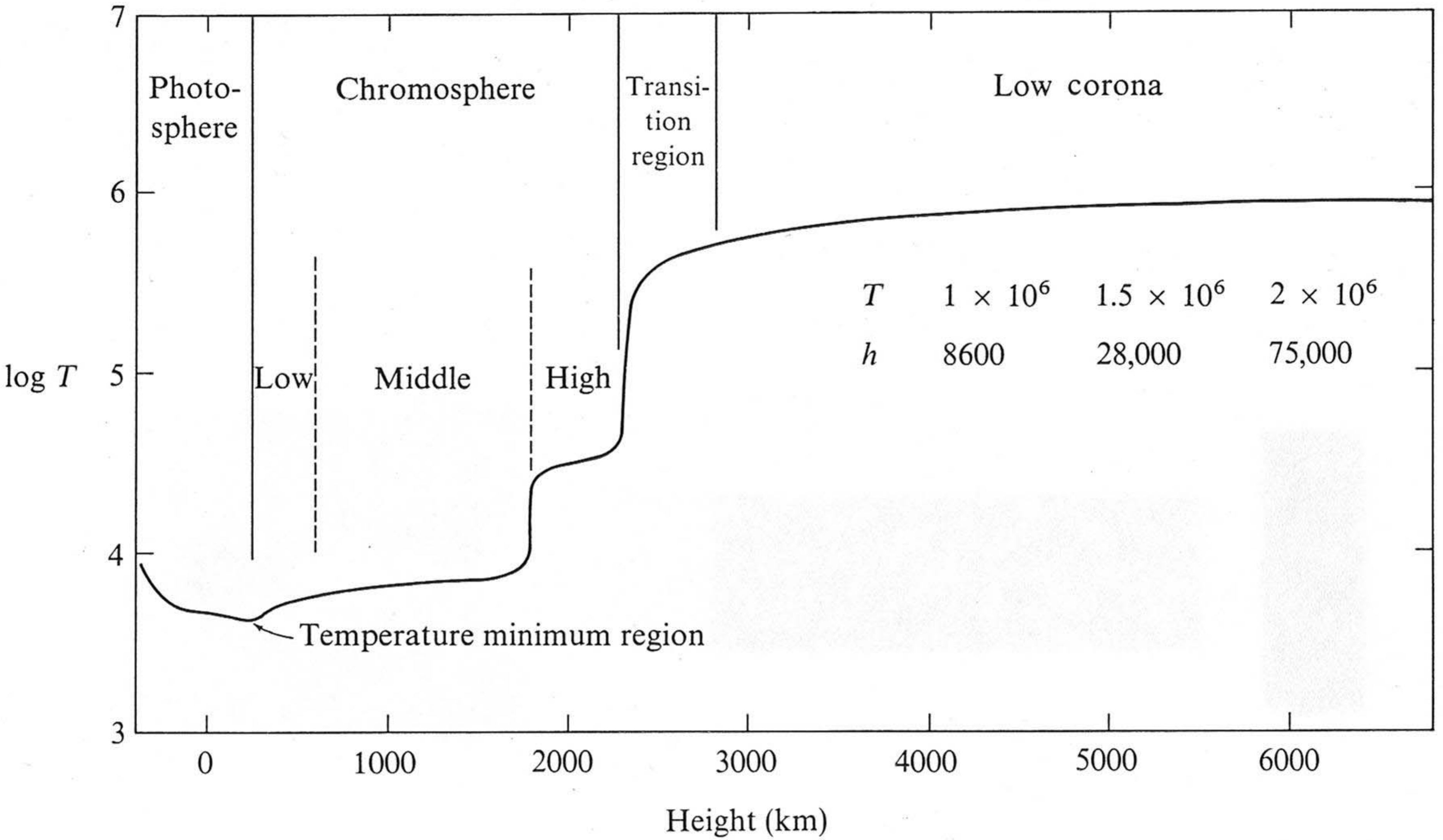
The solar atmosphere

The Sun's atmosphere

- Solar atmosphere generally described by multiple layers. From bottom to top: photosphere, chromosphere, transition region, 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

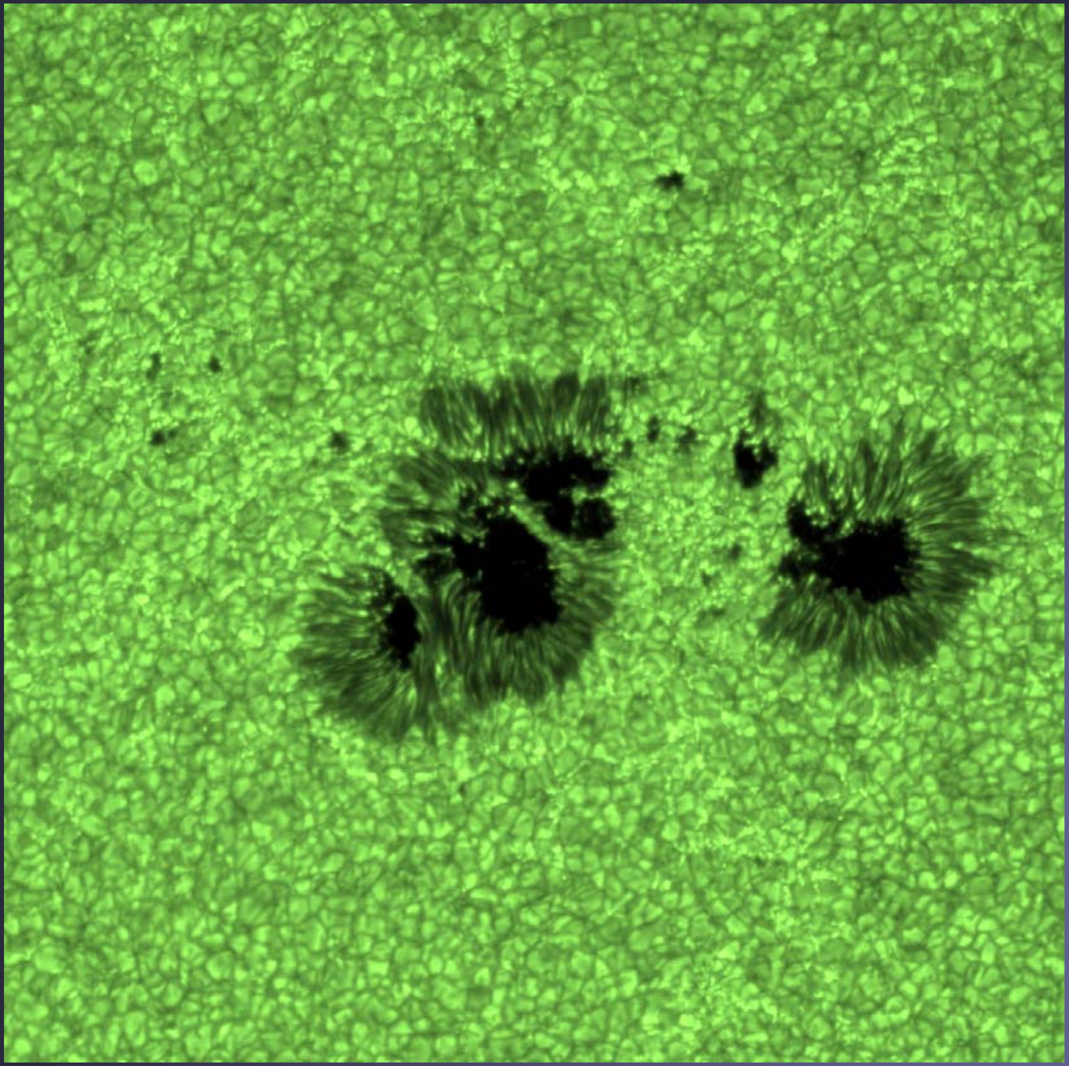
1-D stratification



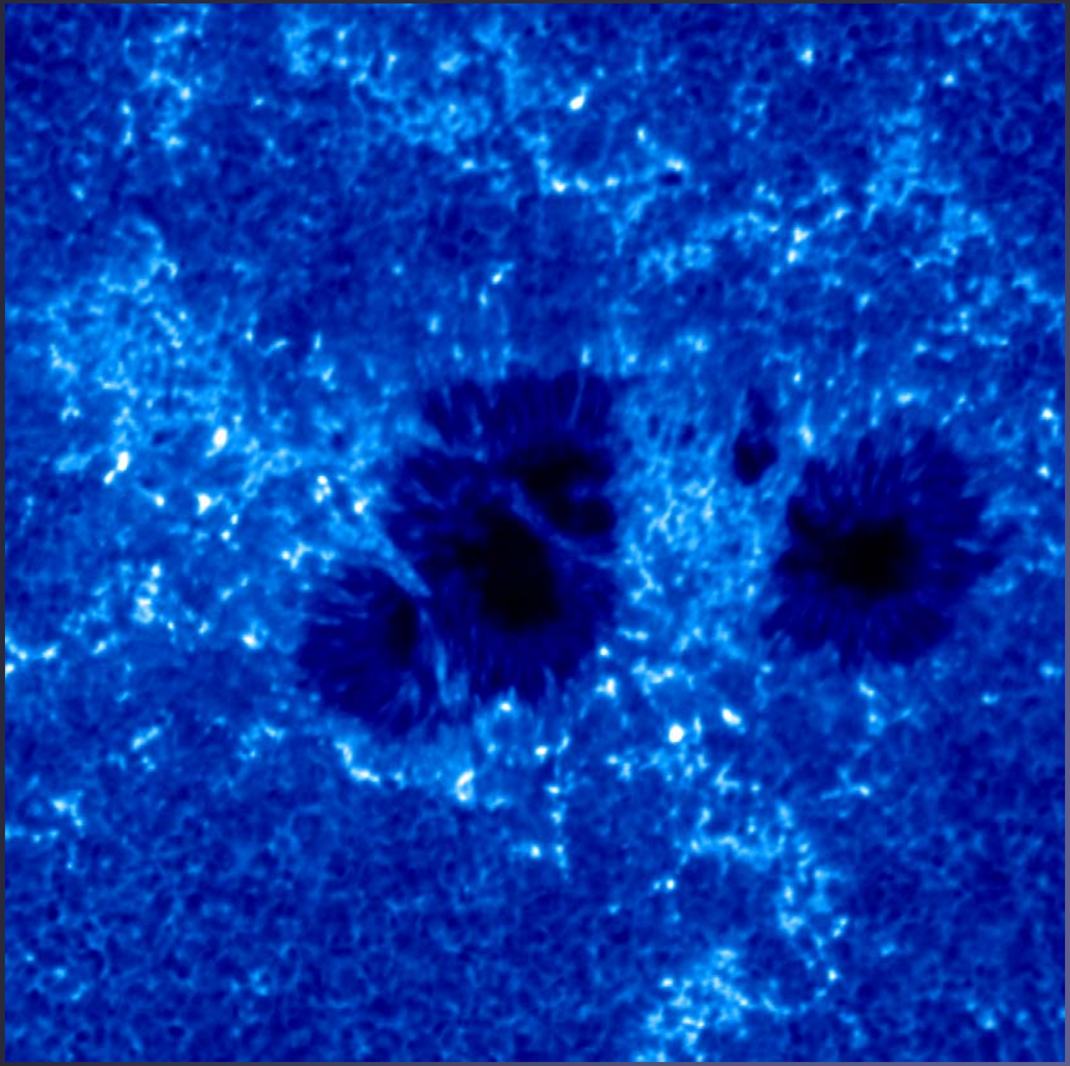
How good is the 1-D approximation?

- 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 particular part of the real Sun.

