6.1 Introduction

This chapter describes the various components of a radio telescope and outlines the various detection mechanisms for the radiation.

6.2 Antennas

6.2.1 Introduction

The elements of a standard radio telescope are the reflector, feed, transmission line and receiver. We will first discuss the reflector, which collects power from an astronomical source and provides directionality. The terms antennas and reflectors are often used interchangeably. However, there is a difference -- an antenna is a device that couples the waves in free space to the confined waves in a transmission line while reflectors concentrate the radiation. The reflector or antenna has two purposes, first they collect power and second they provide directionality. The power collected by an antenna is approximately given by

$$P = S_\nu A \Delta \nu$$

where S_ν is the flux density at the earth from some astronomical source, A is the area of the antenna and $\Delta\nu$ is the frequency interval or bandwidth of the measured radiation. So, the larger antennas collect more power. The antenna also has the capability of discriminating the signals coming from different directions in space.

6.2.2 Diffraction and reciprocity

The operation of antennas, and telescopes in general, are governed by electromagnetic theory and diffraction theory plays an important role. In order to understand this, one first needs to know the reciprocity theorem. This theorem states that the telescope operates the same way whether it is receiving or transmitting radiation. So the response pattern of an antenna that is receiving radiation is the same as the pattern produced when the same antenna is transmitting. A schematic of a response pattern of an antenna is given in Figure 6.1.

