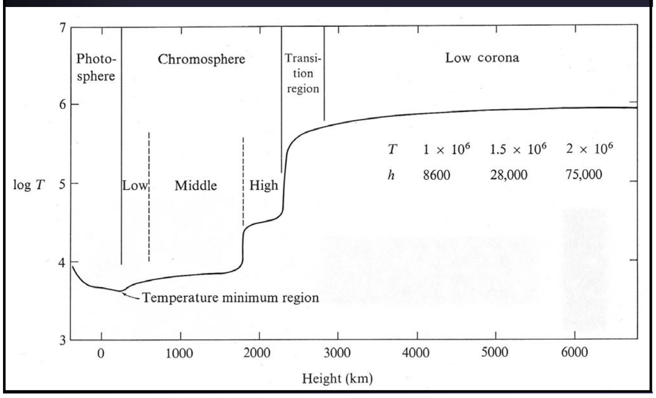
The solar atmosphere

The Sun's atmosphere

- The solar atmosphere is generally described as being composed of multiple layers, with the lowest layer being the photosphere, followed by the chromosphere, the transition region and the corona.
- In its simplest form it is modelled as a single component plane-parallel atmosphere.
- Density drops exponentially: $\rho(z) = \rho_0 \exp(-z/H_\rho)$ (for isothermal atmosphere). T=6000K ➔ $H_\rho \approx 100\text{km}$
- Mass of the solar atmosphere ≈ mass of the Indian ocean (≈ mass of the photosphere)
- Mass of the chromosphere ≈ mass of the Earth's atmosphere

1-D stratification



How good is the 1-D approximation?

- 1-D models reproduce extremely well large parts of the spectrum obtained with low spatial resolution (see spectral synthesis slide)
- However, any high resolution image of the Sun shows that its atmosphere has a complex structure (as seen at almost any wavelength)
- Therefore: 1-D models may well describe averaged quantities relatively well, although they probably do not describe any part of the real Sun at all.

The photosphere

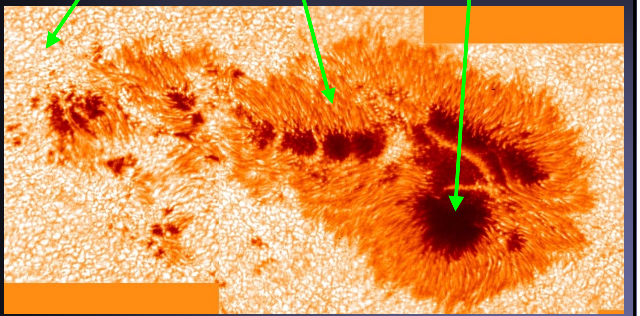
- The photosphere extends between the solar surface and the temperature minimum, from which most of the solar radiation arises.
- The visible, UV (λ> 1600Å) and IR (< 100μm) radiation comes from the photosphere.
- 4000 K < T(photosphere) < 6000 K
- T decreases outwards ➔ $B_\nu(T)$ decreases outward ➔ absorption spectrum
- LTE is a good approximation
- Energy transport by convection and radiation
- Main structures: Granules, sunspots and faculae

The Sun in White Light

(limb darkening removed)

MDI on SOHO

Sunspots



H. Schleicher, KIS/VTT, Obs. del Teide, Tenerife

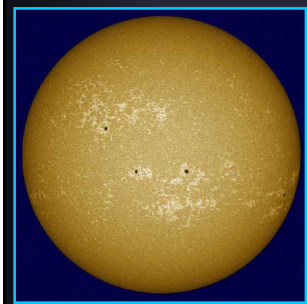
Granulation

- Physics of convection and the properties of granulation and supergranulation have been discussed earlier, so that we can skip them here.

Chromosphere

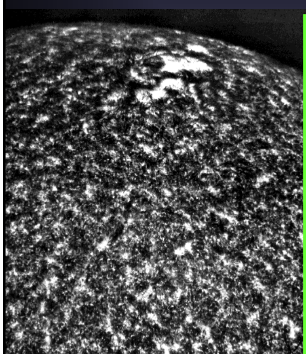
- Layer lying just above the photosphere, at which the temperature appears to be increasing outwards (classically forming a temperature plateau at around 7000 K)
- Assumption of LTE breaks down
- Energy transport mainly by radiation and waves
- Assumption of plane parallel atmosphere is very likely to break down as well.
- Strong evidence for a spatially and temporally inhomogeneous chromosphere (gas at T<4000K is present beside gas with T>8000 K)

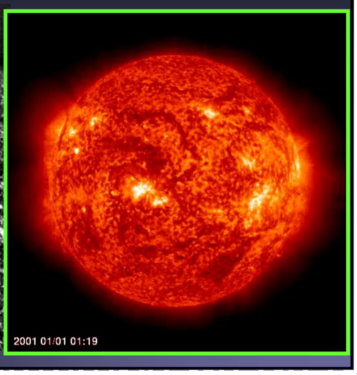
Chromospheric structure



7000 K gas Ca II K

5 10^4 K gas (EIT He 304 Å)



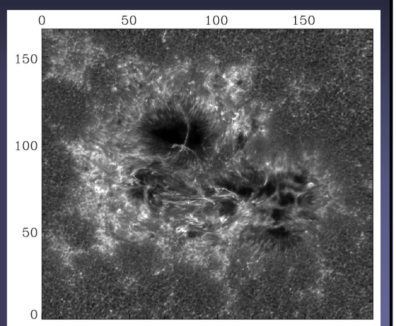


Ca K core, B. Gillespie, NSO

Chromospheric structure II

- The chromosphere exhibits a very wide variety of structures. E.g.,
  - Sunspots and Plages
  - Network and internetwork (grains)
  - Spicules
  - Prominences and filaments
  - Flares and eruptions

DOT Ca II K core chromosphere / DOT G-band photosphere



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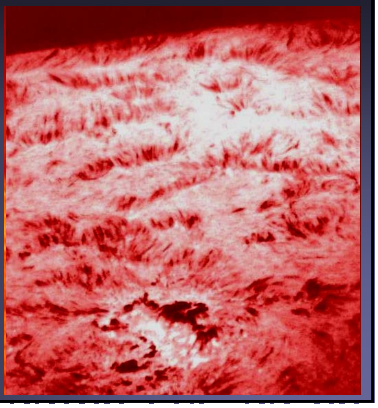

Chromospheric structure

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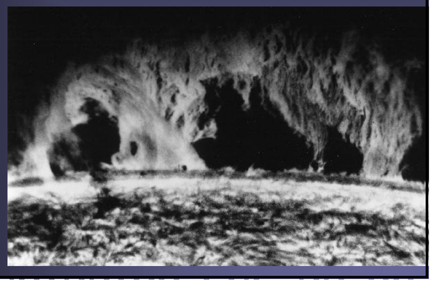
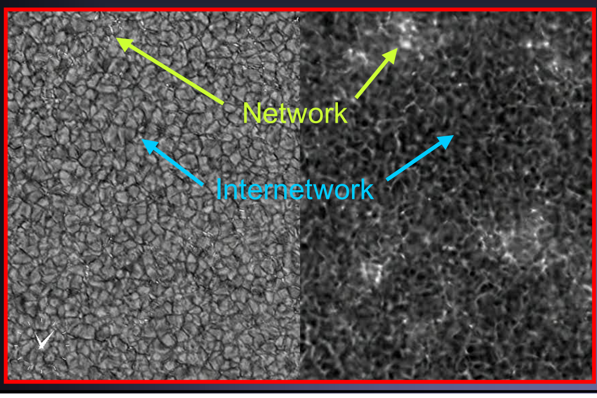
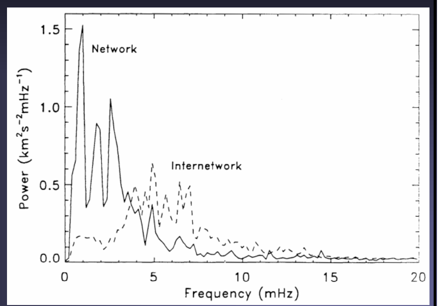


Chromospheric dynamics (DOT)



Chromospheric dynamics

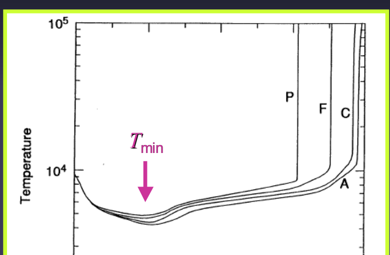
- Oscillations, seen in cores of strong lines
- Power at 3 min in Inter-network
- Power at 5-7 min in Network



Lites et al. 2002

Models: the classical chromosphere

- Classical picture: plane parallel, multi-component atmospheres
- Chromosphere is composed of a gentle rise in temperature between T_{min} and transition region.

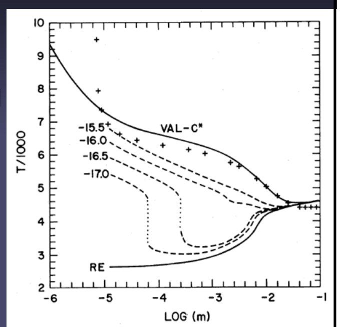


Fontenla et al. 1993

Need to heat the chromosphere

- Radiative equilibrium, RE: only form of energy transport is radiation & atmosphere is in thermal equilibrium.
- VAL-C: empirical model
- Dashed curves: temp. stratifications for increasing amount of heating (from bottom to top).

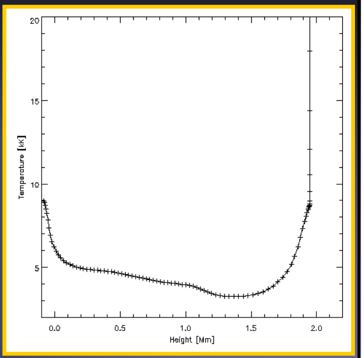
→ Mechanical heating needed to reproduce obs.