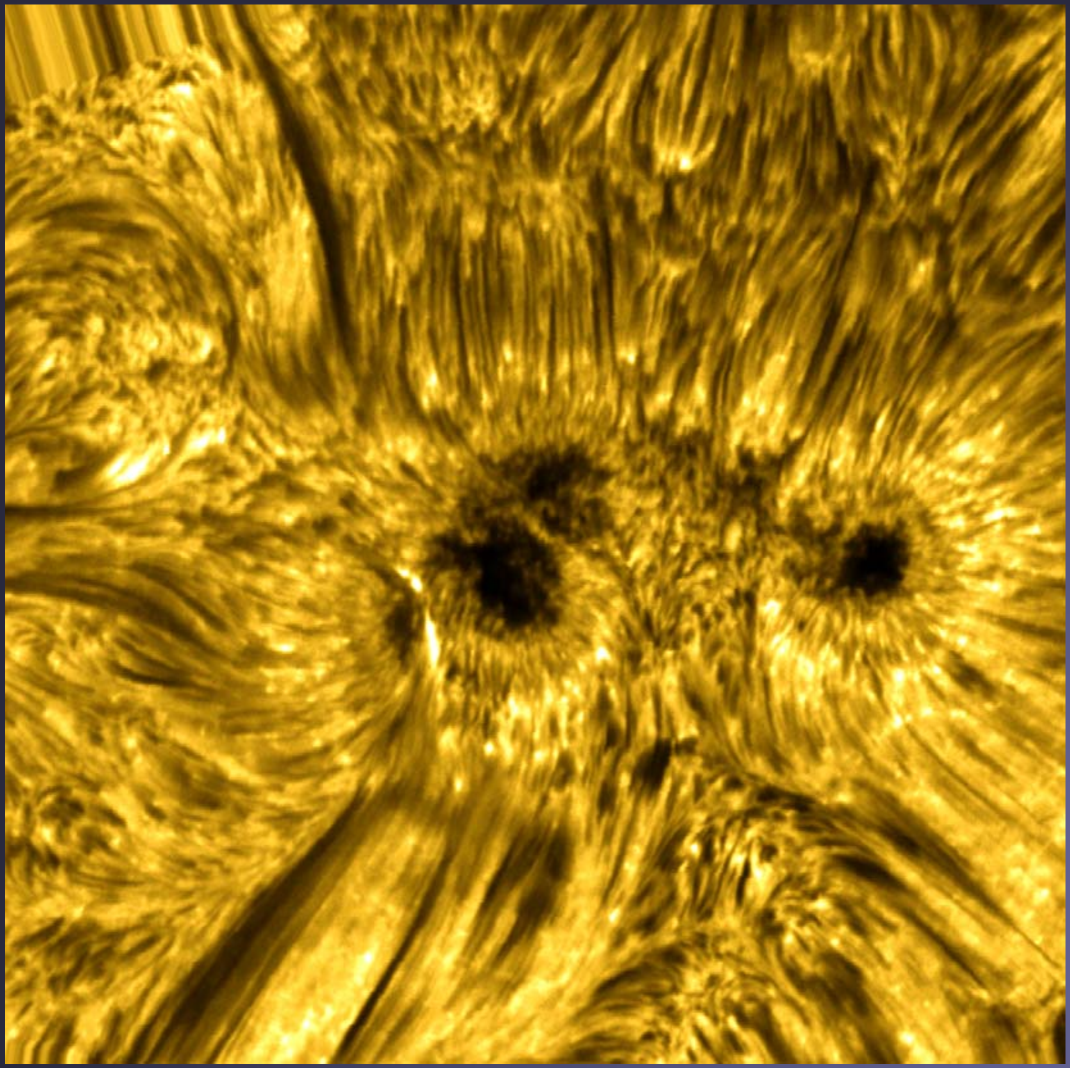
Photosphere



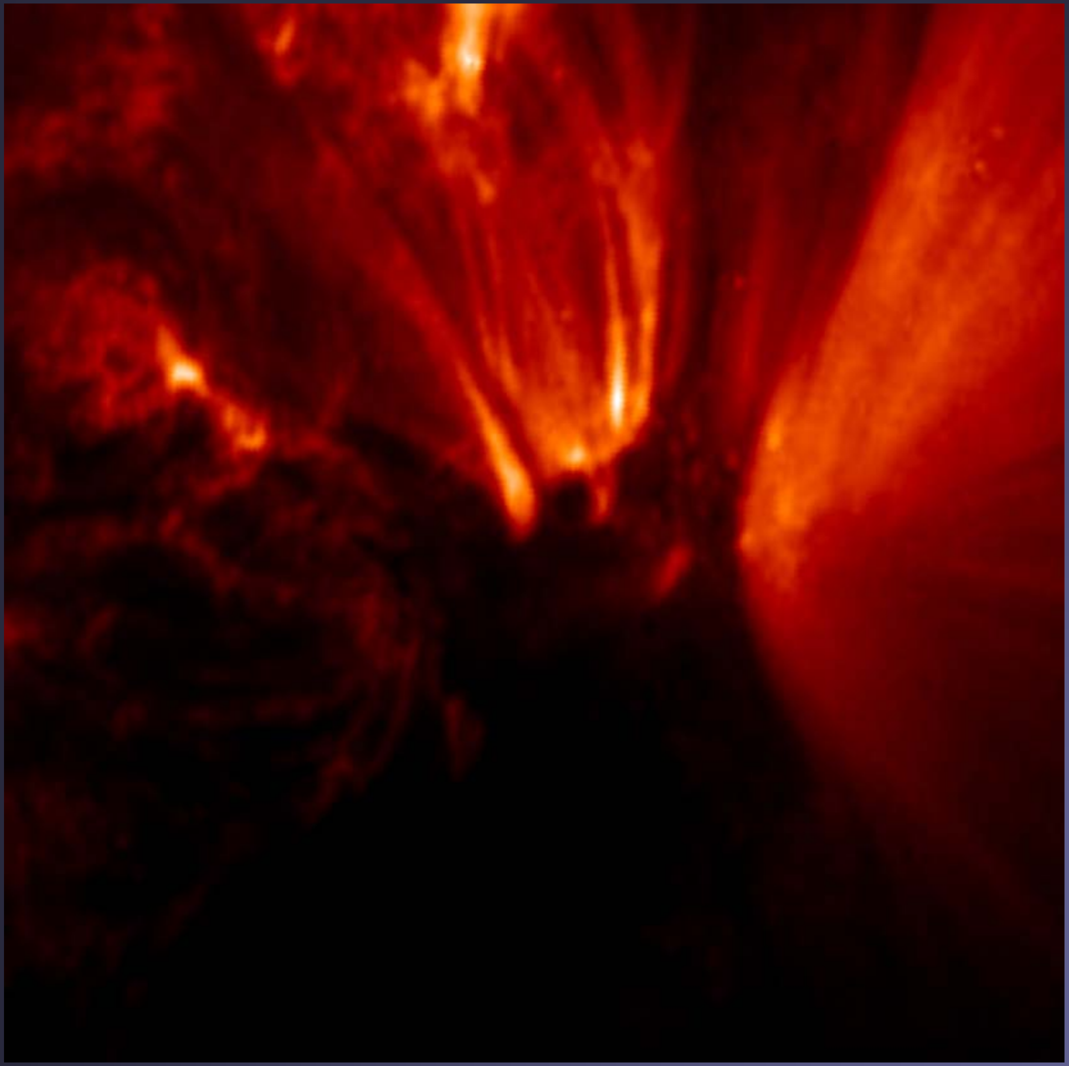
Lower chromosphere



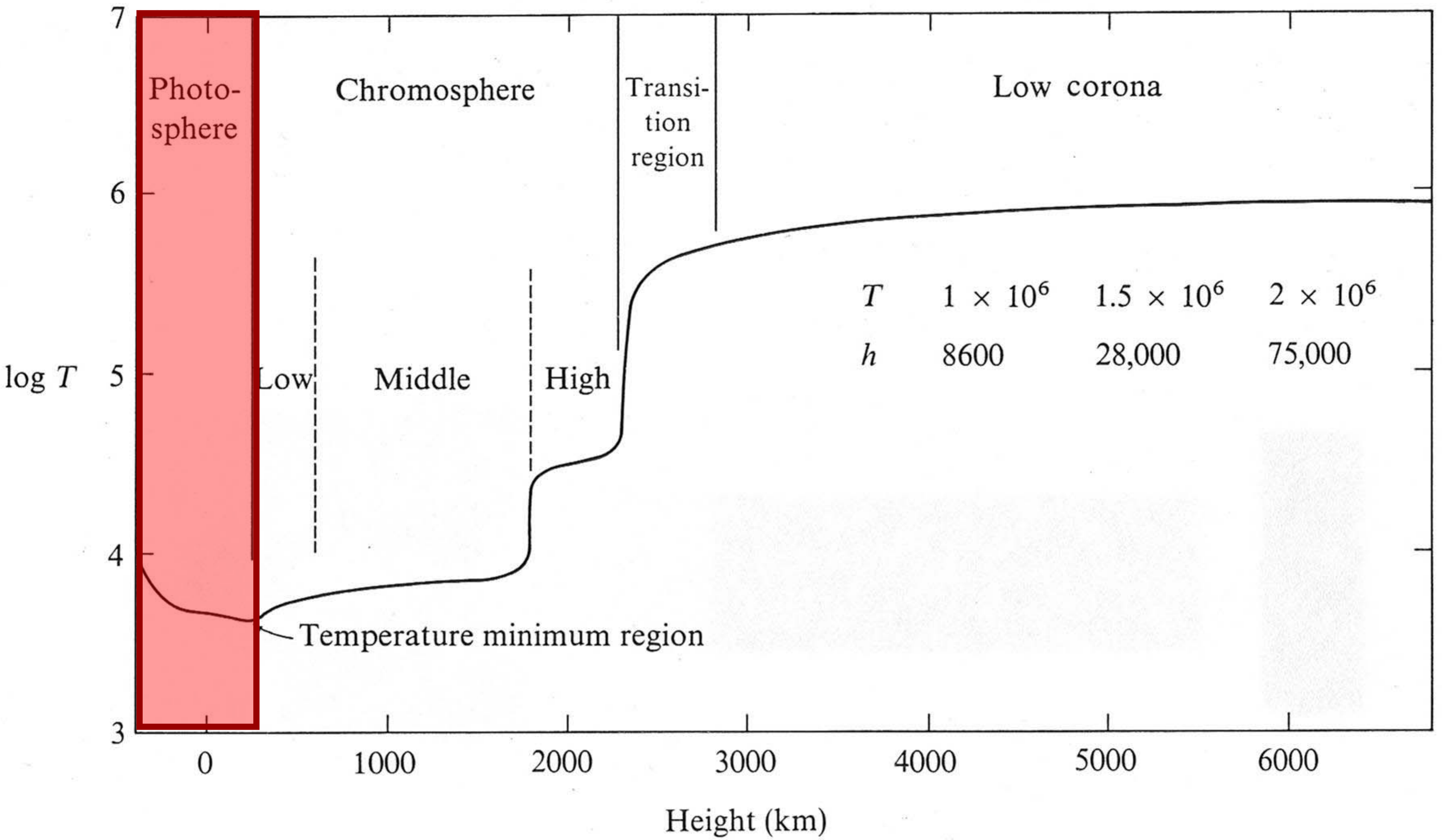
Upper chromosphere



Corona



Photosphere



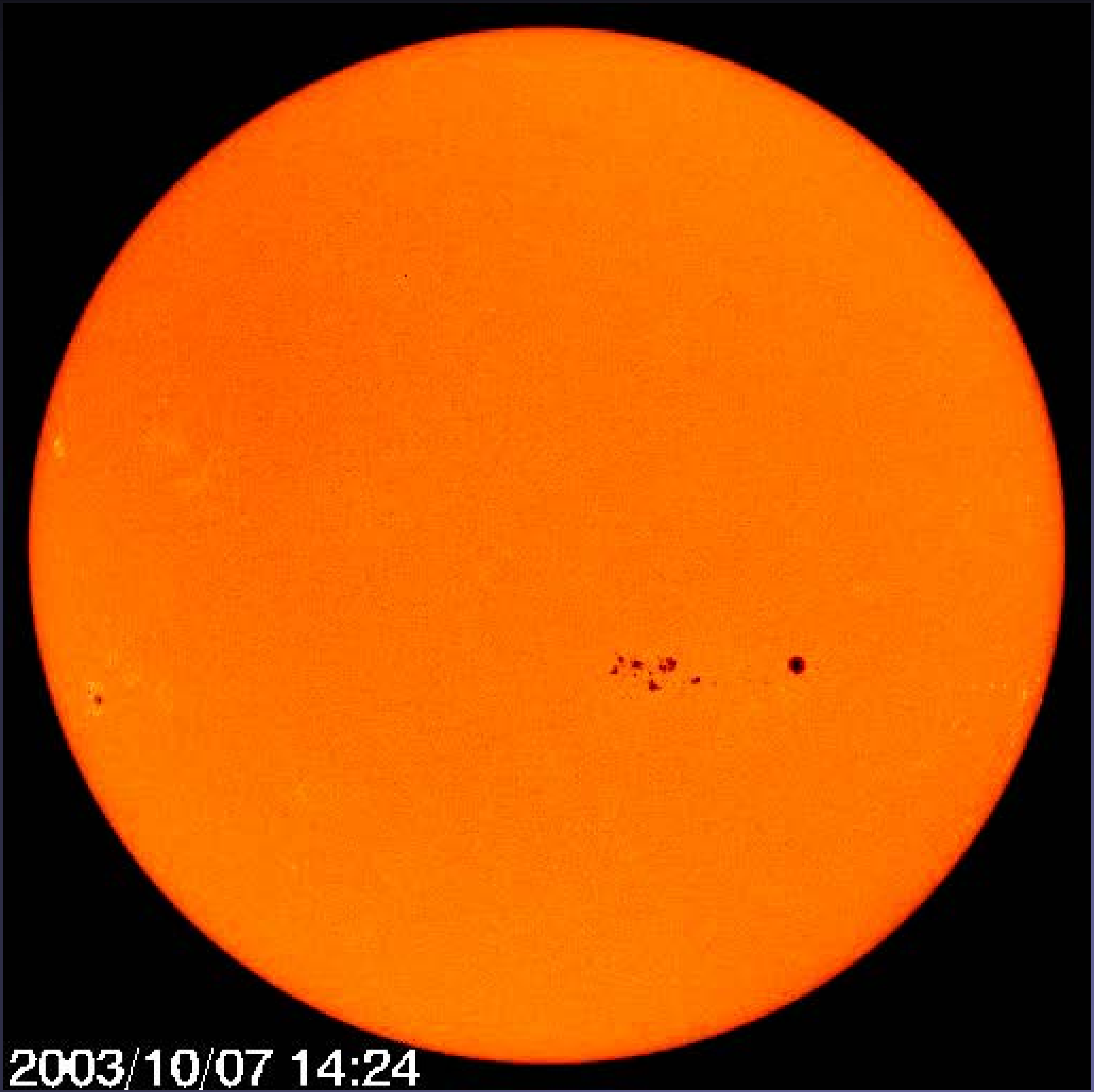
The photosphere

- Photosphere extends between solar surface and temperature minimum (400-600 km)
- It is the source of most of the solar radiation. 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 → photosphere produces an absorption spectrum
- LTE is a good approximation
- Energy transport by radiation (and convection)
- Main structures: Granules, sunspots and faculae