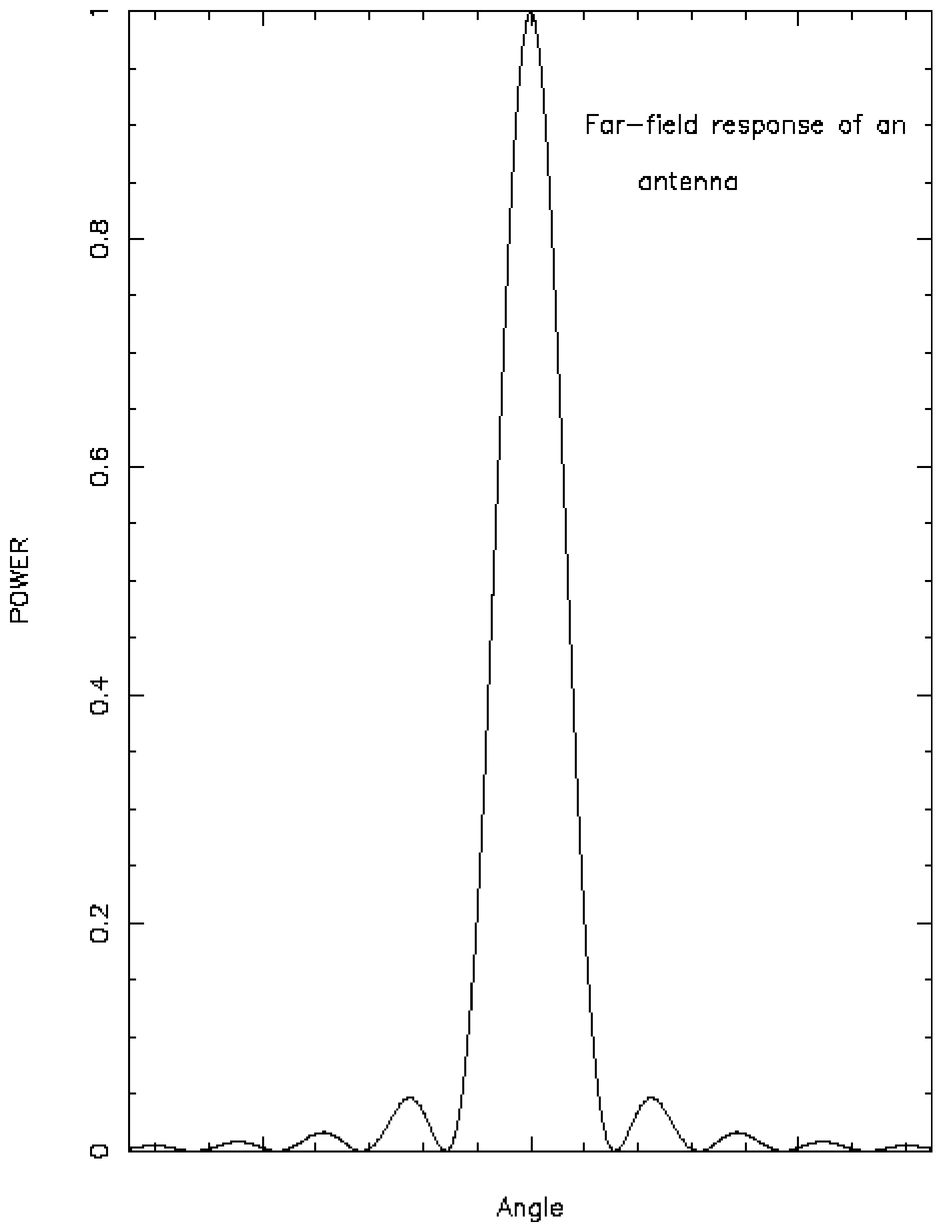

Figure 6.1

A telescope's response pattern is then the same as the far-field diffraction pattern produced by the aperture. In general, when radiation of wavelength λ passes through an aperture of diameter *D*, the radiation diffracts into a beam with angular size $\theta \sim \lambda / D$. At large distances (or the far-field response) the pattern is given by Fraunhofer diffraction theory and the pattern looks like that in Figure 6.1, where the beamwidth is the full width at half power of the main beam. The beamwidth θ is also a measure of the directivity of the antenna. A more precise statement that can be made (and will not be derived here) is that the angular pattern of the electric field in the far-field is the Fourier transform of the electric field distribution across the aperture.

6.2.3 Parabolic Antennas

Parabolic antennas (or reflectors) are common to both radio and optical astronomy. The reflector focuses plane waves to a single point, or in other words, converts plane waves into converging spherical waves. In a radio telescope these spherical waves are then coupled to a transmission line using a feed horn, which is a horn antenna. The feed horn can be placed at the prime focus or at a secondary focus using a Cassegrain design (Figure 6.2).

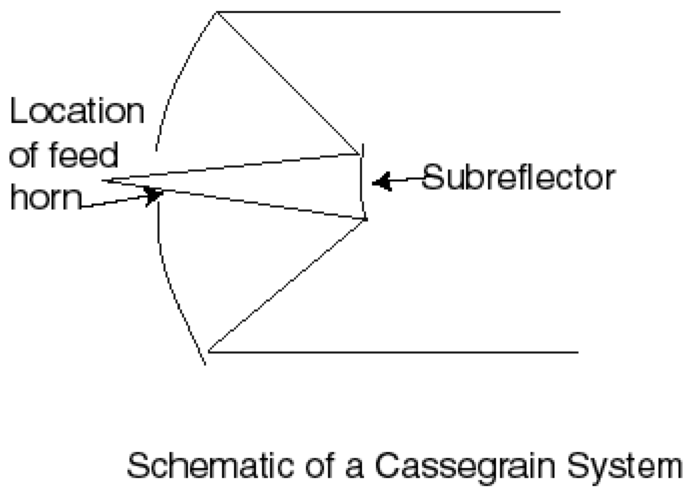

Figure 6.2

Most radio telescopes have a Cassegrain design since placing the feed horn at the prime focus will block more of the surface. Small telescopes (such as the SRT) have a prime-focus arrangement in which the reflector is illuminated by the feed placed at the focal point on the axis of the parabola. The geometry of a parabola is given by

$$y = x^2/(4F)$$

where y is the distance from plane, x is the distance from the vertex, and F is the focal length as shown in Figure 6.3.

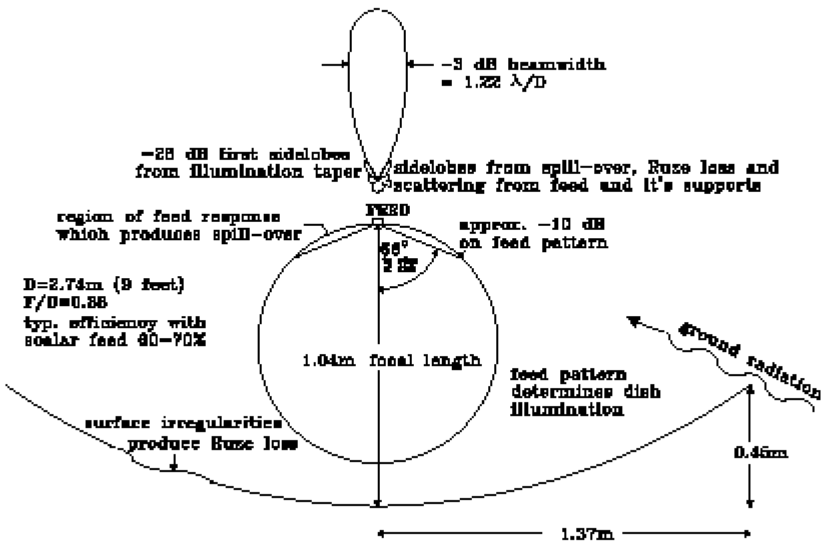

Figure 6.3
The reflector surface must follow a parabola to within a small fraction of a wavelength. An imperfect surface scatters some signal away from the focus and produces a loss known as the *Ruze* loss after John Ruze, who first derived the expression

where *L* is the loss factor, *d* is the root-mean-square (rms) deviation from a

$$L = \exp(-(4\pi d / \lambda)^2)$$

parabola, and λ is the wavelength. For most random distributions, the rms is about one quarter of the peak-to-valley variations. Figure 6.4 shows the Ruze loss in dB as a function of surface quality.