Anderson & Athay 1993

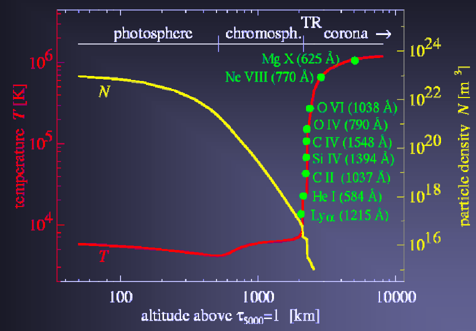
Dynamic models

- Start with piston in convection zone, consistent with obs. of photospheric oscillations
- Waves with periods of ≤3min propagate into chromosphere
- Energy conservation ($\rho v^2/2$ = const.) & strong ρ decrease ➔ wave amplitudes increase with height: waves steepen and shock
- Temp. at chromospheric heights varies between 3000 K and 10000 K



Carlsson & Stein

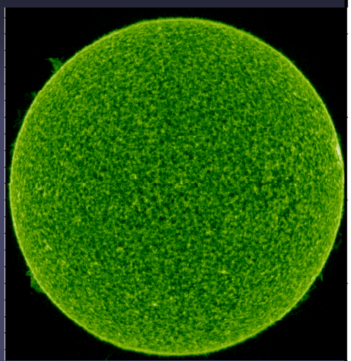
Transition Region



Semi-empirical 1D-models of solar atmosphere: steep increase of T in transition region (TR): < 100 km thick

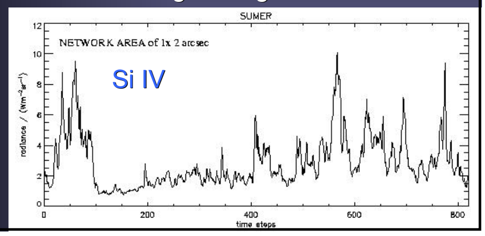
TR spatial structure

- Lower transition region (T<5 10^5 K) shows structure very similar to chromosphere, with network, plage etc.
- C IV (10^5 K) imaged by SUMER
- In upper transition region structures are more similar to corona



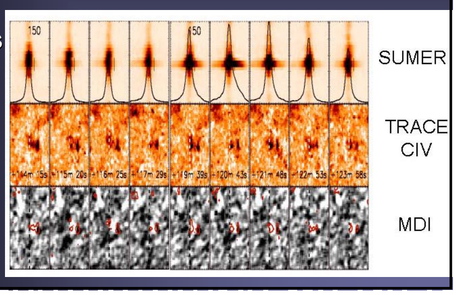
TR dynamic phenomena: blinkers

- Brightness variability in the (quiet) transition region is larger than in any other layer of solar atmosphere
- Typical brightening: blinkers
- Occur everywhere, all the time. Last for minutes to hours. How much of the brightening is due to overlapping blinkers?
- 1 time step ≈ 1 minute



Explosive events

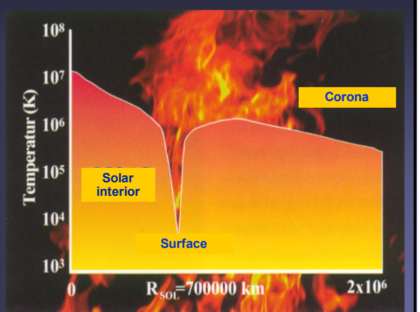
- Broadenings of TR spectral lines at 1-3 10^5 K
- Typical “normal” line width 20 km/s, in explosive event: up to 400 km/s. Cover only a few 1000 km and last only a few minutes
- Typically a few 1000 present on Sun at any given time



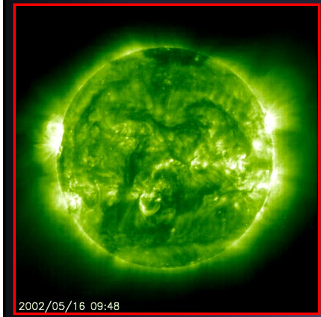
The Solar Corona

While the surface is about 6,000 K, the temperature in the corona reaches about 2 million K.

What causes this rapid increase in temperature is still one of the big mysteries in solar physics.



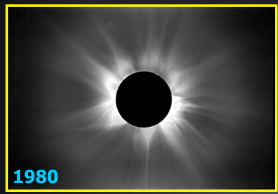
The Hot and Dynamic Corona

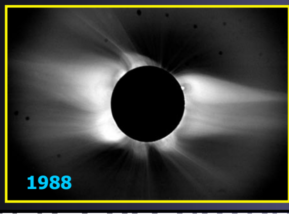


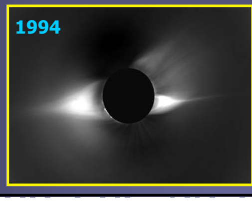
EUV Corona: 10^6 K plasma
(EIT/SOHO 195 Å)

White light corona
(LASCO C3 / SOHO, MPAe)

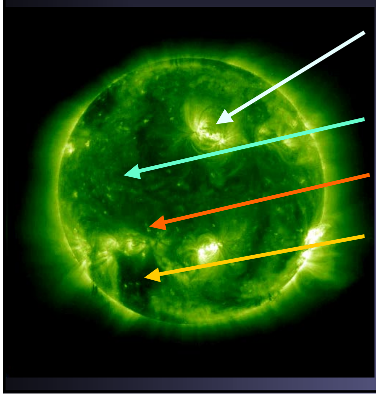
Solar corona during eclipses







Coronal structures

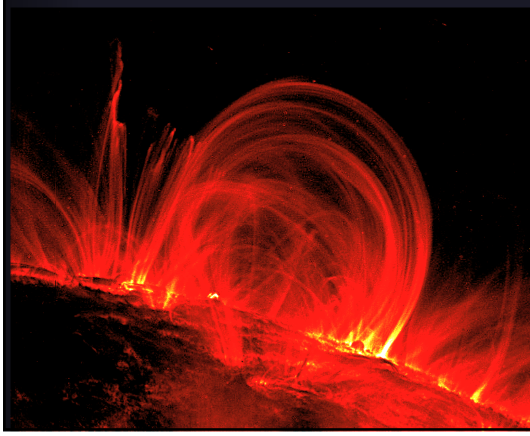


- Active regions (loops)
- Quiet Sun
- X-ray bright points
- Coronal holes
- Arcades

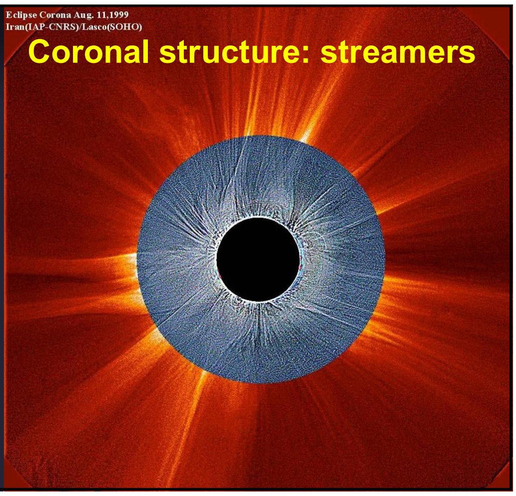
Fe XII 195 Å
(1.500.000 K)
17 May - 8 June 1998

Coronal structure: active region loops

TRACE, 1999

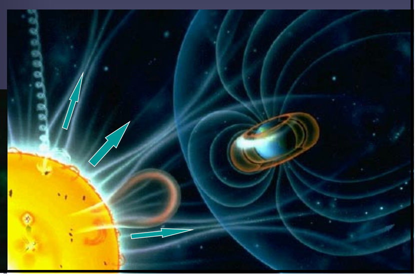


Coronal structure: streamers

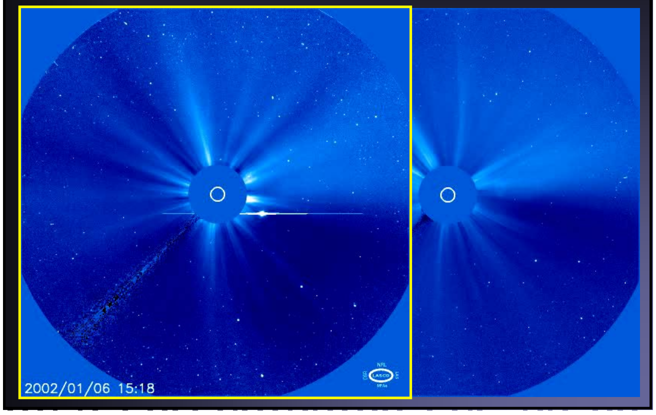


The solar wind

A constant stream of particles flowing from the Sun's corona, with a temperature of about a million degrees and with a velocity of about 450 km/s. The solar wind reaches out beyond Pluto's orbit, with the heliopause located roughly at 100-120 AU



Comets and the solar wind



Solar wind characteristics at 1AU

Fast solar wind
- speed > 400 km/s
- $n_p \approx 3\ cm^{-3}$
- homogeneous
- B≈ 5 nT = 0.0005G
- 95% H, 4% He
- Alfvenic fluctuations
- Origin: coronal holes coronal

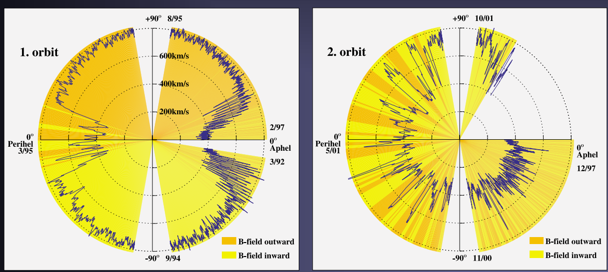
Slow solar wind
- <400 km/s
- ≈ $8\ cm^{-3}$
- high variability
- B < 5 nT
- 94% H, 5% He
- Density fluctuations
- Origin: in connection with streamers

Transient solar wind
- speed from < 300 km/s up to >2000 km/s
- Variable B, with B up to 100 nT (0.01G)
- Often very low density
- Sometimes up to 30% He
- Often associated with interplanetary shock waves
- Origin: CMEs