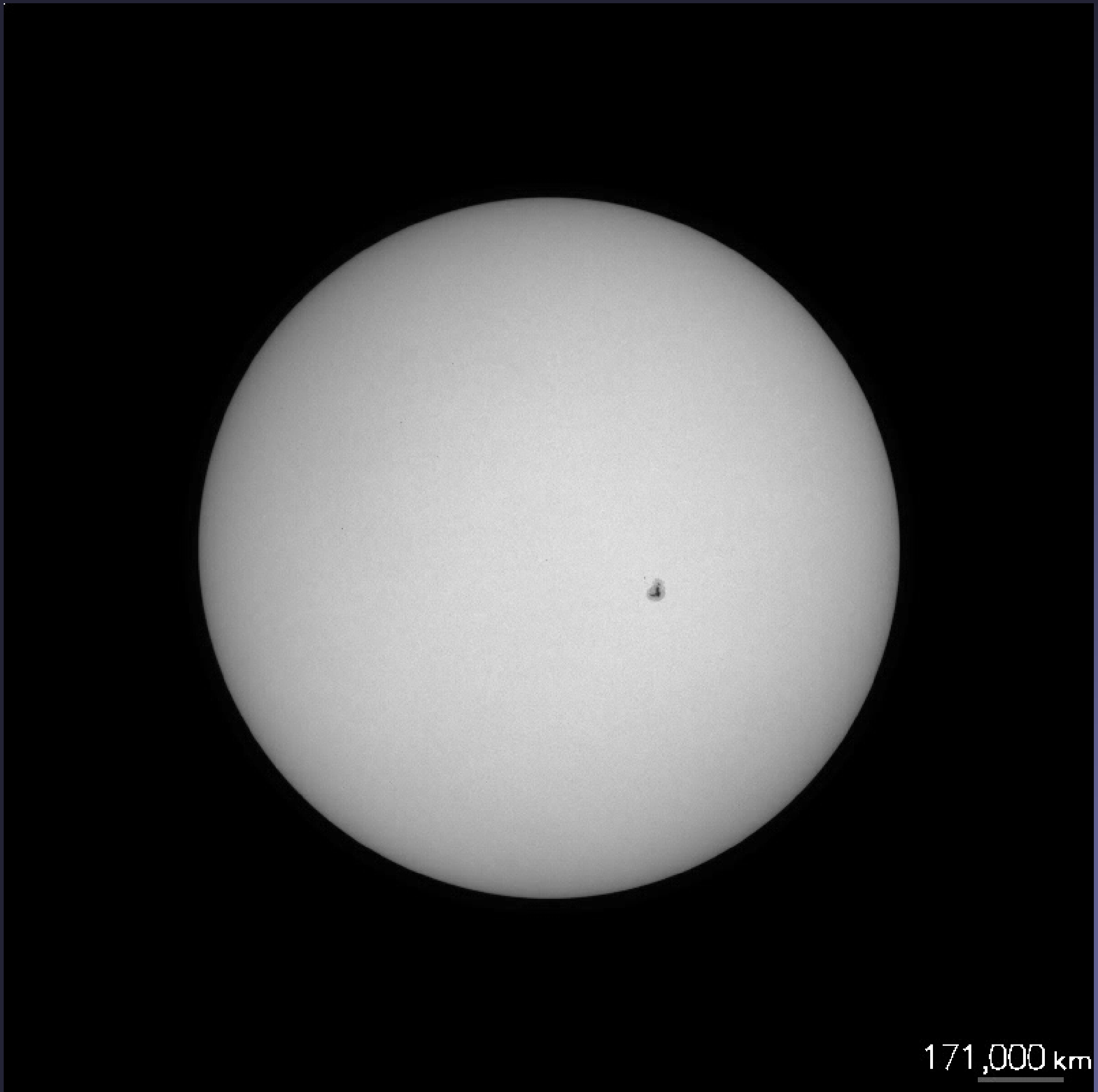
The Sun in White Light

(limb darkening removed)

MDI on SOHO



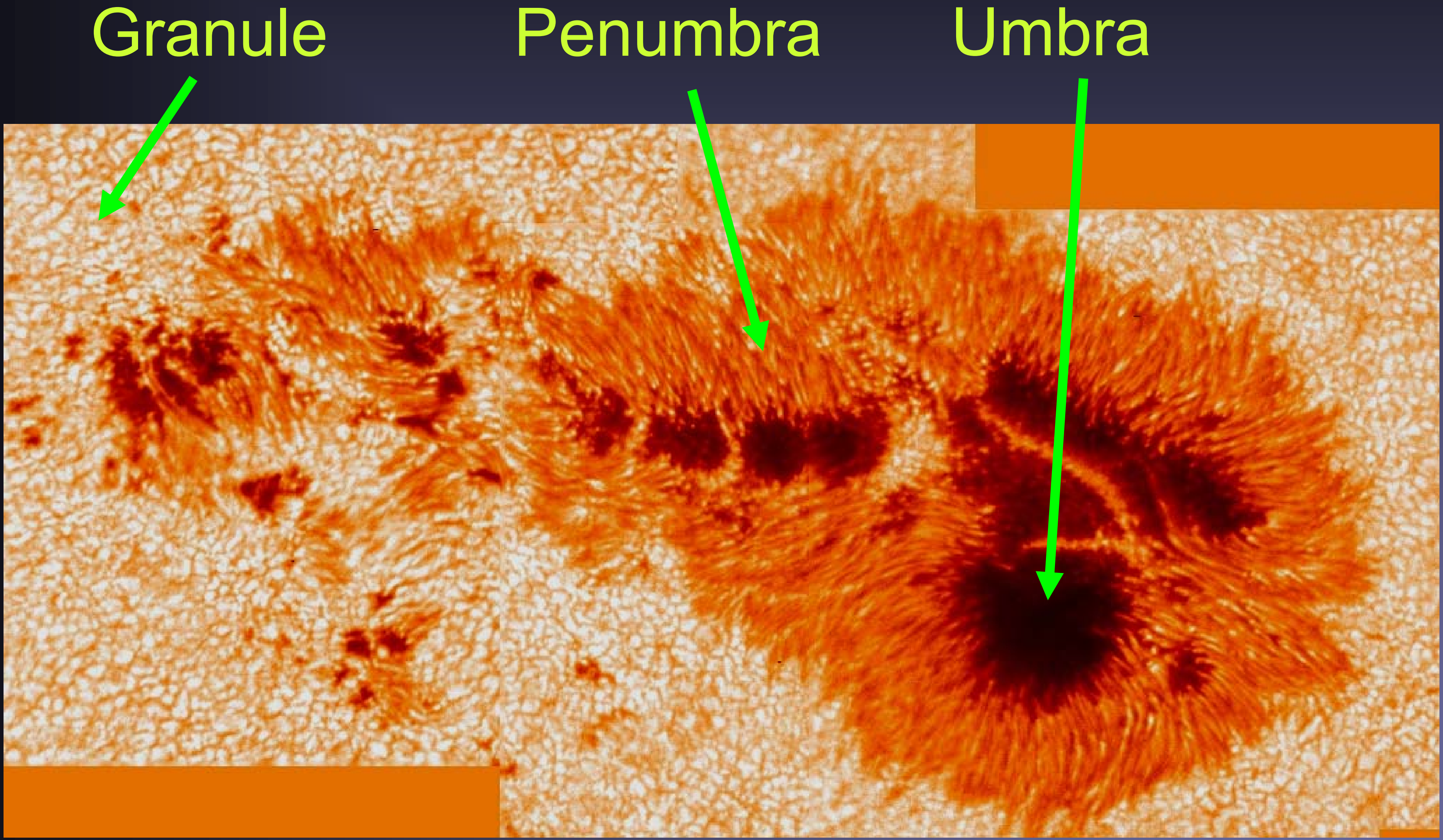
The photosphere: a boring place?



Photospheric structure: Granulation

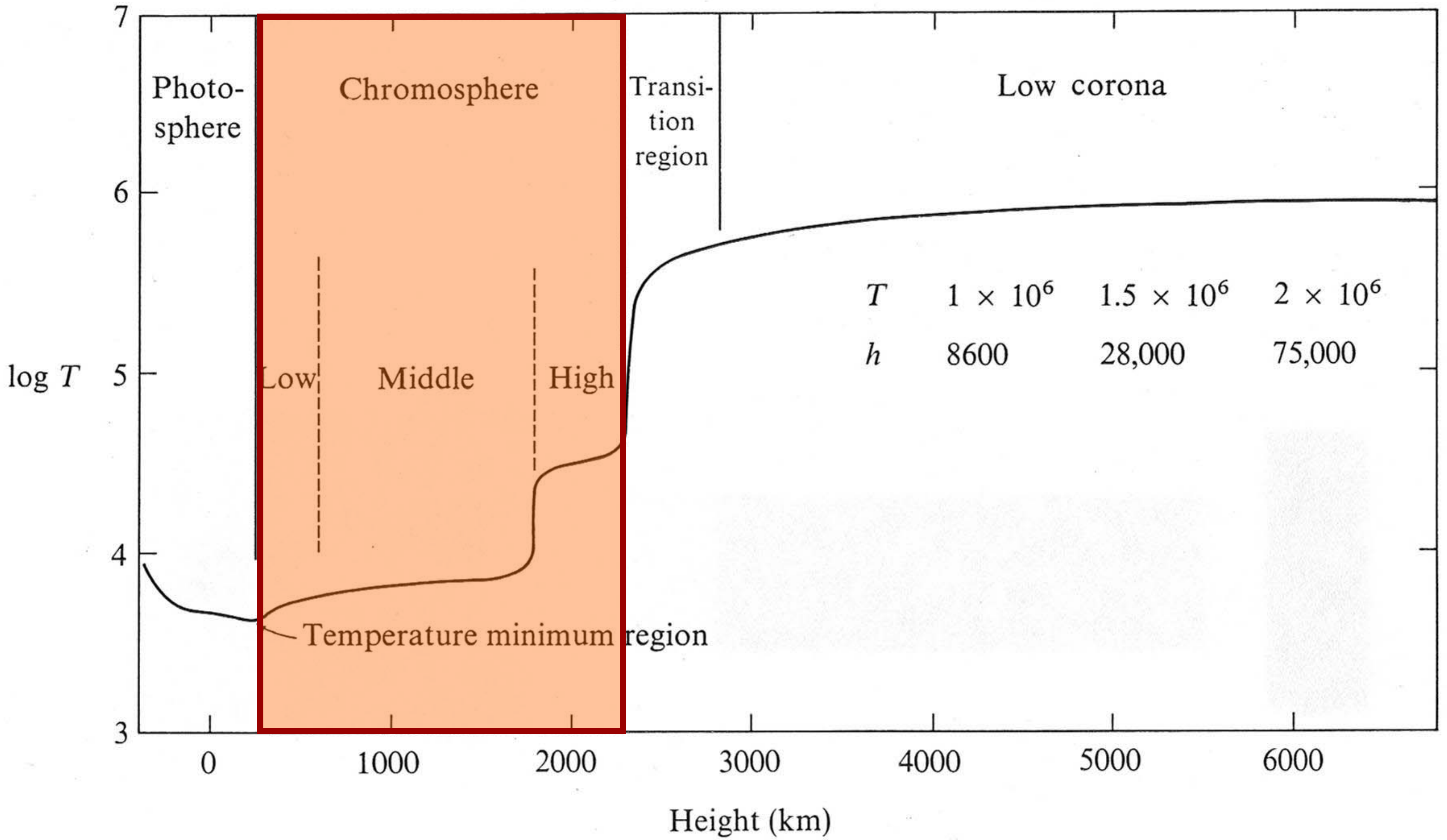
- Physics of convection and the properties of granulation and supergranulation have been discussed earlier, so that we can skip them here.

Photospheric structure: Sunspots



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

Chromosphere

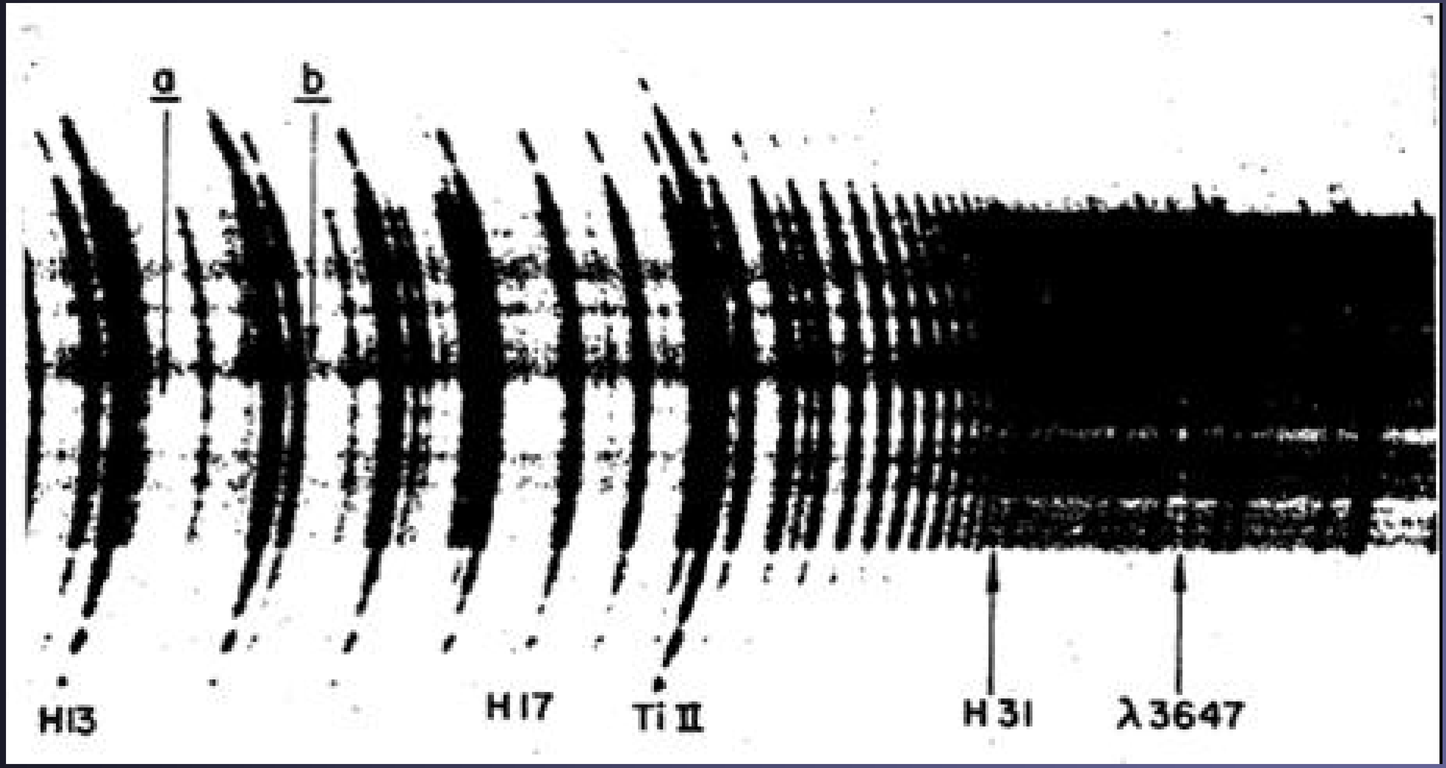


Chromosphere

- Layer just above photosphere, at which temperature appears to increase outwards (classically forming a temperature plateau at around 7000 K)
- Strong evidence for a spatially and temporally inhomogeneous chromosphere (gas at $T<4000$K is present beside gas with $T>8000$ K)
- Assumption of LTE breaks down
- Assumption of plane parallel atmosphere very likely breaks down (i.e. radiative transfer in 3-D needed)
- Energy transport mainly by radiation and waves

