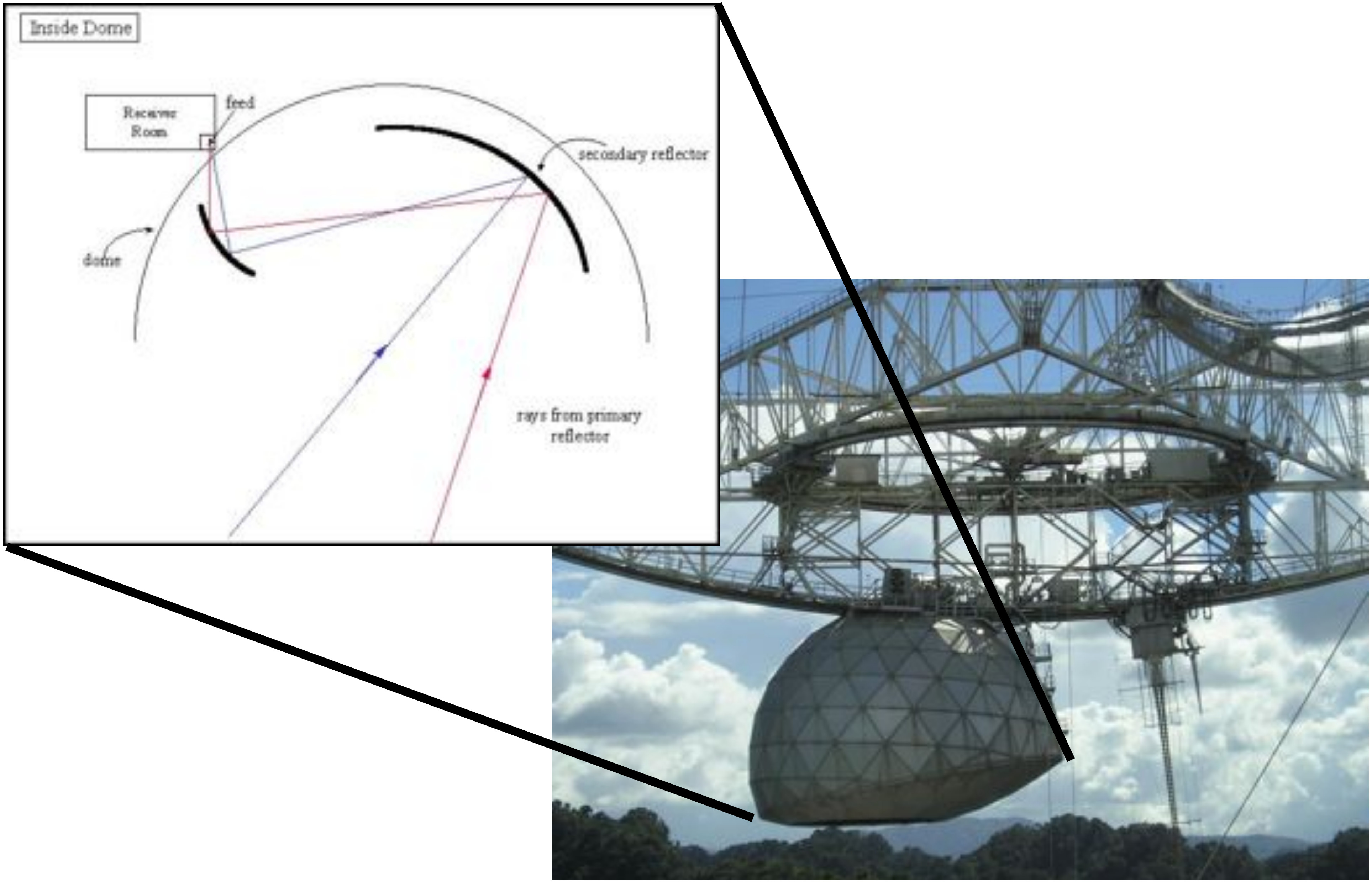
Looking down on the surface from the focus box:



Arecibo is built in a karst depression. The Gregorian secondary hangs on cables that are supported by 3 large towers:



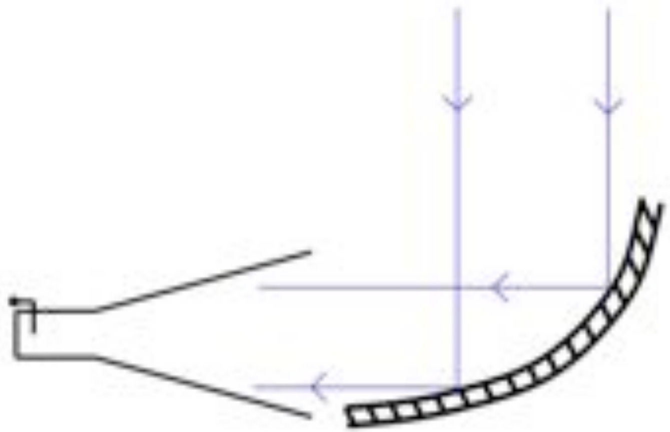
The large secondary feeds a tertiary reflector which in turn feeds a receiver room that has a broad range of receivers.:



Arecibo is an impressive telescope - huge scale:

Even more impresive is what lies under the dish itself: another world... with its own road and river!

Horn antennas:

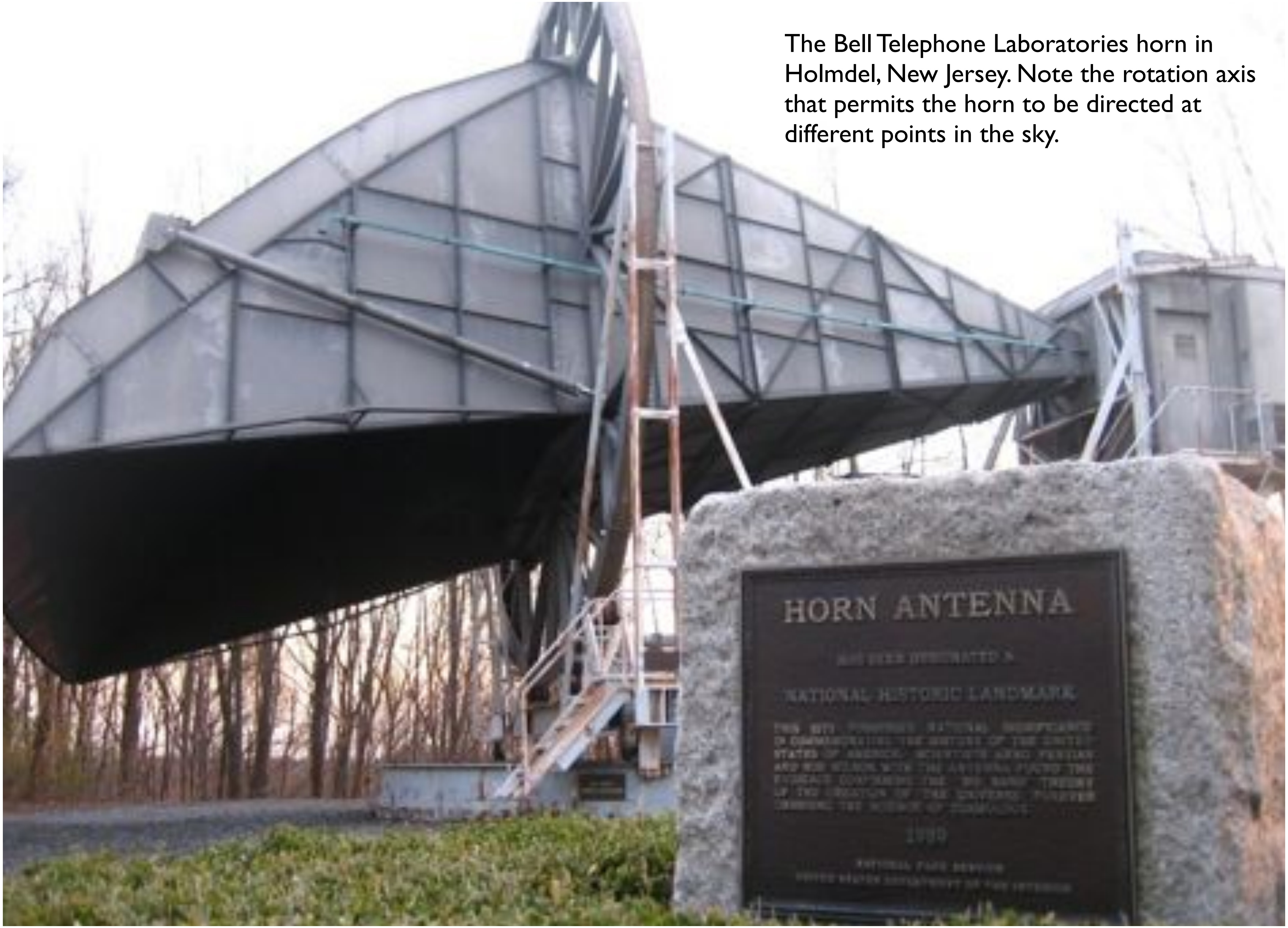


The reflecting "ear" reflects the incoming radio waves towards a horn or bare dipole.

The Horn Antenna combines several ideal characteristics: it is extremely broad-band, has calculable aperture efficiency, and the sidelobes are so minimal that scarcely any thermal energy is picked up from the ground. Consequently it is an ideal radio telescope for accurate measurements of low levels of weak background radiation.

A very famous example is the horn antenna located at Bell Telephone Laboratories in Holmdel, New Jersey, used by Penzias and Wilson to detect the relic radiation of the big bang.

Horn antennas have many practical applications - they are used in short-range radar systems, e.g. the hand-held radar "guns" used by policemen to measure the speeds of approaching or retreating vehicles.



The Bell Telephone Laboratories horn in Holmdel, New Jersey. Note the rotation axis that permits the horn to be directed at different points in the sky.

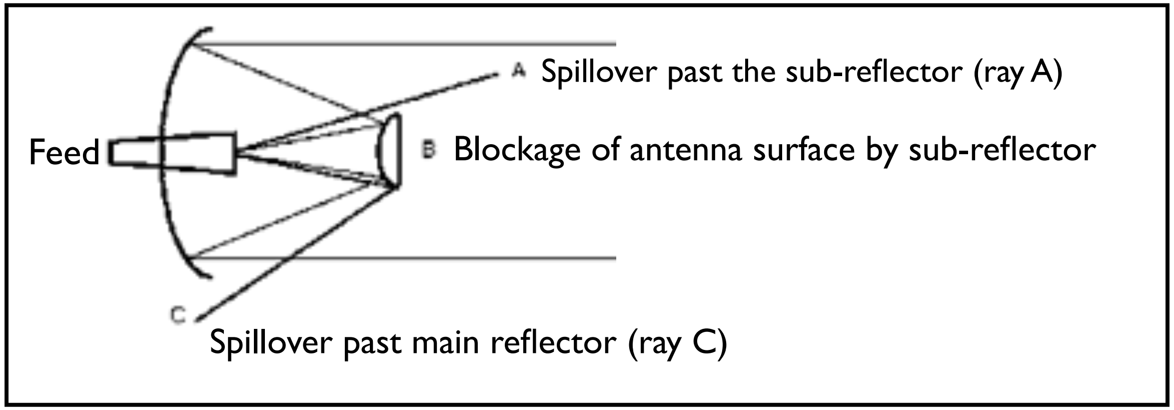
Antenna Performance

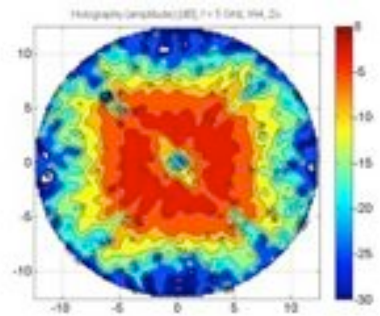
The antenna aperture efficiency η = $\frac{\text{Power collected by feed}}{\text{Power incident on antenna}}$

There are many different potential loss factors: $\eta = \eta_{sf}\eta_{bl}\eta_{sp}\eta_{t}\eta_{misc}$ [6]

η ~ 0.65 <==

Surface efficiency ~ 0.9	Aperture blockage efficiency ~ 0.9	Feed spillover efficiency ~ 0.9	Feed illumination efficiency ~ 0.9	Misc. - other minor losses e.g feed mis-match.





Feed does not illuminate all of antenna surface equally

Antenna surface efficiency

According to the Ruze (1966) formula, the surface efficiency of a paraboloid is well described by:

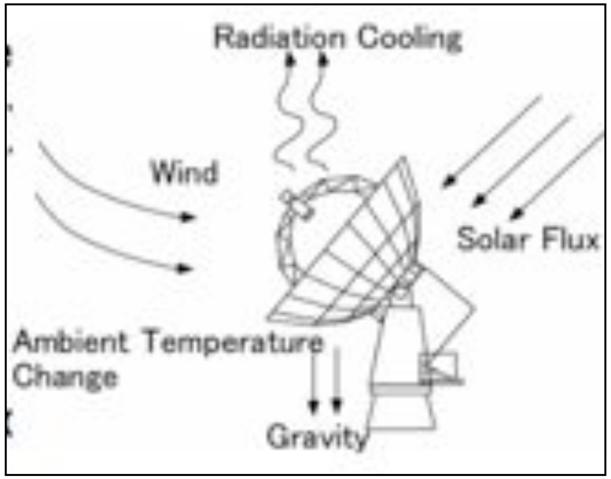
$$\eta_{sf} = e^{-(4\pi\sigma/\lambda)^2} \quad [7]$$

where sigma is the r.m.s. error in the surface of the antenna.

Or re-arranging: $\frac{\sigma}{\lambda} = \frac{1}{4\pi}\sqrt{-ln(\eta_{sf})}$

e.g. For a surface efficiency of 0.7 (typical target value), the required surface error (r.m.s.) is ~ lambda/20.

==> at 7 mm (43 GHz) the surface accuracy must be ~ 350 micron.



==> many different forces acting on an antenna and its surface...

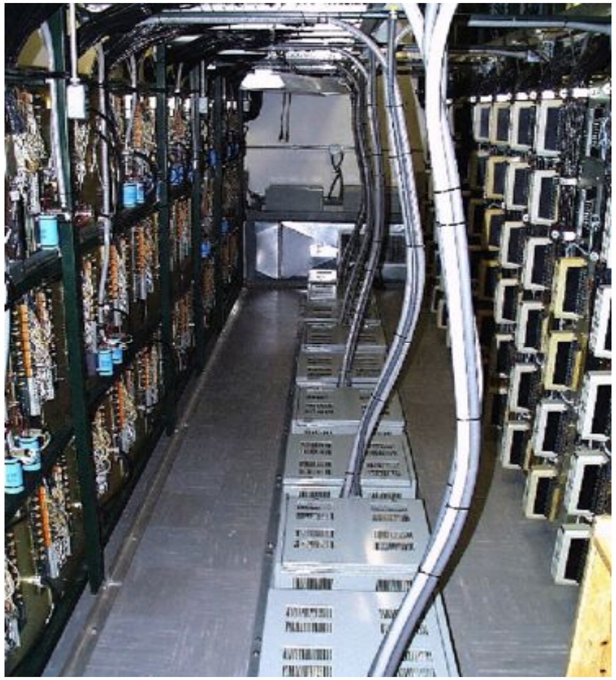
Appreciating the scale of large radio telescopes....

How can we possibly achieve a 350 µm accuracy *(the thickness of three human hairs)* – over a 100 metre diameter surface – an area equal to 2 football fields!

==> "active surface".

GBT Surface has 2004 panels
average panel rms: 68µm.

More than 2000 precision actuators are located under the surface panel corners

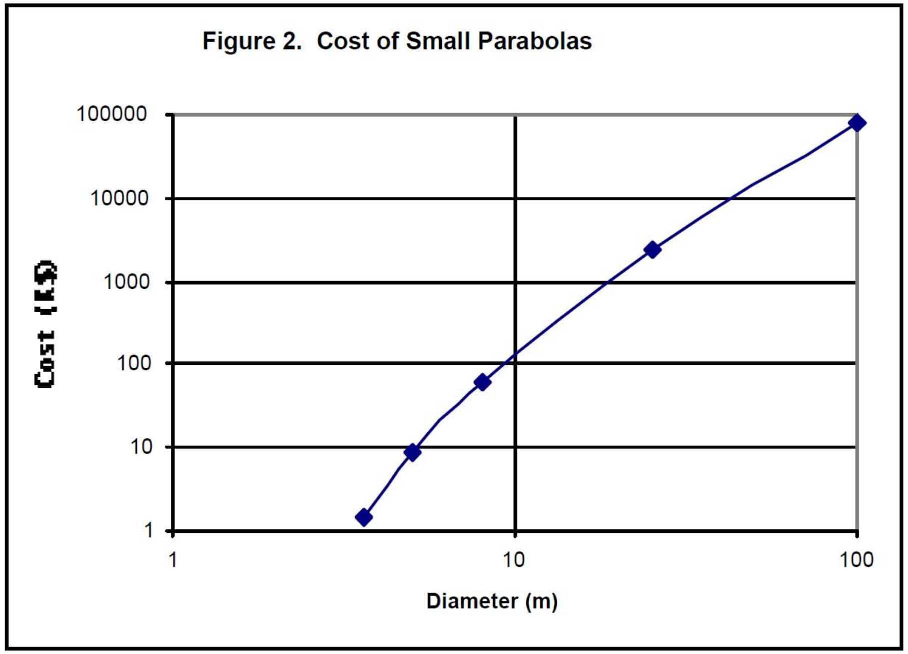


Actuator Control Room (left): 26000 control and supply wires terminate in this room!

How big can parabolic radio telescopes be ?

As the size (diameter) of a radio telescope increases, the gravitational and wind loads on the structure become difficult to manage. The worst problem is the problem of surviving a storm. The degree of wind distortion between paraboloids of different diameters (D) scales as D^3.

The cost of antennas also scales roughly as D^3.



Telescopes like the Jodrell Bank Mark V (right) with a diameter of ~ 305 metres (1970), will probably always remain in model form!

Aside: Power ratios (decibels)

In radio astronomy, decibels are often used to quantify changes in signal level as they pass through the antenna system.

We define power gain (e.g. by an amplifier) as:

$$Gain_{dB} = 10\log(\frac{P_{out}}{P_{in}})$$

We define power loss (e.g. cables) as:

$$Loss_{dB} = 10\log(\frac{P_{in}}{P_{out}})$$

Power gain (in terms of voltages) as:

$$Gain_{dB} = 20\log(\frac{V_{out}}{V_{in}})$$

Absolute Power relative to 1 watt:

$$Power_{dBW} = 10\log(\frac{P}{Watt})$$

Absolute Power relative to 1 milliwatt:

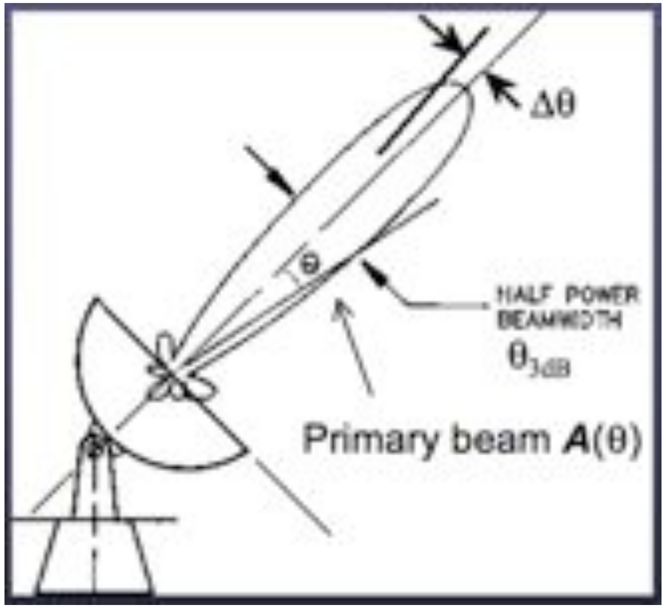
$$Power_{dBm} = 10\log(\frac{P}{mWatt})$$

dB invented by Alexander Graham Bell at Bell Labs.

e.g. a 3dB attenuation of signal power = 50% of signal power being lost.

N.B. a useful feature of dB, is that the total gain or losses associated with a number of inter-connected components is simply the sum of the gains and losses in dB of the individual components [log(abc)=log (a)+log(b)+log(c).

Antenna pointing errors $\Delta\theta$



WSRT pointing errors are typically < 30 arcseconds (or about θ_{3db} /10) at 8 GHz.

The typical goal is: $\Delta\theta < \theta_{3db}/20$

where θ_{3db} is the FWHM of the main lobe of the antenna beam.

N.B. If the antenna moves $\theta_{3db}/20$ off the true pointing centre, this will result in < 1% loss of intensity for a source located on the central axis of the beam.

However, a source located at θ_{3db} will see a 10% loss! This can badly affect the quality of a radio source image towards the edge of the field

At higher frequncies (~ 20 GHz) pointing checks are often made on nearby bright sources to update the pointing model (offsets).

Typically pointing becomes more difficult at higher frequencies and with larger antennas (i.e. smaller primary beams).

Antenna servo performance

The speed at which an antenna can move from one part of the sky to another is also an important performance factor.

This is required for:

(i) observing efficiency

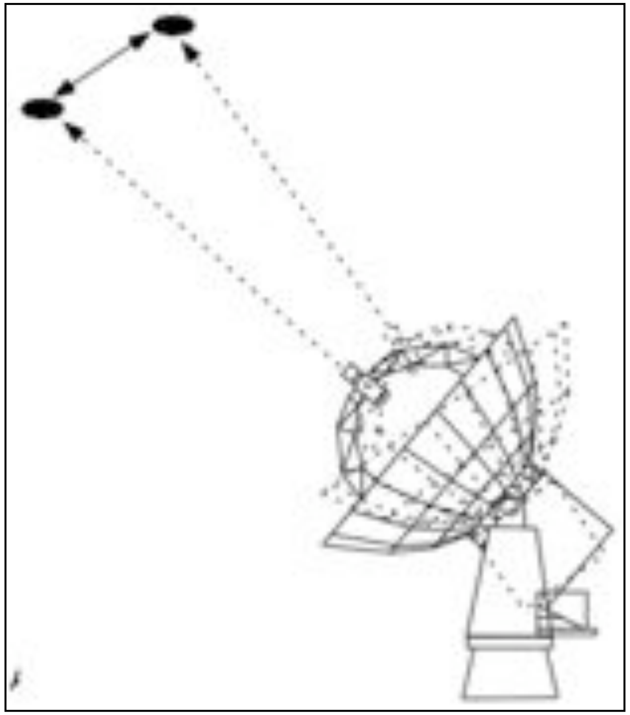
(ii) calibration (e.g. fast switching between nearby sources in the case of phase-referencing).

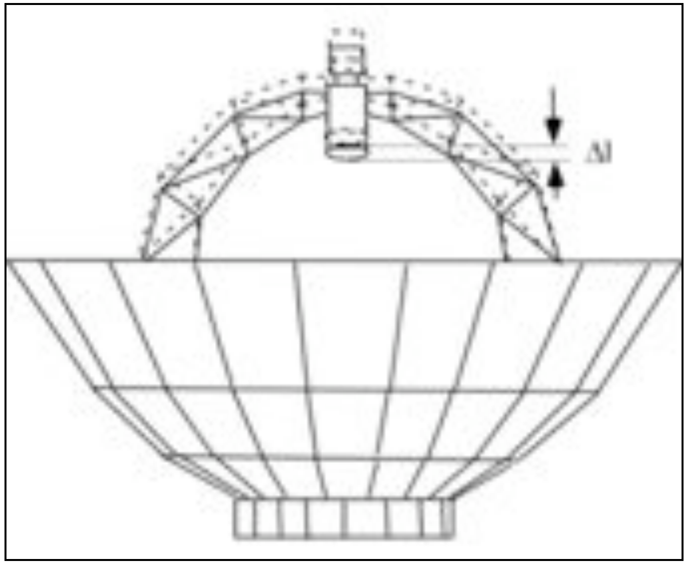
Typical driving rates of modern antennas (e.g. VLBA):

- 90 deg/min in azimuth; 30 deg/min in elevation;
- Settle time ~ 2 secs;
- Time to accelerate to full speed ~ 2 secs.

A rigid structure is important as this:

- minimises "settle time" - the time it takes for antenna to firmly settle on source.
- maintains the optical geometry of the telescope - important for "phase referencing"





e.g. Lovell telescope (right) old, heavy structure is not stiff - leads to over-shooting, long slew and settling times etc.

e.g. ALMA antennas - v. stiff and highly accurate pointing for THz obs.

Antenna Gain and Performance

The Flux Density is equal to the Planck function, "specific intensity" I(v), integrated over solid angle:

$$S = \int I(\nu) d\Omega \qquad (\mathrm{units: Wm^{-2}Hz^{-1}}) \qquad [4]$$

$$S = \int 2kT_b\nu^2/c^2 d\Omega = 2k\nu^2/c^2 \int T_b d\Omega \qquad [5]$$

i.e. the Flux density is just the brightness temperature integrated over the source.

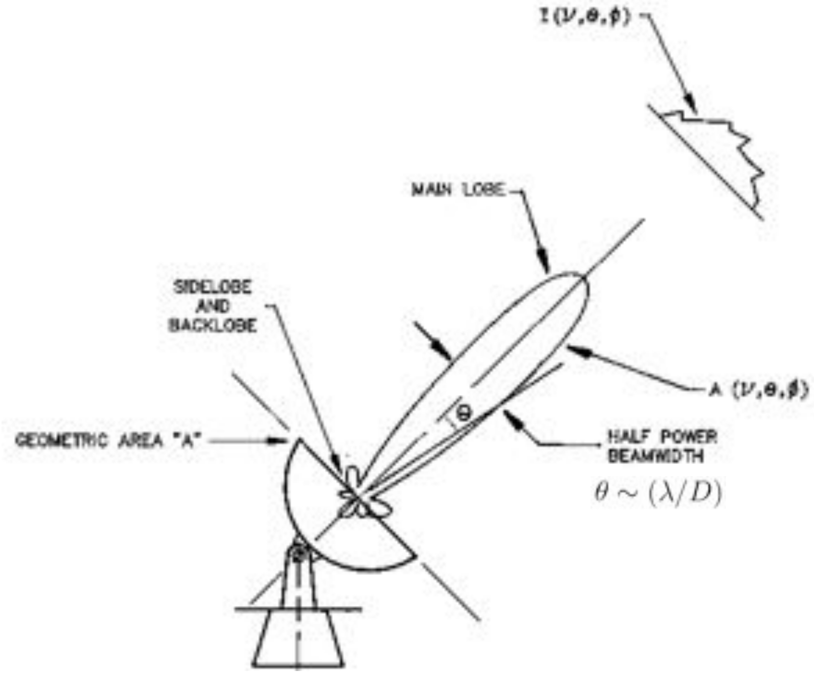
If the source size is equal to the beam size on the sky, the integrated flux density of the source is:

$S = (2kT_b/\lambda^2)\ \Omega_a$ [6] where Ω_a is the solid angle of the beam.

The "peak flux density" is sometimes quoted - this is measured in Jy per beam or Jy/beam.

The angular response of a parabolic antenna with aperture size, D, observing at a wavelength lambda, is diffraction limited and focused into a cone of solid angle:

$$\Omega_A = \frac{\pi}{4}\theta^2 = \frac{\pi}{4}\frac{\lambda^2}{D^2} \sim \frac{\lambda^2}{D^2} \quad [8]$$



Note if we substitute this into eqn [6] we get:

$$S' = \frac{2kT}{\lambda^2}\left(\frac{\lambda}{D}\right)^2 = 2KT/D^2 \quad [9]$$

$$S = 2\frac{P}{\eta A d\nu}$$

Noting that the power (P) into the receiver is given by eqn [3], by equating with [9] we can write:

$$P = kTdv$$ T is known as the antenna temperature, usually denoted T_A. [10]

Eqn[10] is equivalent to the power associated with a resistor placed in a thermal bath at a temperature T - the so-called Johnson-Nyquist formula.