3-D structure of the Solar Wind: Variation over the Solar Cycle

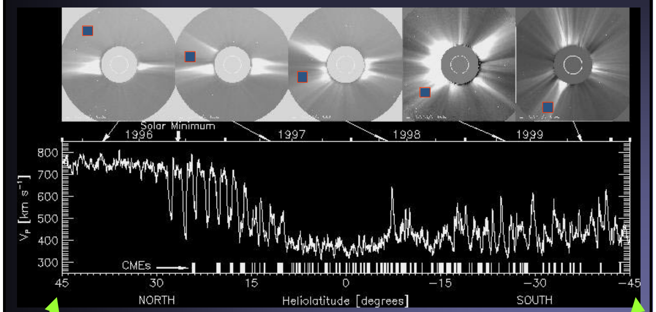
1st Orbit: 3/1992 - 11/1997 declining / minimum phase

2nd Orbit: 12/1997 - 2/2002 rising / maximum phase



Ulysses SWICS data   Woch et al. GRL

Coronal Shape & Solar Wind: Ulysses Data & 3D-Heliosphere



Activity minimum   Activity maximum

Parker's theory of the solar wind

- Basic idea: dynamic equilibrium between hot corona and interstellar medium. Mass and momentum balance equations:

$$\frac{d}{dr}(\rho r^2 v) = 0$$

$$v\frac{dv}{dr} = -\frac{1}{\rho}\frac{dP}{dr}\frac{GM}{r^2}$$

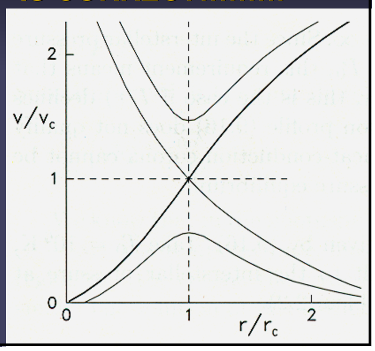
➔Parker's Eq. for solar wind speed (isothermal atmosphere)

$$\frac{1}{v}\frac{dv}{dr}(v^2 - c_S{}^2) = \frac{2c_S{}^2}{r} - \frac{GM}{r^2}$$

Parker's solar wind solutions

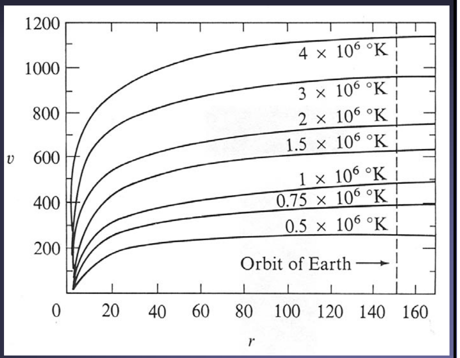
CHECK WHY ONLY ONE SOLN IS CORRECT!!!!!!!!!

- Parker found 4 families of solar wind solutions
- 2 not supported by Obs. (supersonic at solar surface)
- 1 does not give sufficient pressure against the interstellar medium.
- Correct solution must be thick line in Fig.

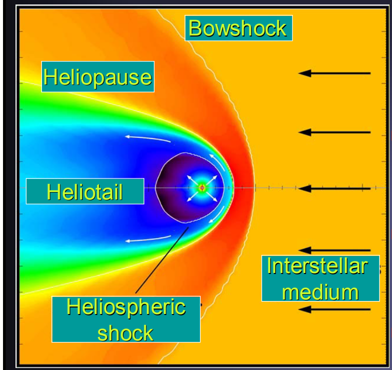


Solar wind speed

- Speed of solar wind predicted by Parker's model for different coronal temperatures (simple, isothermal case; no magnetic field)



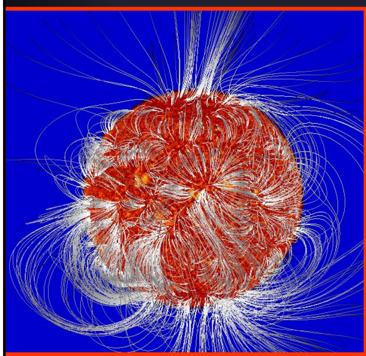
The Heliosphere

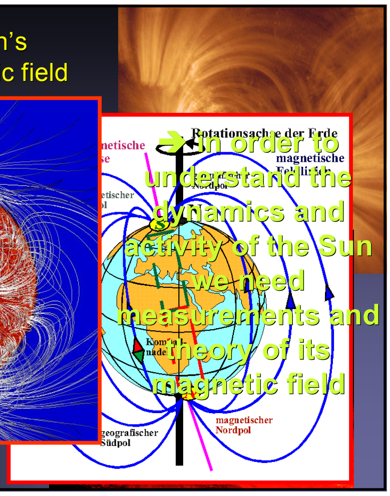


- Heliosphere = region of space in which the solar wind and solar magnetic field dominate over the instellar medium and the galactic magnetic field.
- Bowshock: where the interstellar medium is slowed relative to the Sun.
- Heliospheric shock: where the solar wind is decelerated relative to Sun
- Heliopause: boundary of the heliosphere

Magnetic Field

The source of the Sun's activity is the magnetic field

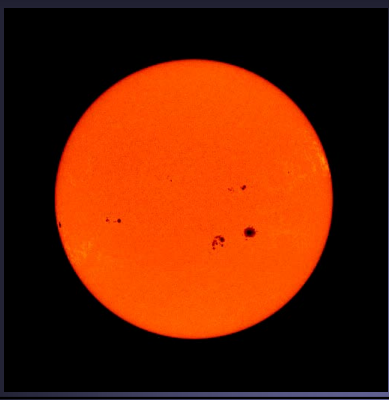




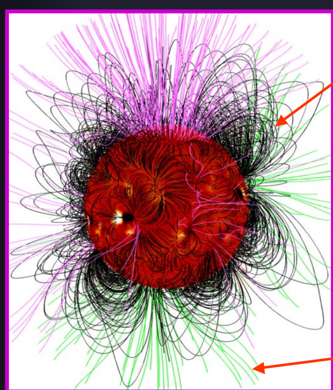
In order to understand the dynamics and activity of the Sun we need measurements and theory of its magnetic field

Thomas Wiegelmann

Correlation of field with brightness



Open and closed magnetic flux



Closed flux: slow solar wind

Most of the solar flux returns to the solar surface within a few $R_\odot$ (closed flux)

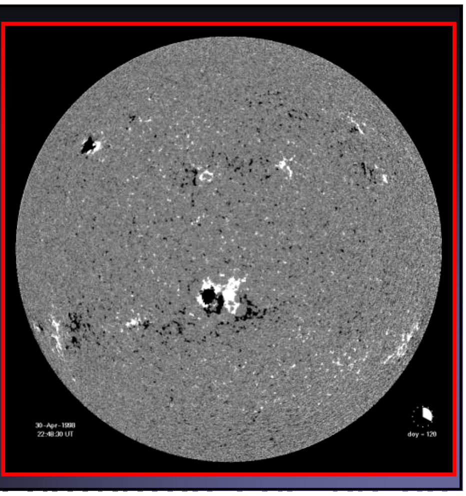
A small part of the total flux through the solar surface connects as open flux to interplanetary space

Open flux: fast solar wind

Measured Magnetic Field at Sun's Surface

Month long sequence of magnetograms (approx. one solar rotation)

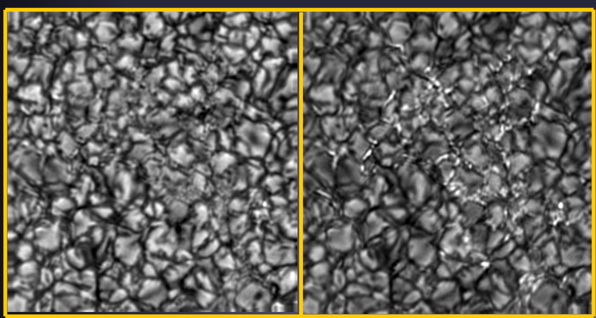
MDI/SOHO May 1998



Methods of magnetic field measurement

- Direct methods:
  - Zeeman effect ➔ polarized radiation
  - Hanle effect ➔ polarized radiation
  - Gyroresonance ➔ radio spectra
- Indirect methods: Proxies
  - Bright or dark features in photosphere (sunspots, G-band bright points)
  - Ca II H and K plage
  - Fibrils seen in chromospheric lines, e.g. Hα
  - Coronal loops seen in EUV or X-radiation

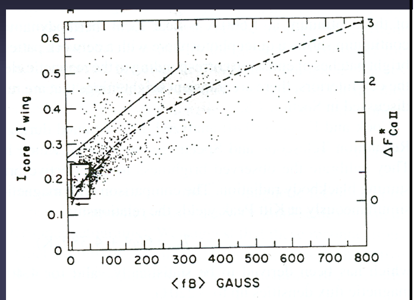
Example of proxies: Continuum vs. G-band



Continuum G-band

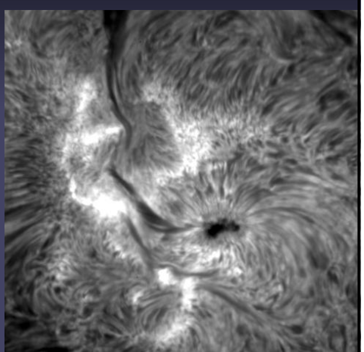
Ca II K as a magnetic field proxy

- Ca II H and K lines, the strongest lines in the visible solar spectrum, show a strongly increasing brightness with non-spot magnetic flux.
- The increase is slower than linear
- Magnetic regions (except sunspots) appear bright in Ca II: Ca plage and network regions



Hα and the chromospheric field

- Hα images of active regions show a structure similar to iron filing around a magnet. Do they (roughly) follow the field lines?
- Relatively horizontal field in chromosphere?
- Note spiral structure around sunspot.