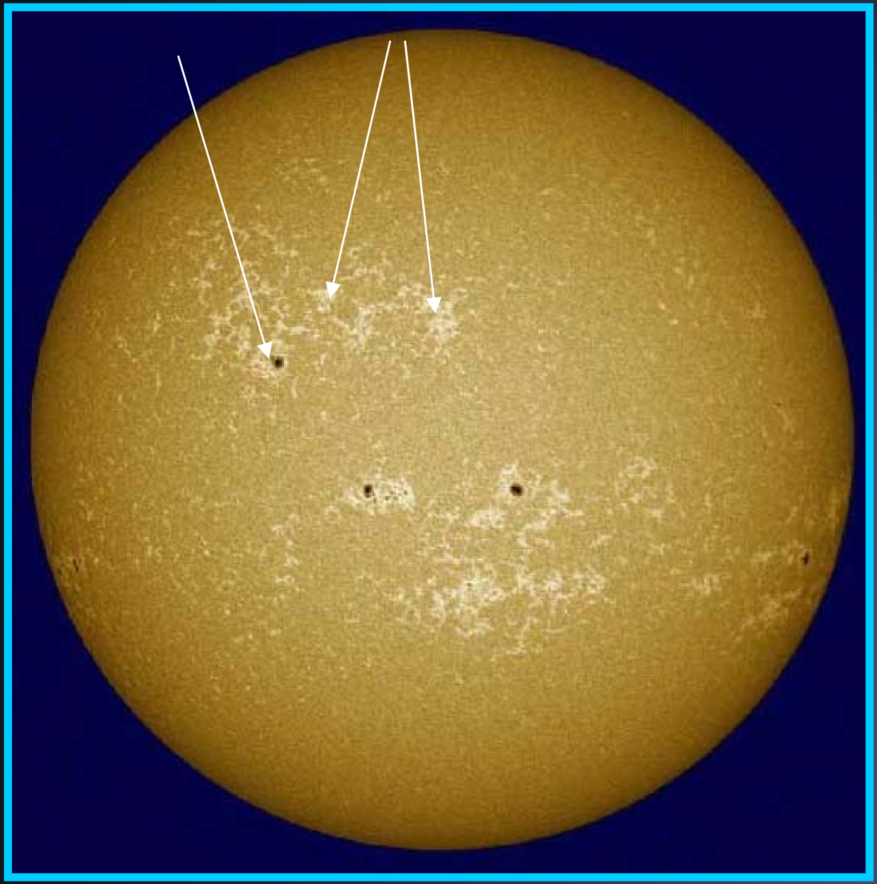
Discovery of Chromosphere

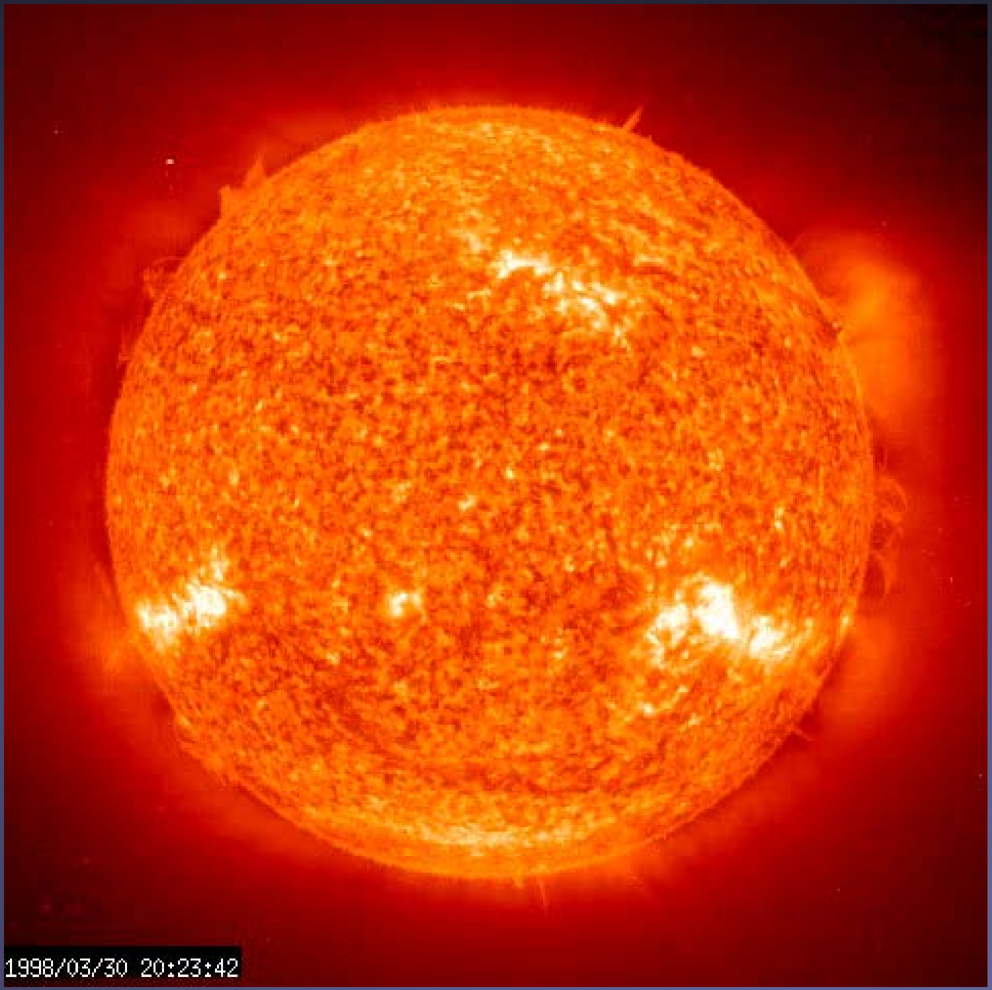
- Flash spectrum seen seconds before and after totality



Chromospheric structure

Spots plages



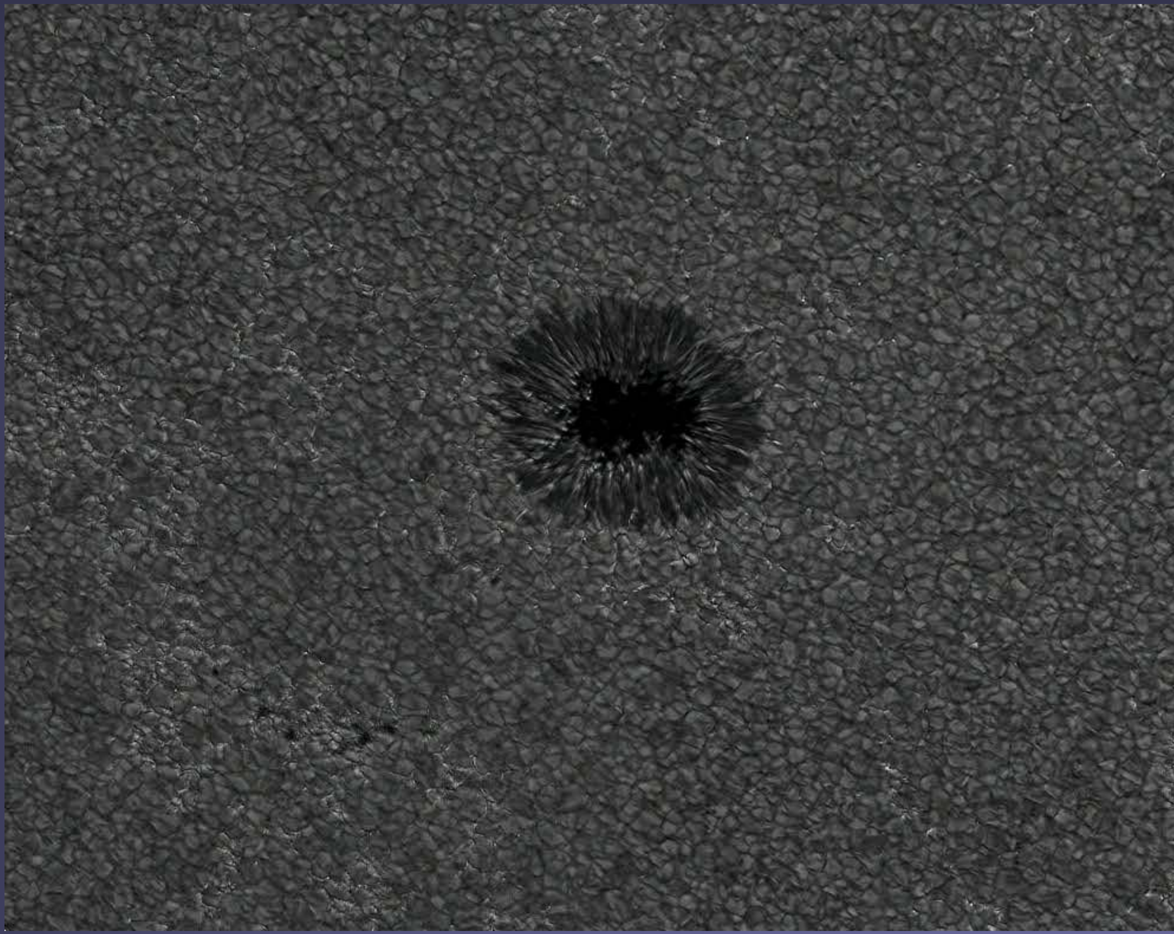


7000 K gas Ca II K

$5\ 10^4$ K gas (EIT He 304 Å)

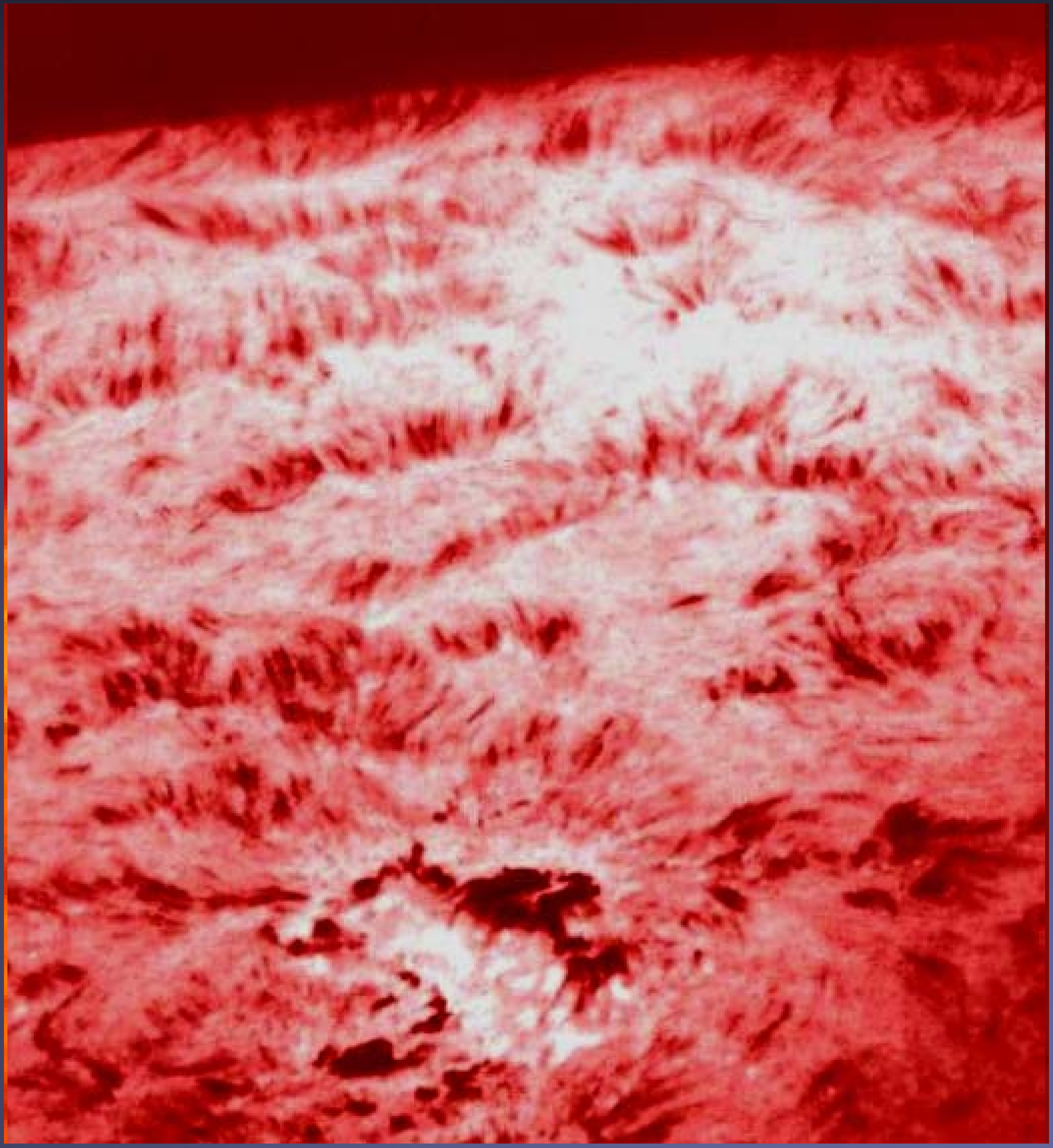
Chromospheric structure III CHANGE!!!