The electrons in the resistor undergo random thermal motion, and this random motion causes a current to flow in the resistor. On average there are as many electrons moving in one direction as in the opposite direction, and the average current is zero. The power in the resistor however depends on the square of the current and is not zero. The power is well approximated by the Nyquist formula:

$$P=kTd\nu$$ where k is the same Boltzmann constant as in the Planck law.

In analogy with this, if a power P is available at an antenna's terminals the antenna is defined to have an antenna temperature of

$$T_A = P/(kd\nu)$$

Note that T_A is not the physical temperature of the antenna!

JULY, 1928 *PHYSICAL REVIEW* *VOLUME 32*

THERMAL AGITATION OF ELECTRIC CHARGE IN CONDUCTORS*

By H. Nyquist

JULY, 1928 *PHYSICAL REVIEW* *VOLUME 32*

THERMAL AGITATION OF ELECTRICITY IN CONDUCTORS

By J. B. Johnson

e.g. A 25-metre telescope, observing a 100 milliJansky (mJy) radio source measures an antenna temperature, T_A, of 0.023 Kelvin!

An isotropic antenna is a (mythical) antenna that collects (or radiates) energy uniformly in all directions. The gain of an antenna is defined as G,

$$G = \frac{\text{Power radiated in a specific direction}}{\text{Power radiated in that direction by an isotropic radiator}}$$

Note that the average gain of any antenna is always 1 if calculated over all angles. The gain of an antenna can also therefore be written as:

$$G = \frac{\text{Solid angle subtended by a sphere}}{\text{Solid angle of antenna beam}} \quad \text{or} \quad G = \frac{4\pi}{\Omega_A} = \frac{4\pi D^2}{\lambda^2} \sim \frac{4\pi A_e}{\lambda^2} \quad [11]$$

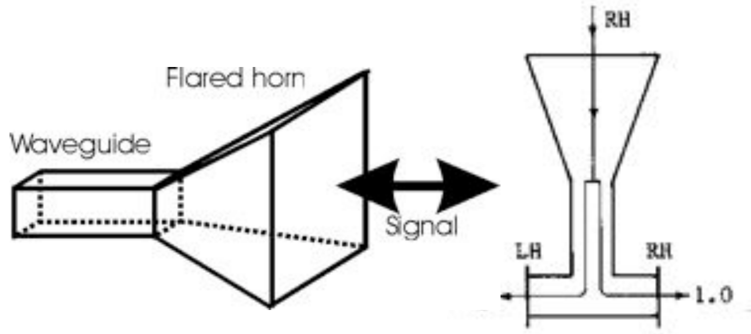
N.B. A_e is the effective area , and is always smaller than the geometric area (see eqn 6).
$A_e = \eta A$ where η is typically in the range of 0.4 - 0.8.

Note that from [11] we can also write:

$$A_e \sim \lambda^2/\Omega_A \quad [12]$$

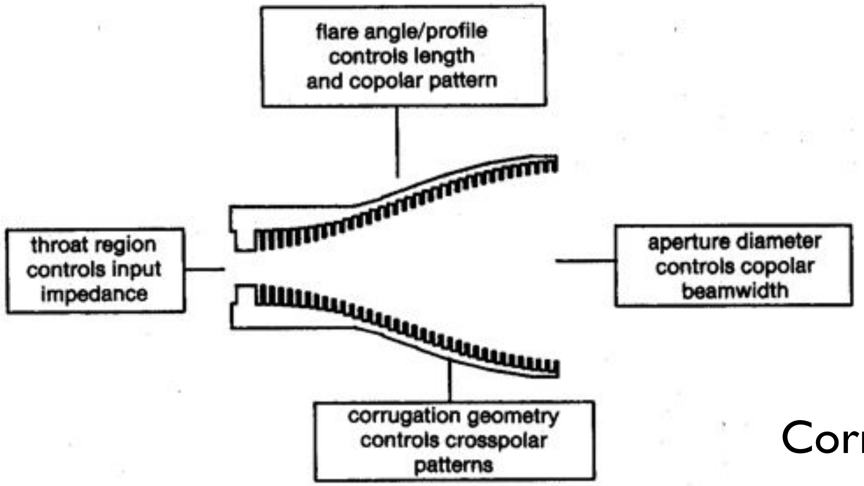
...an interesting fact to bear in mind w.r.t. simple dipoles

A *feedhorn* is the front-end of a waveguide that gathers the e-m signals at or near the focal point, and 'conducts' or guides them to a polarisor that splits the signals into opposite (circular) polarisations (e.g. into independent RH and LH channels - see right).



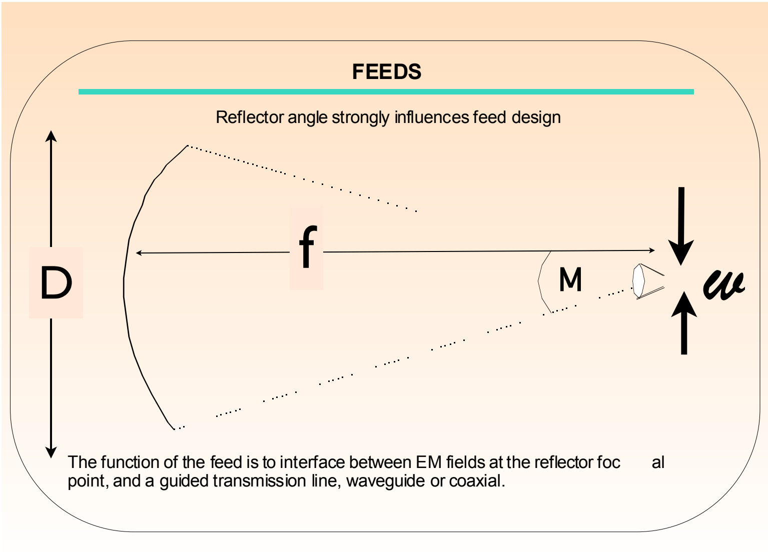
The feedhorn's interior is corrugated in order to increase the surface impedance, so that the wave does not set up voltages in the surface material, but is channelled into a dipole at the end of the horn.

The feedhorn (or "feed") is designed to evenly illuminate the antenna surface. The angle subtended by the reflector as seen by the feed strongly influences the types of feeds which may be used, and the details of their design.



Corrugated horns (above) are the most common.

Generally, the more narrow the angle (M), the larger the feed in units of wavelengths.



Since M ~ D/f (where f is the focal length of the paraboloid), the feed aperture diameter (w) is given by:

$$w \sim \frac{\lambda}{D} f \quad [13]$$

If the f/D ratio is low, say 0.25 to 0.35 then the feed will be close to the dish and needs to spread its power over a wide angle to efficiently illuminate the dish. The feed diameter therefore needs to be small. Note that if the f/D is 0.25 (see earlier slide on parabolic reflectors) the feed is level with the dish aperture, and requires a coverage of 180 degree which is difficult!

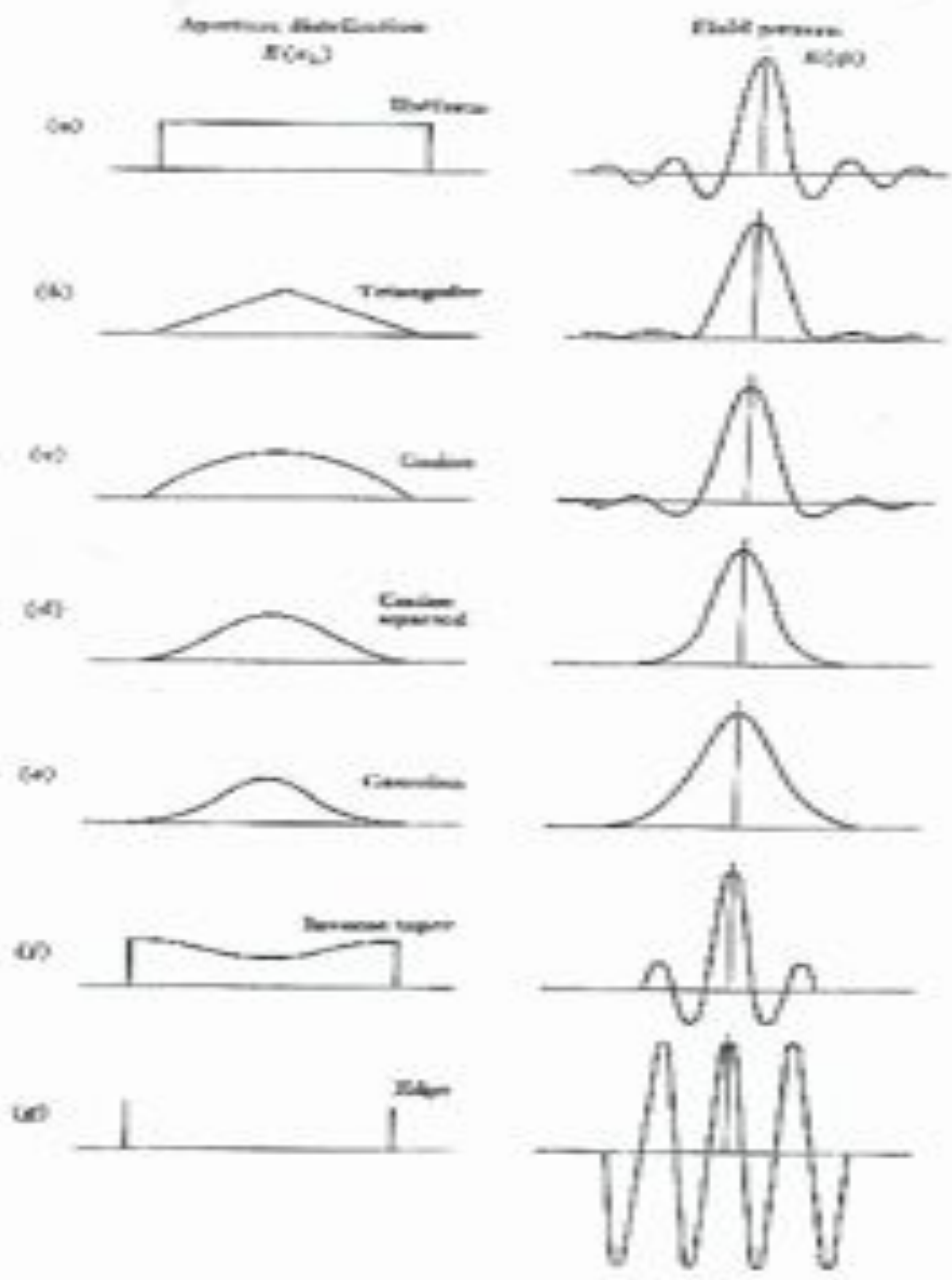
If the f/D is large like 0.75 then the feed will be further away from the dish and needs to spread its power into a narrower angle. The feed needs to be of a larger diameter however.

The illumination of the antenna surface by the feed is usually not uniform

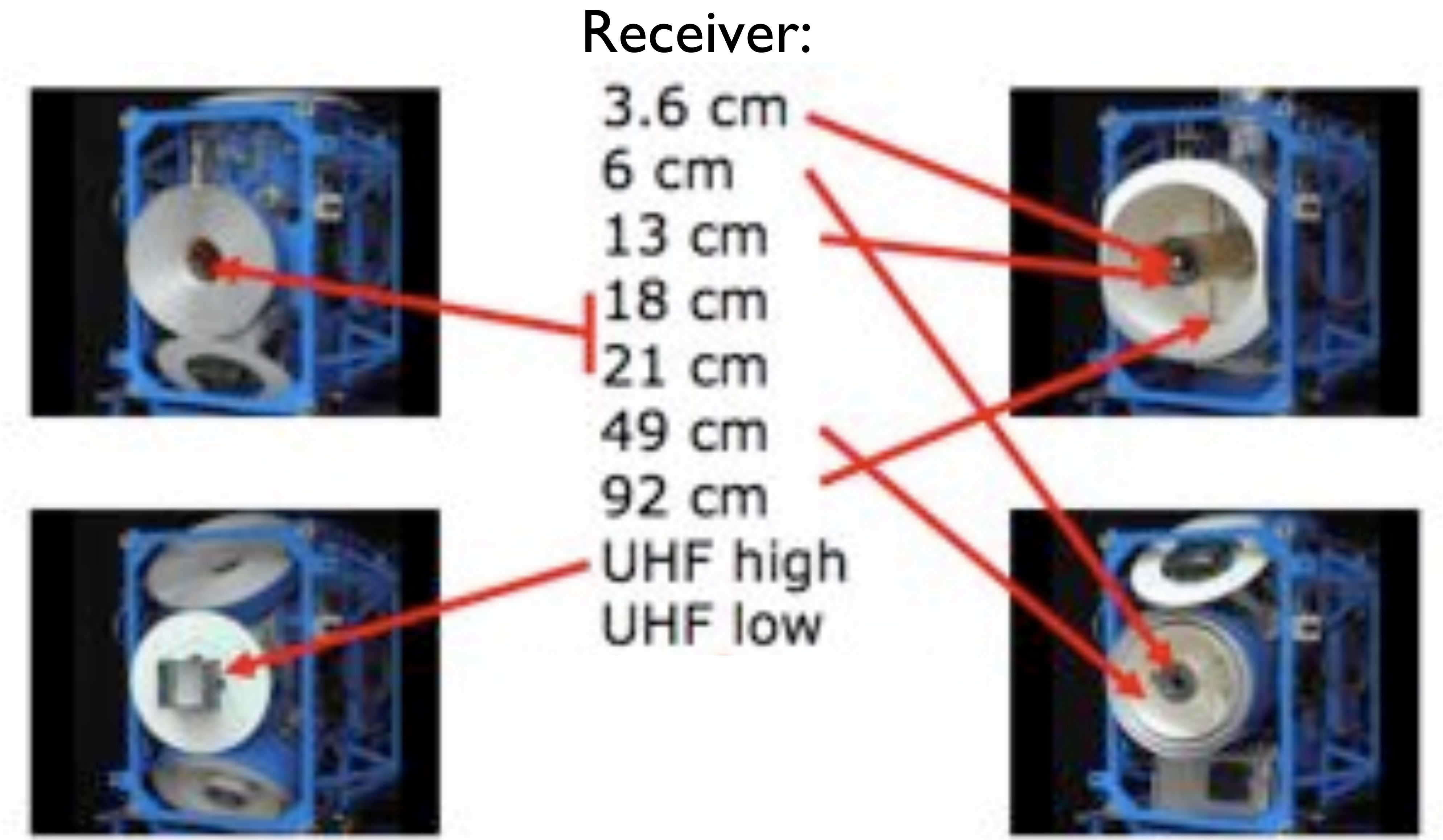
Feeds are usually designed to under illuminate edges of the dish - in order to avoid spillover from the ground.

Such a design produces a larger beam but smaller side-lobes. Cases (b), (c), (d), (e) - right.

Over illumination of the edges results in a narrower beam (better resolution) but high sidelobes. Cases (f) and (g) - right.

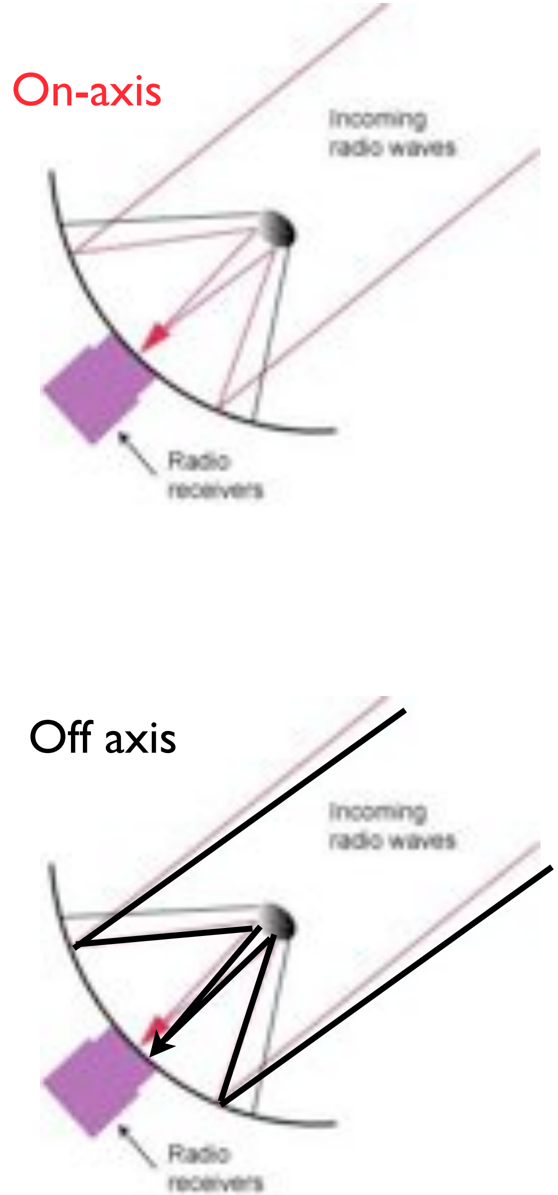


WSRT Multi-frequency front end (MFFE) system (longer wavelengths ==> larger feed - see eqn 13):

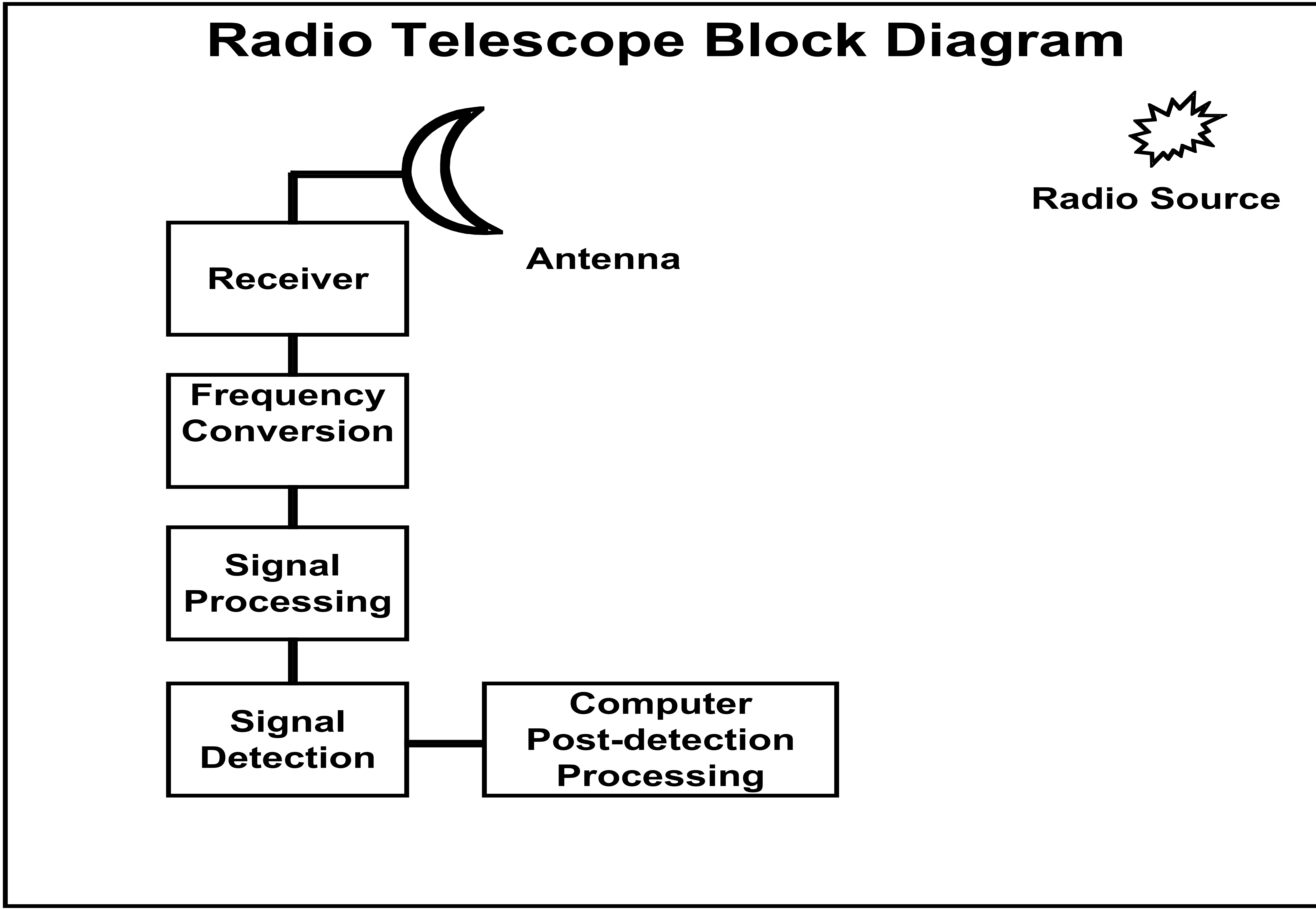


The VLA uses a rotating turret to position each of its feeds (receivers) slightly off-axis. Leads to some calibration problems.

Gregorian Receiver Room



Overview of this lecture



Radio astronomy receiver systems

Two most common types of receivers in radio astronomy: (i) heterodyne receivers and (ii) bolometer receivers.

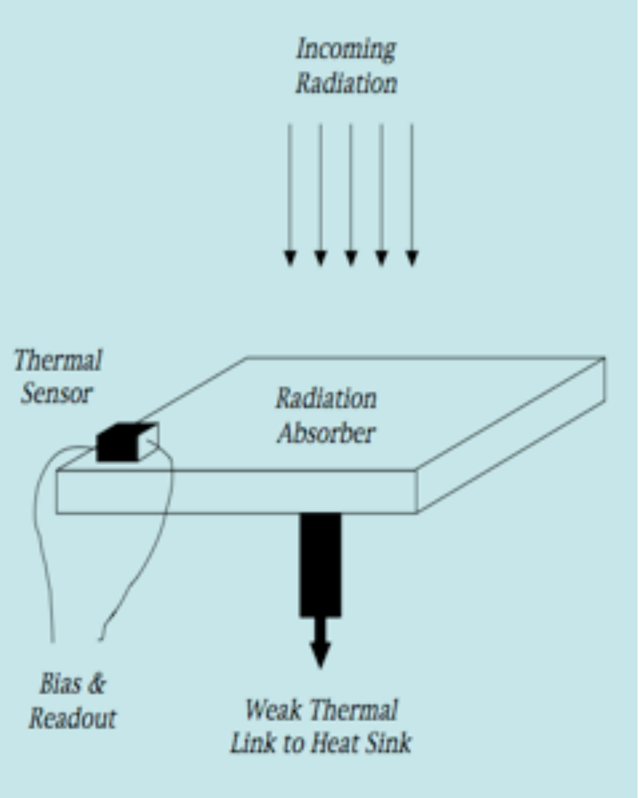
(i) *Heterodyne receivers* - sensitive to the incoming electric field; frequency of the received signal is converted down to a lower frequency by a precise reference signal (mixer) generated locally in the receiver system.

Heterodyne receivers are used at metre, centimetre, millimetre and submillimetre wavelengths. We will study this type of receiver in more detail later.

Example in daily life : Cell phone, WiFi antennas. FM and AM radio ; Example in the laboratory : spectrum analyzer

(ii) *Bolometric receiver* – sensitive to thermal-electrical effects; incoming photons are directly detected, heat is generated and the total power (resistance) changes due to material temperature changes.

Bolometers only record the *intensity* of light but over a very broad range of wavelengths (large bandwidth) e.g. over an entire atmospheric "window" - 310 GHz to 370 GHz. Used exclusively at high (sub-mm) wavelengths to do photometry. Usually cooled to milli-Kelvin level to ensure they are limited only by sky background. N.B. No (or very limited via filters) spectral resolution capability.



General properties and parameters of any detector system

Good detectors preserve the information contained in the incident e-m disturbance or photon stream. Relevant parameters include:

(i) Quantum efficiency - fraction of photons converted into a signal

(ii) Noise - the uncertainty in the output signal - hopefully dominated by statistical fluctuations that are due to the number of photons producing the signal - and free of systematic effects so that longer integrations produce improved noise levels

(iii) Dynamic range - the maximum variation in the signal over which the detector is sensitive and over which no information is lost (e.g. via saturation effects)

(iv) Number and physical size of the pixels (imaging elements) the detector can use simultaneously

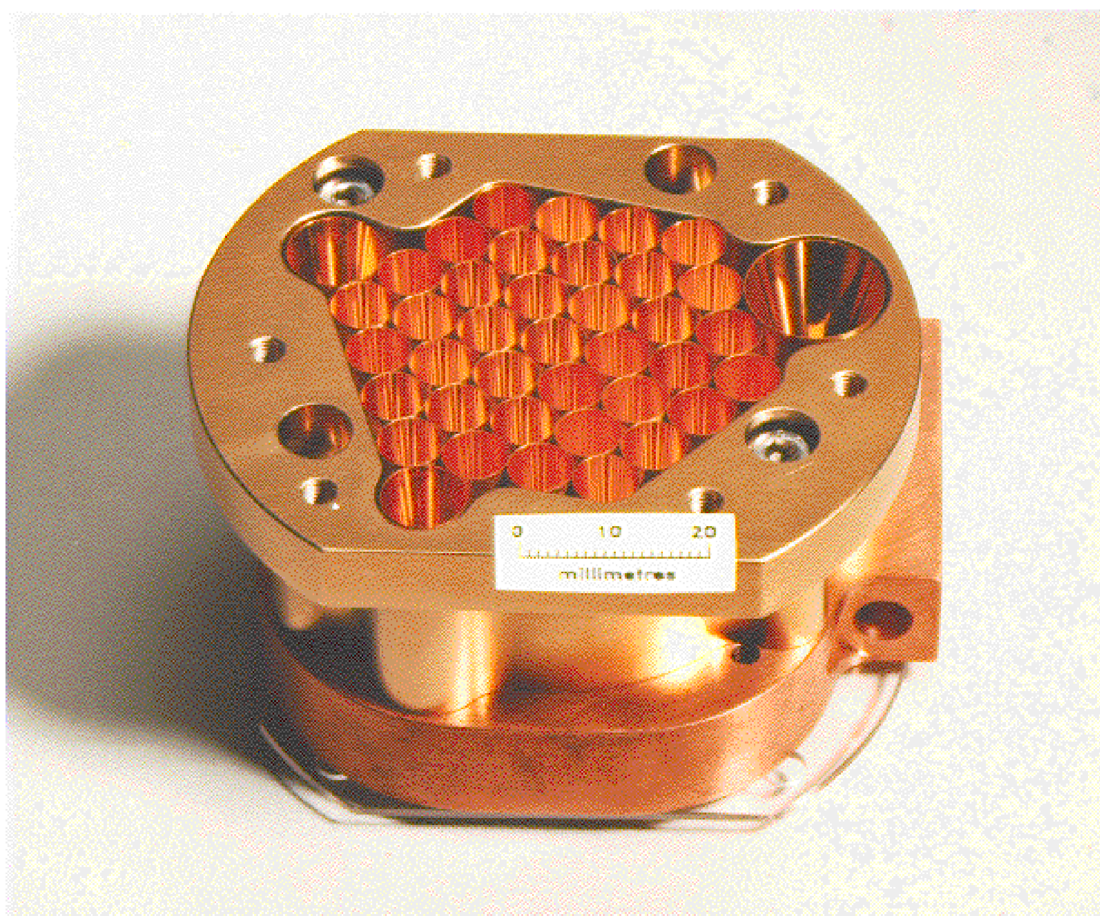
(v) Time response (temporal resolution) of the detector - minimum interval in time over which the detector can distinguish changes in the intensity of the incoming radiation field

(vi) Spectral response - frequency range over which the detector is sensitive to incoming radiation

(vii) Spectral resolution - smallest frequency interval over which the detector is sensitive to incoming radiation.