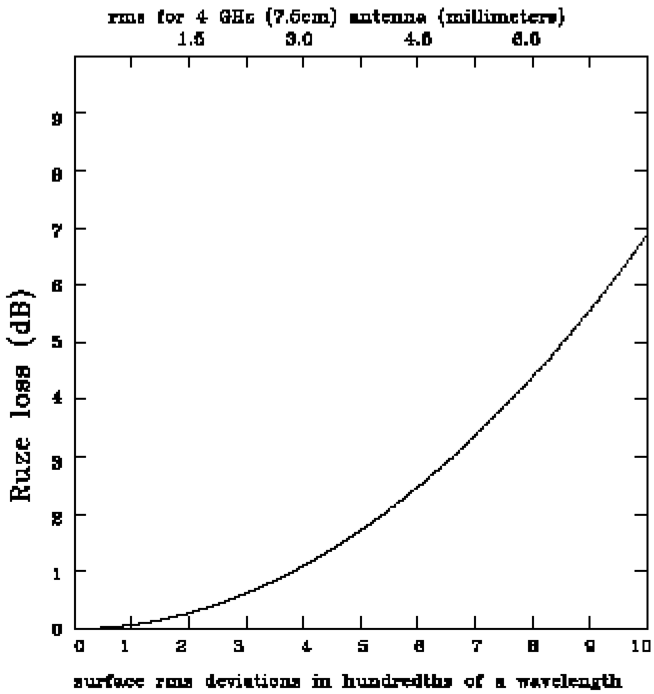

Figure 6.4

The angle subtended by the reflector as seen by the feed is determined by the ratio of focal length to diameter or *F/D* of the dish. Most dishes have a *F/D* ratio close to

0.4. A Radio Shack 9-foot satellite TV dish has an *F/D* of 0.38. For this *F/D*, the edge of the dish is about 64° out as seen by the feed. The feed should ideally be an antenna with a uniform beam that illuminates only the reflector surface. The efficiency in this case would be close to 100%. In practice, a good feed provides about 60 to 70% efficiency, so that the gain of this 9-foot antenna is 39.6 dB at 4.1 GHz. A very popular feed design is a *scalar* feed, which consists of a probe in a circular waveguide surrounded by choke rings as illustrated in Figure 4. The beam of the feed, which usually tapers down by about 10 dB at the edge of the dish, can be adjusted to some extent by the choice of opening size and location of the choke rings. Figure 5 shows the effect on efficiency of varying this taper. The beamwidth of a dish illuminated with such a scalar feed is approximately

$$\theta = 1.22\lambda/D$$

6.2.4 Gain

Radio and radar engineers normally speak about antennas in terms of their gain in dB referred to a half-wave dipole (dBd) or referred to an ideal isotropic antenna (dBi). A half-wave dipole has a gain of 2.15 dBi. Radio astronomers prefer to talk of size and efficiency or effective collecting area The gain, *G*, of an antenna relative to isotropic is related to its effective collecting area, *A*, by

$$G = 4\pi A/\lambda^2$$

where λ is the wavelength. The gain is also related to the directivity of the antenna: An antenna with a smaller beam will have a higher gain. If we think of the antenna as a transmitter, as we can do owing to reciprocity, then if the transmitted energy is confined to a narrow angle, the power in this direction must be higher than average in order for the total power radiated in all directions to add up to the total power transmitted.

To achieve an effective area or aperture of many square wavelengths (gains of more than, say, 26 dBd), a parabolic reflector is the simplest and best approach. For long wavelengths, for which an antenna with more than 26 dBd would have enormous dimensions, other approaches are more appropriate. As radio amateurs doing Moon-bounce know, it is hard to beat an array of Yagi antennas for simplicity and minimum wind loading. A single 20-wavelength-long Yagi can give a gain of 20 dBd. Stacking 2 Yagis adds 3 dB and another 3 dB for every doubling. The effective aperture of a 20 dBd Yagi, however, is only 13 square wavelengths, so that stacking 50 Yagis to get 37 dBd of gain at, say, 4 GHz doesn't make much sense when you can do as well with a 9-foot-diameter dish of 60% efficiency. At UHF frequencies around 400 MHz, the choice between Yagis and a dish is not so clear.

With the availability of excellent LNAs, optimizing the antenna efficiency is less important than optimizing the ratio of efficiency to system noise temperature or gain over system temperature, G/T_s. This means that using a feed with low sidelobes and slightly under-illuminating the dish may reduce T_s by more than it reduces G and so improve sensitivity.

6.2.5 Spillover

With advent of superb Low Noise Amplifiers (LNAs), the antenna noise is also a very important performance parameter along with the gain or equivalent effective aperture. Antenna noise originates from the sky background, ohmic losses, and ground pickup or *spillover* from sidelobes. While the sky noise is fundamental, the losses and sidelobes can be made small by a good design. Sky noise is frequency dependent but never gets any lower than the cosmic 3-K background. The minimum is near 1.4 GHz, where Galactic noise has declined and atmospheric attenuation, due primarily to the water line at 22 GHz, is still low. The lowest system noise achievable is about 18 K. At the NASA deep space network (DSN), where every fraction of a dB of performance is worth $millions, the S-band (2.3 GHz) system budget is approximately 3 K from cosmic background plus 7 K spillover plus 3 K atmospheric plus 5 K LNA. At 408 MHz, galactic noise will dominate, and system noise will be at least 50 K increasing to over 100 K toward the Galactic center. Figure 6.5 shows the relative system-noise contributions as a function of frequency.