- 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



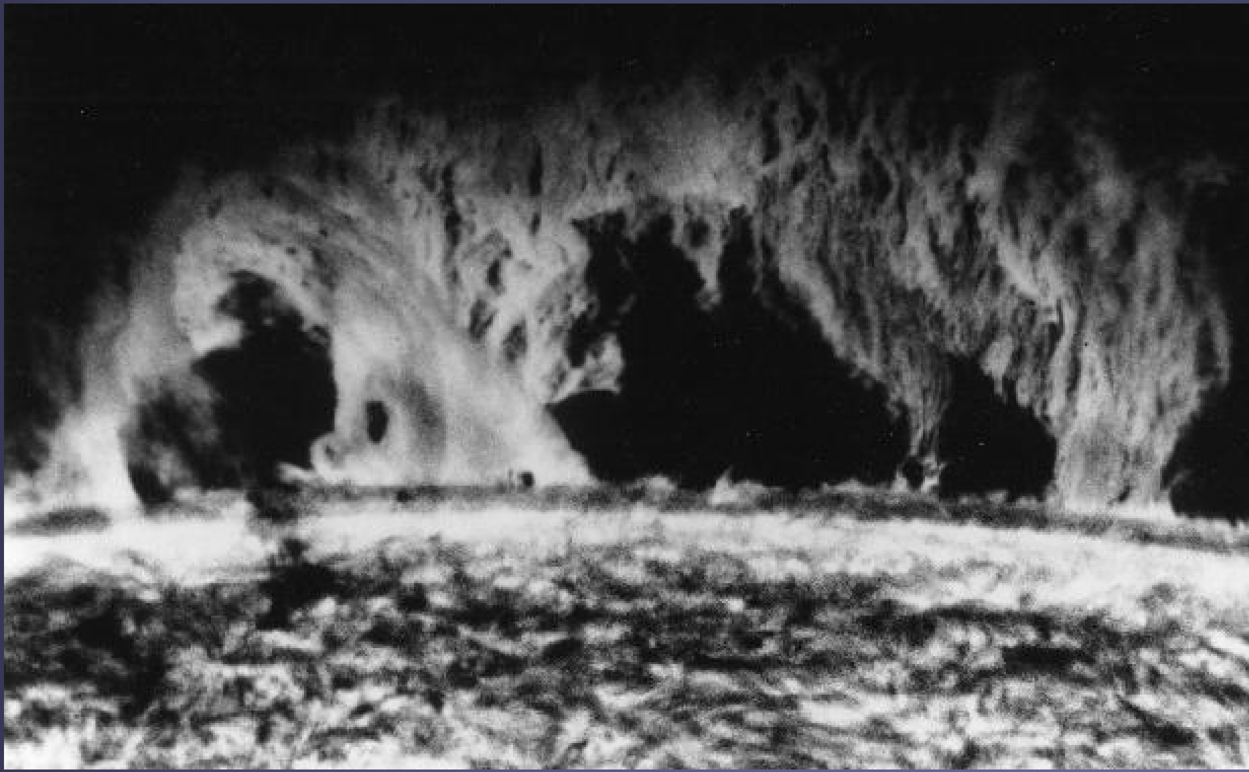
Chromospheric structure

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



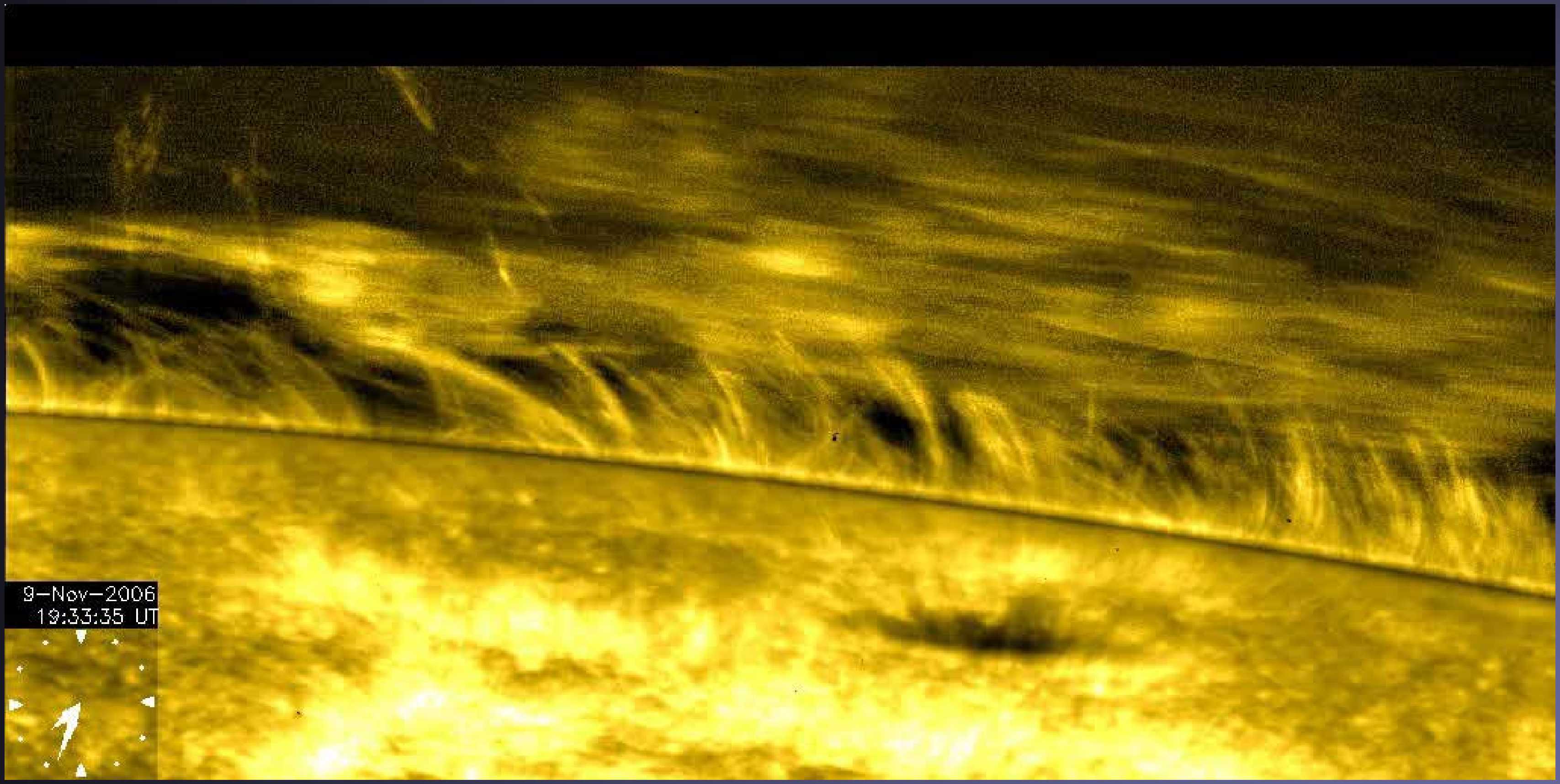
Chromospheric structure

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

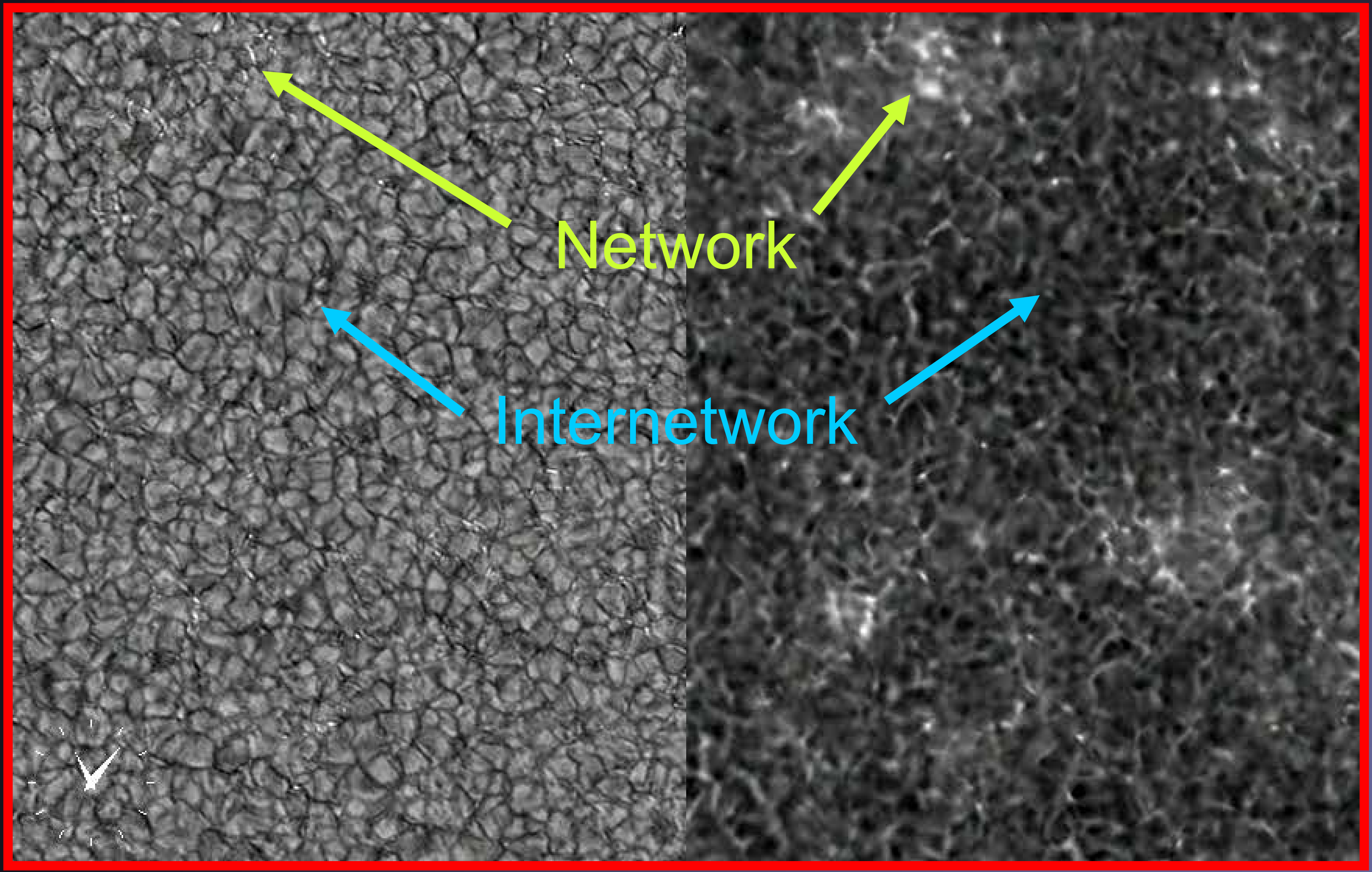


Chromospheric structure

- Spicules
- Prominences and filaments

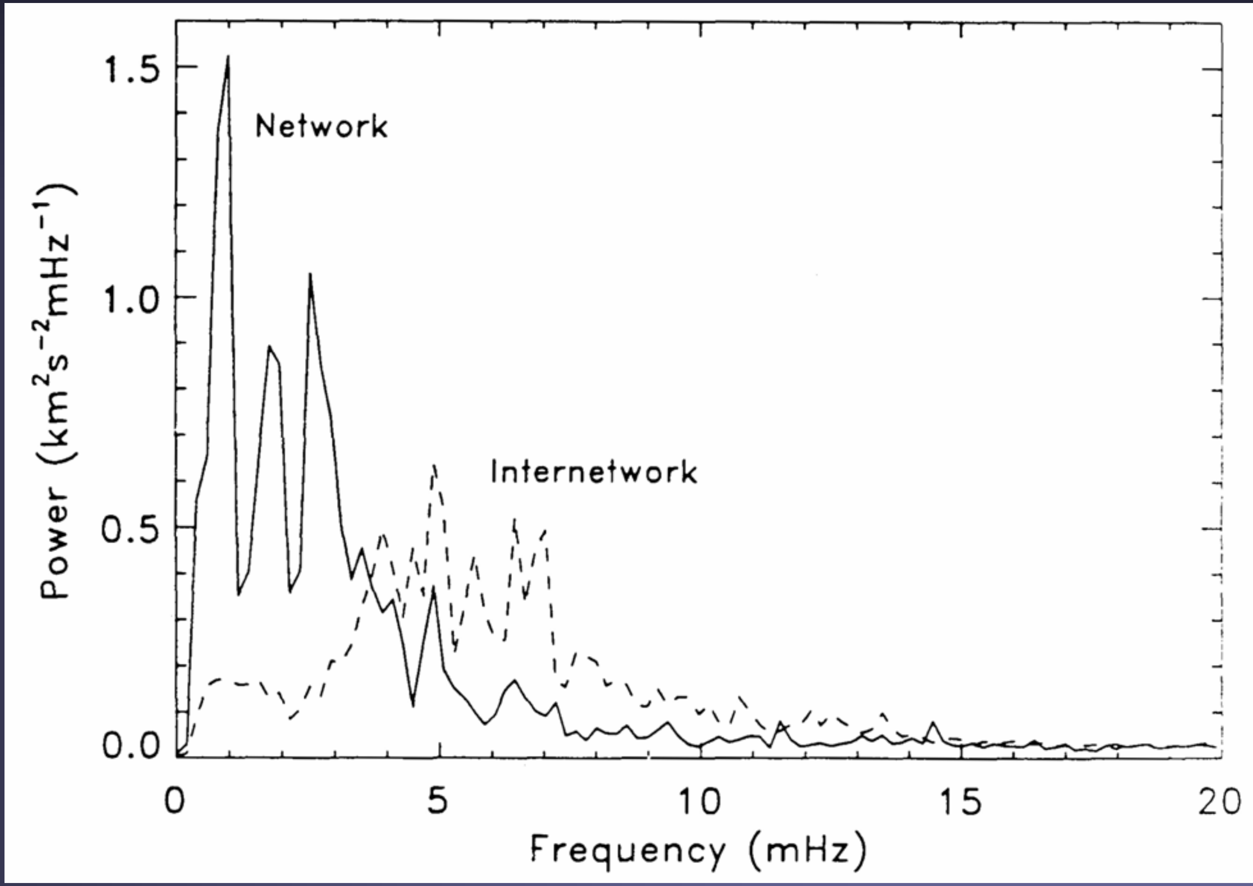


Chromospheric dynamics



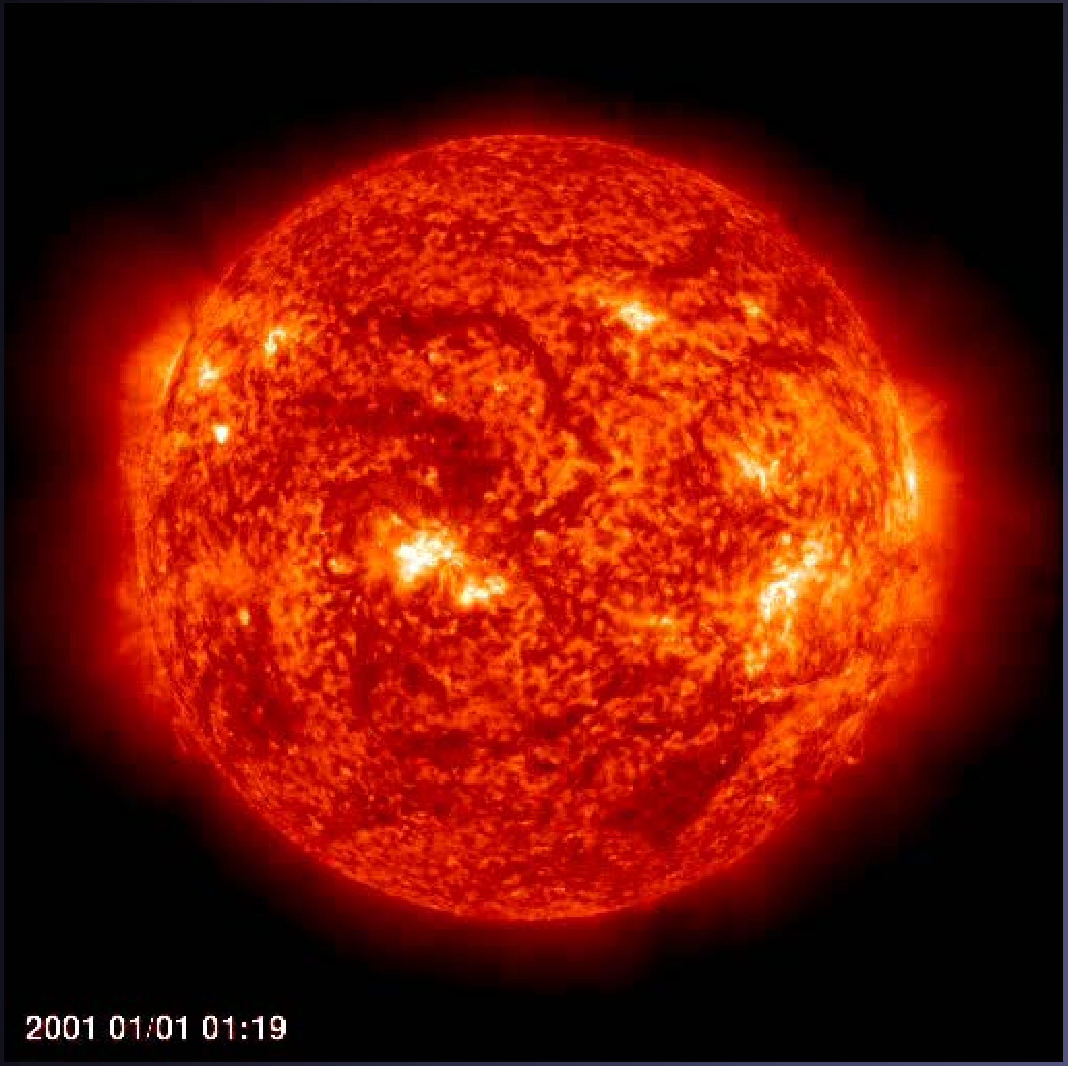
Chromospheric dynamics

- Oscillations, seen in cores of strong lines