Until recently most bolometers had a single detecting element and imaging was performed by a raster process in which telescope moves over a celestial object one pixel at a time! Recently, the first bolometer arrays have been constructed greatly improving survey speed.

SCUBA@JCMT: 450 / 850 um, 91/37 pixels

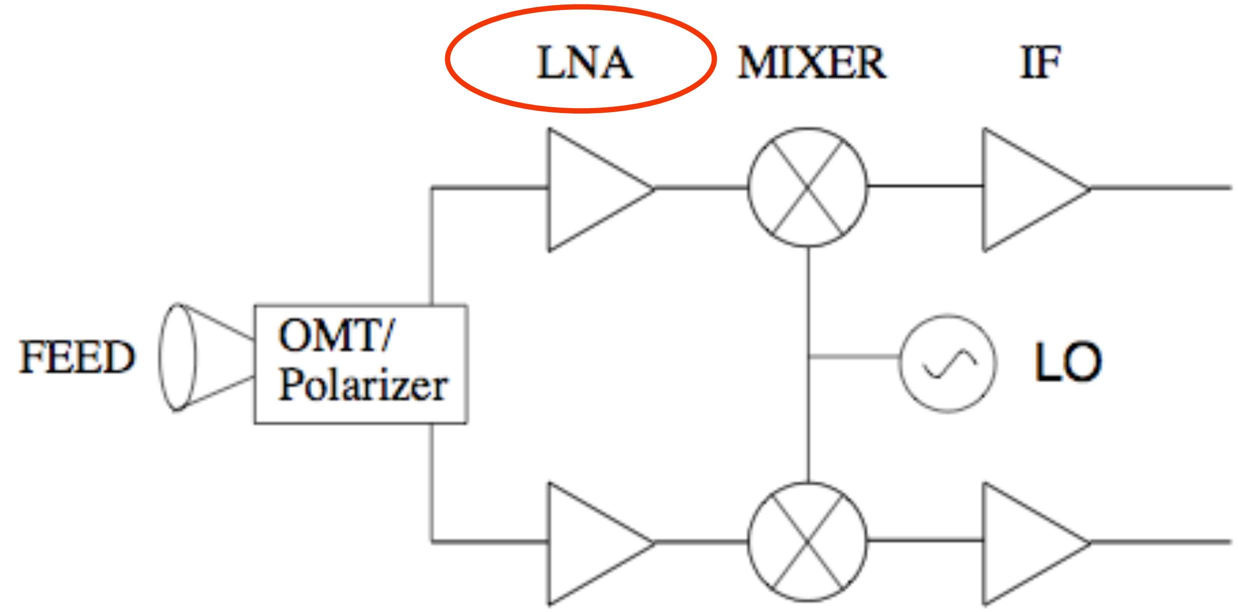


SCUBA-2: 10000 pixels, for JCMT.

SCUBA-2

Heterodyne receiver system

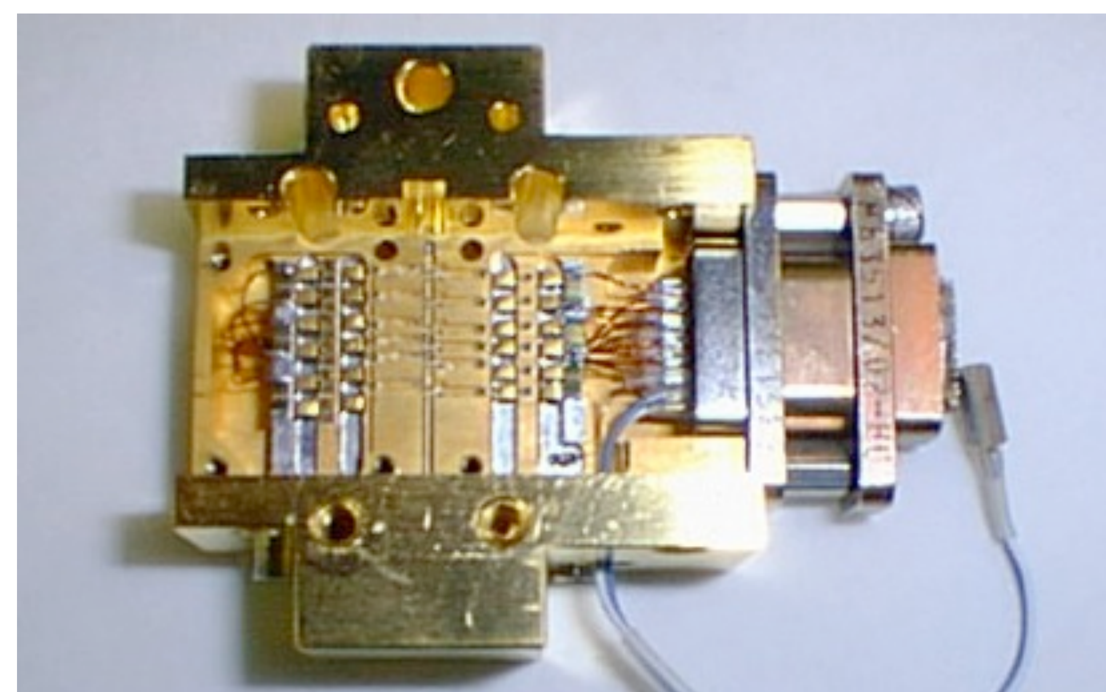
A typical heterodyne radio astronomy receiver system (right). The receiver amplifies the incoming signal from the feed, filters the signal and “down-converts” it to a lower frequency where it can be more easily sampled or detected.



Already covered role of feed and polariser. Some feeds inherently detect and separate polarisations, other types require orthomode transducers (OMT - see above) to separate channels.

LNA = Low Noise Amplifier - amplifies the incoming signal - LNAs are usually cryogenically cooled to minimise the noise they also add to the overall system.

At ν < 10 GHz use cooled gallium arsenide (GaAs) HFET (heterostructure field-effect transistor) amplifiers; at higher frequencies indium phosphide (InP) HFETs have superior performance.



Example of a LNA used by WMAP CMBR satellite.

Thermal noise (Johnson noise) exists in all electronic components and results from the thermal agitation of free-electrons. The noise is typically "white noise" (flat power response with frequency).

In electronics, the noise temperature is a temperature (in Kelvin) assigned to a component such that the noise power delivered by the noisy component is given by,

$$P \sim kTd\nu. \qquad [12]$$

The noise contributions of the various components in a receiver system are usually independent (uncorrelated) and the total noise in a receiver system (T_{RX}) can be estimated by summing all the individual contributions.

The total system temperature, T_{SYS} , is noise from the whole system and includes the antenna Temperature (noise from the sky background, atmosphere, losses in the feed, spillover from the ground) plus the noise from the receiver system itself:

$$T_{SYS} = T_A + T_{RX} = T_{sky} + T_{atm} + T_{loss} + T_{spill} + T_{RX} + \ldots \qquad [13]$$

At centimetre, millimetre wavelenghts, T_{RX} dominates the system noise temperature.

When Penzias & Wilson (of the CMB) made their measurements, they found:

T_{atm} = 2.3 +/- 0.3 K,

T_{loss} = 0.9 +/- 0.4 K,

T_{spill} < 0.1 K.

And they expected $T_{sky} \sim 0$.

So looking straight up, they expected to measure T_A,

$T_A = 2.3 + 0.9 + 0.1 + 0 = 3.2$ K.

What they found was T_A = 6.7 Kelvin!

The excess was the CMB and Galactic emission.

Bell lab advert (right) - 1963 - 3 years before the CMB was detected - and featuring the Penzias & Wilsons horn antenna.

Typical values of contributions to Tsys at cm wavelengths.

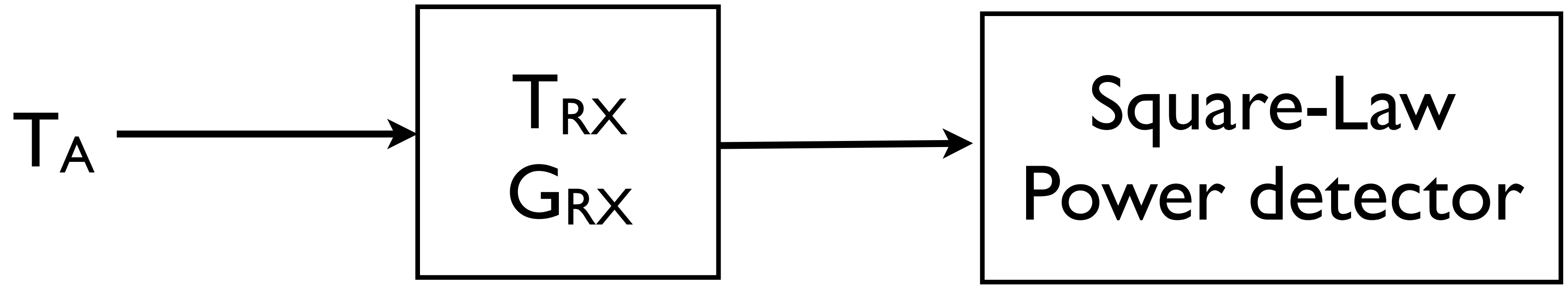
l	T_{sky}	T_{atm}	T_{spill}	T_{loss}	T_{cal}	T_{rx}	T_{sys}
92 cm	25	3	15	7	5	70	125
20 cm	3	3	14	8	2	30	60
6 cm	3	3	7	5	2	30	50
3.6 cm	3	2	5	2	2	16	30
2 cm	3	8	6	13	6	80	116
1.3 cm	3	17	6	21	7	100	154

T_{sky} is ~ 2.7 Kelvin (the CMB signal) at cm wavelengths, but at lower frequencies, radio emission from the Milkyway becomes increasingly strong (T_{sky} ~ 2000 Kelvin at 70 MHz).

T_{atm} is ~ 3 Kelvin at cm wavelengths but this increases as one moves to higher frequencies e.g. T_{atm} ~ 20 Kelvin at 1 cm.

Receiver Performance

Simple receiver system:



Consider an incoming signal with a noise temperature of T_A, the receiver output is:

$$P = k dv G(T_A + T_{RX}) \quad [14]$$

where:

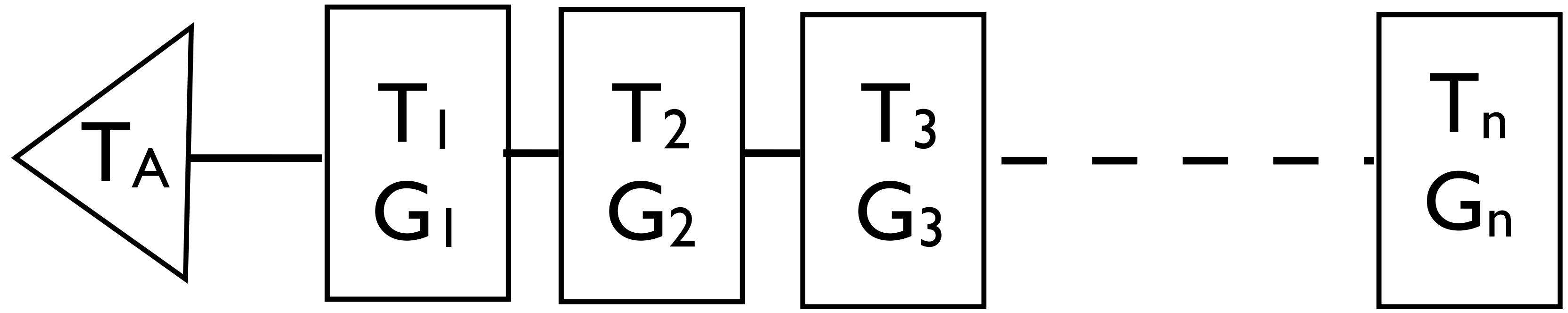
P = the output power in Watts
dv= Bandwidth (Hz)
G = receiver Gain
T_A = the antenna temperature;
T_{RX} = the receiver noise temperature;
k = Boltzmann's constant, 1.38E-23 Joule/Kelvin

e.g. If dv=50MHz, and the receiver Gain is 100db, and T_{sys} is 40K, then the output power of the receiver is:

1.38E-23 x 50E6 x 1E10 x 40 = 1.6E-6 =2.76E-4 Watts.

This is a fairly typical power output in radio astronomy. The result is often expressed in dbm: e.g. 0.27 mW (-5.6 dbm).

Consider a multi-stage receiver system:



Receivers usually have more than 1 amplification stage and also include "passive" devices e.g. cables. Some "lossy" components may gave $G < 1$.

In any case, the power at the output of the receiver given above is just:

$$P = kdv\,(T_A G_1 G_2 G_3 \ldots G_n + T_1 G_1 G_2 G_3 \ldots G_n + T_2 G_2 G_3 \ldots G_n + T_3 G_3 \ldots G_n + \ldots)$$

where T_A is the antenna temperature; T_n is the noise temperature of each stage in the reciever chain, and G is the gain.

This expression can be re-written as:

$$P = kdvG_1 G_2 G_3 \ldots G_n (T_A + T_1 + T_2/G_1 + T_3/G_1 G_2 + \ldots + T_n/G_1 G_2 G_3 \ldots G_{n-1}) \quad [15]$$

Or $P = kdvG(T_A + T_{RX})$ where T_{RX} is the overall receiver noise temperature (see [14]).

For a general receiver system with n stages, the total system temperature of the receiver is (see from eqn [14] and [15]):

$$T_{RX} = T_1 + T_2/G_1 + T_3/G_1G_2 + ... + T_n/G_1G_2G_3... G_{n-1} \quad [16]$$

N.B. the implication is that the noise properties of the first amplifier (or indeed component) come to dominate the overall system temperature! It's therefore important that the first amplifier is as good as possible. State-of-the art noise values ~ 10 Kelvin.

All components in a receiver generate noise - even "passive" devices such as transmission lines, couplers, cables, etc.

These components are not perfect and attenuate the signal. The Gain, G, of such a device is always less than 1, and the associated "Loss Factor" or attenuation, L, is defined as

$L = 1/G > 1$

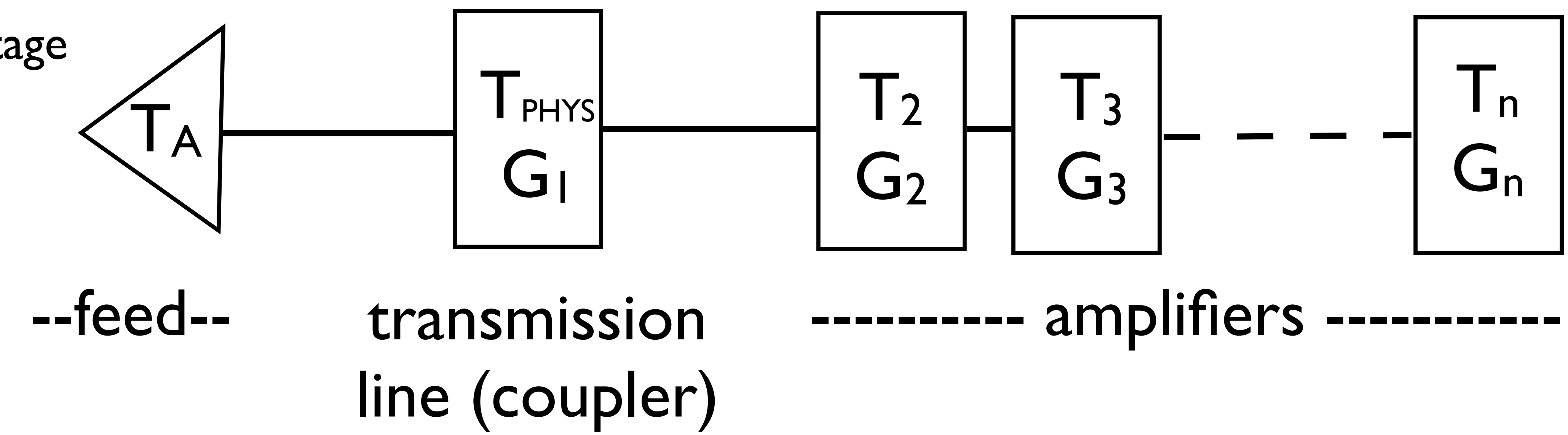
In addition, "lossy" components also add their own noise to the system:

$T_L = T_{phy} (L-1)$

where T_{phy} is the physical temperature of the lossy device, usually room temperature for cables etc.

Note that if this lossy component is used before the 1st stage of amplification, its noise characteristics are also important in terms of achieving good noise performance.

Another multi-stage receiver system:



In the scenario above, the transmission line introduces a system loss (G <1), acting as an attenuator to the signal. Noting that L=1/G and the form of equation [16] the receiver noise temperature is given by:

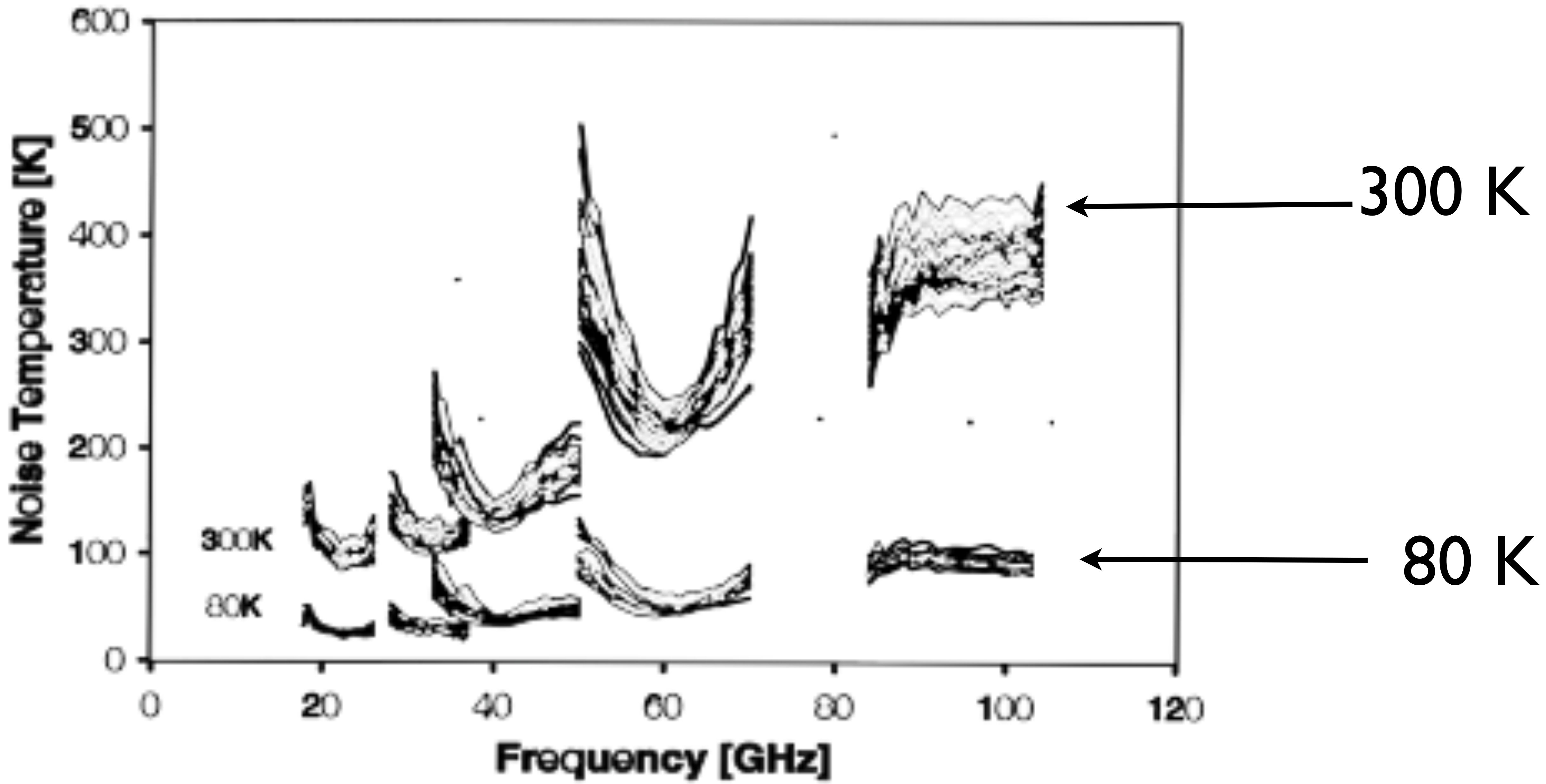
$$T_{RX} = (L-1)T_{phys} + LT_2 + LT_3/G_2 + \ldots LT_n/G_2G_3\ldots G_{n-1} \qquad [17]$$

e.g. if the room-temperature transmission line above has a loss of 0.4dB, and there are 2 stages of amplification with noise temperatures/gains of 45K/13dB and 140K/13 dB, and assuming the rest of the receiver has a noise temperature of 800K, then...... the overall receiver noise temperature is (via [17]):

$$T_{RX} = (1.1-1) \times 290 + (1.1 \times 45) + (1.1 \times 140/20) + 1.1 \times 800/(20\times 20) =$$

$$29.0 + 49.5 + 7.7 + 2.2 = 88 \text{ Kelvin}$$

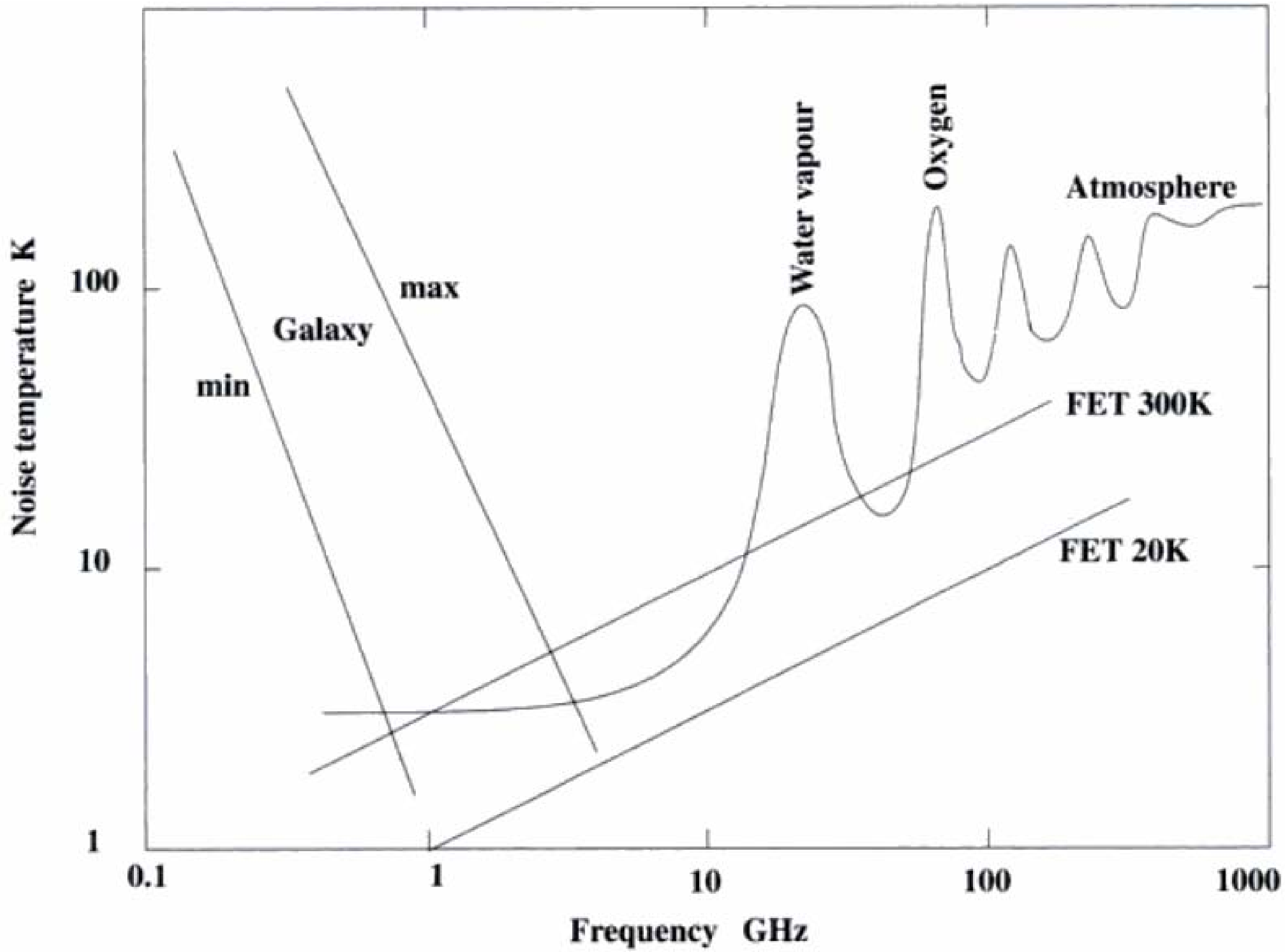
Note in the previous calculation that the noise temperature of the first stage (including lossy components) is all important. Note also that the loss term ahead of the first amplifier contributes to every following stage, not just the first, so it should be as small as possible. Often , the front-end of most radio astronomy receivers will be cooled by liquid helium.



Data courtesy M. Pospieszalski of NRAO Central Development Laboratory

Above: Measurements of 5 different LNA batches at 300 Kelvin (room temperature) and 80 Kelvin. The noise typically increases with rising frequency and cooling temperature.

Many receivers boast noise figures that are comparable or below the sky background. Note that at low frequencies (< ~1 GHz) there is not much point in cooling receivers - relatively cheap (i.e. noisy!) amplifiers can be used.



Field Effect Transistor (FET) receivers, cooled and uncooled.

Telescope Performance - figures of merit

The sensitivity (or "gain") of an telescope is often measured in K/Jy, and usually changes with observing frequency. This is also called the telescope DPFU (Degree per Flux Unit).

e.g. a typical 25-metre telescope has a DPFU of 0.1 K/Jy. i.e. when a 1 Jy source enters the beam of a telescope, T_A rises by 0.1 Kelvin.

Another measure of sensitivity is the SEFD (System equivalent Flux Density).