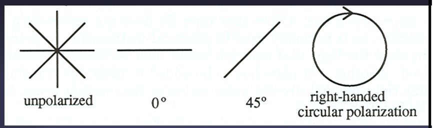
Zeeman diagnostics

- Direct detection of magnetic field by observation of magnetically induced splitting and polarisation of spectral lines
- Important: Zeeman effect changes not just the spectral shape of a spectral line (often subtle and difficult to measure), but also introduces a unique polarisation signature

➔ Measurement of polarization is central to measuring solar magnetic fields.

Polarized radiation

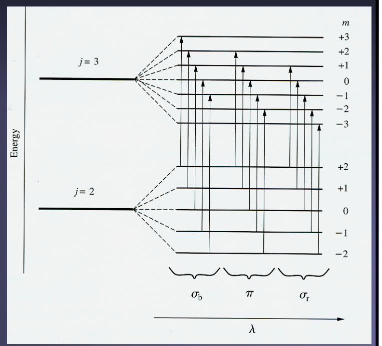
- Polarized radiation is described by the 4 Stokes parameters: I, Q, U and V



- I = total intensity = $I_{\text{lin}}(0^o) + I_{\text{lin}}(90^o) = I_{\text{lin}}(45^o) + I_{\text{lin}}(135^o) = I_{circ}(right) + I_{\text{circ}}(\text{left})$
- $Q = I_{\text{lin}}(0^o) - I_{\text{lin}}(90^o)$
- $U = I_{\text{lin}}(45^o) - I_{\text{lin}}(135^o)$
- $V = I_{\text{circ}}(\text{right}) - I_{\text{circ}}(\text{left})$
- Note: Stokes parameters are sums and differences of intensities, i.e. they are directly measurable

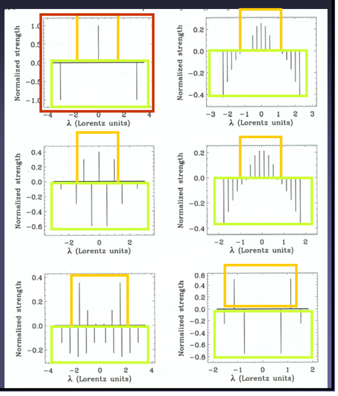
Zeeman splitting of atomic levels

- In the presence of a B-field a level with total angular moment J will split into $2J+1$ sublevels with different M.
- $E_{J,M} = E_J + \mu_0 g M_J B$
- Transitions are allowed between levels with $\Delta J = 0,\pm1$ & $\Delta M = 0$ (π), ±1 (σ_b, σ_r)
- Splitting is determined by Lande factor g : $g(J,L,S) = 1+(J(J+1)+S(S+1)-L(L+1))/2J(J+1)$

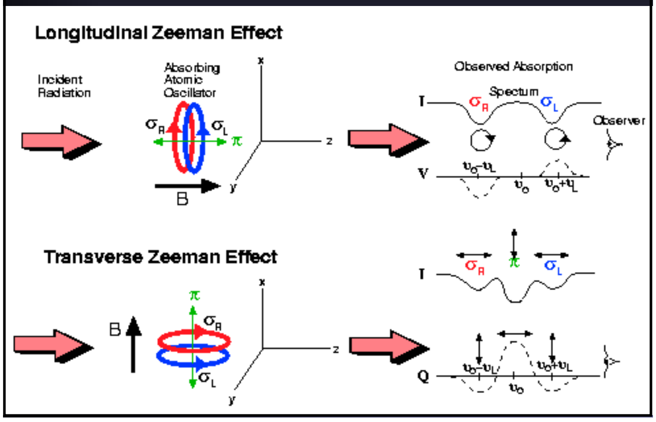


Splitting patterns of lines

- Depending on g of the upper and lower levels, the spectral line shows different splitting patterns
- Positive: π components: $\Delta M=0$
- Negative: σ components: $\Delta M=\pm1$
- Top left: normal Zeeman effect (rare)
- Rest: anomalous Zeeman effect (usual)



Polarization and Zeeman effect



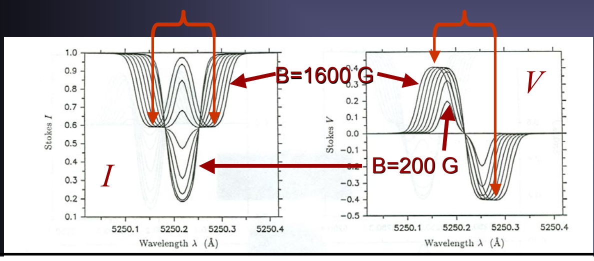
Effect of changing field strength

Formula for Zeeman splitting (for B in G, λ in Å):

$\Delta\lambda_H = 4.67\ 10^{-13}\ g_{\text{eff}} B \lambda^2$ [Å]

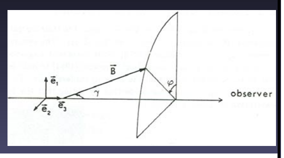
g_{eff}=effective Lande factor of line

For large B: $\Delta\lambda_H$ = Δλ between σ-component peaks

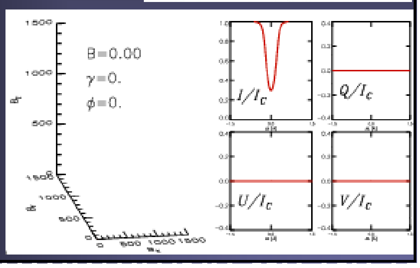


Dependence on B, γ, & φ

- $I \sim \kappa_\sigma(1+\cos^2\gamma)/4 + \kappa_\pi \sin^2\gamma/2$
- $Q \sim B^2 \sin^2\gamma \cos 2\varphi$
- $U \sim B^2 \sin^2\gamma \sin 2\varphi$
- $V \sim B \cos \gamma$



- Q, U: transverse component of B
- V: longitudinal component of B

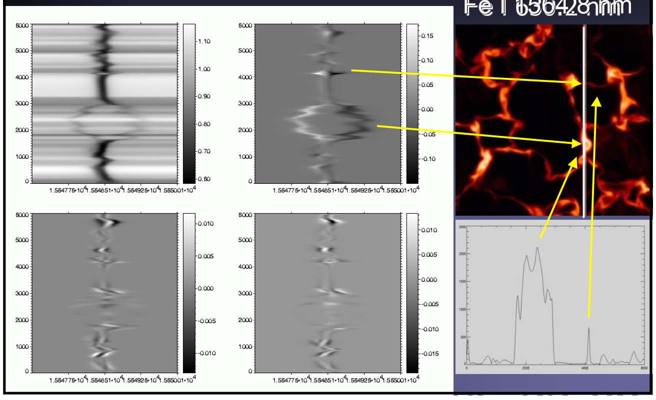


Juanma Borrero

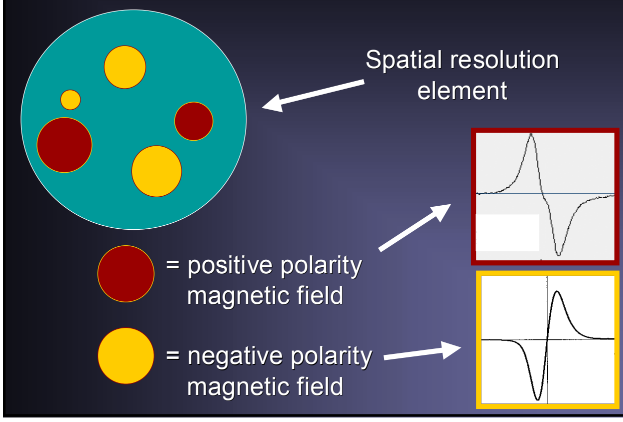
Zeeman polarimetry

- Most used remote sensing of astrophysical (and certainly solar) magnetic fields
- Effective measurement of field strength if Zeeman splitting is comparable to Doppler width or more: $B > 200$ G … 1000 G (depending on spectral line) ➔ works best in photosphere
- Splitting scales with λ ➔ works best in IR
- Sensitive to cancellation of opposite magnetic polarities ➔ needs high spatial resolution

Effect of wavelength of spectral line

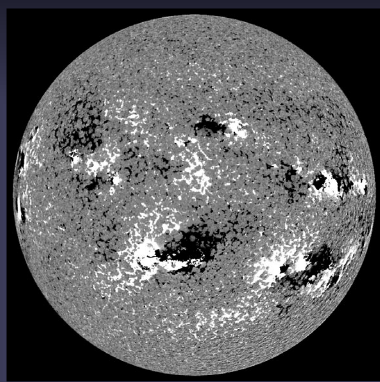


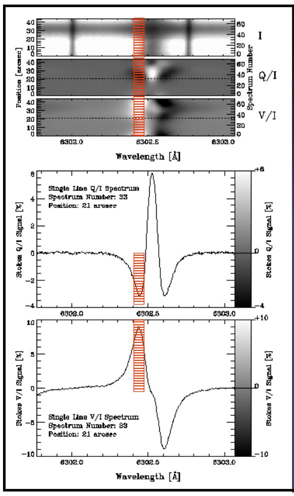
Cancellation of magnetic polarity



Magnetograms

- Magnetograph: Instrument that makes maps of (net circular) polarization in wing of Zeeman sensitive line.
- Example of magnetogram obtained by MDI
- Conversion of polarization into magnetic field requires a careful calibration.

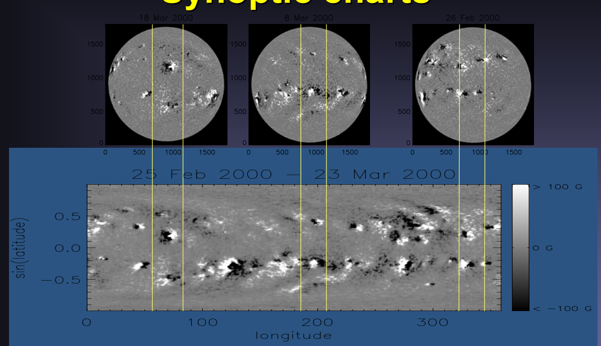




What does a magnetogram show?