- Power peaks: oscillations, not turbulence
- Power at 3 min in Inter-network
- Power at 5-7 min in Network



Lites et al. 2002

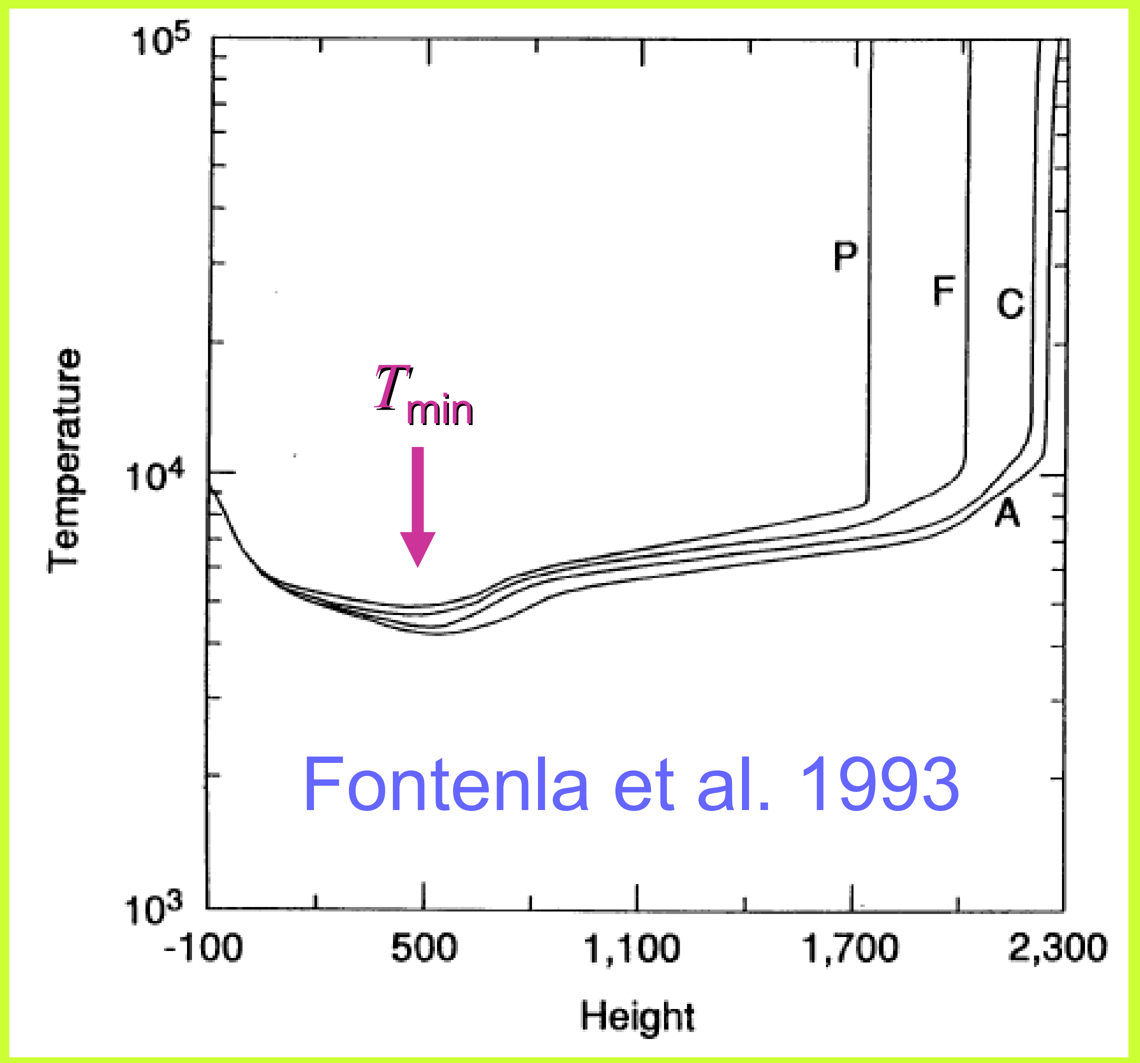
1 year of EIT: upper chromosphere



- He II 304 Å with EIT/SOHO
- samples gas at ≈50 kK (with contribution also from corona: coronal holes are visible)

Models: the classical chromosphere

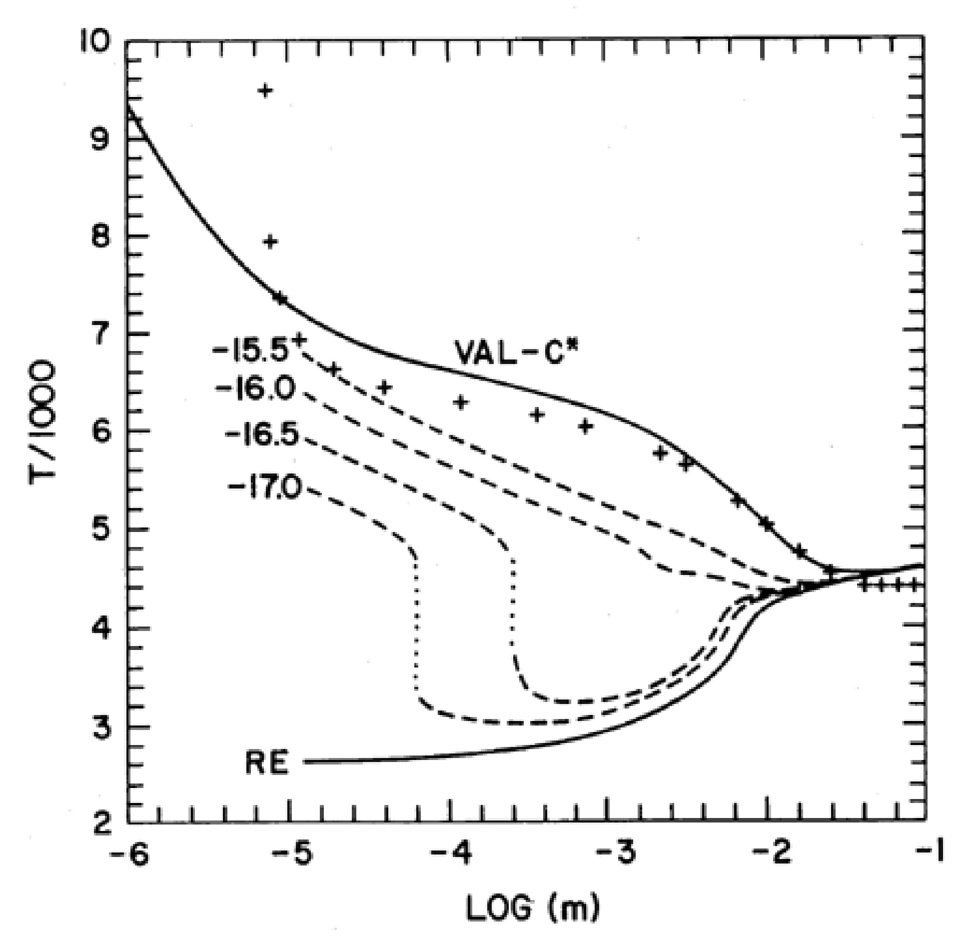
- Classical picture: plane parallel, multi-component atmospheres
- Chromosphere is composed of a gentle rise in temperature between T_{min} and transition region.



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.