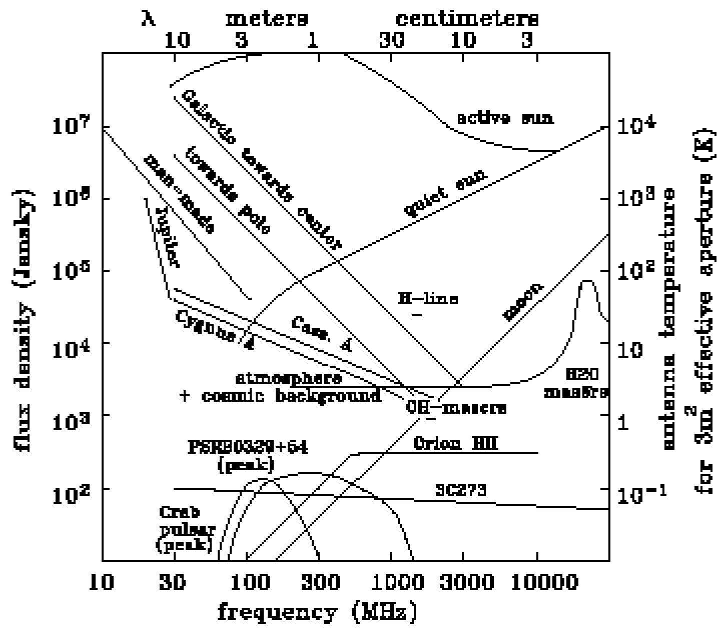

Figure 6.5

6.3 Receivers

6.3.1 Introduction

Radio telescope receivers filter and detect radio emission from astronomical sources. In most cases the emission is incoherent radiation whose statistical properties do not differ either from the noise originating in the receiver or from the background radiation that is coupled to the receiver by the antenna. In addition, these signals are extremely weak, so amplifiers have to be constructed in order to increase the signal to a detectable level.

Figure 6.6 is a schematic of a minimalist receiver for continuum radio astronomy.

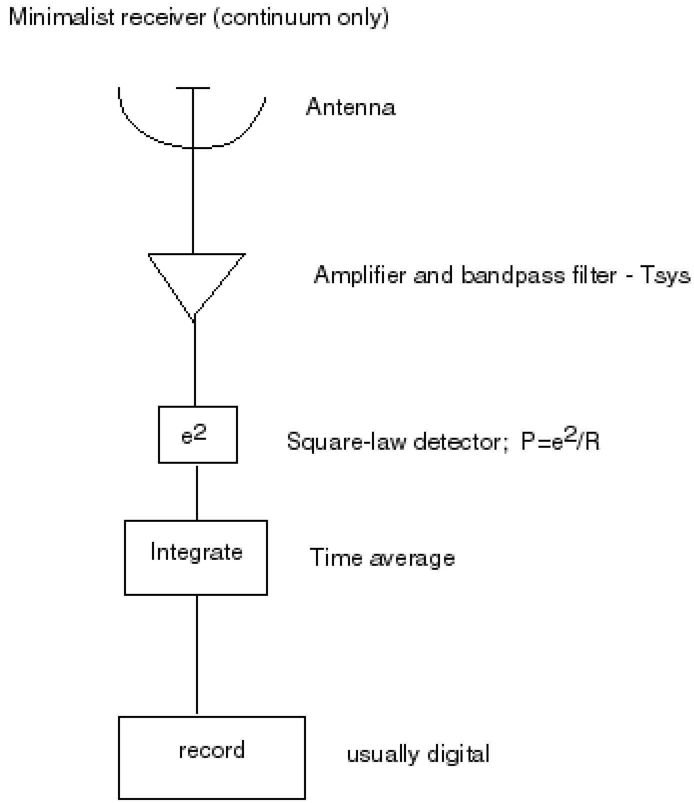

Figure 6.6

After the antenna, the first stage of the receiver, the low-noise amplifier (LNA), is probably the most important component of a radio telescope. Since the signals are so weak, the noise performance of the receiver is crucial, and this leads to extraordinary efforts, such as cryogenic cooling, to reduce noise in the LNA. The noise performance of radio-astronomy receivers is usually characterized by an equivalent system temperature, T_{sys} (in Kelvins), referred to the feed or even to outside Earth's atmosphere. Using temperature units for the system allows direct comparison with source temperatures. Typical system temperatures are ten to a hundred K for centimeter wavelengths or up to several hundred K for millimeter and submillimeter wavelengths. These numbers should drop as technological progress is made.