This is the ratio of the TSYS to the DPFU with units Jy:

$$\text{SEFD} = \text{TSYS}/\text{DPFU} \quad [20]$$

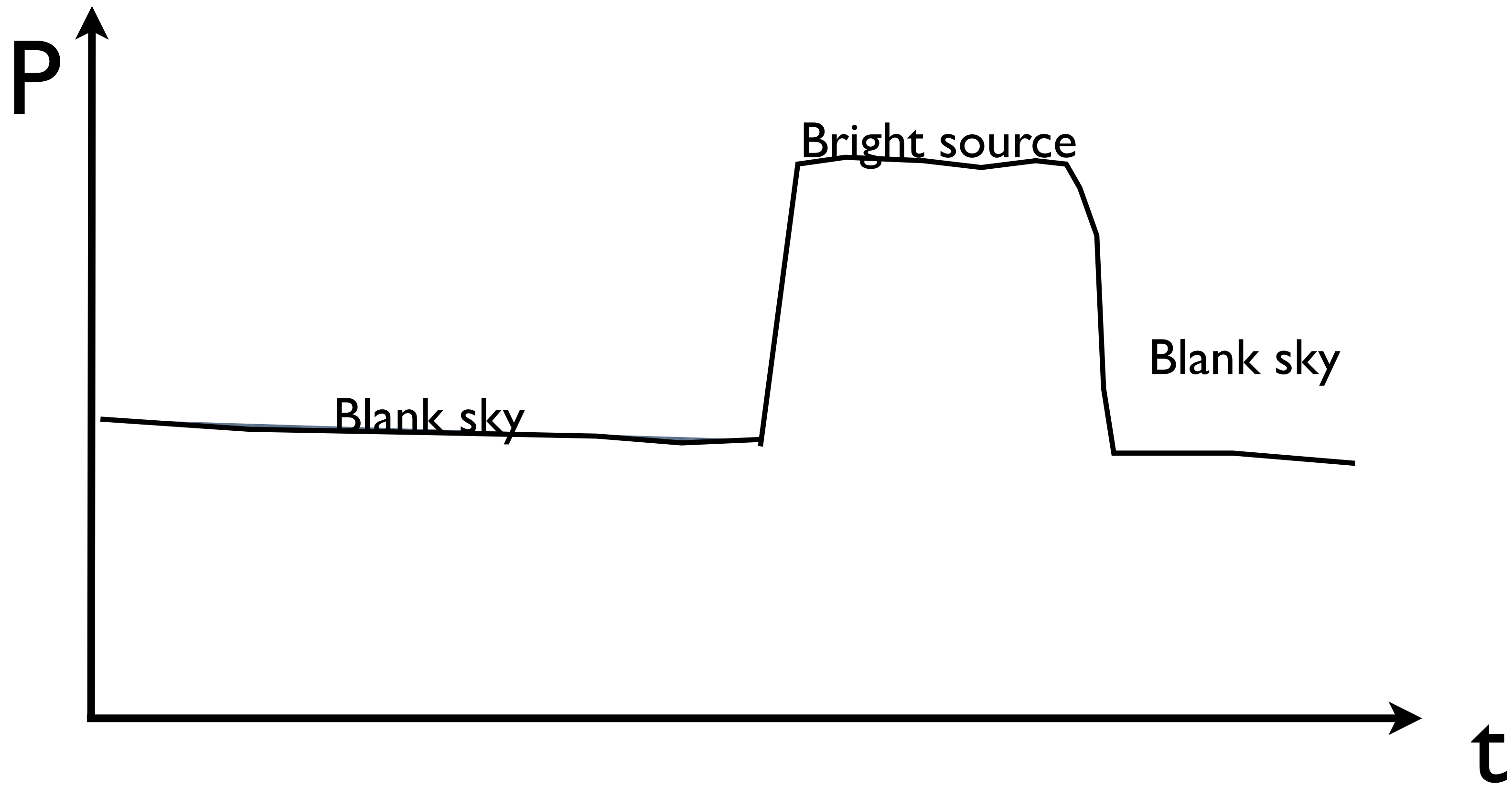
e.g. the Effelsberg 100-m telescope has an DPFU of 1.4 K/Jy and a TSYS (at 18 cm) of about 30 Kelvin. Think of the SEFD as the total system noise (power) of the telescope. For Effelsberg the SEFD is ~ 20 Jy.

For comparison the SEFD of a single WSRT antenna is ~ 300 Jy. and for a 6-metre ATA telescope is ~ 6000 Jy. At the other end of the scale, the Arecibo telescope has an SEFD of 3Jy.

N.B. In the case of the SEFD - "small is beautiful"; in the case of the DPFU - "big is beautiful".

Measuring the SEFD of a real radio telescope (in practice!)

The SEFD of a radio telescope can be measured by switching the telescope between a very bright source (of known flux density, S) and an empty patch of blank sky.



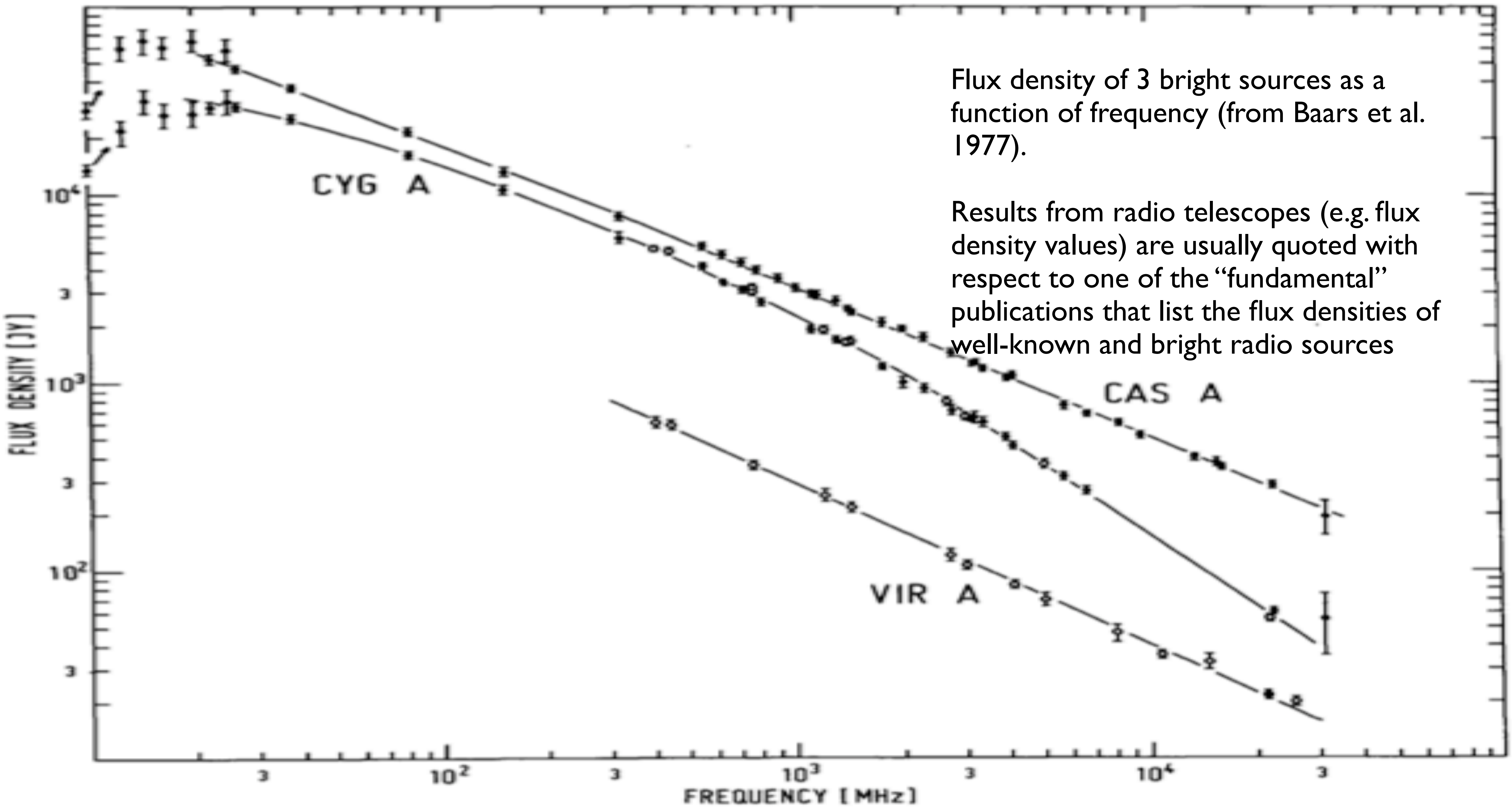
If we then compare the measured power output levels of the receiver then:

$$Y = \frac{P_{on\_source}}{P_{blank\_sky}} = \frac{G\, dv\, (T_{sys}+T_{bright\_src})}{G\, dv\, T_{sys}} = \frac{T_{sys} + T_{bright\_src}}{T_{sys}} = 1 + \frac{T_{bright\_src}}{T_{sys}}$$

Recalling equation:

$$\mathrm{SEFD} = T_{sys}/\mathrm{DPFU} = T_{sys}/(T_{bright\_src}/S) = S/(Y-1)$$

Careful, fundamental flux density measurements of the brightest sources in the sky have been made (e,g. Baars et al. 1977, Ott et al. 1994) using very well calibrated telescopes over a wide range of frequencies:



Flux density of 3 bright sources as a function of frequency (from Baars et al. 1977).

Results from radio telescopes (e.g. flux density values) are usually quoted with respect to one of the "fundamental" publications that list the flux densities of well-known and bright radio sources

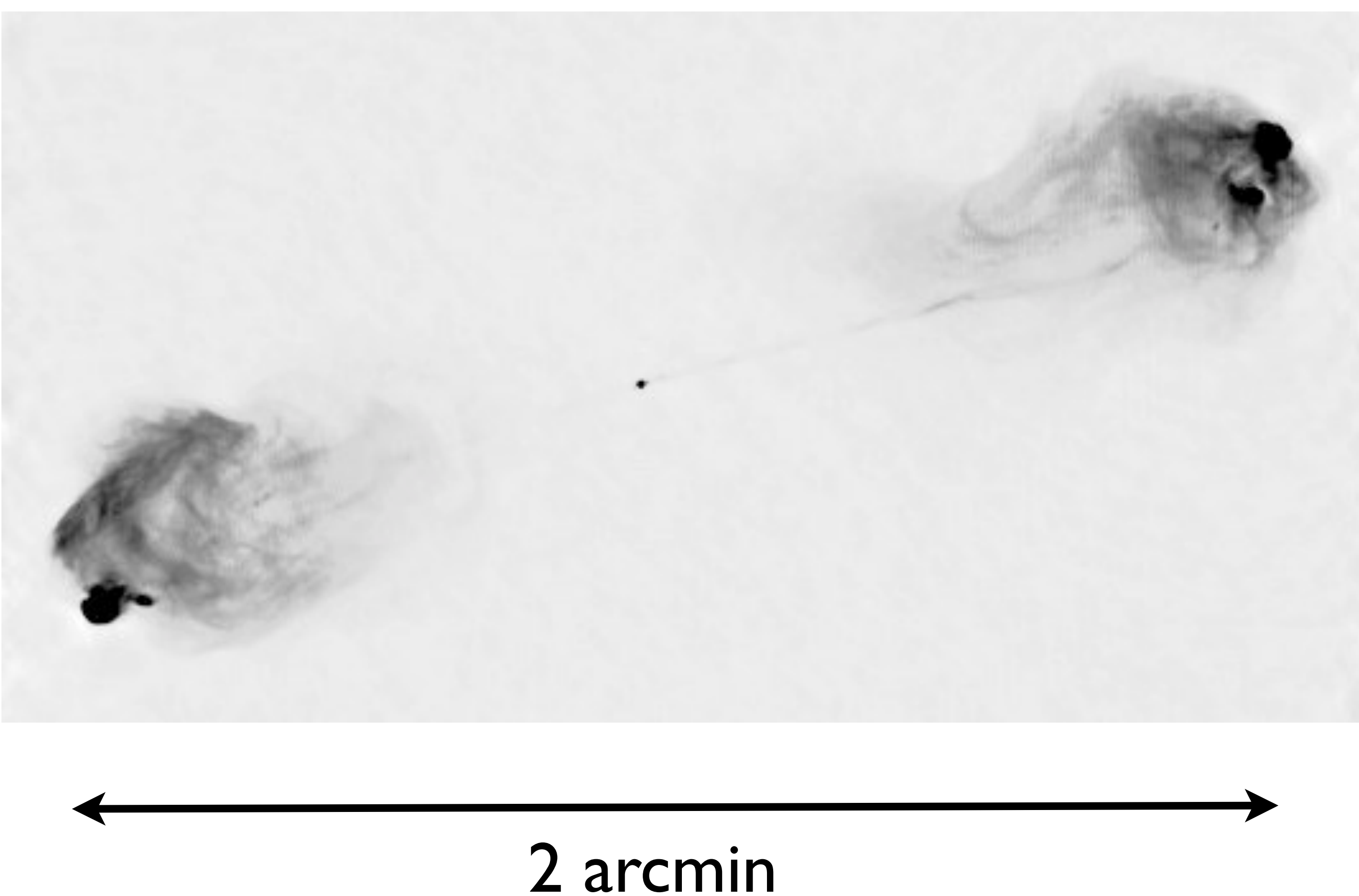
Fig. 1a. The absolute spectra of Cas A and Cyg A and the semi-absolute spectrum of Vir A. Solid symbols are absolute measurements, open symbols relative measurements

The telescopes SEFD should be known in order to place single-dishes and interferometer arrays on an absolute flux density scale.

Ideally SEFD "calibrator" sources should be unresolved, non-variable and bright.

Some sources that are appropriate calibrators for small telescopes, are inappropriate telescopes for large telescopes.

e.g. Cyg-A will be unresolved for a small telescope but extended for a large telescope e.g. Effelsberg. In addition, Cyg-A is so strong that it may itself saturate the receiver of a large telescope, pushing the receiver into the non-linear regime.



At a wavelength of 3.6cm:

D=10 metre => 15 arcmin beam

D= 25 metre => 6 arcmin beam

D=100 metre => 1.5 arcmin beam

Sensitivity of a radio telescope

Recall that a 100 mJy source raises the T_A by only 0.023K. How do we measure this against a system noise which is at least a factor of 1000 greater ?

One way is to average together many telescope samples - since the noise is random it is uncorrelated between one sample and another. The source signal is correlated however. In this way we can "beat down" the noise level.

Consider Nyquist sampling of a passband (dv) over an integration time, τ . In this case we will obtain $2d\nu\tau$ independent samples.

For Gaussian noise, the uncertainty ΔT in the measurement of T_A will reduce with the square-root of the number of samples:

$$\Delta T = \frac{T_{tot}}{\sqrt{(2d\nu\tau)}}$$

e.g. for an integration time of 120 secs, a bandwidth of 20 MHz, and a T_{SYS} (T_{tot}) of 40 Kelvin, the uncertainty ΔT is ~ 0.0006 Kelvin. So a 100 mJy source can be detected with a signal-to-noise ration of ~ 40.

In most cases T_{Tot} is dominated by $T_{SYS.}$ This is not the case when we observe very bright sources such as the Sun - then the source itself dominates. At low frequencies the sky noise begins to dominate - in these cases the requirements on receiver performance (T_{RX}) can be greatly relaxed.

Receiver gain stability

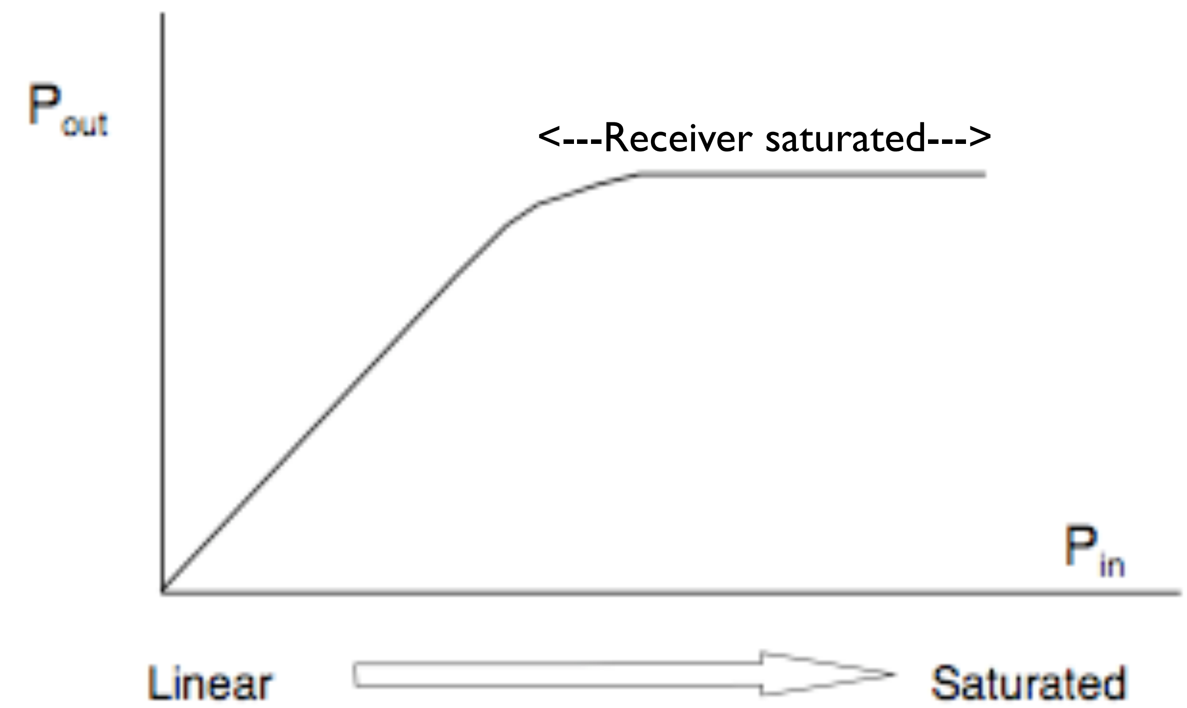
In practice the gain of the receiver changes with time and is not perfectly constant. This is a bit of a problem, the detector cannot distinguish between a change in signal power and a change in the gain of the receiver.

To understand the impact of receiver stability note that if our receiver has a system temperature of 40K, a 1% increase in the gain produces a ΔT of 0.4 Kelvin which is > 600x larger than the sensitivity we could reach (see previous example) using a perfectly stable system.

Note this implies that our receiver has a gain stability << 0.0015%. Receivers must be constructed very carefully with the best cables and components, and be placed in a temperature controlled environment in order to achieve this level of stability.

An important characteristic of a receiver is stability and linearity.

Gain stability is often dependent on maintaining a stable operating environment.



A receiver (amplifier) has a limited operating range in terms of the input power they can handle. At some input power level the response of the receiver becomes non-linear or saturated (sometimes referred to as the receiver being "compressed"). This can happen in the presence of powerful artificial signals e.g. man-made radio frequency interference (RFI).

In severe cases, "notch filters" are used to suppress specific frequency ranges where RFI is prevalent and strong e.g. local point-to-point communication lines, geo-synchronous satellites etc. This is required because when a strong RFI signal overloads a receiver, the receiver is compressed across the entire passband (bandwidth) making it unuseable throughout its entire frequency range.

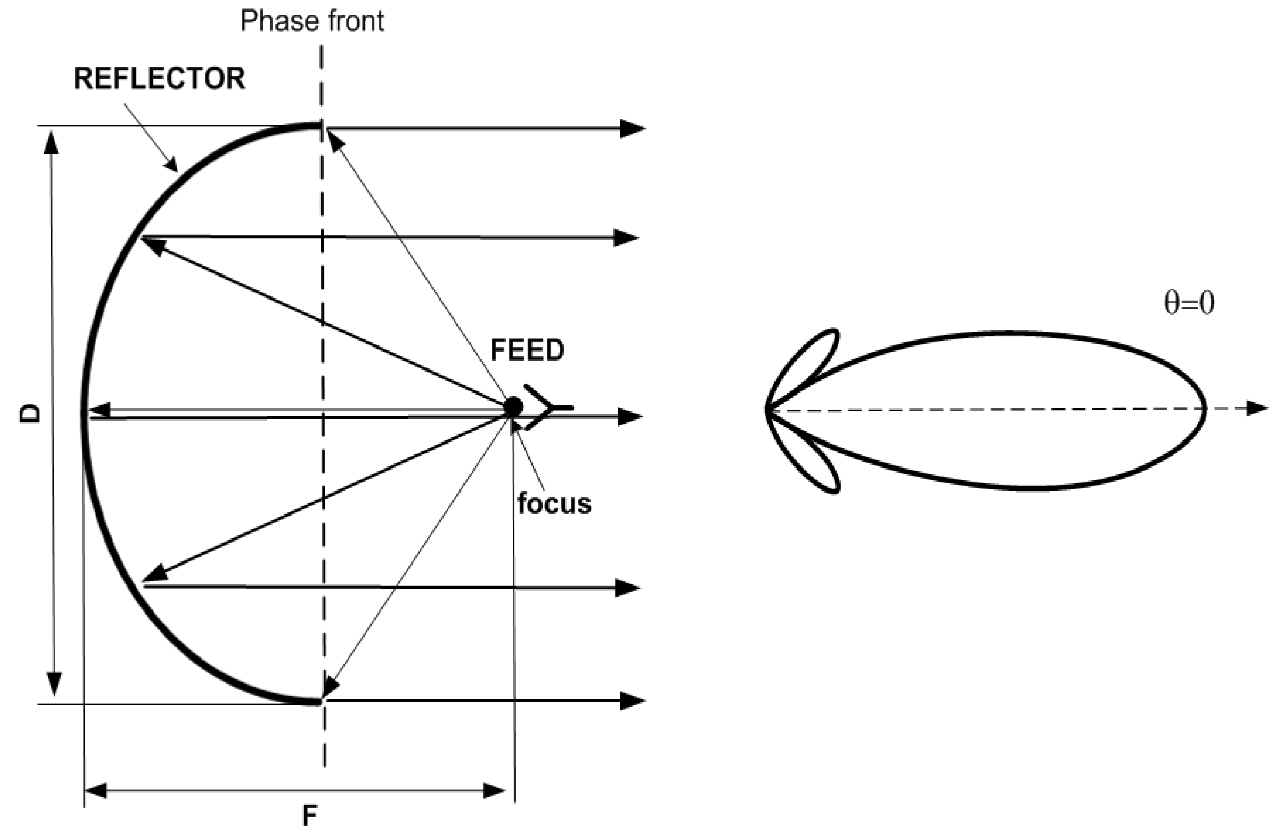
As a passive "service", Radio astronomers are waged in a constant battle with other users (and abusers) of the spectrum. Some protection is provided: e.g. 1421 MHz (neutral hydrogen).

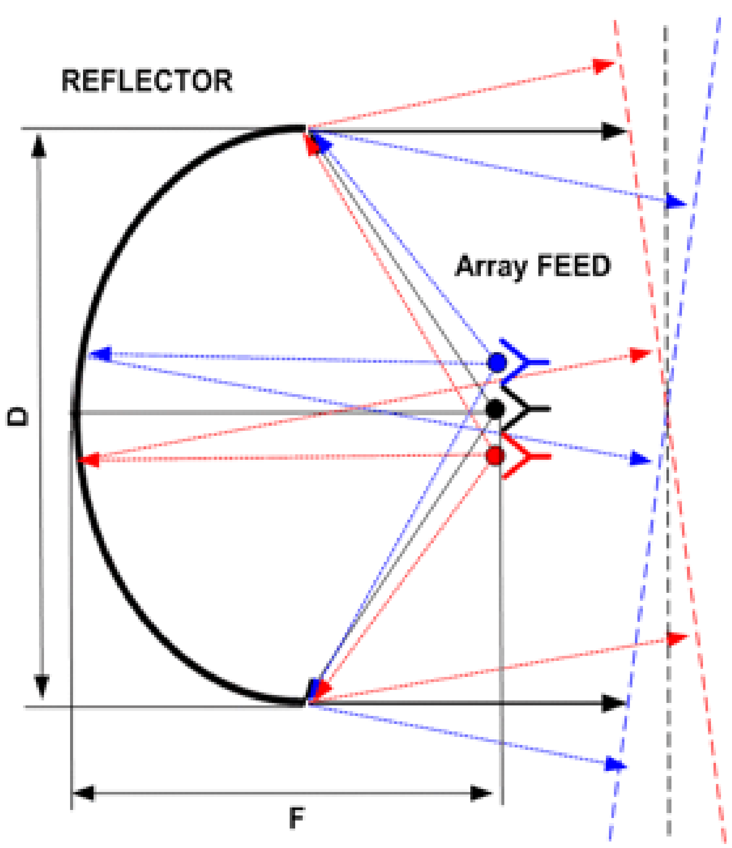
CRAF (right) is the Committee on Radio Astronomy Frequencies and does battle with the multi-national communications industry.

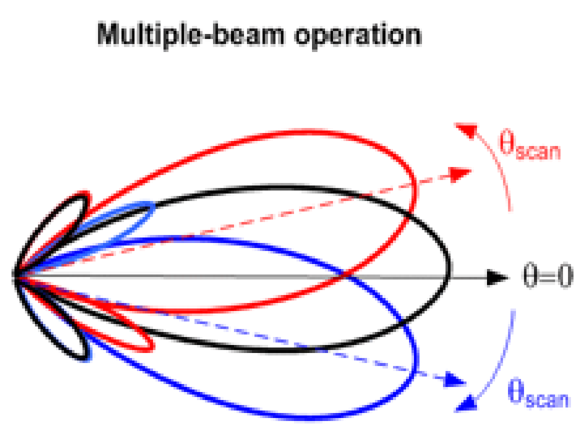
Will radio astronomy still be possible in Europe by 2020 ?

Multi-beam receiver systems

Standard single-beam ("single pixel") system:







Multi-beam system (left) - greatly expands a telescopes field of view - important for survey speed e.g. Parkes 64-metre 13 “beam” (feed) system:



e.g. Arecibo 7 “beam” (feed) system:



Aperture Arrays (AA)

At low frequencies (< ~ 1 GHz) Aperture Arrays make up a large physical aperture by combining (phasing) elements - e.g. (groups of) dipoles - spread over a large physical area.

Effective aperture:

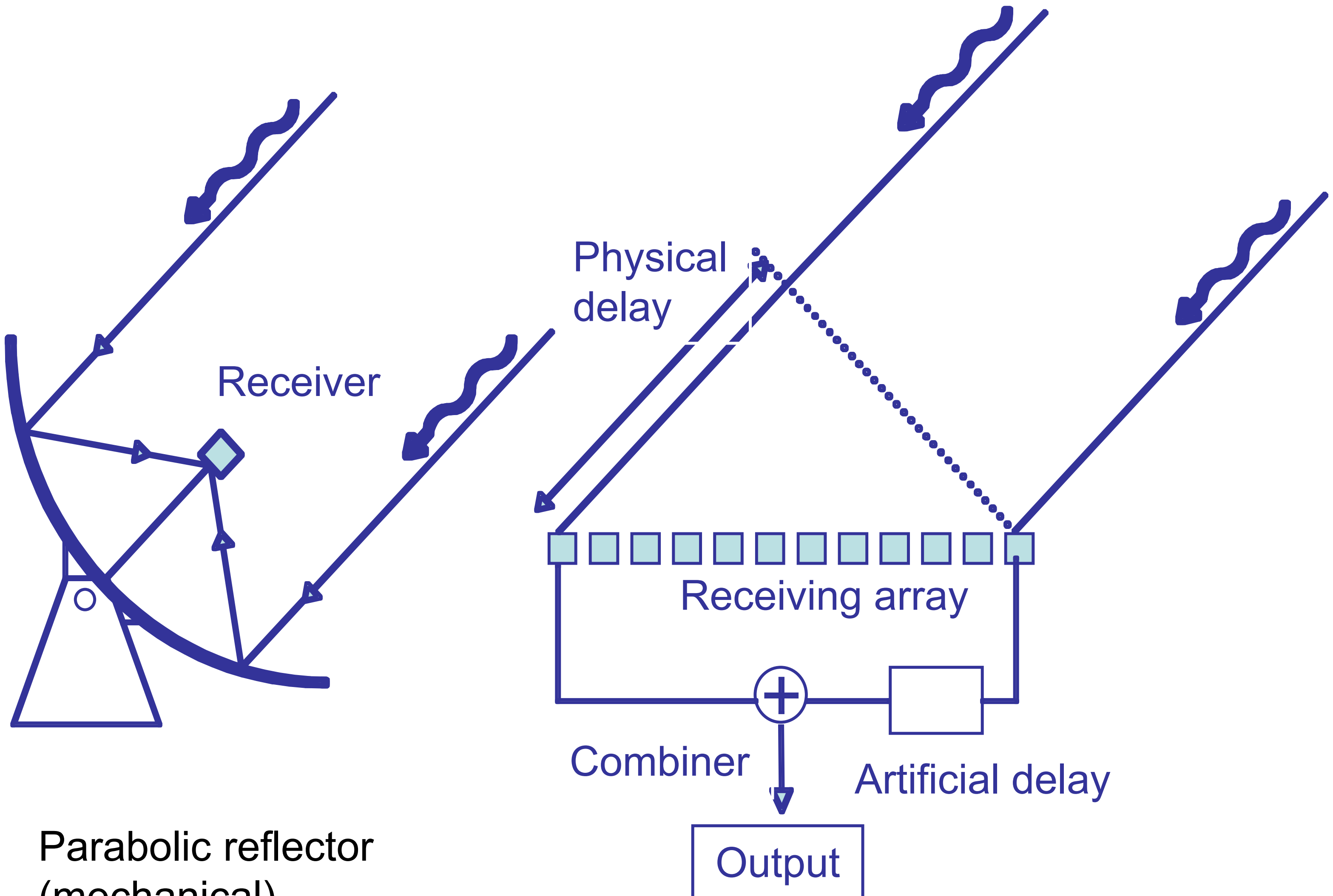
$$A_e \sim \lambda^2/\Omega_A$$

For an idealised isotropic beam pattern, $\Omega_A \sim 2\pi$. For a simple isolated dipole $\Omega_A \sim 3$, and for a dipole in a half-wavelength spaced dense array, $\Omega_A \sim 4$.