- Plotted at left:
  - Top: Stokes I, Q and V along a spectrograph slit
  - Middle: Sample Stokes Q profile
  - Bottom: Sample Stokes V profile
  - Red bars: example of a spectral range used to make a magnetogram. Generally only Stokes V is used (simplest to measure), gives longitudinal component of B.

Synoptic charts



Synoptic maps approximate the radial magnetic flux observed near the central meridian over a period of 27.27 days (= 1 Carrington rotation)

Polarized radiative transfer

- Complication: RTE required for 4 Stokes parameters: Written as differential equation for Stokes vector $\mathbf{I}_\nu = (I_\nu, Q_\nu, U_\nu, V_\nu)$
- Eq. in plane parallel atmosphere for a spectral line (Unno-Rachkowsky equations):

  $\mu\, d\mathbf{I}_\nu / d\tau_c = \mathbf{\Omega}_\nu \mathbf{I}_\nu - \mathbf{S}_\nu$

- $\mathbf{\Omega}_\nu$ = absorption matrix (basically ratio of line to continuum absorption coefficient), $\mathbf{S}_\nu$ = source function vector, τ_c = continuum optical depth.

Polarized radiative transfer II

The absorption matrix

$$\Omega_\nu = \begin{pmatrix} 1+\eta_I & \eta_Q & \eta_U & \eta_V \\ \eta_Q & 1+\eta_I & -\rho_V & \rho_U \\ \eta_U & \rho_V & 1+\eta_I & \rho_Q \\ \eta_V & -\rho_U & -\rho_Q & 1+\eta_I \end{pmatrix}$$

SAY MORE ABOUT MATRIX ELEMENTS!!!!!! SHOW HOW ZEEMAN EFFECT ENTERS INTO THEM, ETC.!!!!!

Polarized radiative transfer III

- The Zeeman effect only enters through $\mathbf{\Omega}_\nu$
- $\mathbf{\Omega}_\nu$ contains besides absorption due to Zeeman-split line $(\eta_I, \eta_Q, \eta_U, \eta_V)$ also magnetooptical effects, such as Faraday rotation (ρ_Q, ρ_U, ρ_V) : rotation of plane of polarization when light passes through B.
- $\mathbf{\Omega}_\nu = \mathbf{\Omega}_\nu(\gamma, \varphi, B)$, i.e. $\mathbf{\Omega}_\nu$ depends on the full magnetic vector (in addition to the usual quantities that the absorption coefficient depends on)

LTE

- In LTE the Unno-Rachkowsky equations simplify since

  $\mathbf{S}_\nu = (B_\nu, 0, 0, 0)$

  Here B_ν = Planck function
- Also, $\mathbf{\Omega}_\nu$ is simplified. The $\eta_I, \eta_Q, \eta_U, \eta_V$ and ρ_Q, ρ_U, ρ_V values only require application of Saha-Boltzmann equations (similar situation as for LTE in case of normal radiative transfer). Each of these quantities is, of course, frequency dependent.

Solution of Unno Eqs

- General solution best done numerically (even formal solution is non-trivial: exponent of matrix $\mathbf{\Omega}_\nu$)
- Simple analytical solutions exist for a Milne-Eddington atmosphere (i.e. for $\mathbf{\Omega}_\nu$ independent of τ_ν and $\mathbf{S}_\nu$ depending only linearly on τ_ν). Particularly simple if we neglect magneto-optical effects
- $I(\mu) = \beta\, \mu\, (1+\eta_I)/\Delta$
- $P(\mu) = \beta\, \mu\, \eta_P/\Delta$, where $P = Q, U,$ or V
- $\Delta = (1+\eta_I)^2 - \eta_Q^2 - \eta_U^2 - \eta_V^2$

  takes care of line saturation
- β is derivative of Planck function with respect to τ_ν .

Basics: magnetic pressure

- Magnetic field exerts a pressure. Pressure balance between two components of the atmosphere, 1 and 2 (Gauss units):

$$\frac{B_1^2}{8\pi} + P_1 = P_2 + \frac{B_2^2}{8\pi}$$

- If, e.g. $B_2 = 0$, then $P_1 < P_2$ and it follows:

→ Magnetic features are evacuated compared to surroundings.

- If $B_2 = 0$ and $T_1 = T_2$, then also $\rho_1 < \rho_2$, so that the magnetic features are buoyant compared to the surrounding gas.
- In the convection zone this buoyancy means that rising field bundles (flux tubes) keep rising (unless stopped by other forces, e.g. curvature forces.

Basics: plasma β

- Plasma β describes the ratio of thermal to magnetic energy density:

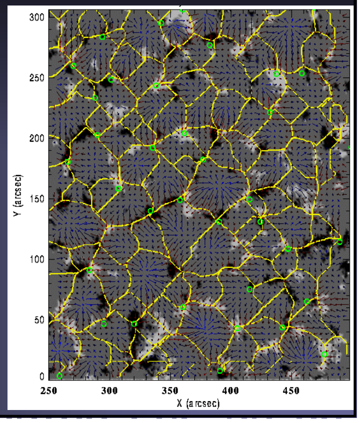
$$\beta = \frac{8\pi P}{B^2}$$

- $\beta < 1$ ➔ Magnetic field dominates and dictates the dynamics of the gas
- $\beta > 1$ ➔ Thermal energy, i.e. gas dominates & forces the field to follow
- β changes with $r/R_\odot$
  - $\beta > 1$ in convection zone
  - $\beta < 1$ in atmosphere, particularly in corona $\beta << 1$

Supergranules and magnetic field

- Magnetogram: black and white patches
- Horizontal velocity: arrows
- Divergence: blue arrows > 0; red arrows: < 0
- Supergranule boundaries: yellow
- Magnetic field is concentrated at edges of supergranules

➔ B swept out by flow



Frozen-in magnetic fields

- Magnetic field is swept to supergranule boundaries ➔ magnetic field is “frozen” into the plasma
- This happens if there are a sufficient number of ionised particles, or equivalently, if the electric conductivity is very high, since charged particles cannot cross field lines (gyration)
- This is the case, even in the cool photosphere of sunspots (only 10^{-4} of all particles are ionized), due to the large number of collisions

➔ If plasma moves perpendicularly to the field, it drags the field with it (or is stopped by the field) and vice versa. Flows parallel to the field are unaffected.

Magnetic field in the convection zone

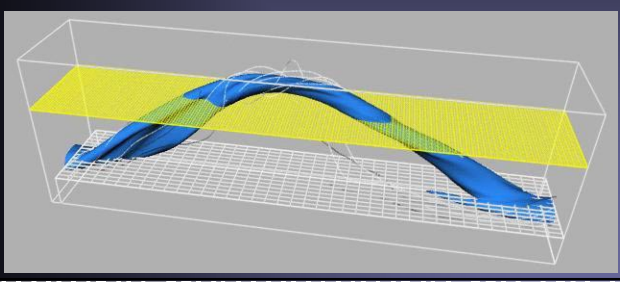
- Magnetic field thought to be produced by a dynamo located near the bottom of the convection zone (e.g. in the overshoot layer below the convection zone).

➔ toroidal flux tubes

- Once field becomes strong enough, it is susceptible to buoyancy (Parker instability)
- A loop-like structure moves towards the solar surface and breaks out.

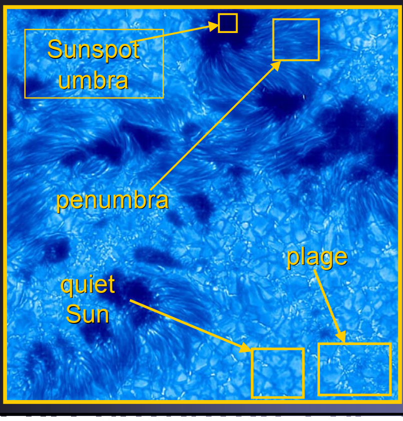
Emergence of a magnetic flux tube

Magnetic field is believed to be generated mainly in the Tachocline near bottom of convection zone.
Due to its buoyancy (see earlier slide; Parker instability), a magnetic field will rise towards the solar surface. At the solar surface it will produce a bipolar active region.



Emergence and evolution of active region: GET BETTER MOVIE!!!!!!!!!!

The photospheric field

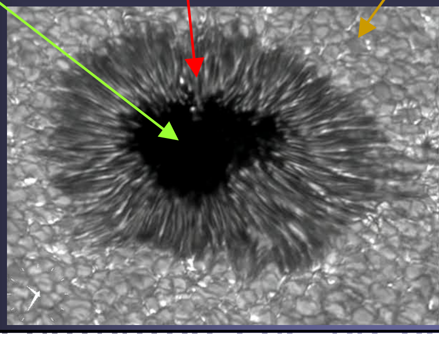


Sunspots

Umbra $T_{eff} \approx 4500$ K

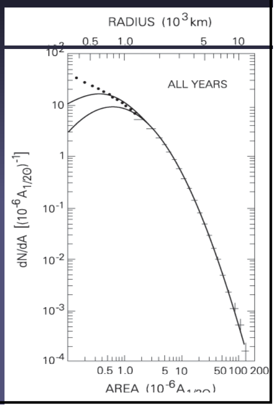
Penumbra $T_{eff} \approx 5500$ K

Granule $T_{eff} \approx 5800$ K



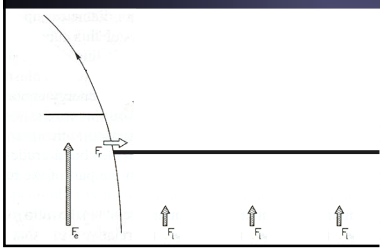
Sunspots, some properties

- Field strength: Peak values 2000-3500 G
- Brightness: umbra: 20% of quiet Sun, penumbra: 75%
- Sizes: Log-normal size distribution. Overlap with pores (log-normal = Gaussian on a logarithmic scale)
- Lifetimes: T between hours & months: Gnevyshev-Waldmeier rule: $A_{max} \sim T$, where A_{max} = max spot area.