A. The noise equations

The usual root-mean-square (rms) noise calculations in radio astronomy are based on

$$\Delta T = \frac{\alpha\gamma T_{sys}}{\sqrt{\beta t}}$$

where T_{sys} is the system temperature in Kelvins, β is the noise bandwidth, which is only approximately equal to the resolution in spectroscopy or the total usable bandwidth in continuum, and t is the total integration time (on plus off) for normal switched observing. Set α to 2 for ordinary single switching where 2 is the product of two $\sqrt{2}$, one for spending half the time on source, another for differencing two equally noisy measurements. The correlation quantization correction γ is approximately 1.16 for Haystack's spectrometer with its modified 3×3 multiplication table. Set γ to 1 for continuum. Then ΔT is the rms fluctuation in the corresponding measurement. The βt in the denominator of this equation is, in effect, the number of samples averaged to make the measurement.

B. Why heterodyne?

Most receivers used in radio astronomy (all receivers used for spectroscopy) employ so-called superheterodyne schemes. The goal is to transform the frequency of the signal (SF) down to a lower frequency, called the intermediate frequency (IF) that is easier to process but without losing any of the information to be measured. This is accomplished by mixing the SF from the LNA with a local oscillator (LO) and filtering out any unwanted sidebands in the IF. A bonus is that the SF can be shifted around in the IF, or alternatively, the IF for a given SF can be shifted around by shifting the LO.

C. Why square-law detectors?

Inside radio-astronomy receivers, the signal is usually represented by a voltage proportional to the electric field as collected by the antenna. But we normally want to measure power or power density. So, at least for continuum measurements and for calibration, we need a device that produces an output proportional to the square of the voltage, a so-called square-law detector, and also averages over at least a few cycles of the waveform.

6.3.2 Extracting weak signals from noise

As mentioned above, radio-astronomy systems usually operate close to the theoretical noise limits. With a few exceptions, signals are usually extremely weak. One such exception is the Sun. Depending on frequency, Solar cycle, antenna size, and system noise temperature, pointing an antenna at the Sun normally increases the received power several fold. Toward other sources, it is not unusual to detect and measure signals that are less than 0.1% of the system noise. The increase in power, measured in K, due to the presence of a radio source in the beam is given by

$$T_a = AF/(2k)$$

where *A* is the effective aperture (m^2) or aperture efficiency times physical aperture, *F* is the radio flux density in watts/m^2/Hz, and *k* is Boltzmann's constant, 1.38×10^{-23} w/Hz/K. The factor of 2 in the denominator is because radio astronomers usually define the flux density as that present in both wave polarizations, but a receiver is sensitive to only one polarization. Radio telescopes use linear or circular polarization depending on the type of observations being made, and with two LNAs and two receivers, one can detect two orthogonal polarizations simultaneously. In order to detect and measure signals that are a very small fraction of the power passing through the receiver, signal averaging or *integration* is used. If the receiver gain were perfectly stable, our ability to measure small changes in signal is given by the noise equation in the previous section. There ΔT is the one-sigma measurement noise.

If the receiver bandwidth is 1 MHz and T_{sys} = 100 K, for example, then we can measure down to 0.013 K in one minute. For a sure detection, we need to see a change of 10 sigma or about 0.1 K change. The receiver gain in practice is seldom exactly constant, and the additional spillover noise and atmospheric noise may also be changing, so it will be difficult at this level to distinguish a real signal from a change in gain or atmospheric noise. There are several solutions to this problem, depending on the type of observing, all of which rely on some way of forming a reference. If we are making spectral-line measurements, the reference is often just adjacent frequencies. If we scan the frequency or simultaneously divide the spectrum into many frequency channels, then the gain or atmospheric noise changes will be largely common to all frequencies and will cancel with baseline subtraction in the final spectrum. In making measurements of broadband or *continuum* radio emission, we usually use a synchronous detection technique known as *Dicke* switching after its inventor Robert Dicke. An example of Dicke switching is the use of a switch to toggle the input of the LNA between two antenna outputs that provide adjacent beams in the sky. If we switch fast enough in this case and take the difference between the power of the two outputs synchronously with the antenna switch, then receiver gain changes will largely cancel. Furthermore, if the two antenna beams are close together on the sky, then changes in the atmospheric noise will tend to be common to both beams and will also cancel. Since we are taking a difference and spending half the time looking at the reference, the ΔT given above will have to be doubled.

Another powerful technique for extracting weak signals from noise is correlation. The radio telescope in this case has two or more receivers either connected to the same antenna, or, more often, two or more separate antennas. The signal voltages are multiplied together before averaging instead of multiplying the signal voltage by itself to obtain the power. With separated antennas, the correlation output combines the antenna patterns as an interferometer, which generates lobes on the sky that are separated in angle by the wavelength divided by the projected baseline between the antennas. Correlation techniques are common in radio astronomy, and they are becoming popular also in communications. Correlation is used, for example, to detect and demodulate spread-spectrum signals as in code-division multiple-access (CDMA) digital cellular telephones.