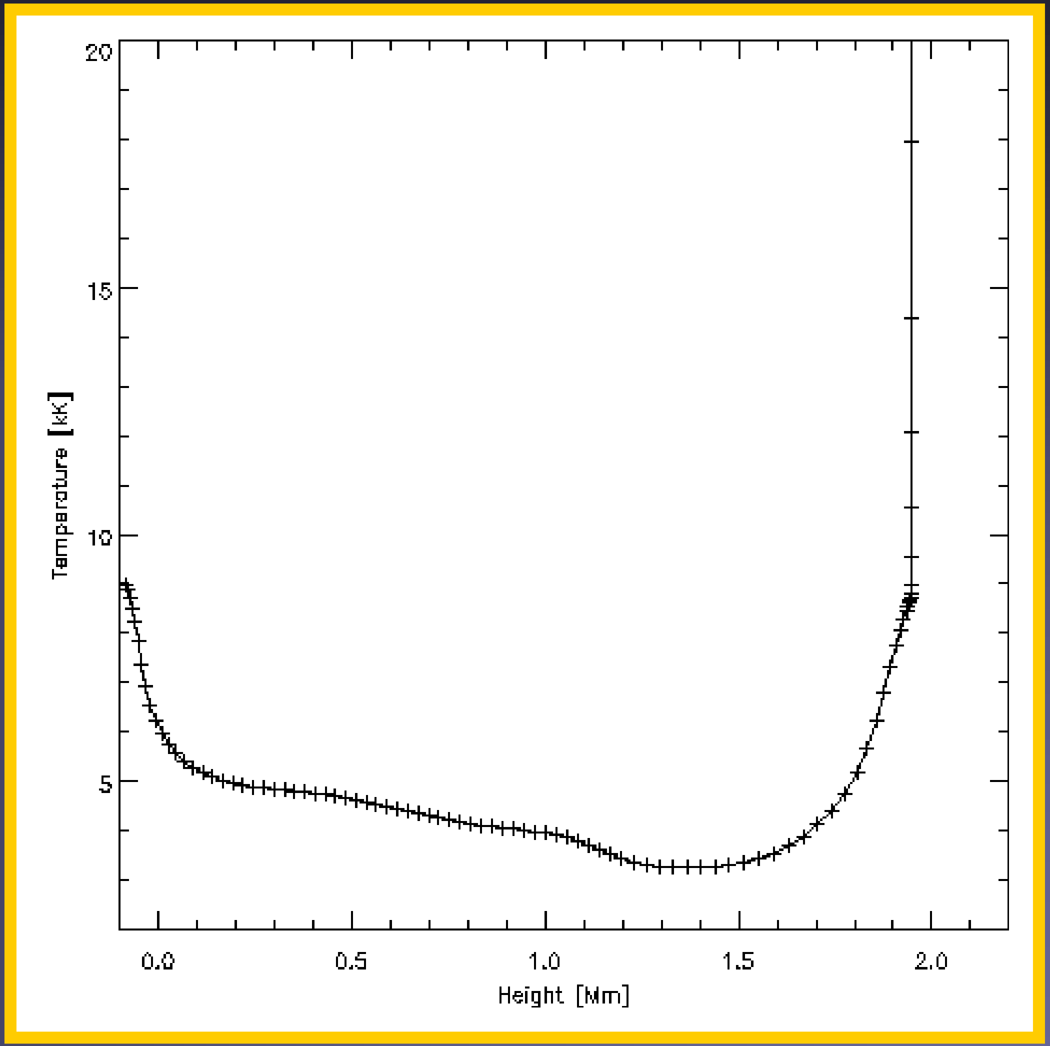


Anderson & Athay 1993

Dynamic models

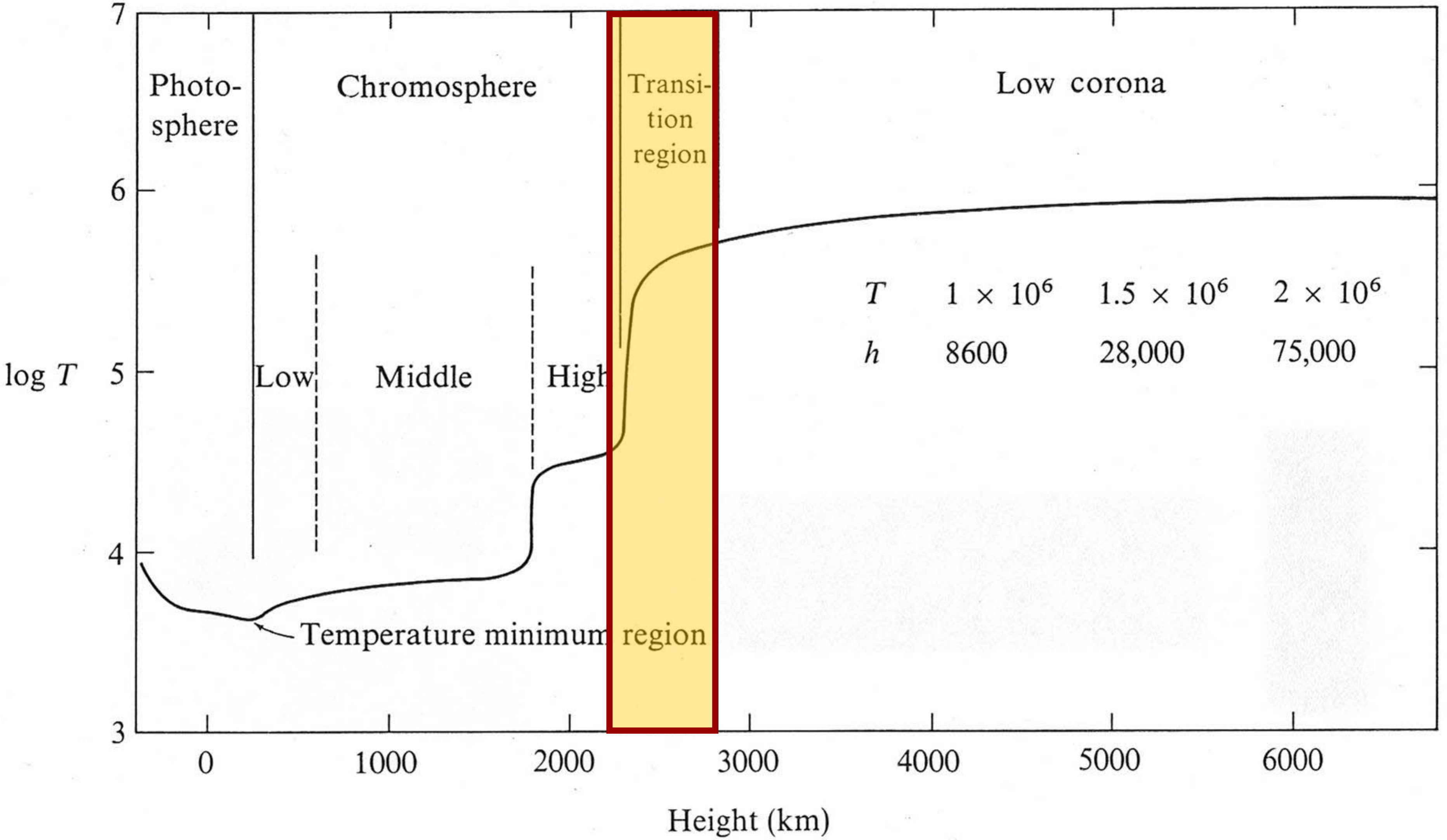
- Start with piston in convection zone, consistent with obs. of photospheric oscillations
- Waves with periods of ≤3min propagate into chromosphere
- Energy conservation ($\rho v^2/2$ = const.) & strong ρ decrease ➔ wave amplitudes increase with height: waves steepen and shock

➔ Temp. at chromospheric heights varies between 3000 K and 10000 K

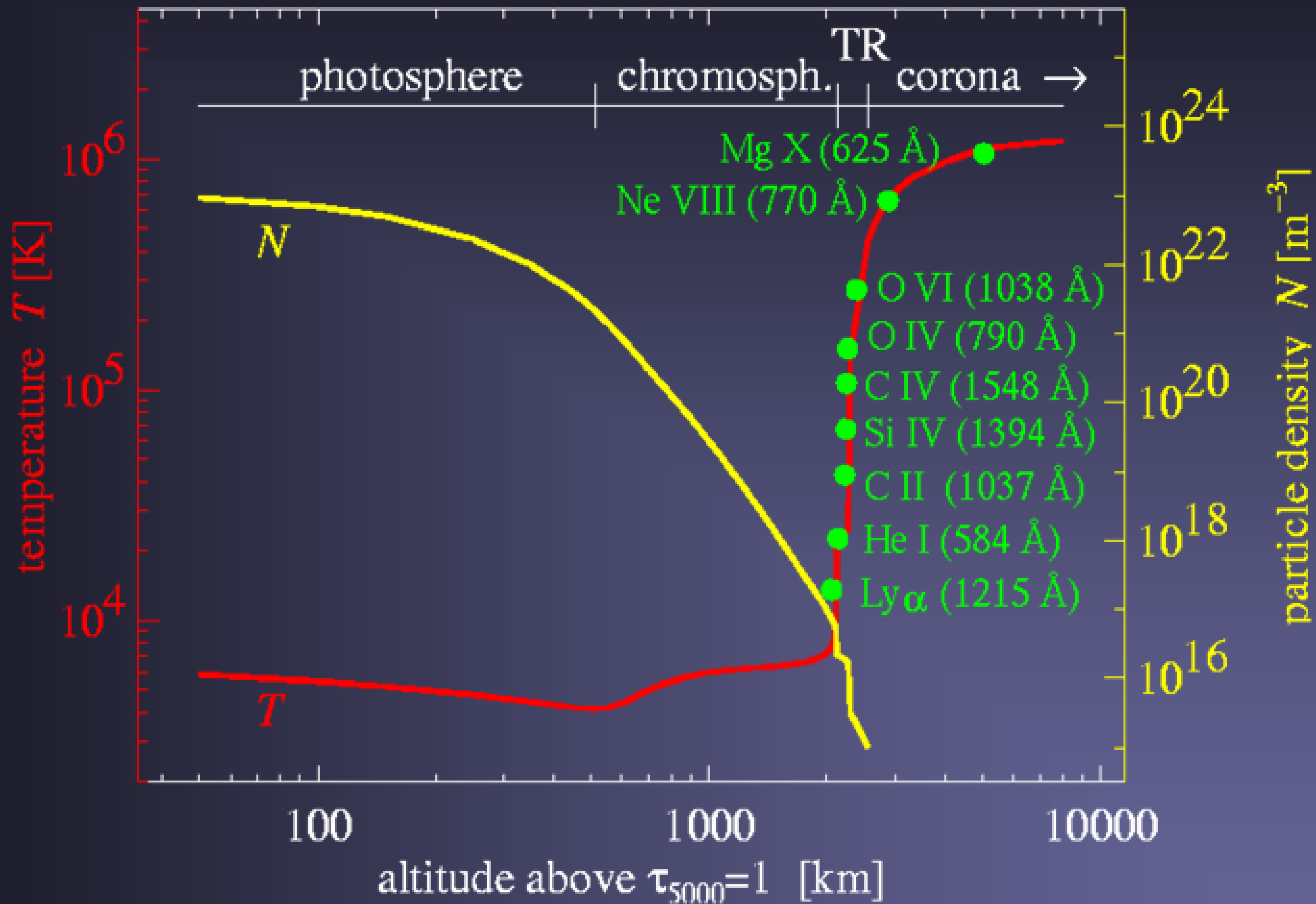


Carlsson & Stein

Transition Region



Transition Region



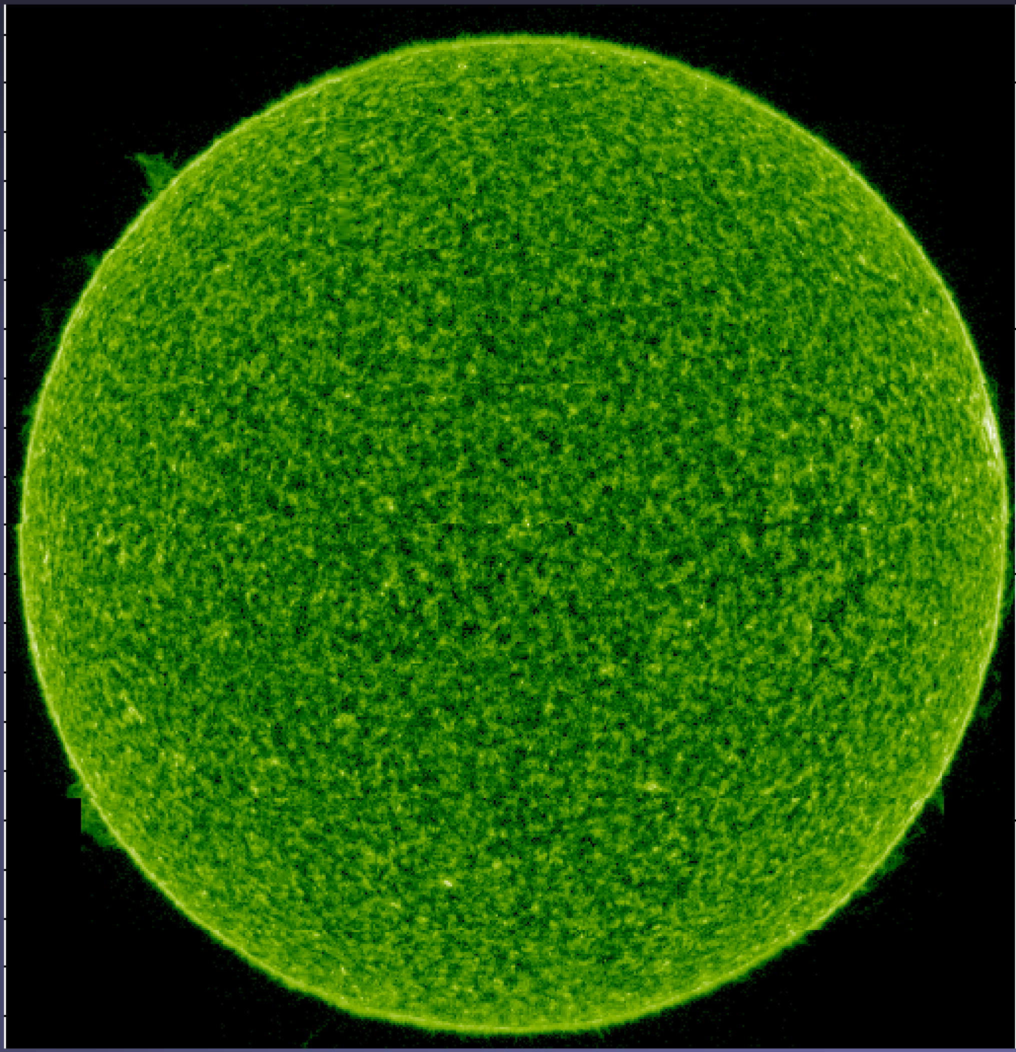
Semi-empirical 1D-models of solar atmosphere: steep increase of T in transition region (TR): < 100 km thick

Transition region properties

- Temperature increases from $5\ 10^4$ K to 1 MK
- Density drops ➔ P_g remains almost constant
- Divided into
  - lower transition region: T< $5\ 10^5$ K. Shows networt structure, similar to Chromosphere
  - upper transition region: T> $5\ 10^5$ K. Shows loop structures, similar to Corona
- Populated by 3 types of structures: footpoints of coronal loops, footpoints of open field lines, cool transition region loops.
- Heating thought to be mainly by heat conduction from corona (for those parts magneticly connected to corona).