Why are sunspots dark?

- Basically the strong nearly vertical magnetic field, not allowing motions across the field lines, quenches convection inside the spot.
- Since convection is the main source of energy transport just below the surface, less energy reaches the surface through the spot ➔ dark



Why are sunspots dark? II

- Where does the energy blocked by sunspots go?
- Spruit (1982) showed: both heat capacity and thermal conductivity of CZ gas is very large

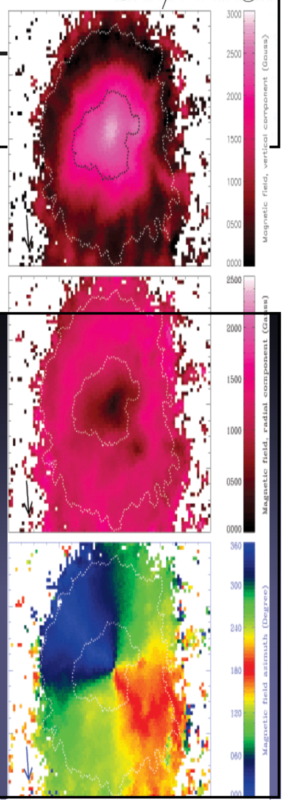
➔ High thermal conductivity: blocked heat is redistributed throughout CZ (no bright rings around sunspots)

High heat capacity: the additional heat does not lead to a measurable increase in temperature

- In addition: time scale for thermal relaxation of the CZ is long, 10^5 years: excess energy is released almost imperceptibly.

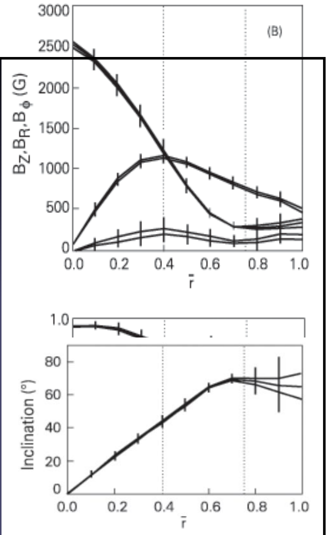
Magnetic structure of sunspots

- Peak field strength ≈ 2000 – 3500 G (usually in darkest, central part of umbra)
- B drops steadily towards boundary, $B(R_{spot}) \approx 1000$ G
- At centre, field is vertical, becoming almost horizontal near R_{spot} .
- Regular spots have a field structure similar to a buried dipole

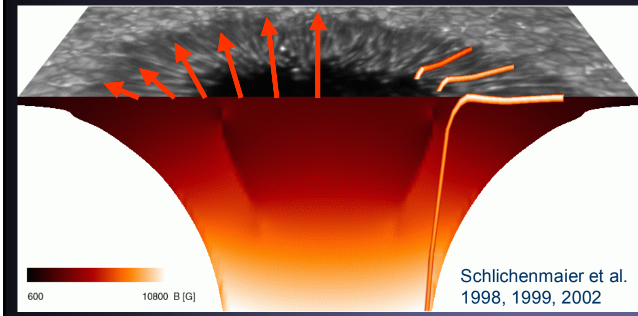


Magnetic structure of sunspots II

Azimuthal averages of the various magnetic field components in a sample of regular (near-circular) medium-sized sunspots.

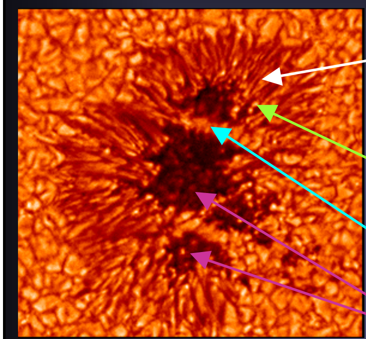


Magnetic structure of sunspots



- Regular on large scales (≈ dipole, B_{max} ≈ 2500 G, for simple spots)
- Extremely complex on small scales (penumbra, subsurface)

Sunspot fine structure



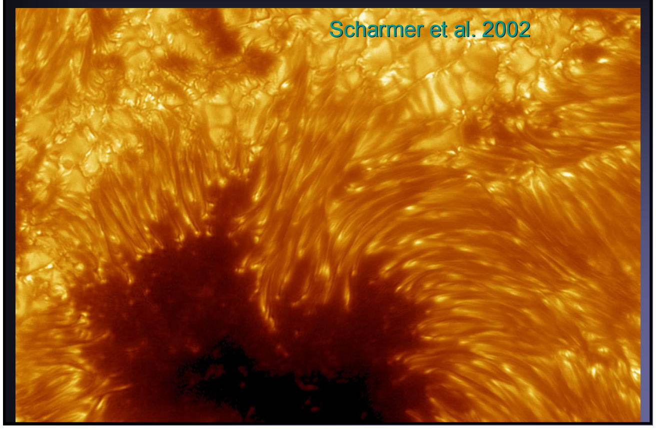
Penumbral filaments (bright and dark)

Penumbral grain (seen to move inward)

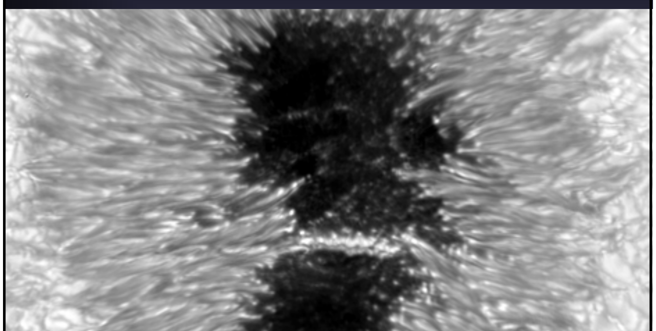
Light bridge

Umbral dot

Highest resolution

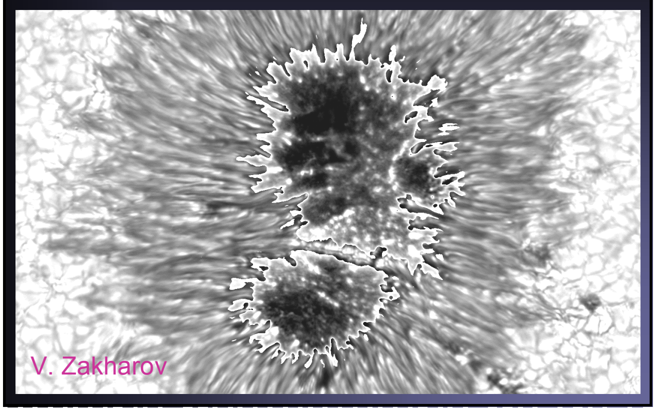


Sunspots

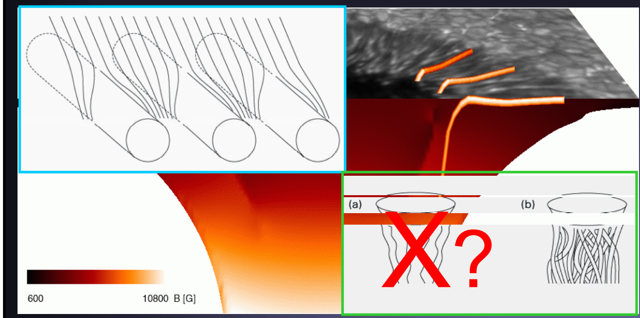


TiO Zakharov 2004

Umbral dots seen in TiO



Magnetic structure of sunspots



Sunspots span too many spatial and temporal scales to be successfully simulated from first principles.

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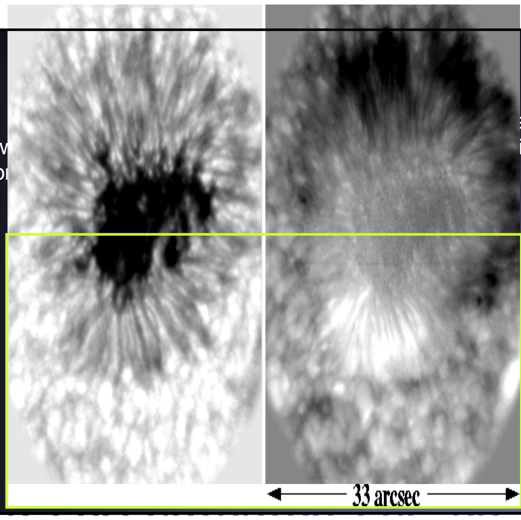
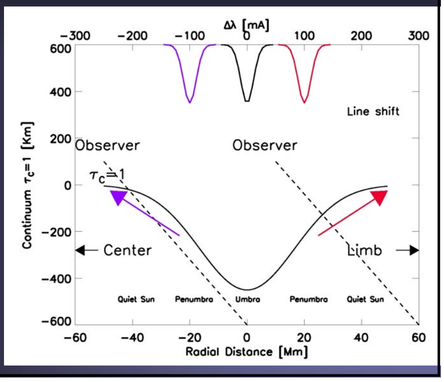
flow



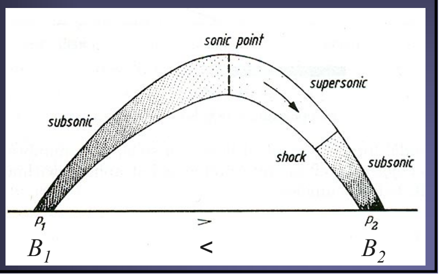
Evershed effect: illustration

- Horizontal outflow of matter.
- Thought to be driven by a siphon flow mechanism
- NEED NEW SLIDE !!!!!!!



Siphon flow model of Evershed effect

- Proposed by Meyer & Schmidt (1968).
- If there is an imbalance in the field strength of the two footpoints of a loop, then gas will flow from the footpoint with lower B to that with higher B.
- Supersonic flows are possible.