The result is that very large collecting areas can be achieved at low-frequencies using large arrays of dipoles arrays.

Reber's 2MHz Square km dipole array.

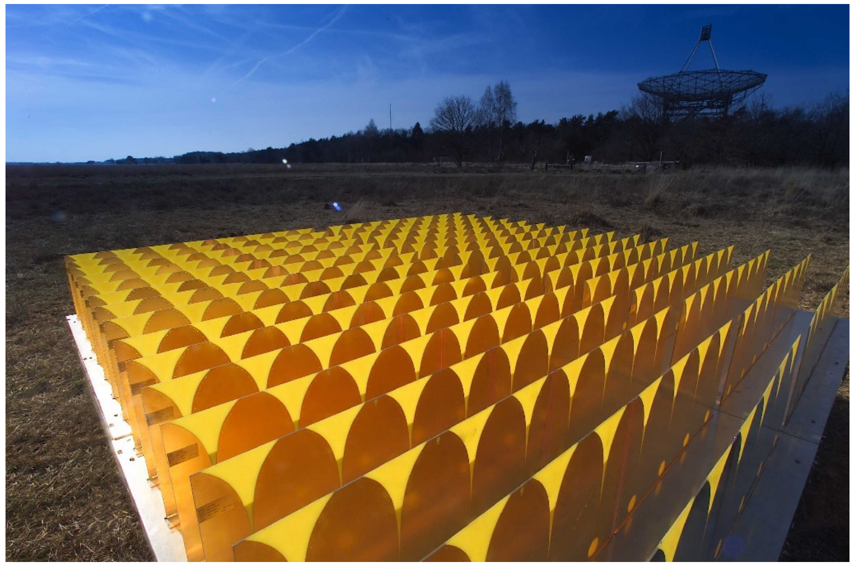
Recently, the combination of simple “front-end” collecting areas (e.g. dipole arrays) and state-of-the-art “back-end” data signal processing/super-computing has led to the development of distributed, low-frequency radio telescopes e.g. LOFAR, MWA, LWA etc.



Parabolic reflector (mechanical)

Aperture Array (electronic)

Test AA at ASTRON



Electronic components are used to add an appropriate delay to each elements signal so that they may be combined “in phase”. The electronics are used to focus the incoming wavefront, replacing the need for a parabolic reflector. This process is known as “beam-forming” - multiple beams can be created, only limited by digital electronics, processing power and costs.

Aperture array approach attempts to replace expensive steel parabolas with cheap off-the-shelf digital electronics and signal processing. This development represents a paradigm shift in radio astronomy technology.

Advantages of aperture arrays:

1) The effective area of an isolated dipole ~ $(\text{lambda})^2$. At low frequencies very large collecting areas can be built relatively cheaply. N.B. the cost of a parabola reflector scales as $D^{2.5\text{-}2.7}$

2) Aperture arrays are *very* flexible: it is possible to generate multiple-beams simultaneously - number of beams is only limited by available electronics (but see point 2 below).

3) Aperture arrays have a LARGE field of view - important for survey science.

4) AAs have no moving parts, and take advantage of commercial investments and Moore's law/ mass production of electronic components.

5) If the data from each dipole is buffered, we can re-point an aperture array telescope at a particular point in the sky in the past ! e.g. before a triggered event occurs.

Aperture array approach attempts to replace expensive steel parabolas with cheap off-the-shelf digital electronics and signal processing.

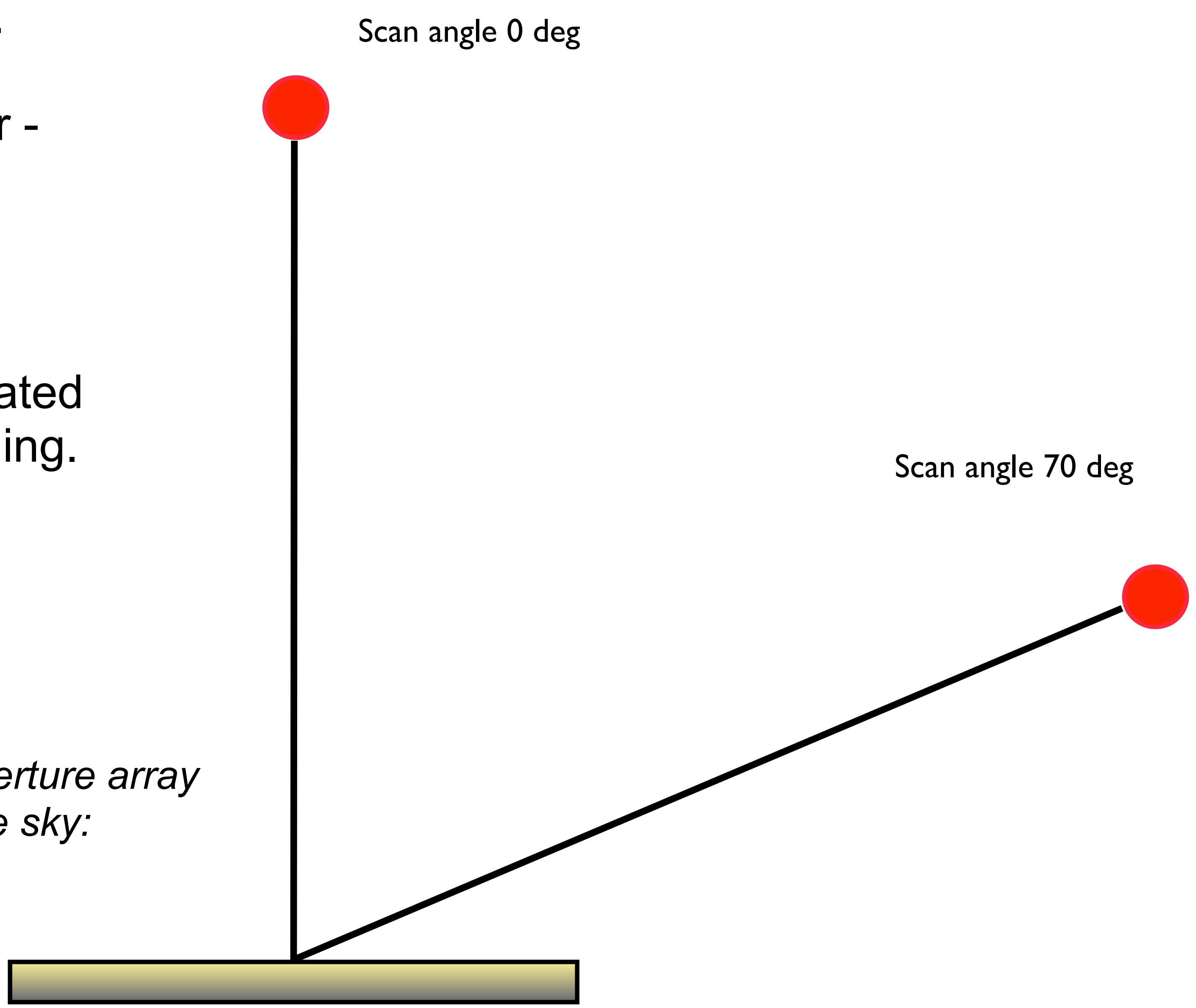
Disadvantages of aperture arrays:

1) Scan angle dependance (“foreshortening”): effective area scales as Cosine (scan angle).

2) Rely on digital electronics - requires power - trade rising costs of steel (for reflectors) with rising energy costs.

3) Still to demonstrate that AA approach can generate the very stable electronically generated beams required for high dynamic range imaging.

Effective area and beam shape of an aperture array changes as the source moves across the sky:

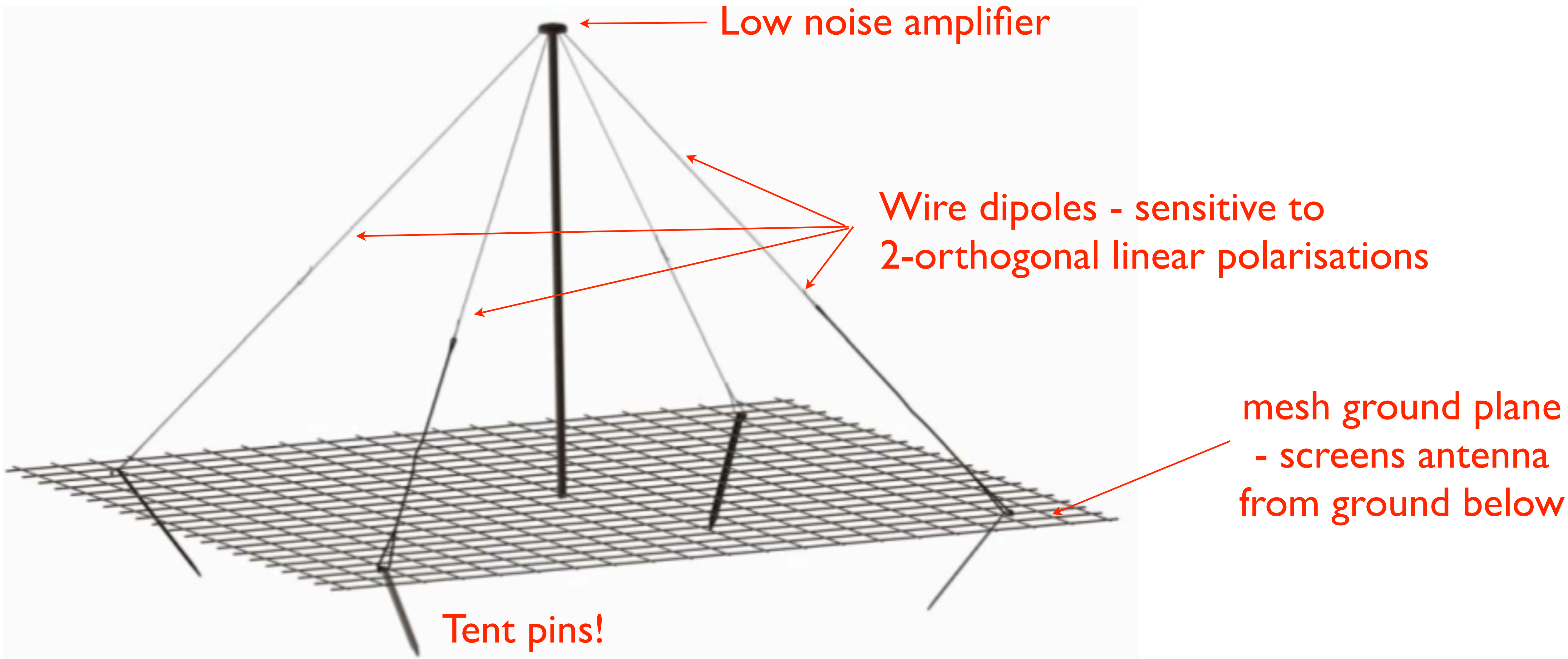


Traditional paraboloid and LOFAR aperture array telescopes side-by-side at Effelsberg, Germany.

LOFAR - an aperture array telescope

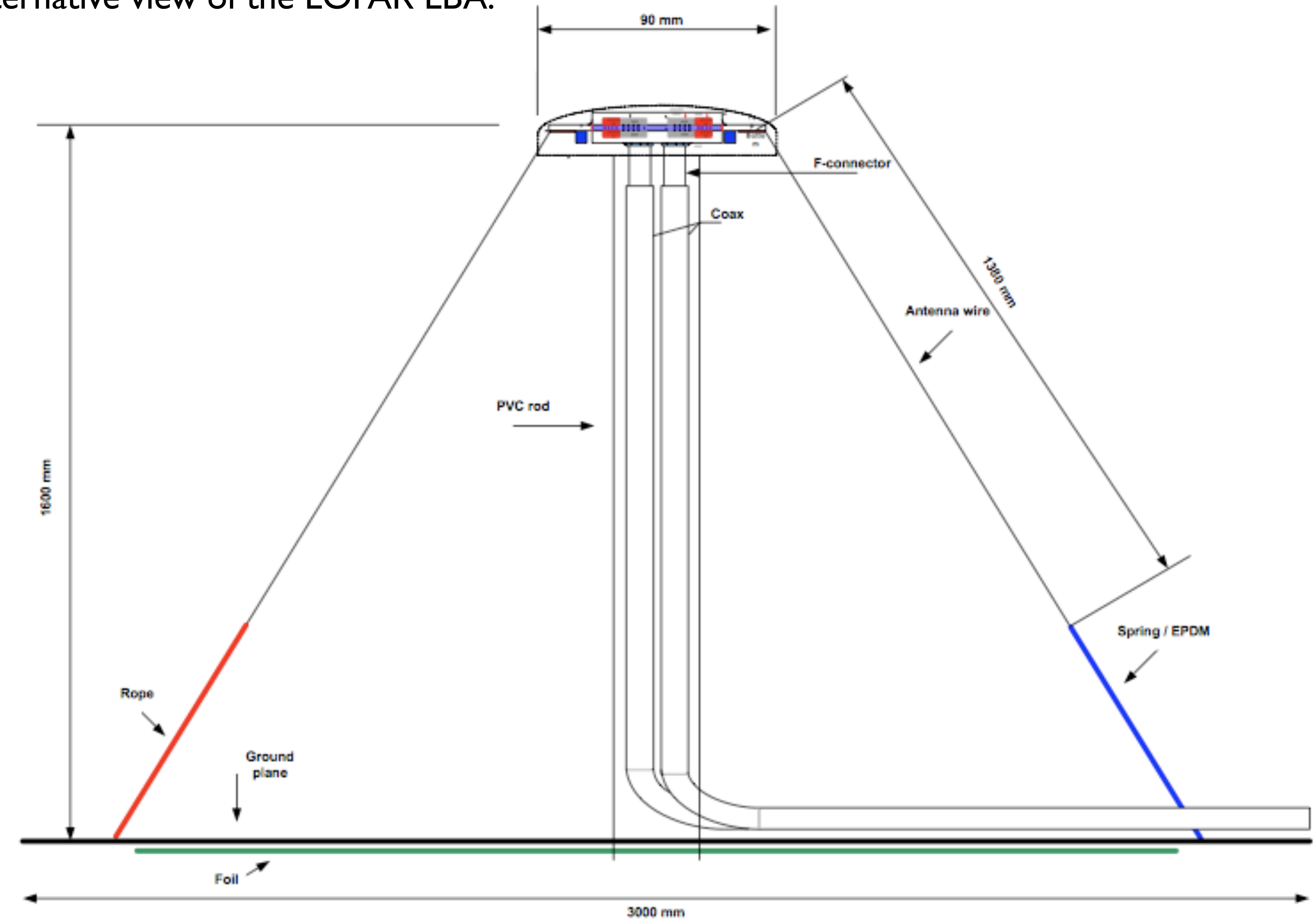
LOFAR uses two types of antenna - Low Band Antenna (LBA - 15-80 MHz) and High Band Antenna (HBA - 120-240 MHz). We focus here on the simplest - the LBA.

The LBA antenna is a simple "inverted V" (sometimes referred to as a "droopy dipole"):



There are 96 of these antennas in each LOFAR "station".

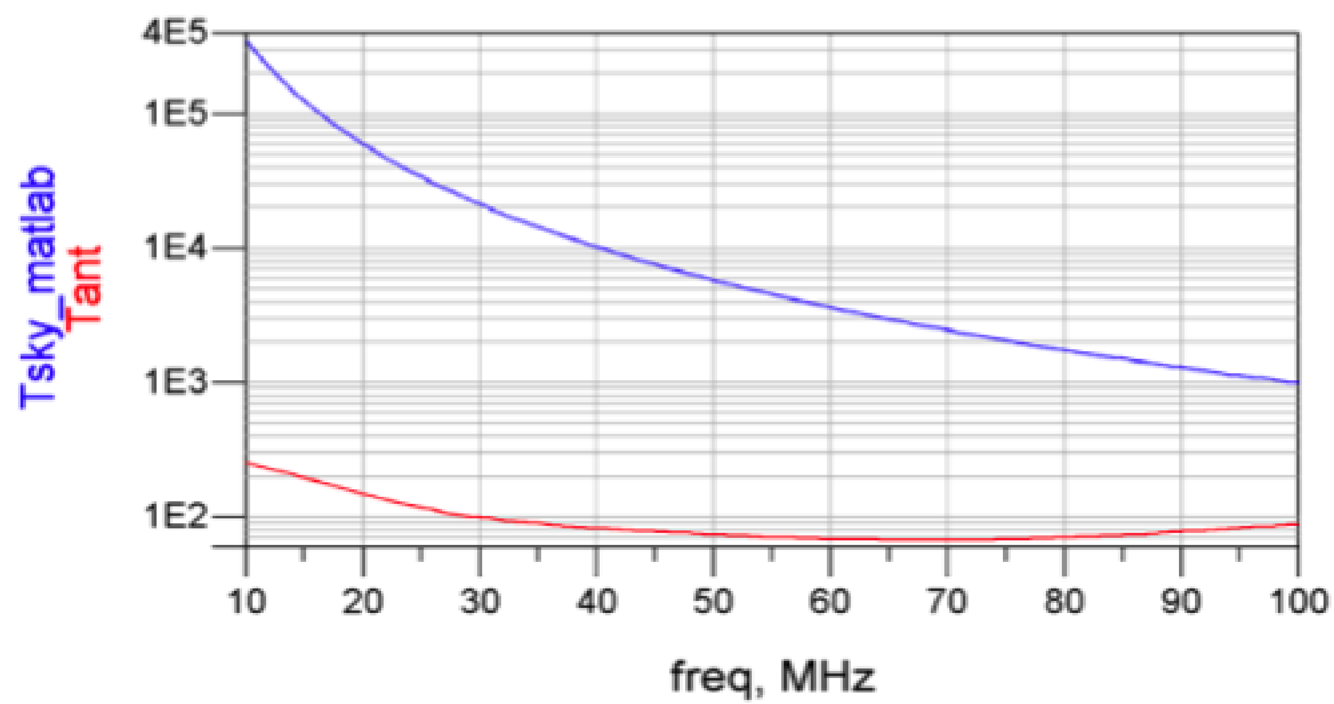
An alternative view of the LOFAR LBA:



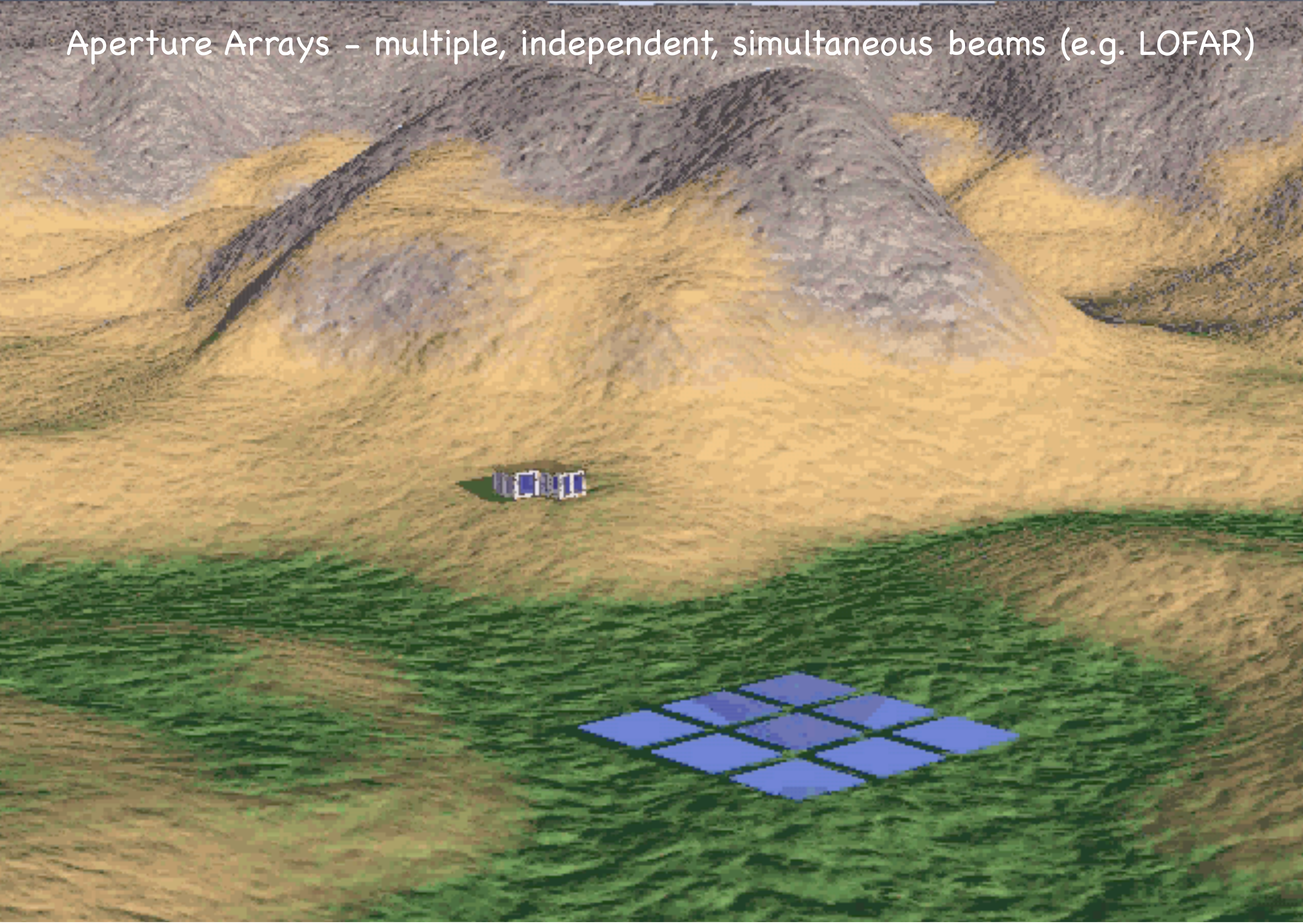
The central PVC rod houses coax-cables that connect the 2 (polarisation) analogue outputs of the Low-noise amplifier (LNA) to the central Receiver Unit (RCU). The coax cables also deliver power to the LNA.

The receiver unit (RCU) converts the signals into the digital domain and splits them up into a set of narrow band (256 kHZ) output signals for each polarisation). These subbands are converted into station beams by a "Beam Former". Multiple beams can be formed. The total bandwidth (including all beams) is 48 MHz. For LOFAR, the resulting multiple beam signals (up to 8) are transported via the LOFAR optical fibre Wide Area Network (WAN) towards the CEntral Processor (CEP) in Groningen.

The LBA is "sky noise limited" i.e. the noise associated with the antenna and receiver system is << the sky noise (see right). Since the sky noise is very bright at low frequencies (mainly due to the galaxy) this is relatively easy. LOFAR can therefore employ very simple low-cost technology, enabling large station arrays to be built cheaply (one LOFAR station < 1M€).