6.3.3 An analog-to-digital converter (ADC)

Since all the final processing of a radiometer output is done with a computer, we need to convert analog voltages from the detector to numbers that can be processed in software. A very accurate and effective ADC is a voltage-to-frequency converter followed by a counter. This ADC provides integrated power with as many bits as are needed to represent the count over the integration interval. If reading the output of the counter at perfectly regular intervals is difficult, then another counter can be used simultaneously to count the constant-frequency output of a crystal oscillator. The integrated power is then proportional to the ratio of the counts from the voltage-to-frequency converter to the counts from the crystal oscillator.

6.3.4 Interference

Radio astronomy is often limited by interference especially at low frequencies. The spectrum is overcrowded with transmitters: Earth-based TV, satellite TV, FM, cellular phones, radars, and many others. Radio astronomy has some protected frequency bands, but these bands are often contaminated by harmonics accidentally radiated by TV transmitters, intermodulation from poorly designed transmitters, and noise from leaky high-voltage insulators and automobile ignition noise. Some of the worst offenders are poorly designed satellite transmitters, whose signals come from the sky so that they effect even radio telescopes that are well shielded by the local terrain. Radio telescopes and their receivers can be made more immune to interference by:

a) Including a bandpass filter following the LNA to prevent interference from being generated inside the receiver by intermodulation.
b) Placing the telescope in a location with as much shielding as possible from the local terrain. Low spots (e.g., valleys) are good for low-frequency radio telescopes because they reduce the level of interference from ground-based transmitters. (Prefer a dry mountain top to reduce atmospheric attenuation at millimeter and shorter wavelengths.)
c) Tracking down interference and trying to reduce it at the source.
d) Designing and using an antenna with very low sidelobes.
e) Using an interferometer and correlation processing, which is far more immune to interference.
f) Using data editing to remove data corrupted by interference.

6.4 Spectrometers and spectroscopy

A source of electromagnetic radiation that is in a solid form, such as the surface of a planet or a small grain of dust in interstellar space, has a very smooth spectrum, that is, the intensity of the emission varies quite slowly with frequency. Such emission is called continuum emission -- the spectrum is a continuous function of frequency without sharp features. In this case there is not much restriction on the bandwidth that can be used to detect the radiation. One can use the largest bandwidth permitted by the radiometer to obtain the highest sensitivity.

However, in the case of atoms and molecules in a gaseous state, the emission is discrete. A gas does not produce continuous emission but rather the emission is over a small range of frequencies. The spectrum consists of narrow "spikes" of emission whose width is determined primarily by the motions of the emitting atoms or molecules.

In this case one needs to have much narrower bandwidths which decreases the sensitivity. In order to detect these spectral lines spectrometers are used.

6.4.1 Scanning filter

To measure spectral-line emission or absorption from molecules or atoms, we need a device to measure power spectra -- a spectrometer. An intuitive method to measure power spectra is to scan a narrow tunable bandpass filter across the frequencies to be measured and record its power output as a function of frequency. A variant of this scheme, actually used in spectrum analyzers, has a fixed filter in the IF that is scanned, in effect, by scanning an LO in a heterodyne configuration.

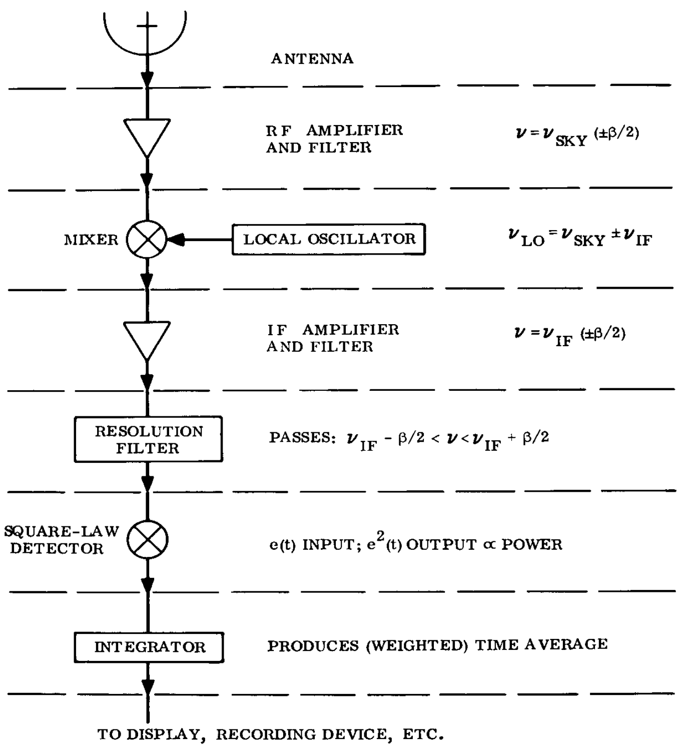

Figure 6.7

Figure 6.7 shows an example of a scanning filter receiver. Variations in the power spectrum narrower than the width of the scanning filter are smoothed over and lost. The width and shape of this filter characterize the spectrometer's resolution. This scheme works but is wasteful because all the information outside the instantaneous position of the filter is ignored.