The Wilson effect

- Near the solar limb the umbra and centre-side penumbra disappear

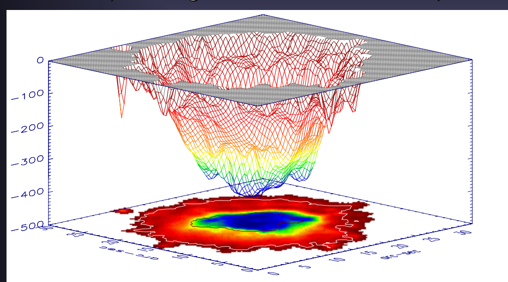
➔ We see 400-800 km deeper into sunspots than in photosphere

- Correct interpretation by Wilson (18th century).

Other interpretation by e.g. W. Herschell: photosphere is a layer of hot clouds through which we see deeper, cool layers: the true, populated surface of the Sun.

Sunspot Wilson depression

Map of Wilson depression (determined from T & B measurements and assumption that sunspot magnetic field is close to potential)



Shibu Mathew

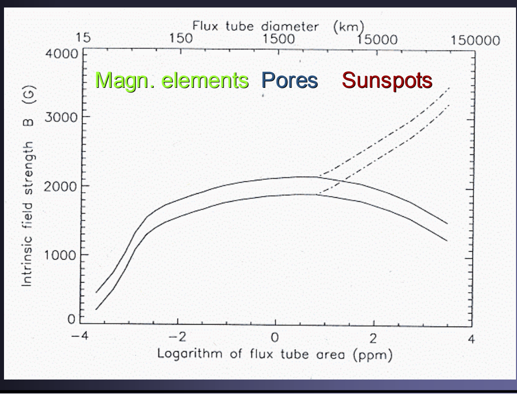
What causes the Wilson depression?

- B square/8 pi means gas pressure lower in spot than outside i.e. density also lower, i.e. fewer atoms to absorb, i.e. opacity also lower
- we see deeper into spot

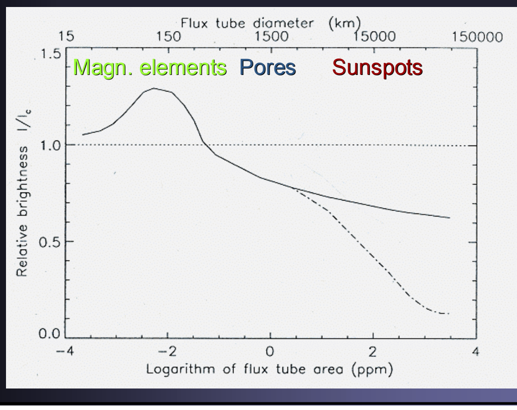
Magnetic elements

- Most of the magnetic flux on the solar surface occurs outside sunspots and pores (=smaller dark magnetic structures).
- These most common magnetic features, called magnetic elements, are small (diameters partly below spatial resolution of 100 km), bright and concentrated in network and facular regions.
- Magnetic elements are usually described by thin magnetic flux tubes (i.e. bundles of nearly parallel field lines).

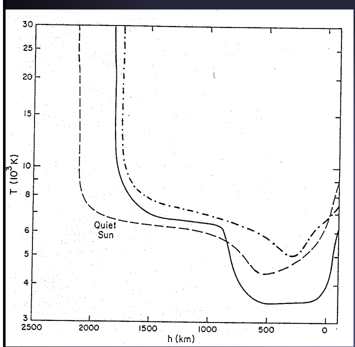
Surprisingly constant field strength



Temperature contrast vs. size

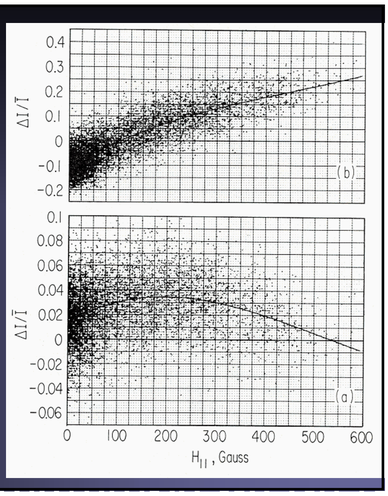


Temperature stratifications of quiet Sun, sunspot, plage

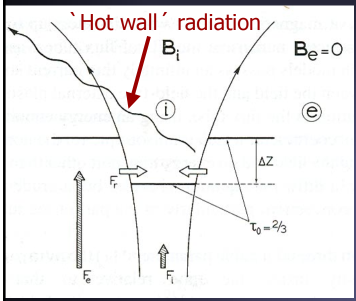


- Dashed line: Quiet Sun atmosphere
- Solid line: sunspot atmosphere
- Dot-dashed line: active region plage atmosphere
- Plage is hottest everywhere in atmosphere
- Sunspot coldest in photosphere, but gets hotter in chromosphere

Increased contrast of magnetic elements in higher layers



Why are magnetic elements bright?

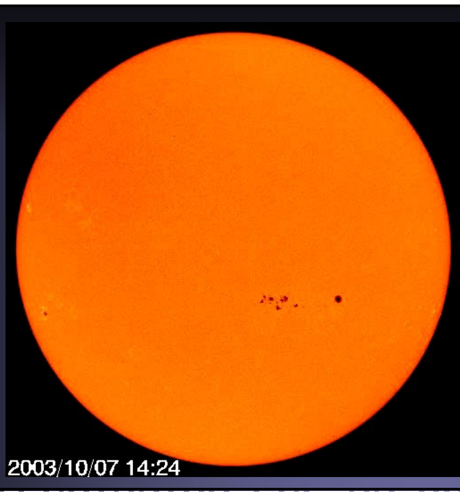


- Quenching of convection
- Partial evacuation
  - → enhanced transparency
  - → heating by `hot walls´
  - → *local flux excess*
- Inflow of radiation wins because the flux tubes are narrow (diameter ~ Wilson depression).
- High heat conductivity
  - → flux disturbance partly propagates into the deep convection zone
  - → Kelvin-Helmholtz time

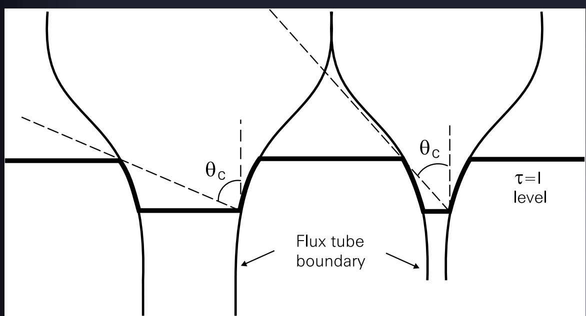
Why are faculae best seen near limb?

The Sun in White Light, with limb darkening removed

MDI on SOHO



Flux tube brightening near limb



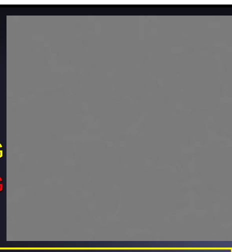
- The flux tubes expand with height (pressure balance
- They appear brightest when hot walls are well seen, i.e. near limb (closer to limb for larger tubes

B_z (Z=0)

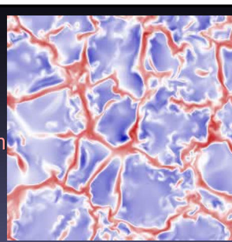
>500G

>1000G

>1500G



6 Mm



v_z (Z=0)

3-D compressible radiation-MHD simulations

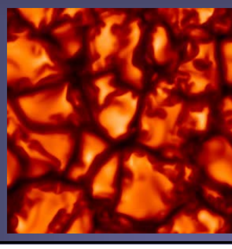
Plage: $B_z(t=0) = 200$ G

Grid Size: 288 x 288 x 100

Vertical extent: 1.4 Mm

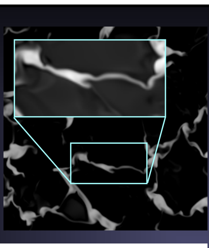
Horizontal extent: 6 Mm

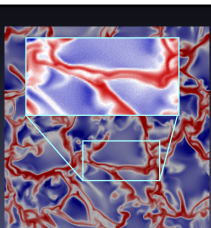
Alexander Vögler et al.



I_c (Z=0)

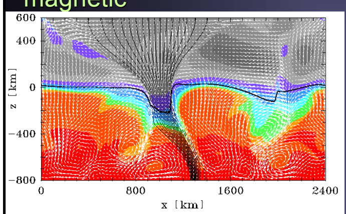
B_z

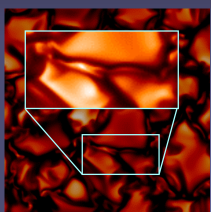




v_z

Details of thin magnetic





I_c

Horizontal cuts near surface level

Vögler et al. 2005

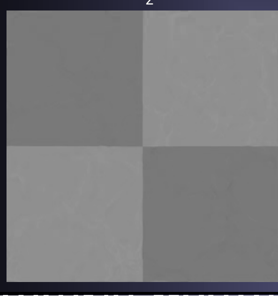
3-D Radiation-MHD Simulations

Alexander Vögler, Robert Cameron, Manfred Schüssler

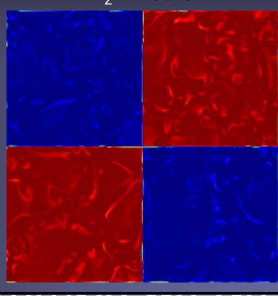
Mixed polarity simulations: diffusion & cancellation of opposite polarities (20 km resolution): $\langle B_{initial} \rangle = 200$G.

B_z

$B_z < 200$ G



6000 km

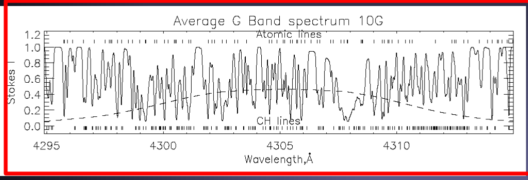


Simulations require computers...