Aperture Arrays - multiple, independent, simultaneous beams (e.g. LOFAR)

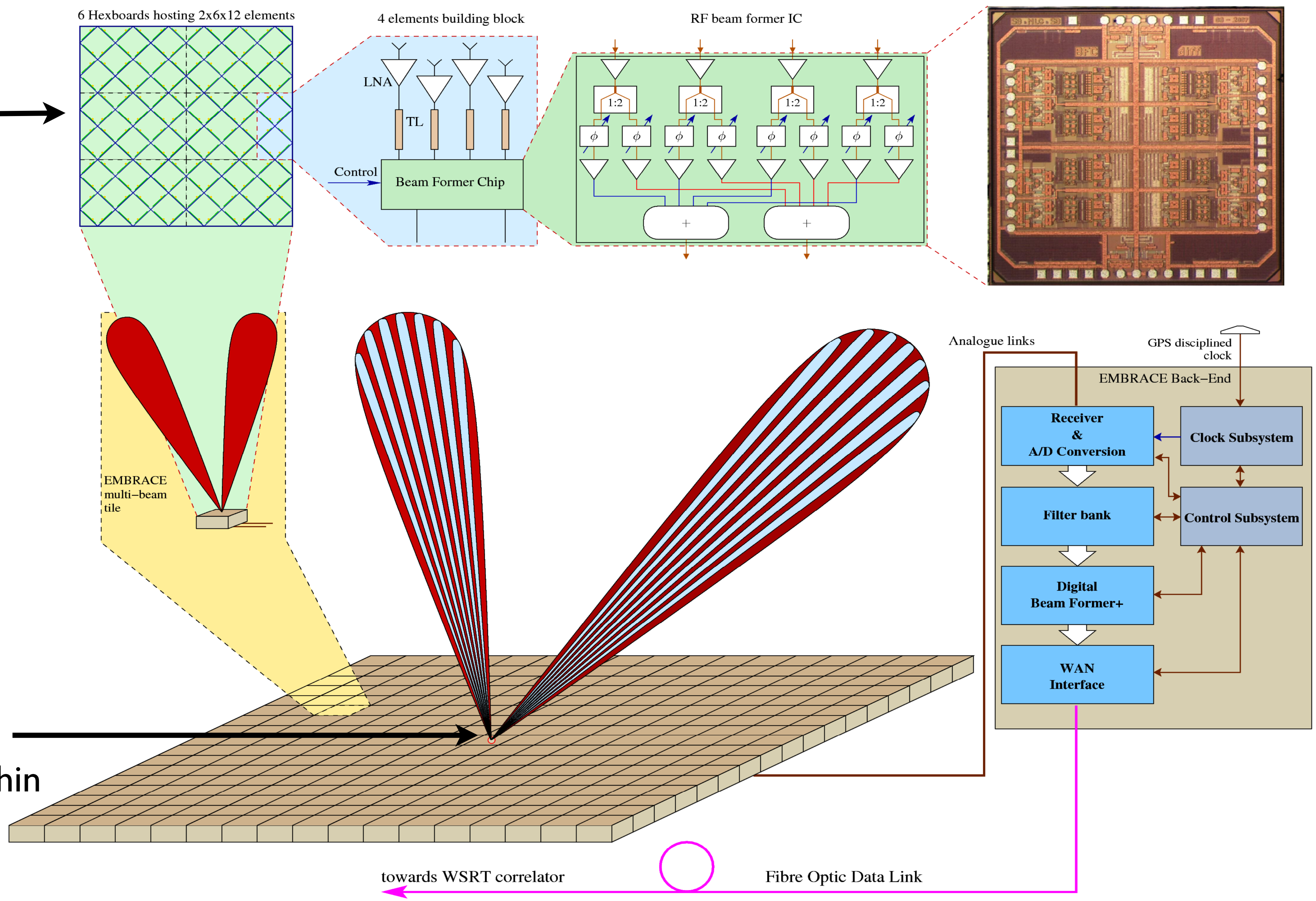


Aperture Array Prototypes

Aperture Array prototypes are currently being developed to operate at higher frequencies than LOFAR (e.g. in the range of 0.5 - 1.5 GHz. The best example is EMBRACE which is a European project led by ASTRON:

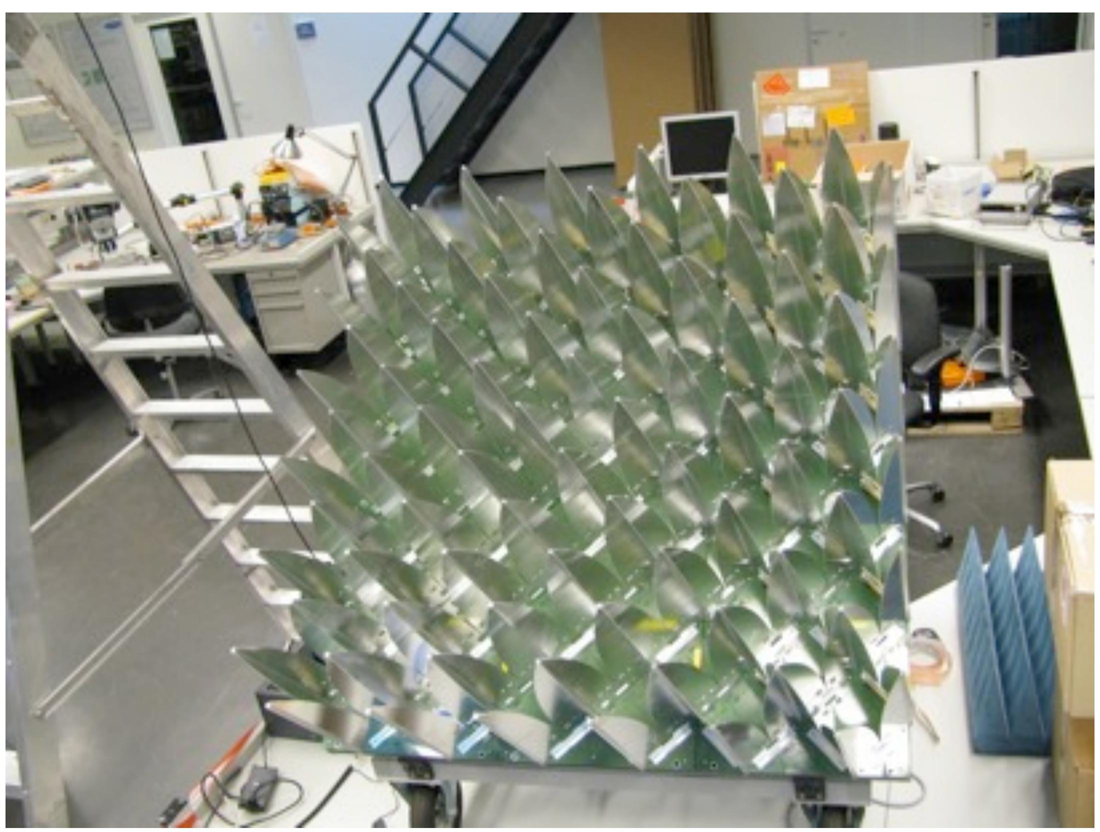
Single tile

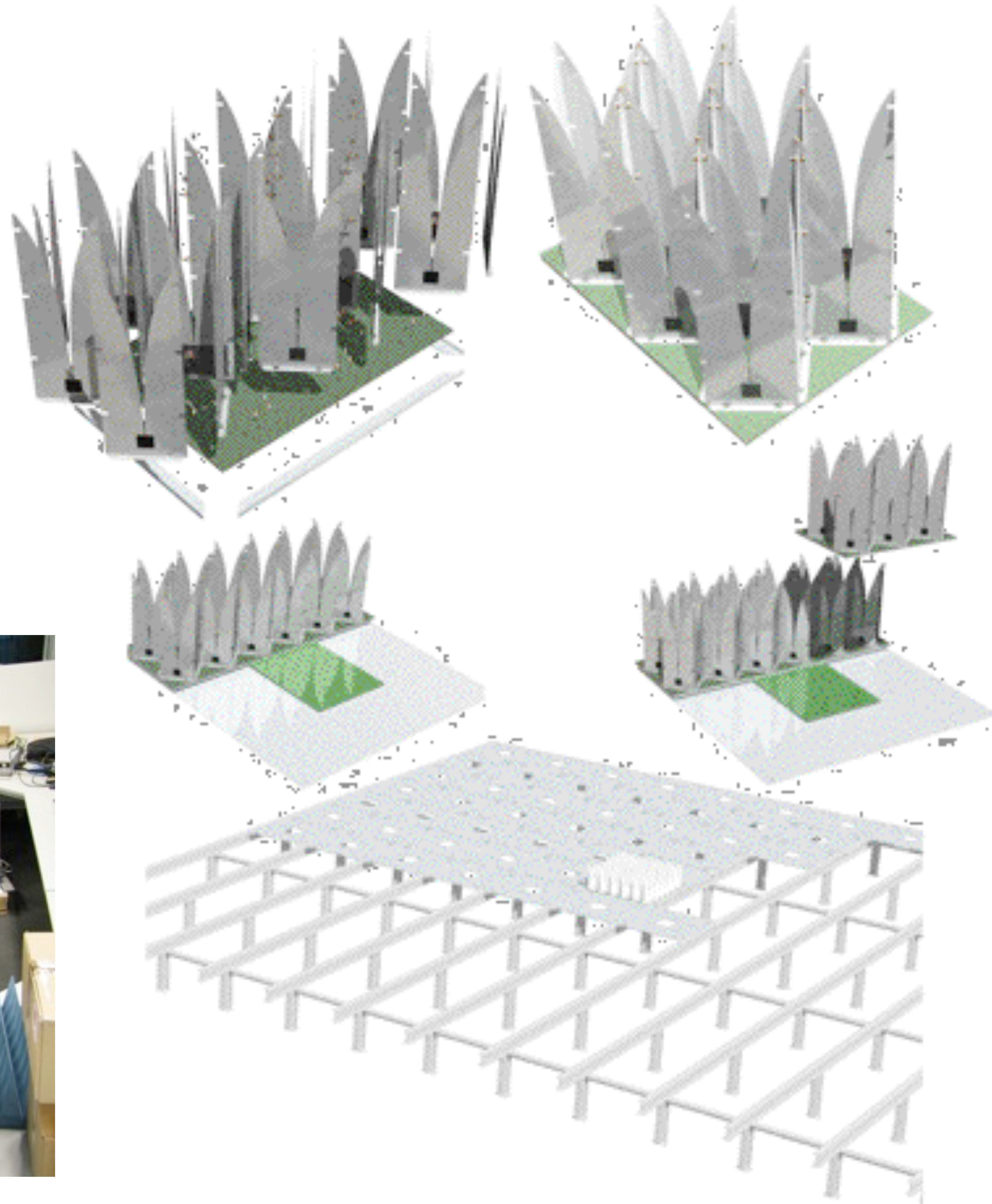
2-independent fields of view - many-beams within each FoV



EMBRACE parameters:

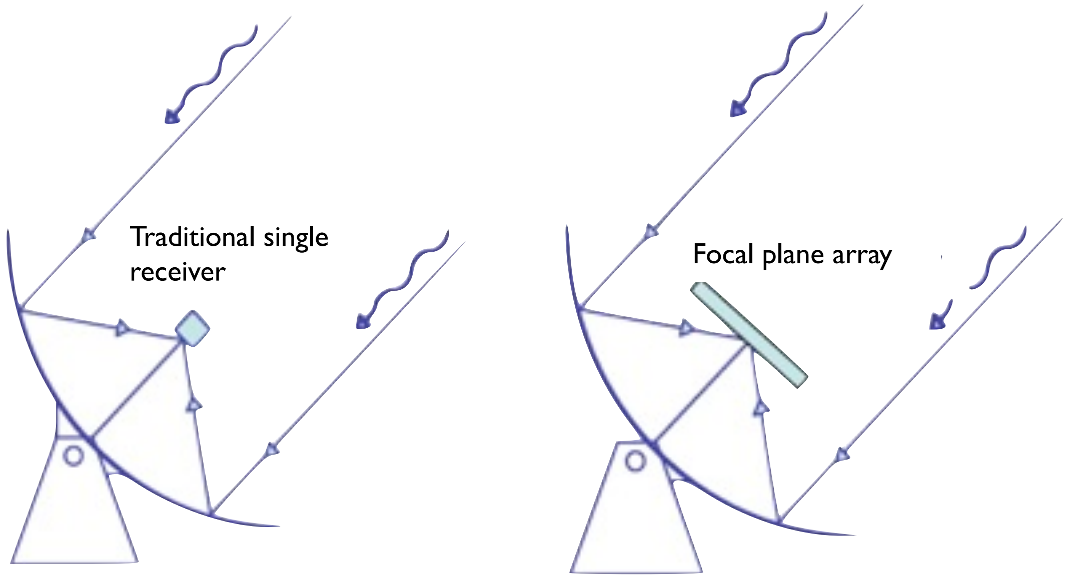
- Freq range: 500-1500 MHz
- Polarisation: single
- Collecting area: 300 sq metres
- Aperture efficiency > 80%
- Electronic scan range +/- 45 deg
- Tsys < 100 Kelvin
- Instantaneous bandwidth ~ 40 MHz
- Dynamic range A/D convertor: 60db

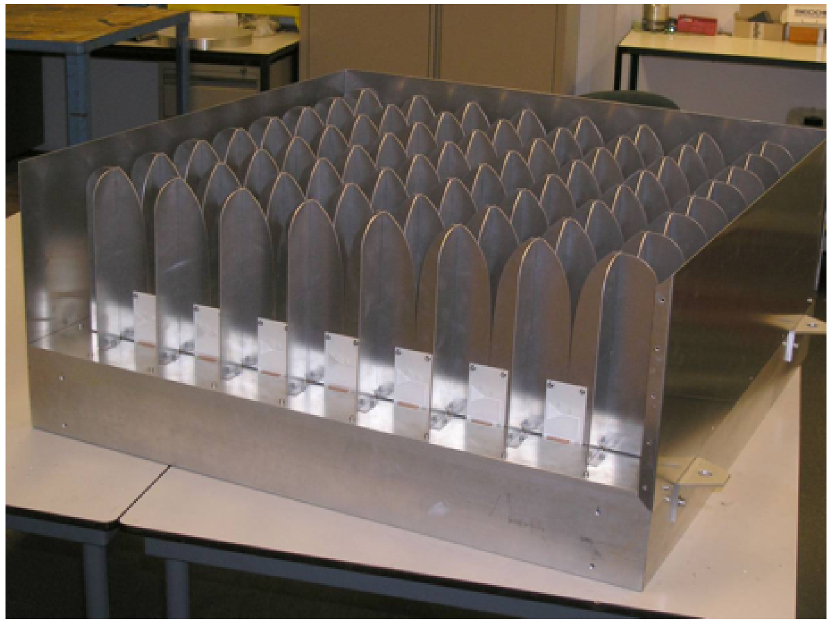




Focal Plane Arrays (FPA)

In Focal Plane Arrays (FPA) systems a 2-dimension array of recievers is created around the focus of a telescope by combining many simple antenna elements together and weighting them optimally:





A focal plane array system in the lab (ASTRON)

The goal is to form many beams on the sky.

This greatly increases the telescopes field of view.

Huge technology step forwards (from a single pixel to a CCDlike camera array)

A focal plane array demonstration system (Digestif) installed on one of the telescopes at WSRT. (ASTRON)

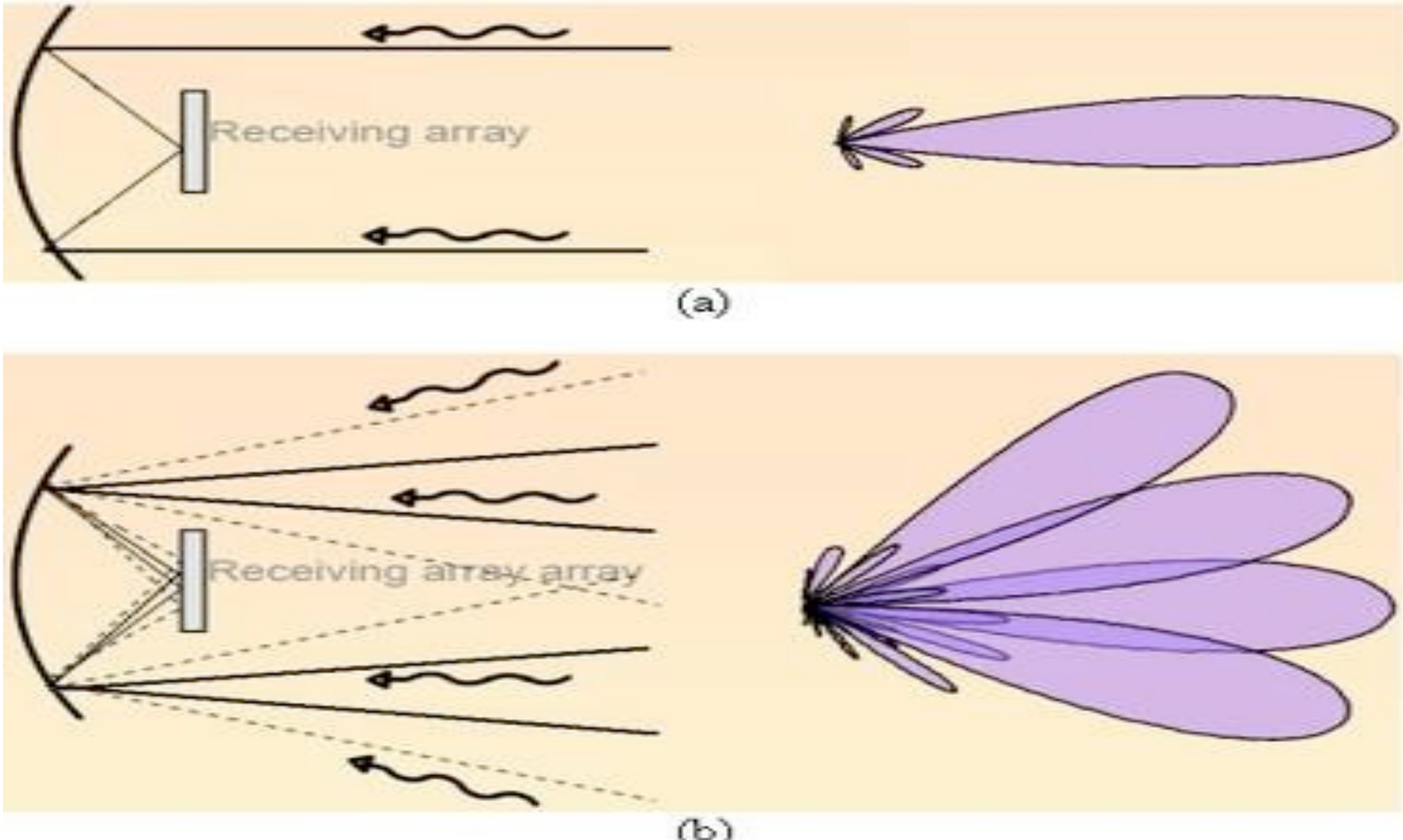
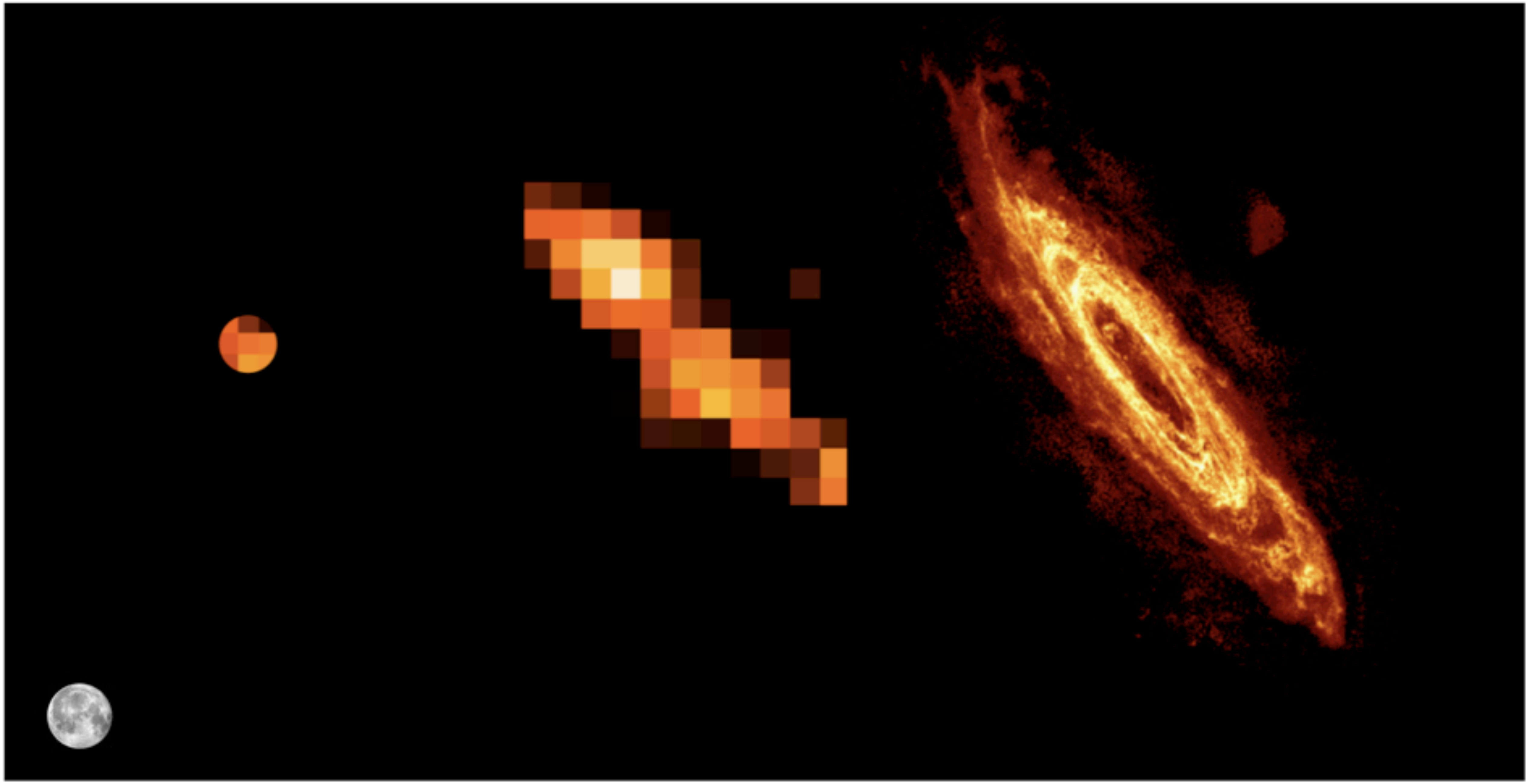


Fig. (a) The radiation pattern of a one-beam telescope, and (b) the radiation pattern of a four-beam telescope by placing a FPAs in the focal plane. With a focal plane array installed, a conventional radio telescope is transformed from a 1-D, single pixel detector to a 2-D multi-pixel camera. Courtesy: U of Birmingham, UK

1 wsrt antenna

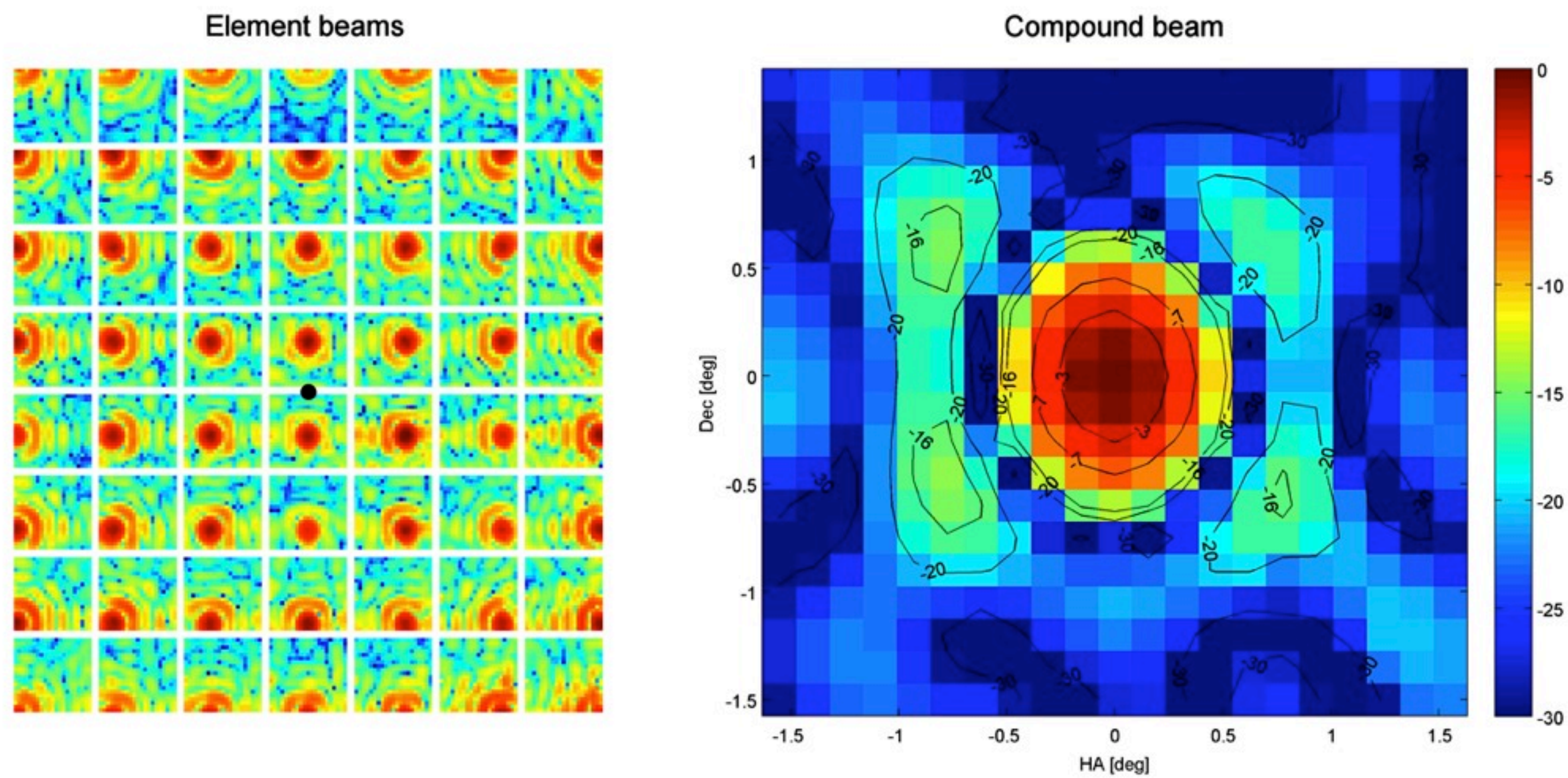
1 wsrt antenna fitted with FPA

14 wsrt antennas with FPA



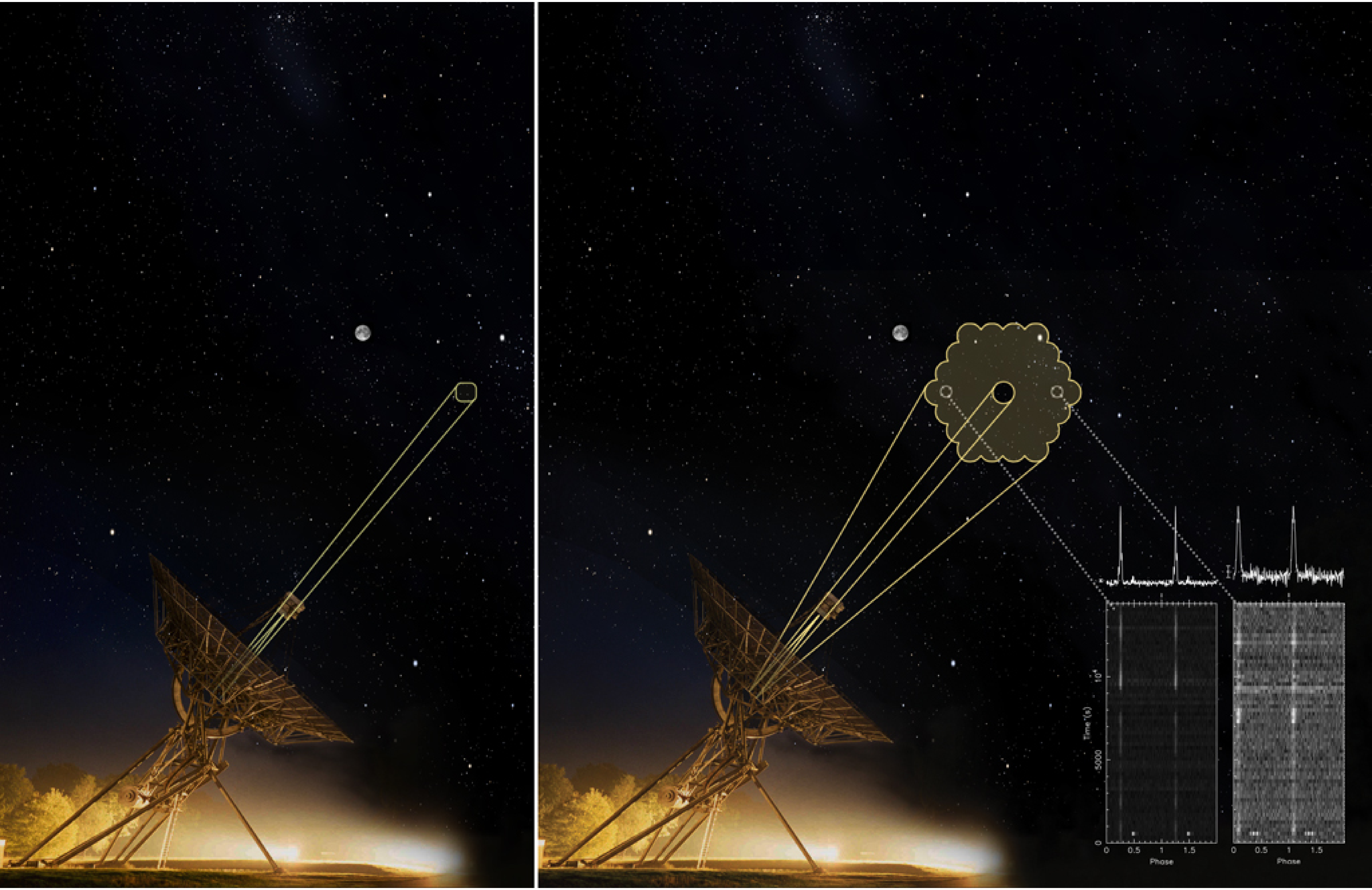
Variable element response

Optimal beam forming: Measured response of beam by combining together many individual FPA elements:



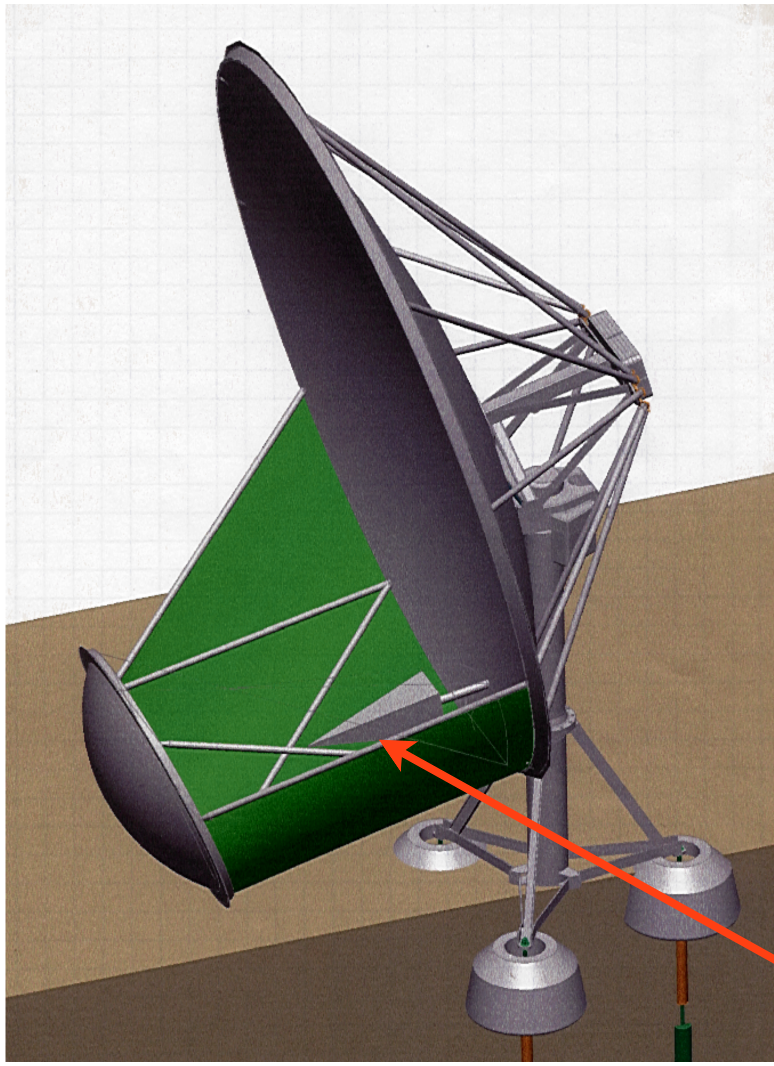
FPAs offer several advantages over classical multi-feed (multi-receiver) systems:

- dense sampling of the focal plane permits the beams to overlap on the sky (difficult to achieve this with bulky multiple horns)
- can optimise beam in a given direction
- can increase anttenna aperture efficiency via precise control of surface illumination



Time (s)
Phase
Phase

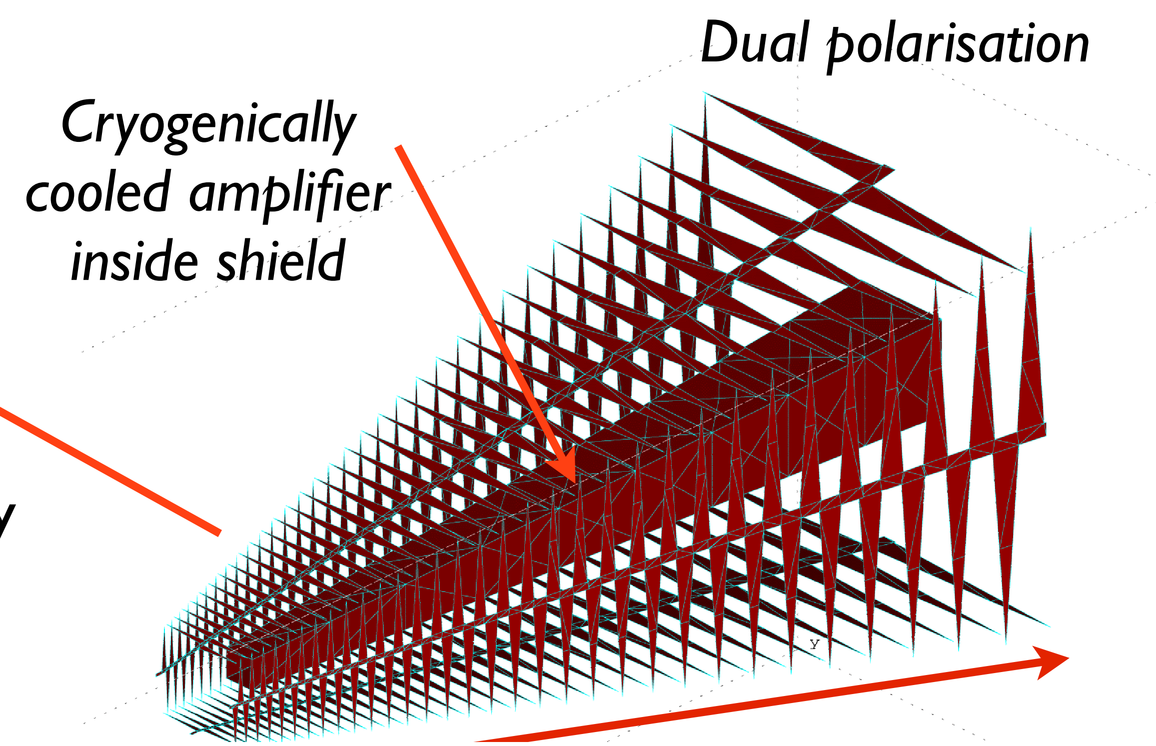
Concepts like the ATA (Allen Telecope Array) emphasise:



(i) cheap, 6 x 7-metre offset Gregorian antenna; 2.4-metre secondary.

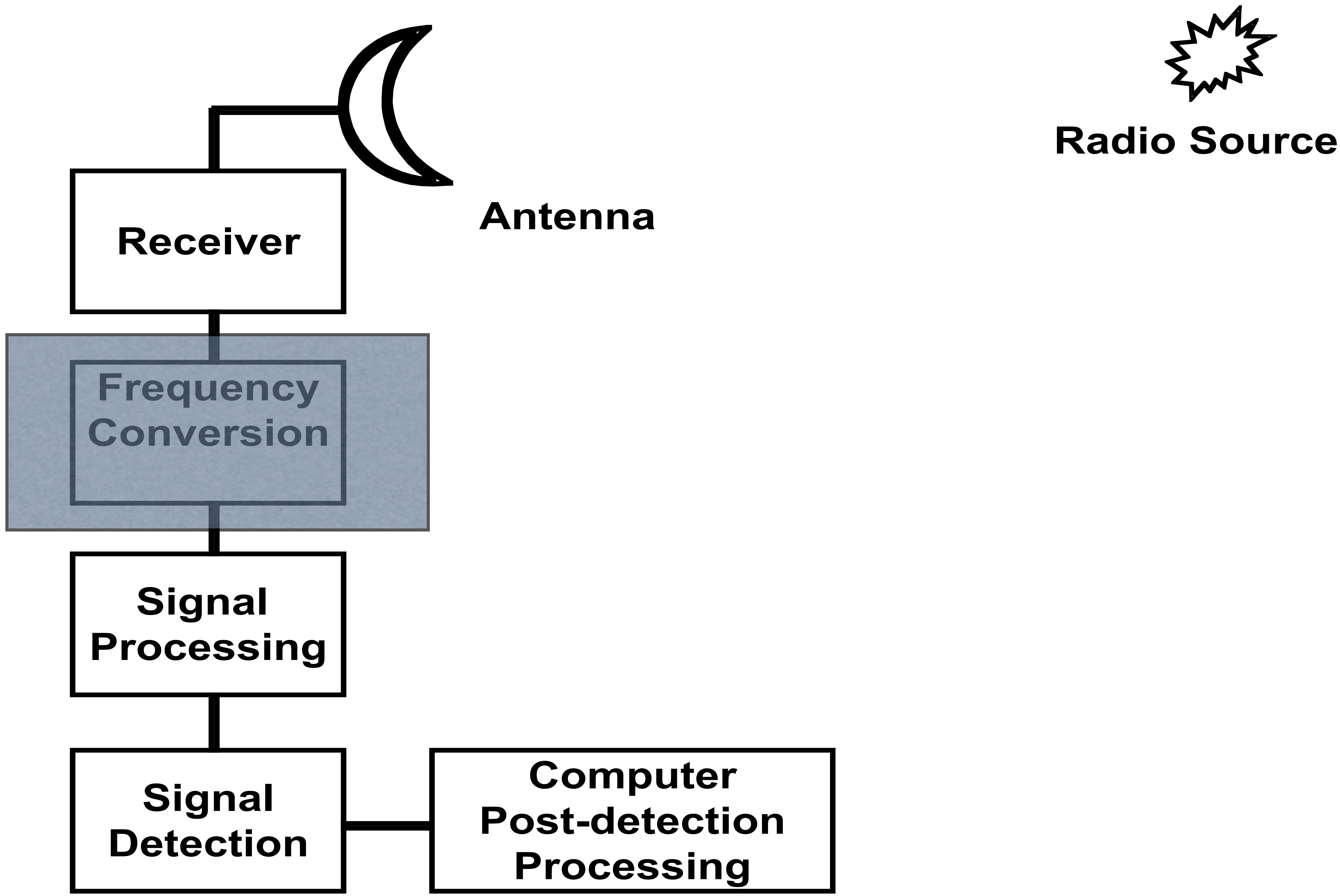
Small antennas also provide huge field of view (5 x size of full moon at 1.4 GHz)

(ii) Arrays like the ATA employ many of these small antennas - for the SKA several thousand would be used - as we shall see this greatly improves the imaging capability of these arrays but also increases the data processing costs.

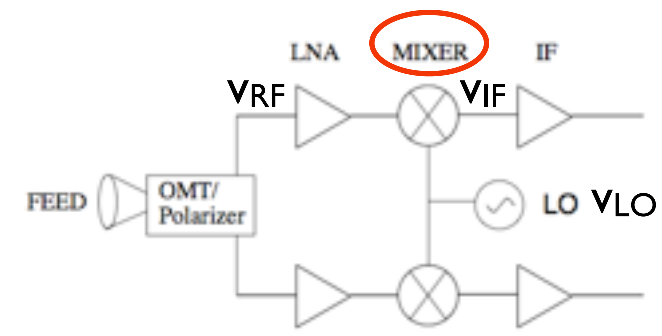


(iii) log-periodic feeds that are fed at the tip and employ cryogenically cooled LNAs. These receivers work over a decade in frequnecy e.g. 0.5-11 GHz for the ATA!

Radio Telescope Block Diagram



Frequency (down) conversion



A typical receiver tries to down-convert the "sky signal" or "Radio Frequency" (or RF) to a lower, "Intermediate Frequency" (or IF) signal

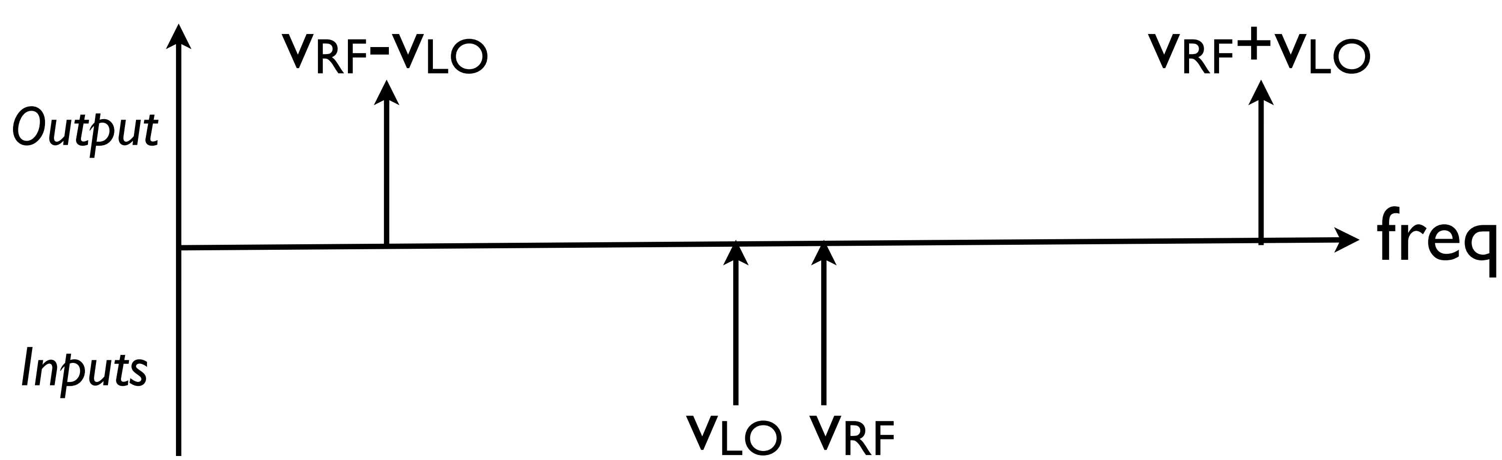
The reasons for doing this include: (i) signal losses (e.g. in cables) typically go as frequency^2; (ii) it is much easier to mainpulate the signal (e.g. amplify, filter, delay, sample/process/digitise it) at lower frequencies.

We use so-called "heterodyne" systems to mix the RF signal with a pure, monochromatic frequency tone, known as a Local Oscillator (or LO) or "mixer".

Consider an RF signal in a band centred on frequency $\nu_{RF,}$ and an LO with frequency ν_{LO}, these can be represented as two sine waves with angular frequencies w and w_o:

$$\nu_{IF} = \nu_{RF}\nu_{LO} \sim sin(\omega t)sin(\omega_o t) = 1/2(cos(\omega - \omega_o)t + cos(\omega + \omega_o)t)$$

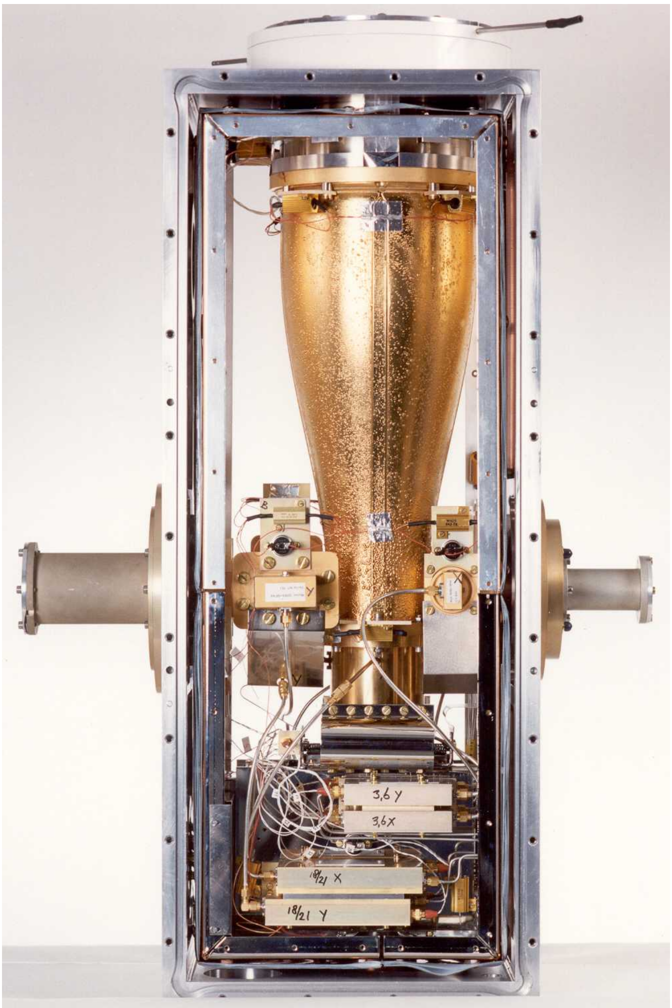
– Difference frequency – *– Sum frequency –*



The higher frequency component ("sum frequncy" $\nu_{RF}+\nu_{LO}$) is usually removed by a filter that is included in the LO electronics. Hence the process of down-conversion, takes a band with centre frequency ν_{RF} and converts it to a lower (difference) frequency, $\nu_{RF}-\nu_{LO}$.

Above: a typical LO used at cm wavelengths. These are sometimes locked to precise frequency standards (e.g. H-masers for VLBI) to maintain coherence.

Below: 6cm receiver system at wsrt



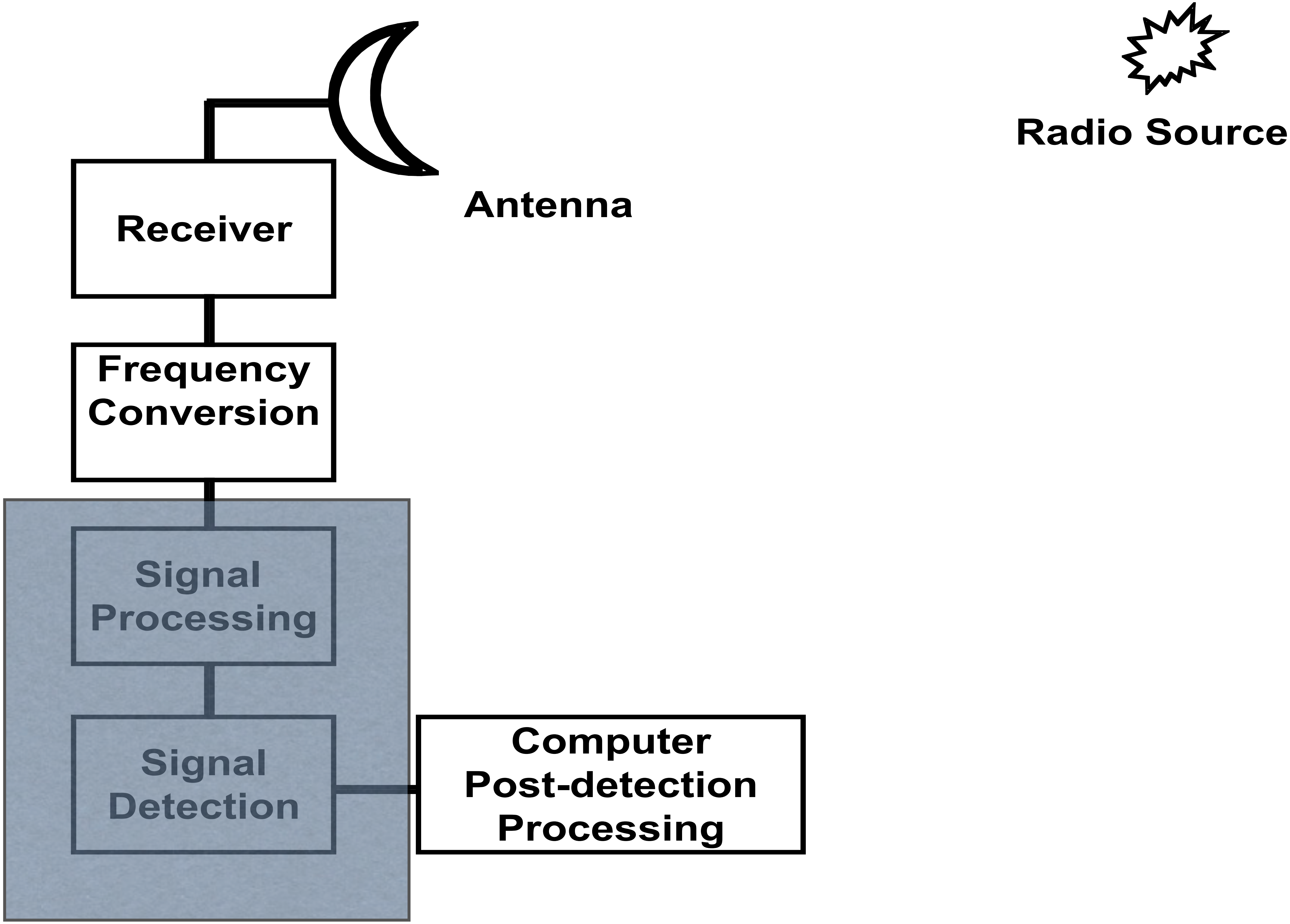
The mixer signal products preserves the noise characteristics of the input RF (sky) signal, but they contain an arbitrary phase-shift due to the unknown phase of the LO.

Usually there will be several mixers and frequency conversions in a receiver system. Eventually one edge of the frequency band reaches 0 Hz, known as a "base-band" or "video" signal.

Note that by changing the LO frequency, the sky frequency also changes. By tuning the LO you can bring a certain sky frequency (e.g. HI or other spectral lines) into the observing passband.

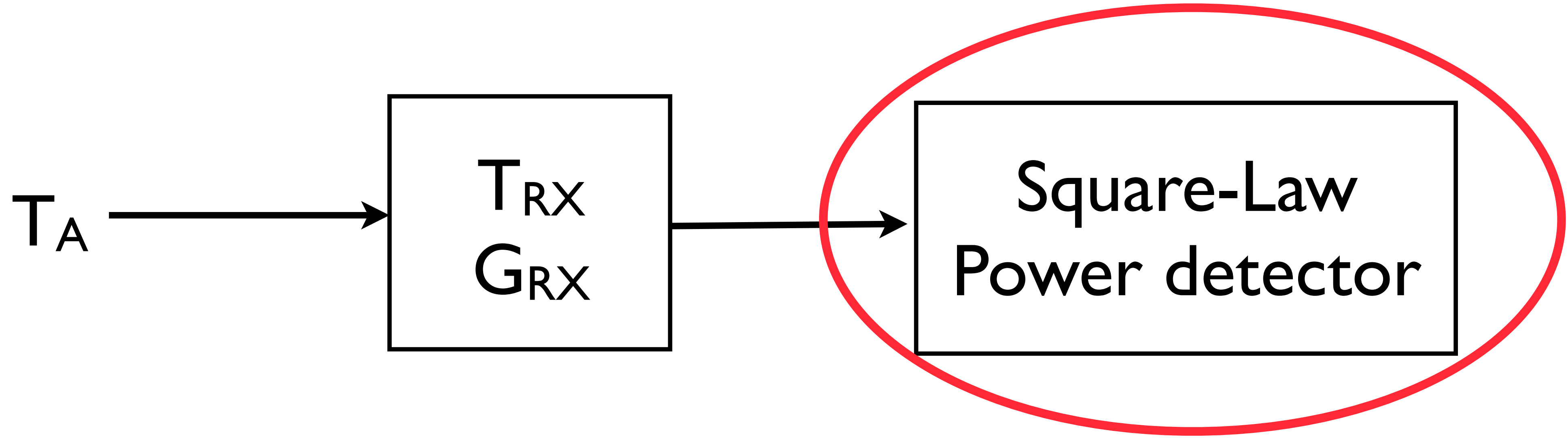
At high frequencies (e.g. millimetre wavelengths), down-conversion occurs before amplification.

Radio Telescope Block Diagram



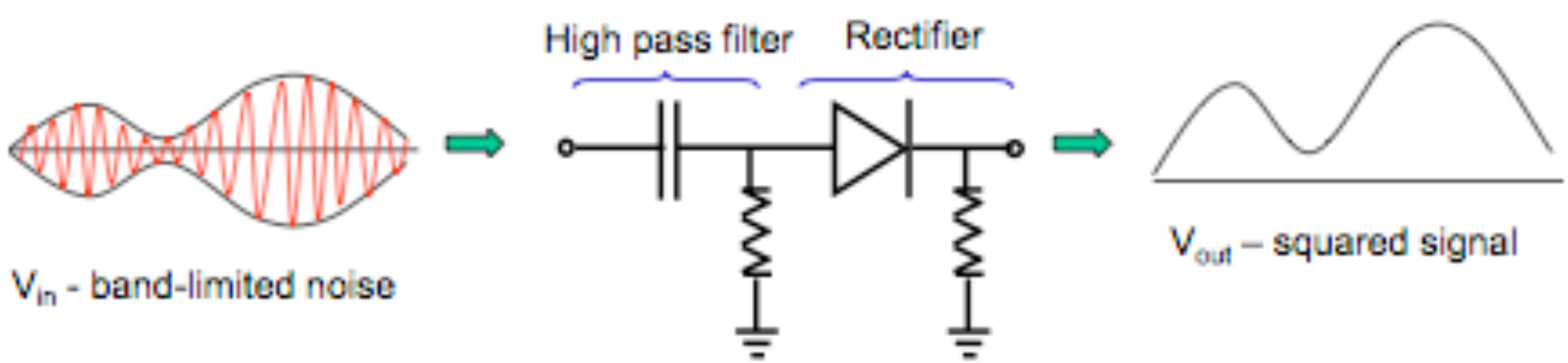
Square-Law detector

Simple receiver system:

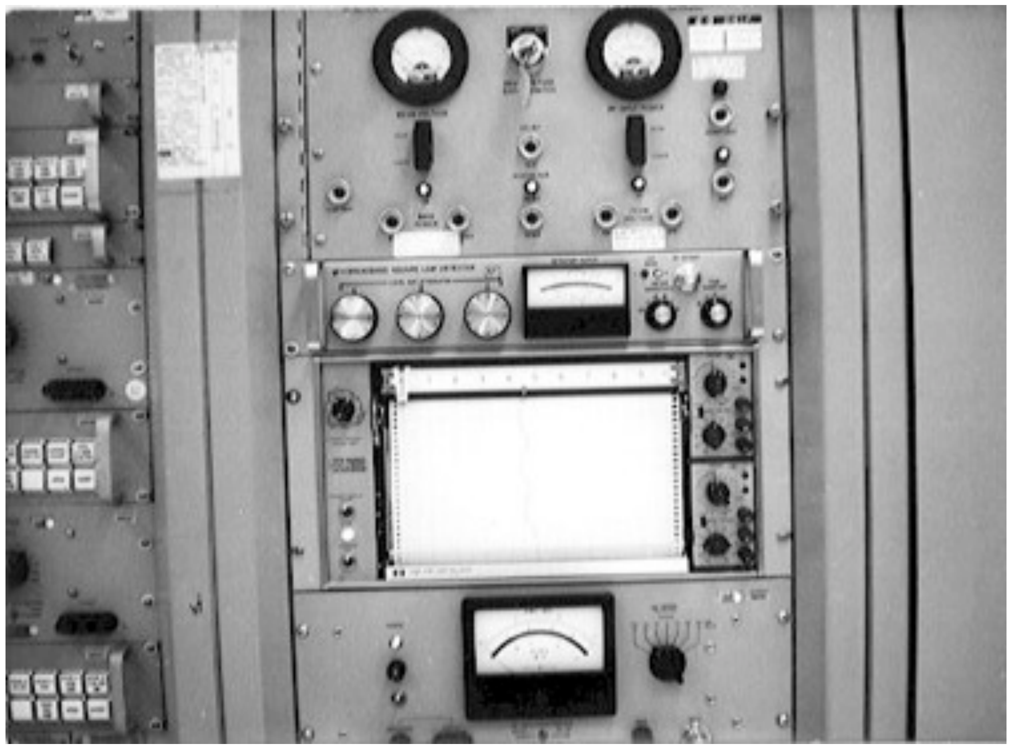


Since radio astronomy signals have the characteristics of white noise, the voltage induced in the receiver output alternates positively and negatively about zero volts. Any measurement of the Voltage expectation value or time average will read zero (e.g. hooking up a receiver to a DC voltmeter will not measure any signal).