6.4.2 Comb of filters -- filter bank

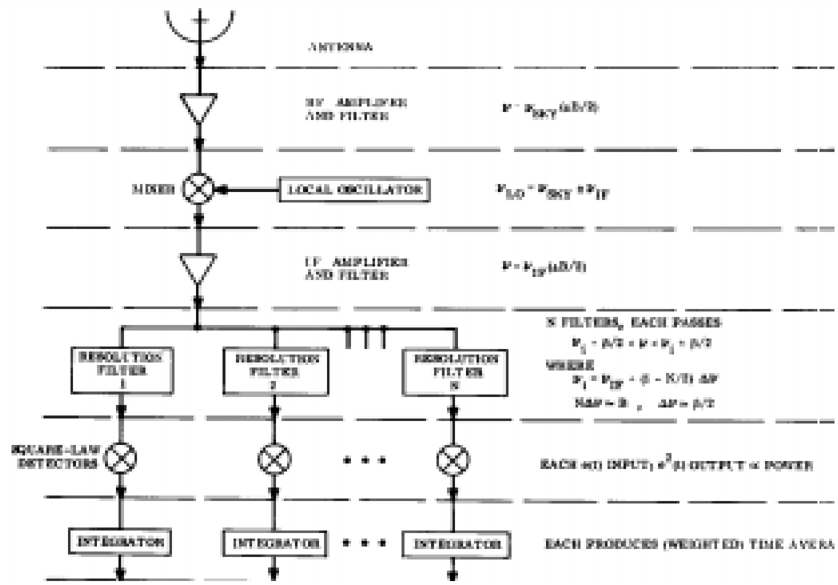

Figure 6.8

A significant improvement in observing efficiency results from having a comb or bank of bandpass filters placed side-by-side in frequency and recording all their outputs simultaneously. Figure 6.8 shows an example of such a filter bank. Choosing filter shapes and spacings for such a spectrometer is, however, not intuitive. The popular almost-square filters placed just touching, for example, give spectra that are difficult to interpret whenever spectral features are comparable to the filter widths. With today's technology, filter banks are expensive and troublesome compared to various digital alternatives.

6.4.3 Autocorrelations and Fourier transforms

A. Why?

Some authors define the power spectrum to be the Fourier transform of the autocorrelation of the voltage and then show that this definition accords, at least approximately, with the intuitive scanning filter. If autocorrelations are done for lags, τ, up to some maximum, $\tau_{\max}$, and Fourier transforms are done with no weighting, then, except for noise considerations, the resulting power spectra will be the same as would be obtained with a scanning filter whose shape is $\sin(2\pi\nu\tau_{\max})/(2\pi\nu\tau_{\max})$. If the spectra to be measured are band limited, perhaps by a preceding lowpass filter, then, by the Nyquist theorem, autocorrelations can be done at uniform finite lag steps, $\tau_S = 1/(2\nu_{\max})$, where $\nu_{\max}$ is the maximum frequency of this band. This sampling corresponds to two data per cycle of $\nu_{\max}$. The resulting spectra are also band limited because they contain no lags above $\tau_{\max}$, which is the Nyquist limit. The band-limited spectra from an autocorrelation spectrometer are, then, smooth continuous functions of frequency, and they can be specified at a finite set of evenly spaced points provided that these points are no farther apart than a Nyquist step, $\nu_{step} = 1/(2\tau_{\max})$. The number of these frequency steps is the same as the number of lag steps, namely $2\nu_{\max}\tau_{\max}$, which sometimes leads to confusion. Smooth continuous spectra can be obtained from finite sets of points by convolution by the same $\sin(2\pi\nu\tau_{\max})/(2\pi\nu\tau_{\max})$.

The point of using autocorrelations to get power spectra and of quantizing both autocorrelations and spectra is to allow these operations to be done digitally. The autocorrelations are usually done in hardware, the Fourier transforms in software. This is usually a significant simplification compared to a filter bank.