G-band Spectrum Synthesis

G-Band (Fraunhofer): spectral range: 4295-4315 Å

contains many temperature-sensitive molecular lines (CH)

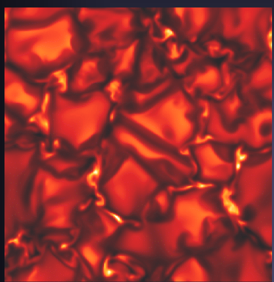


For comparison with observations, we define as G-band intensity the integral of the spectrum obtained from the simulation data:

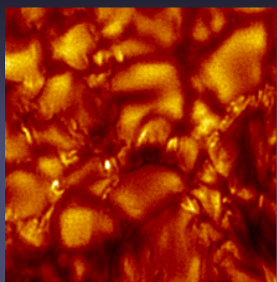
$$I_G = \int_{4295\,A}^{4315\,A} I(l)\,dl$$

Shelyag et al. 2004

G-band: Simulation vs. Observation



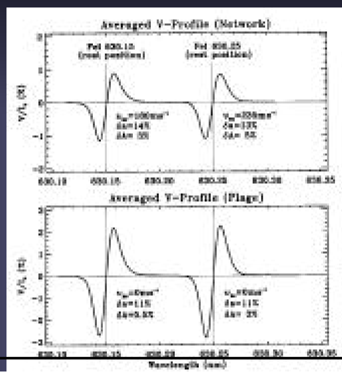
Simulation (20 km resolution)
Schüssler et al. 2003
Shelyag et al. 2004



Observation (100 km resol.)
(SST, La Palma)
Scharmer et al. 2002

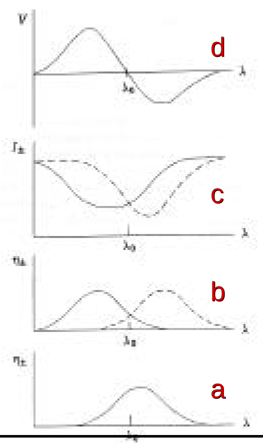
Stokes V asymmetry

- Stokes V profiles observed in quiet Sun and in active region plage are asymmetric: typically blue wing has larger area 'A' and amplitude 'a' than red wing

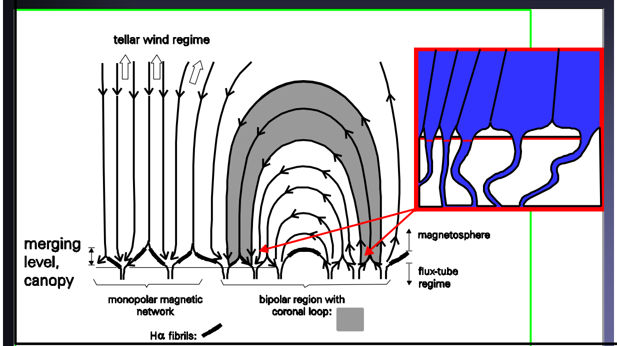


Producing Stokes V asymmetry

- Consider 2-layered atmosphere:
  - Bottom layer a: v but no B
  - Top layer b: B but no v
- Note the importance of line saturation for producing asymmetric V profile.

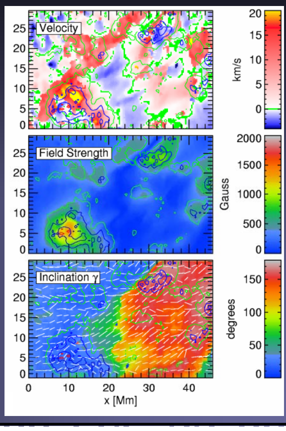


Flux Tubes, Canopies, Loops and Funnels



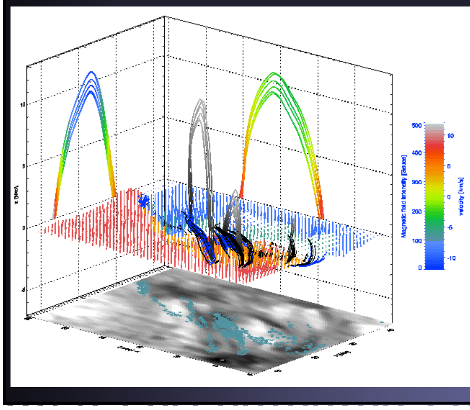
Measurement of *B* at coronal base

- Previously: Magnetic vector only known at solar surface. However, magnetic field has main effect in corona. Exception: radio observations give |B| but low resolution.
- Now: Direct measurement of full magnetic vector at base of corona & in cool loops possible
- Measurement using He I 10830 Å (TIP, VTT, Tenerife) & simple inversion code



Solanki et al. 2003, Lagg et al. 2004

Structure of Magnetic Loops



Magnetic loops deduced from measurements of He I 10830 Å

Stokes profiles in an emerging flux region.

Left projection: Field strength

Right projection: Vertical velocity

Andreas Lagg

Magnetic field extrapolations: Force free and potential fields

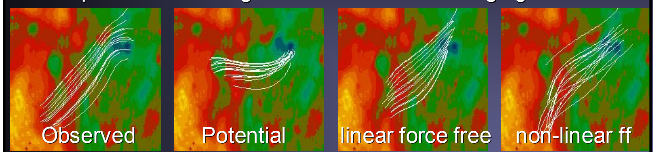
- General problem in solar physics: Magnetic field is measured mainly in the photosphere, but it makes the music mainly in the corona.

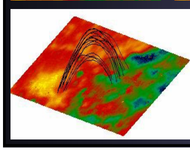
➔Either improve coronal field measurements or extrapolate from photospheric measurements into the corona.

- If $\beta << 1$ then we can neglect the influence of the gas on the field: the field is force-free. Considerable simplification of the computations
- If we further assume that there are no currents, the computations become even simpler (potential field).

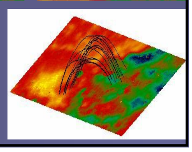
Testing Magnetic Extrapolations

- Non-linear force-free fields reproduce the loops reconstructed from observations better than the linear force-free ones and far better than potential field extrapolations.
- Loops harbour strong currents while still emerging.

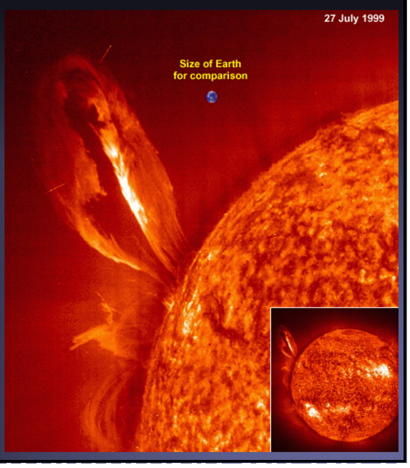




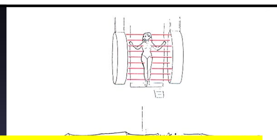
Wiegelmann et al. 2004



Prominences



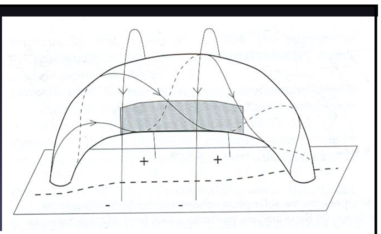
Kippenhahn's magnetic circus

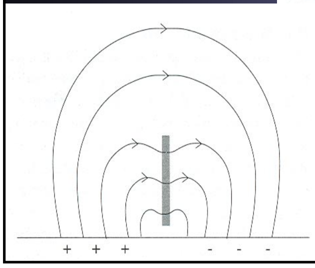


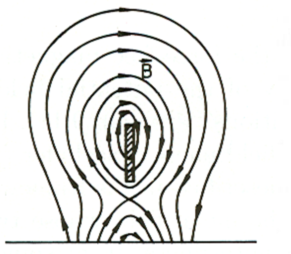
Problem: prominence material is dense and cool and high in the sky ➔ It must be supported against gravity. Obvious supporting structure: magnetic field. However, magnetic field must be curved upwards to keep the material from flowing down along the field lines. Different solutions have been proposed.

Prominence models

Kippenhahn-Schlüter (below), Kuperus-Raadu (below right) and flux tube (right)

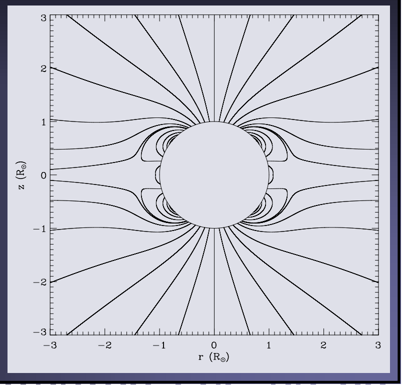






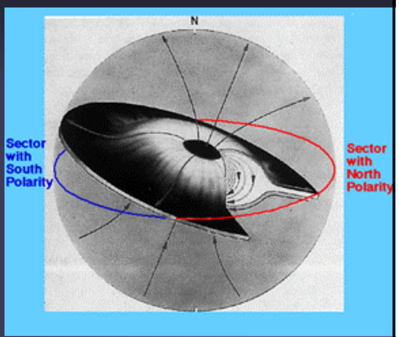
Large scale magnetic structure of the quiet Sun

- At large scales dipolar component of the magnetic field survives, since multipoles ➔ $B \sim r^{-n-1}$, where n=2 for dipole, n=3 for quadrupole, etc.
- Closer to sun ever higher order multipoles are important



Solar current sheet at activity minimum

- At activity minimum solar magnetic field is like a dipole, whose field lines are stretched out by the solar wind.
- Field lines with opposite polarity lie close to each other near equator: euatorial current sheet.
- If dipole axis inclined to ecliptic: magnetic polarity at Earth changes over solar rotation.



Heliospheric current sheet and Parker spiral

- Since solar wind expands radially beyond the Alfven radius (where the energy density in the wind exceeds that in the magnetic field) and the Sun rotates (i.e. the footpoints of the field), the structure of the field (carried out by the wind, but anchored on rotating surface) shows a spiral structure.