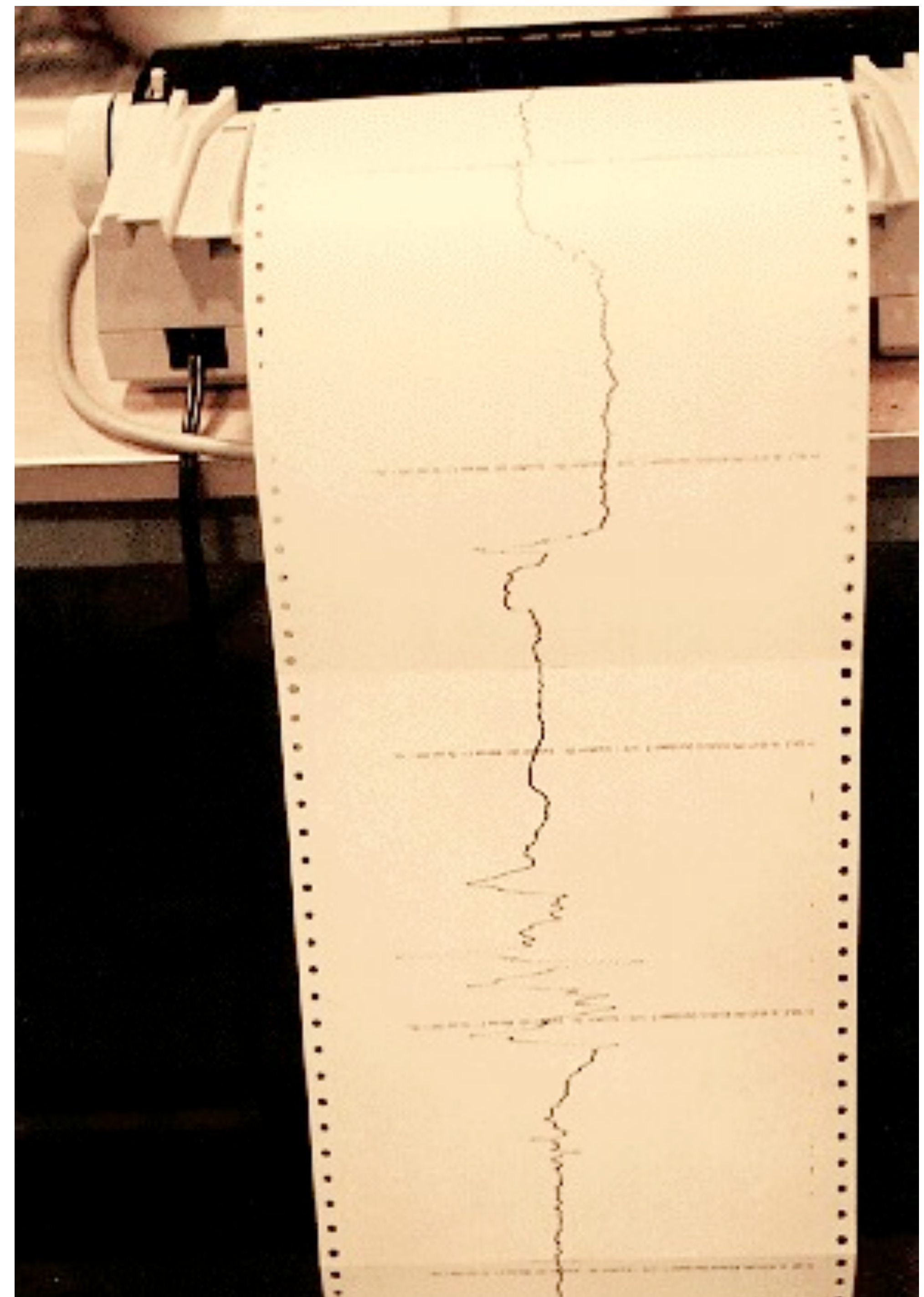
What is needed is a non-linear device ($V_{out} = AV_{in}^2$) that will only measure the passage of the signal in one preferred direction (either positive or negative) i.e. we must incorporate a semiconductor diode into our measuring system.



The simplest detectors that radio astronomers use are so-called "Square-Law Detectors". In these systems the DC component of the diode output is proportional to the square of the AC input voltage i.e. proportional to the power of the incoming signal.



Old Square-law detector and chart recorder system as used by NASA (DSN)



In the crudest systems of yesteryear, this DC voltage can be used to drive a penchart recorder. With this kind of system Pulsars were discovered!

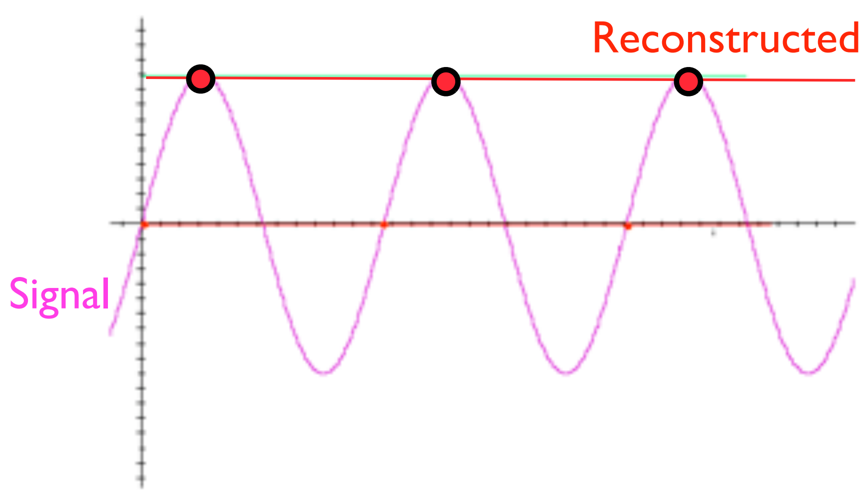
Online data sampling

Square-law detectors are not used these days. The receiver produces a varying analogue output voltage that is usually digitised and stored for further (offline) processing.

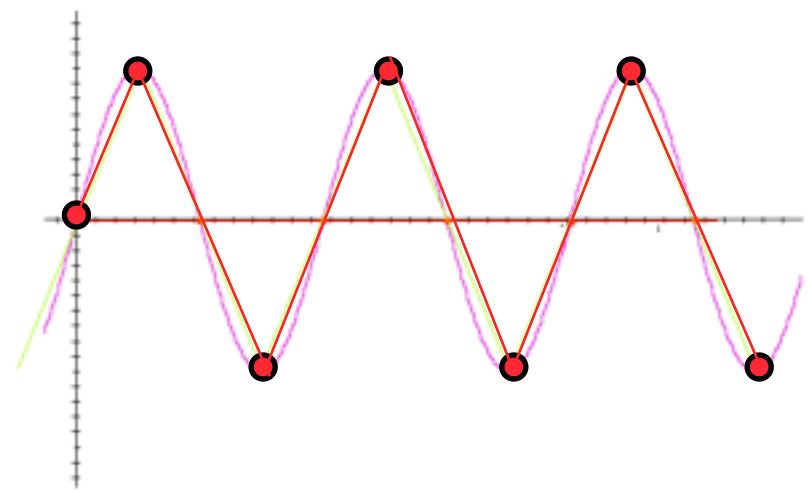
How often must be sample the signal ?

Consider the following sine wave:

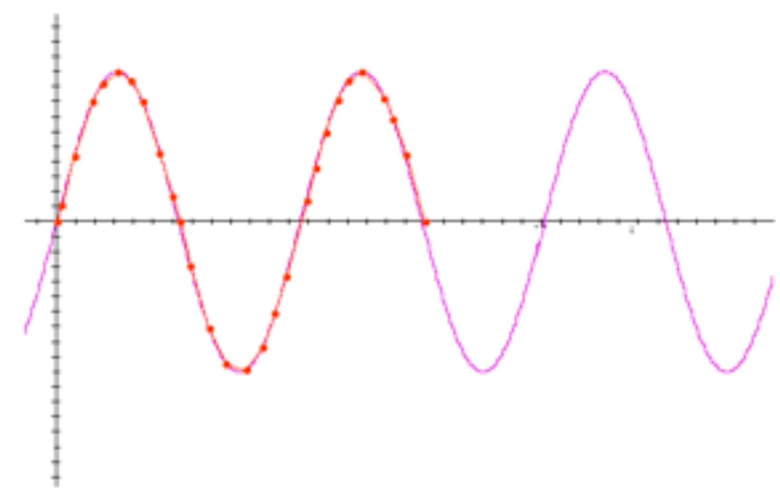
If we sample once per cycle time (period) we would consider the signal to have a constant amplitude.



If we sample twice per cycle time (period) we get a saw-tooth wave that is becoming a good approximation to a sinusoid.



For lossless digitisation we must sample the signal at least twice per cycle time.

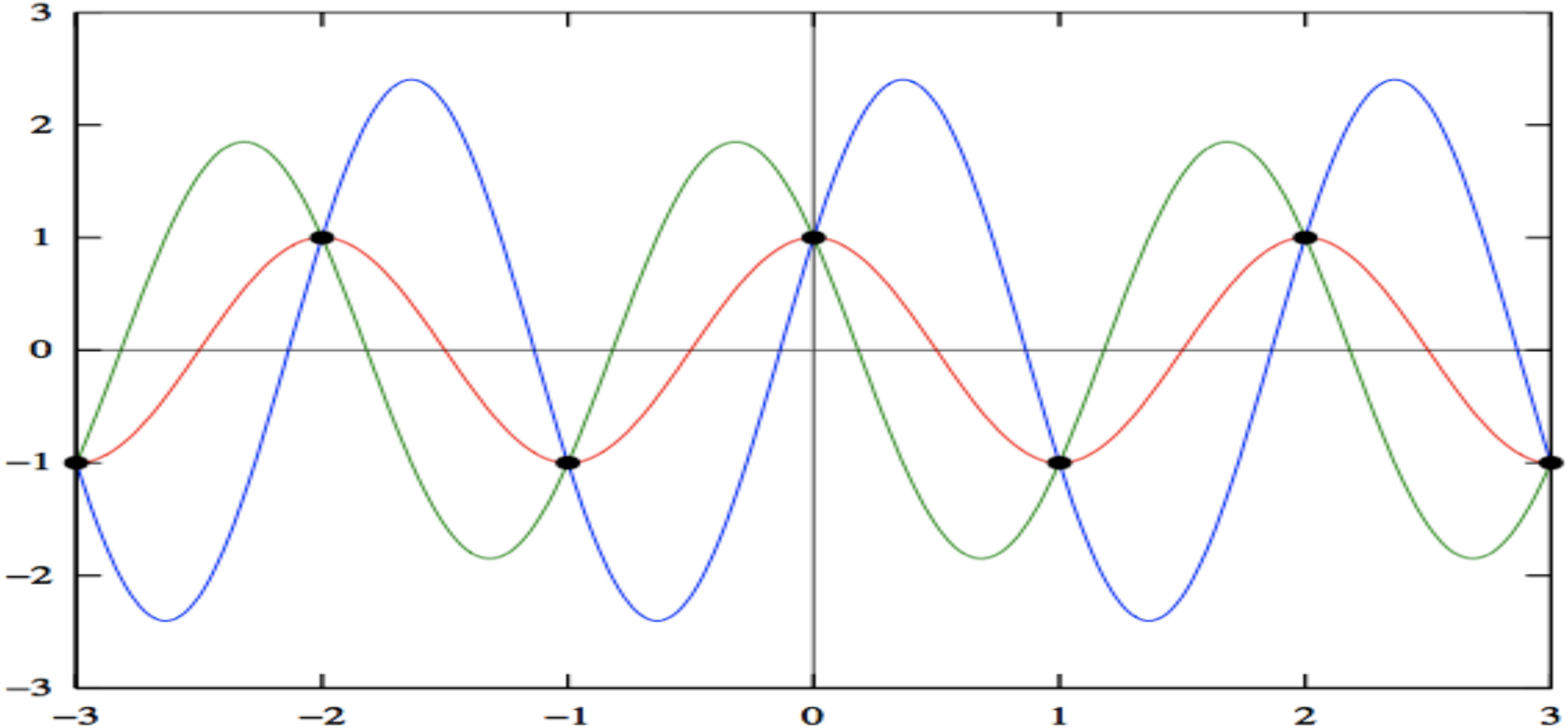


Nyquist's sampling theorem states that for a limited bandwidth signal with maximum frequency f_{max}, the equally spaced sampling frequency f_s must be greater than twice the maximum frequency f_{max}, i.e. $f_s > 2\ f_{max}$ in order for the signal to be uniquely reconstructed without aliasing.

The frequency $2f_{max}$ is called the Nyquist sampling rate.

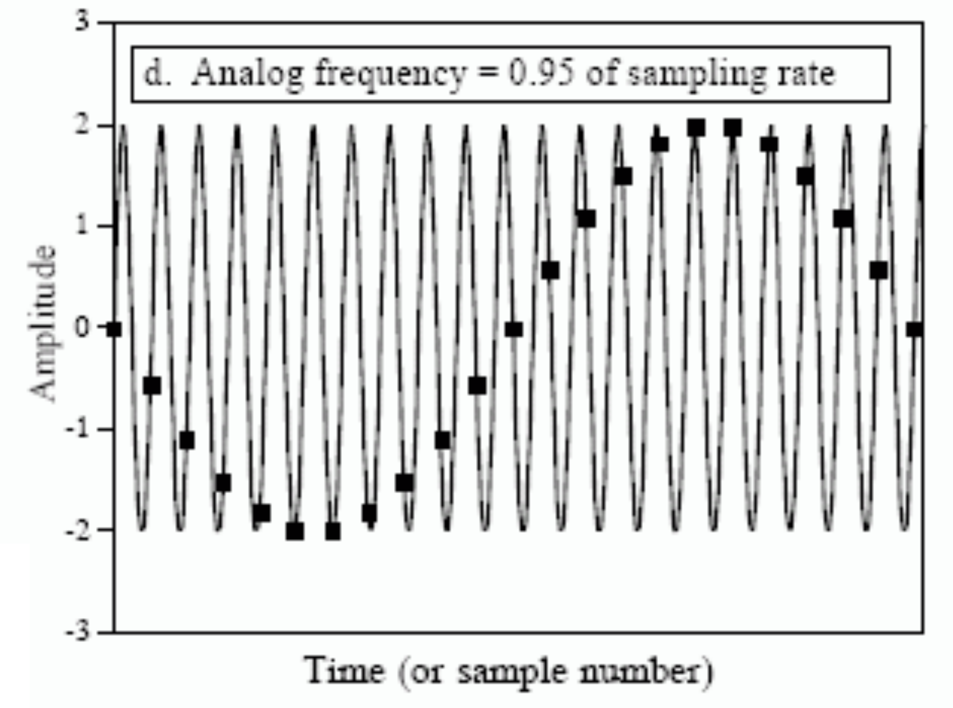
e.g. If a reciever system provides a baseband signal of 20 MHz, the signal must be sampled 40E6 times per second.

Note that strictly speaking, the sampling frequency (rate) must be strictly greater than the Nyquist rate ($f_s > 2 f_{max}$) of the signal to achieve unambiguous representation of the signal. In the pathlogical case where the signal contains a frequency component at precisely the Nyquist frequency, then the corresponding component of the sample values cannot have sufficient information to reconstruct the signal.

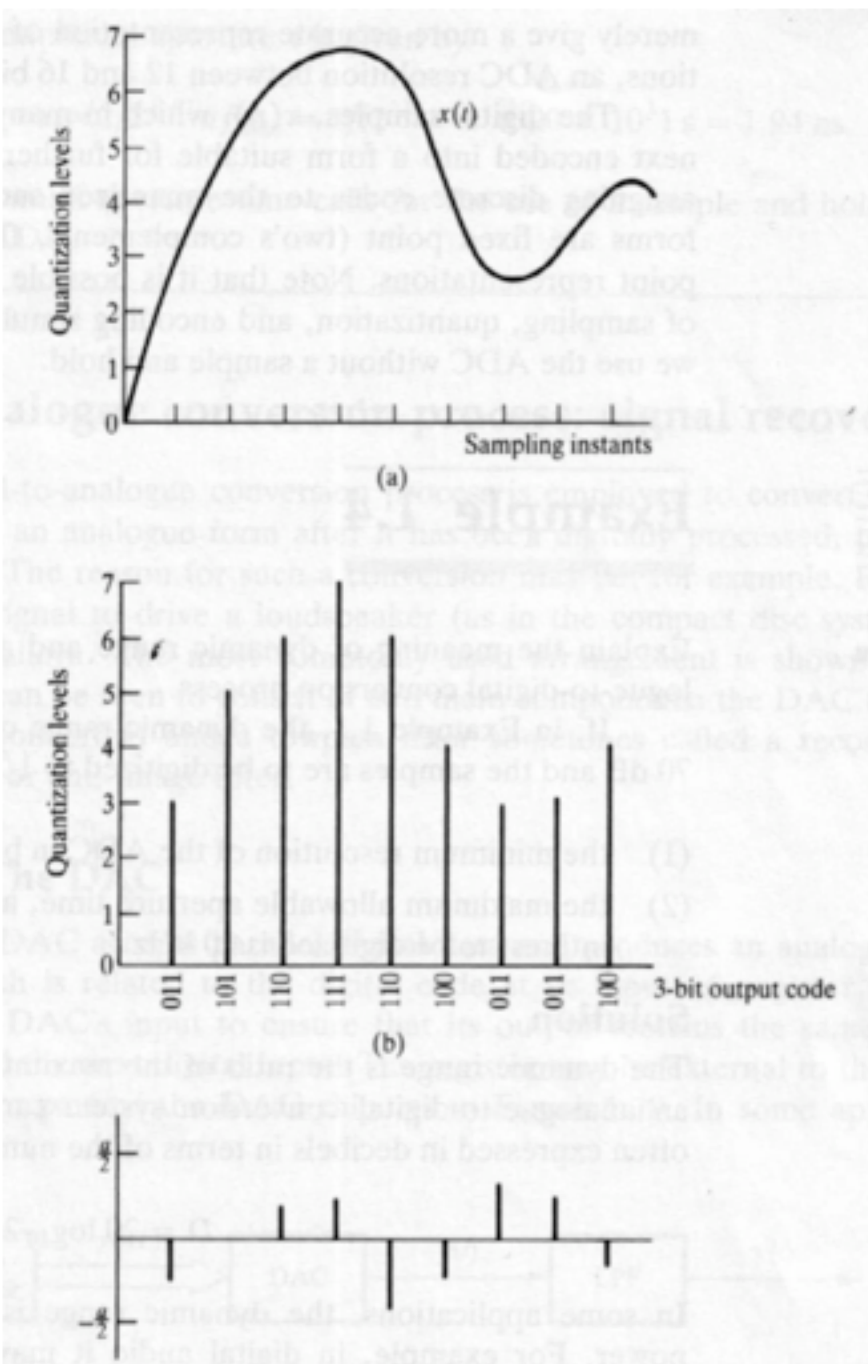


A family of sinusoids at the critical frequency, all having the same sample sequences of alternating +1 and −1. That is, they all are aliases of each other, even though their frequency is not above half the sample rate. Note that strictly speaking, the sampling frequency (rate) must be strictly greater than the Nyquist.

Sampling at less than the Nyquist rate leads to aliasing of the original signal (right).



Since the final processing of radio astronomy data takes place via digital computers, devices known as *Analogue to digital converters* (ADC) are used to sampled at regular intervals the voltage signal from the receiver.



The sampling frequency employed is usually the Nyquist rate or sometimes the signal may be over-sampled.

The number of bits (referred to as the quantisation) used for one value of the discrete signal sets the accuracy of the signal magnitude.

Left: a 3-bit (9 level) quantisation is used [b] in order to characterise the original signal [a]. The errors (or residuals) are shown in [c].

Somewhat surprisingly, even low levels of quantisation result in a relatively modest degredation in signal-to-noise, at least in the case where the signals are not strong:

No. of Bits	Relative performance
1	64%
2	81%
3	88%
infinity	100%

2-bit samples are commonly used in current radio telescope systems. As can be seen from the table, the degredation of the signal-to-noise for 2-bit sampling is much less than that of 1-bit sampling, the value achieved is 0.88 of an ideal system. A larger number of bits can be used but the point of diminishing returns is rapidly reached and the compute burden begins to rise with for very little real gain.

The table (left) assumes Nyquist sampling. Some modest gains can be made by also increasing the sample rate.

This analysis is correct assuming we are sampling signals with a limited range of power. The process of quantisation is inherintely non-linear, and in the presence of strong signals (such as RFI) a larger number of bits is required to characterise the wide range of signal strength.

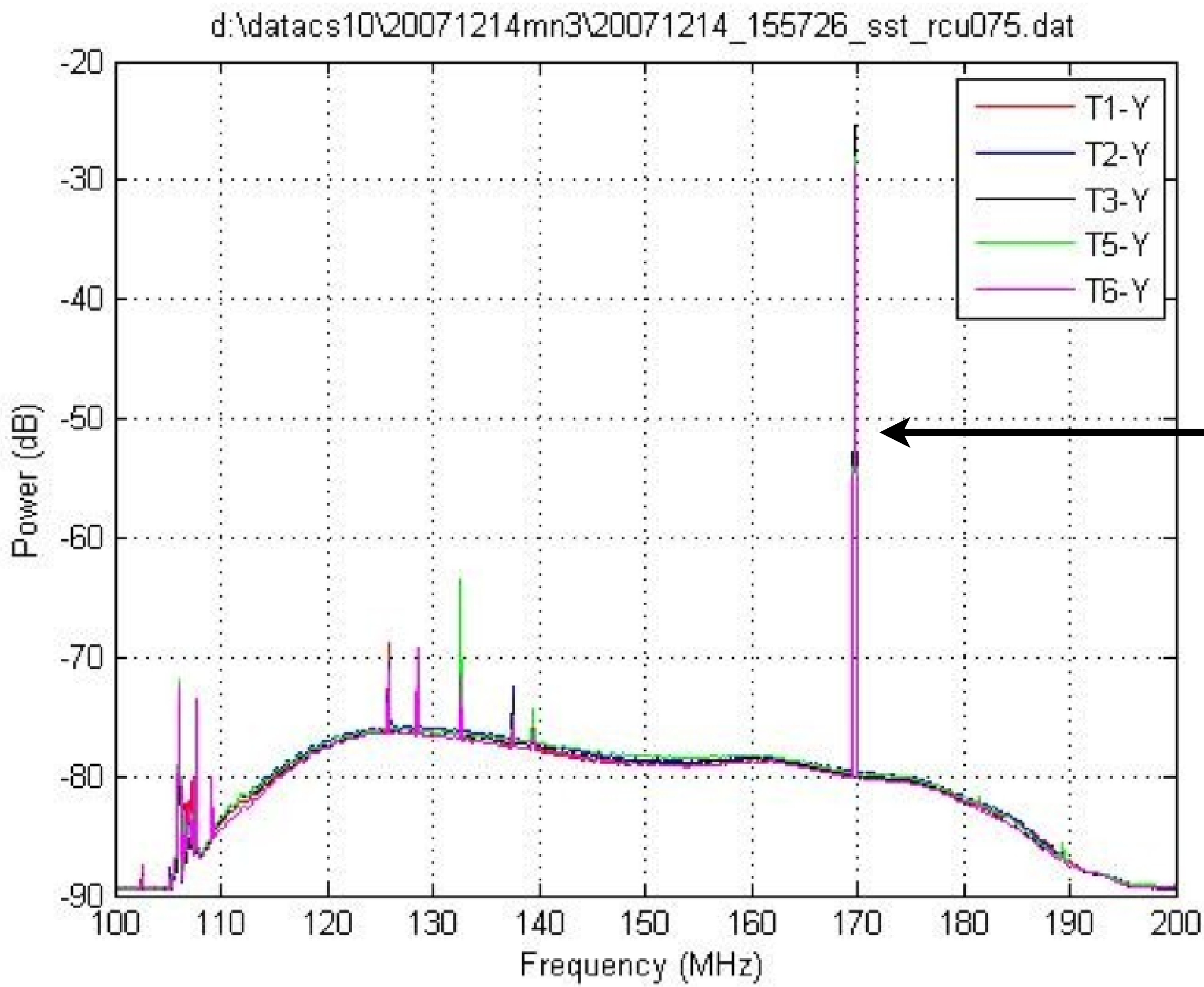
e.g. LOFAR uses 12-bit samples... leading to very large data rates!

e.g. LOFAR operates with a bandwidth of 48 MHz. With Nyquist sampling, each LOFAR station generates 48E6 x 2 x 12 bits ~ 1.1 Gbit per second per polarisation product.

RFI signals can produce signals that are 80-100 dB (8 to 10 orders of magnitude) above the receiver power nosie level. Since they are usually very narrow-band, this input power (P) gets diluted across the observing band to ~ 40-50 dB.

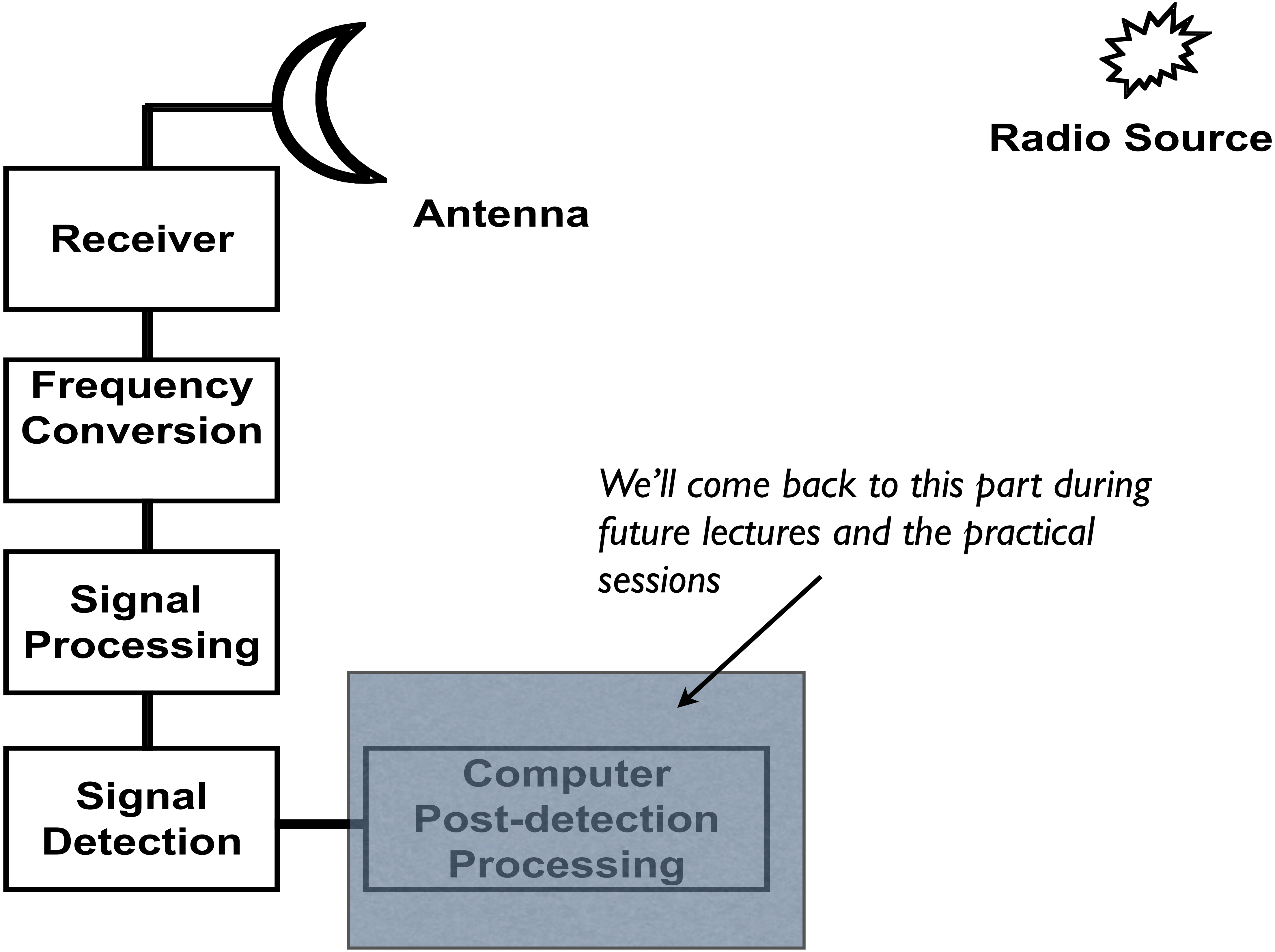
An ADC with "N-bit" sampling permits us to measure a range of voltage (V) of 2^N levels. In term of power the range is 2^{2N} (since $P \sim V^2$).

e.g. a 10-bit system can measure a range of power spanning 2^{20} ~ 60dB.



100 - 200 MHz spectrum of a LOFAR High Band Antenna. An RFI signal > 50dB above the receiver noise level is clearly present.

Radio Telescope Block Diagram



Sources

- Figures & slides this lecture: Garrett

http://www.astron.nl/astrowiki/doku.php?id=UvA_MSc_RadioAstronomy_2013

Questions?