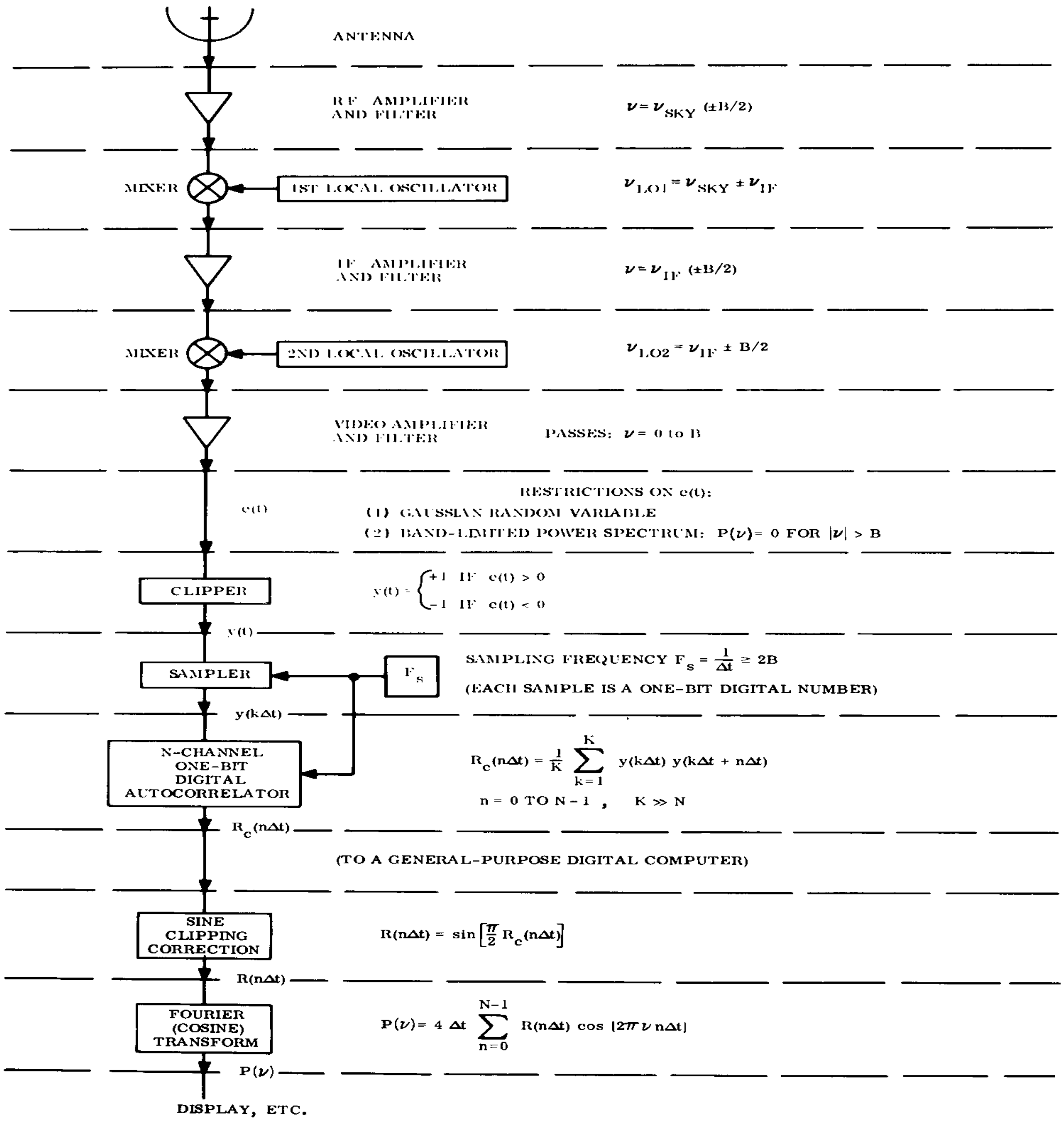

Figure 6.9

B. Multibit vs one bit

A further simplification is possible because of the nature of the signals to be measured. For Gaussian random noise, autocorrelations and power spectra can be computed from one-bit (just sign) samples of the voltage. Figure 6.9 shows an example of such a configuration. The price to pay for this simplification is about 31% additional noise and a little more computer arithmetic to correct the one-bit autocorrelations before the Fourier transform. Haystack now uses 1.5-bit (i.e., three-level) sampling, which adds about 16% additional noise (γ in the noise equation above) and is only a little more complex than one-bit.

6.4.4 Switching schemes and baselines -- Why switch?

One of the most persistent and difficult problems in spectral measurements in radio astronomy involves the difficulty of obtaining good flat baselines -- the parts of spectra with no signal. The corresponding baseline problem with continuum measurements involves stable measurements off source (cold sky) to subtract from on-source measurements. There are numerous instrumental effects that contribute to bumpy baselines, and many of the effects are a substantial percentage of the system temperature and are typically much larger than the signals to be measured. To reduce the severity of such instrumental effects, almost all radio-astronomy measurements are made using one of several possible switching schemes. The ideal switching scheme would have the source itself turn off and on synchronously with a prescribed periodicity and with nothing else changing. Then the difference between signal and comparison is precisely the desired measurement. Provided that T_s does not change between on and off, one can show that noise is minimized by spending half the time on source, half off.

Among practical switching schemes, we can move the antenna pointing on and off the source either by actually moving the antenna or by offsetting the effective pointing by moving the feed or its image. But this works only if the source is confined in angle. Or we can move the source in and out of the passband by moving the LO frequency. But this works only if the source is confined in frequency. Or we can switch the input to the receiver alternately from the feed to an absorber or load. This last scheme is less good because many of the instrumental effects to be ameliorated are in the feed and beyond.

Switching schemes are needed for either continuum or spectral measurements. A price to be paid for switching is increased noise. There are two contributions to the added noise: The receiver spends only half the time looking at the signal, and the result is the difference between two equally noisy measurements. The result is twice the noise compared with not switching for the same integration time (this is α in the noise equation above), or four times the integration time to achieve the same noise.