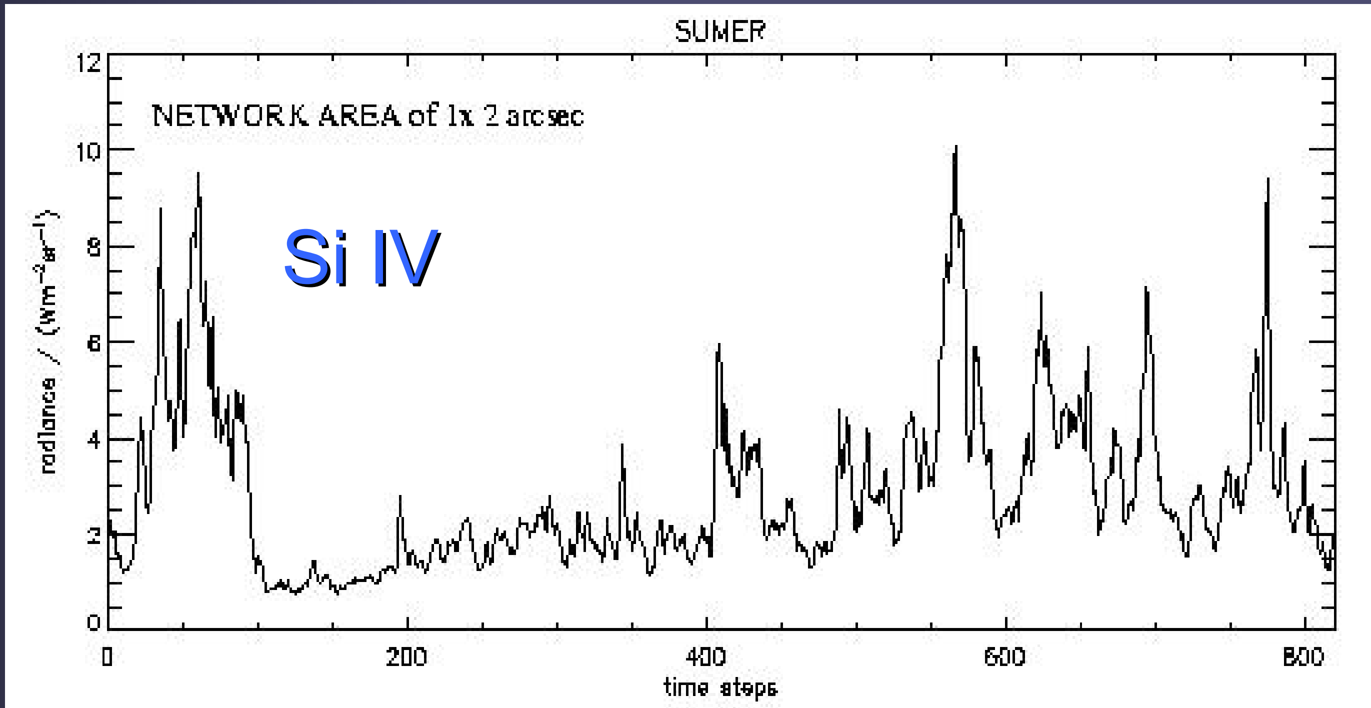
TR spatial structure

- 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



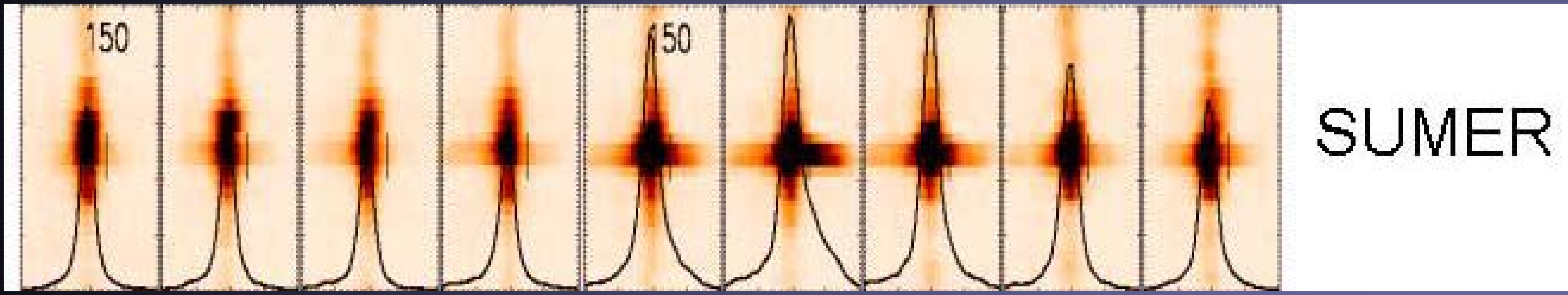
TR dynamic phenomena: blinkers

- Brightness variability in Quiet-sun 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 is 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



Corona

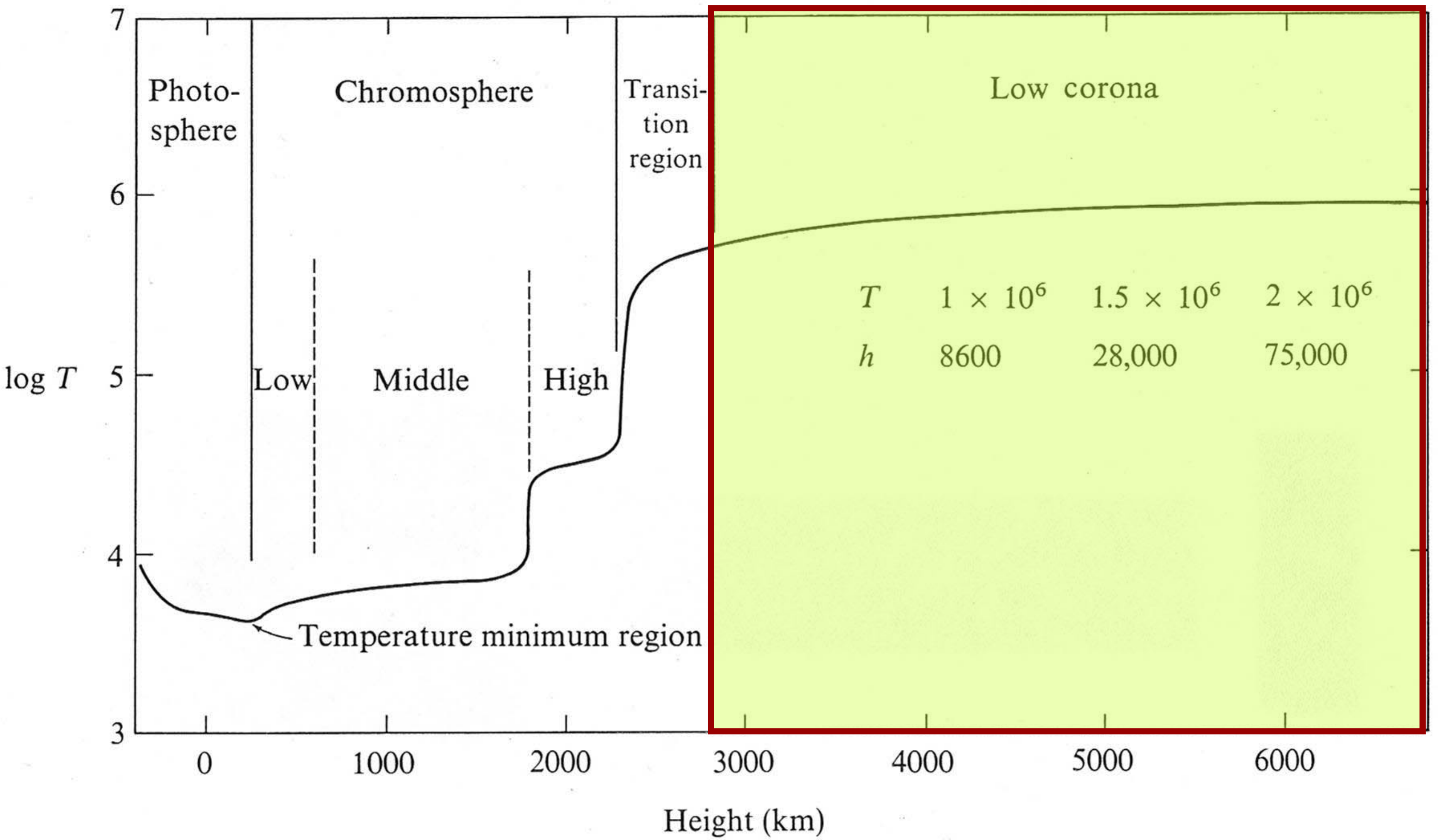
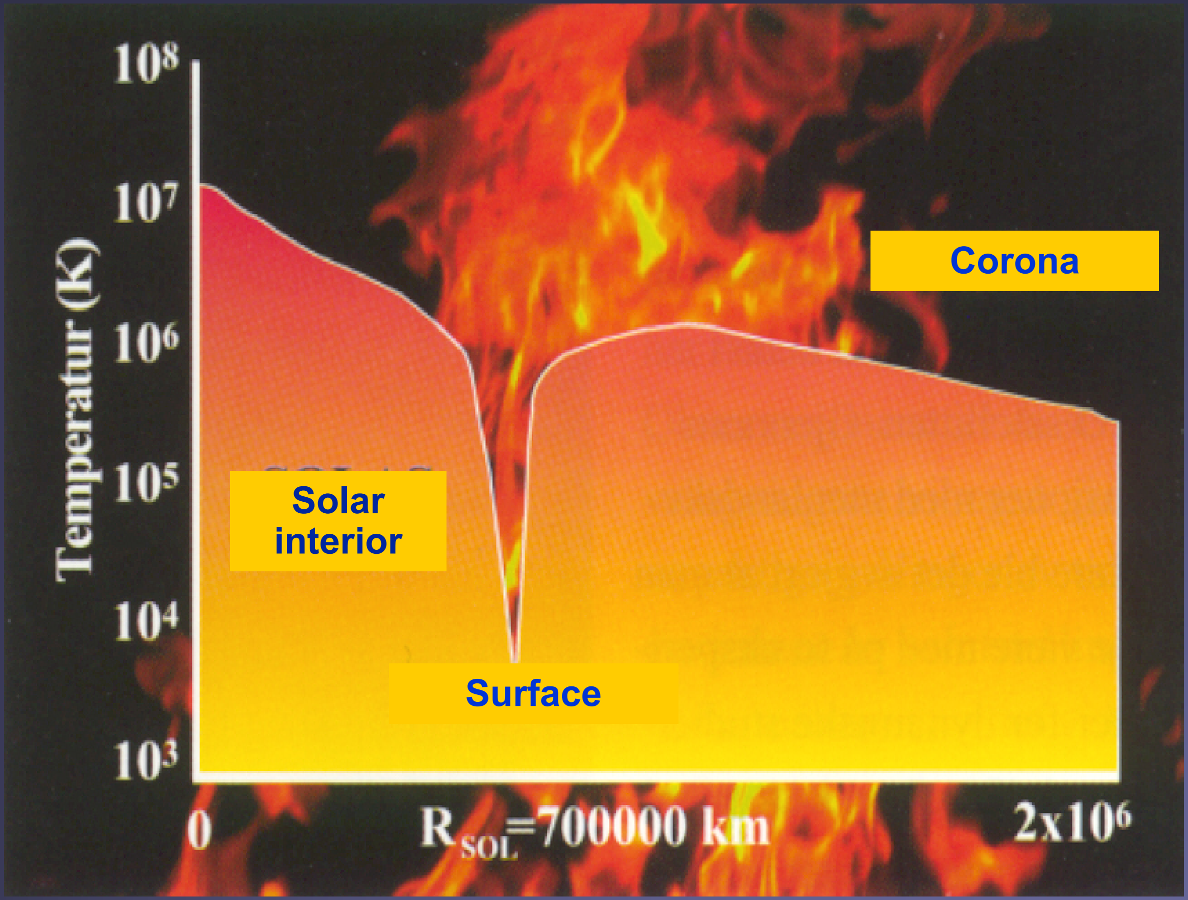


Photo-
sphere
Chromosphere
Transi-
tion
region
Low corona
Low
Middle
High
T 1 × 10^6 1.5 × 10^6 2 × 10^6
h 8600 28,000 75,000
Temperature minimum region
log T
7
6
5
4
3
0
1000
2000
3000
4000
5000
6000
Height (km)

The Solar Corona

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

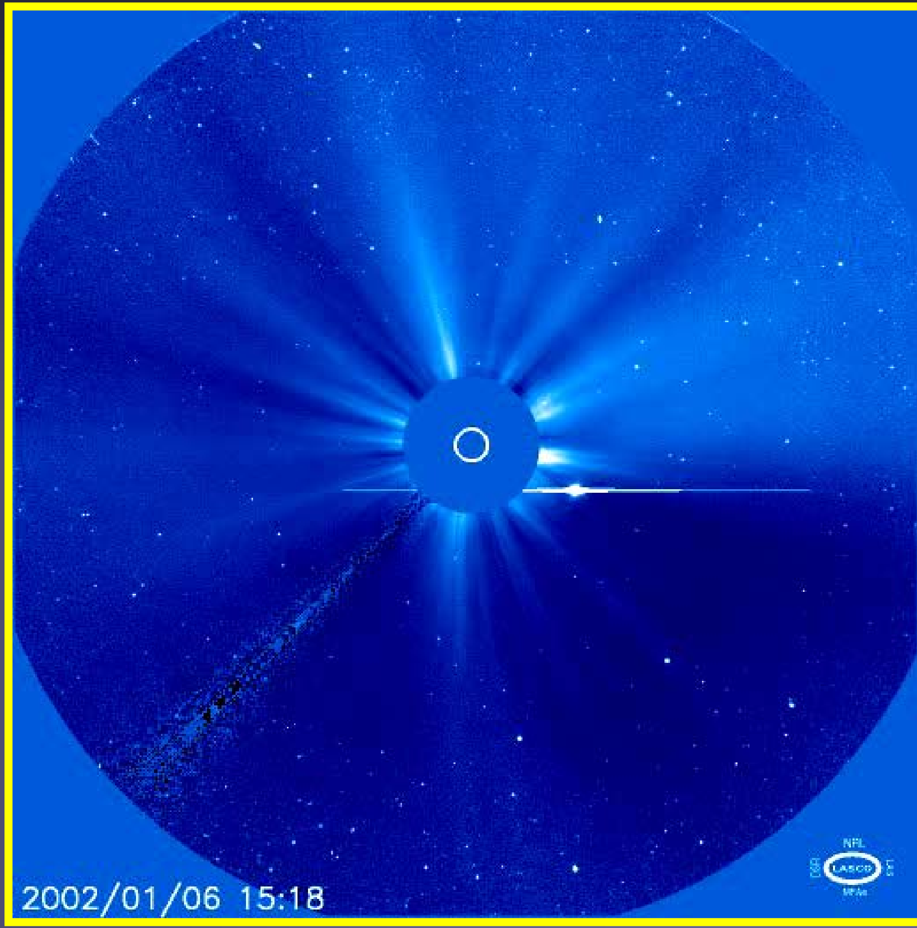
What causes this rapid increase in temperature is still one of the big mysteries in solar physics.



The Hot and Dynamic Corona

Corona during an Eclipse

EUV Corona: Plasma at >1 Mio K (EIT 195 Å)



Artificial eclipse (LASCO C3 / SOHO, MPS)

Eclipse corona

- K corona: Inner portion of sun's corona, having a continuous spectrum caused by electron scattering (Thomson scattering)
- F corona: Outer portion of solar corona, consisting of sunlight scattered from interplanetary dust between sun and earth. Also known as Fraunhofer corona: shows Fraunhofer lines.
- L corona: Emission line corona (forbidden lines). Negligible contribution to coronal brightness

Coronal brightness

- Total visible flux from corona
  - Activity maximum: $1.5\ 10^{-6}\ F_{\odot} = 0.66$ full Moon
  - Activity minimum: $0.6\ 10^{-6}\ F_{\odot} = 0.26$ full Moon
- Visibility: during an eclipse corona typically extends for 4 solar radii until its brightness drops below sky brightness levels

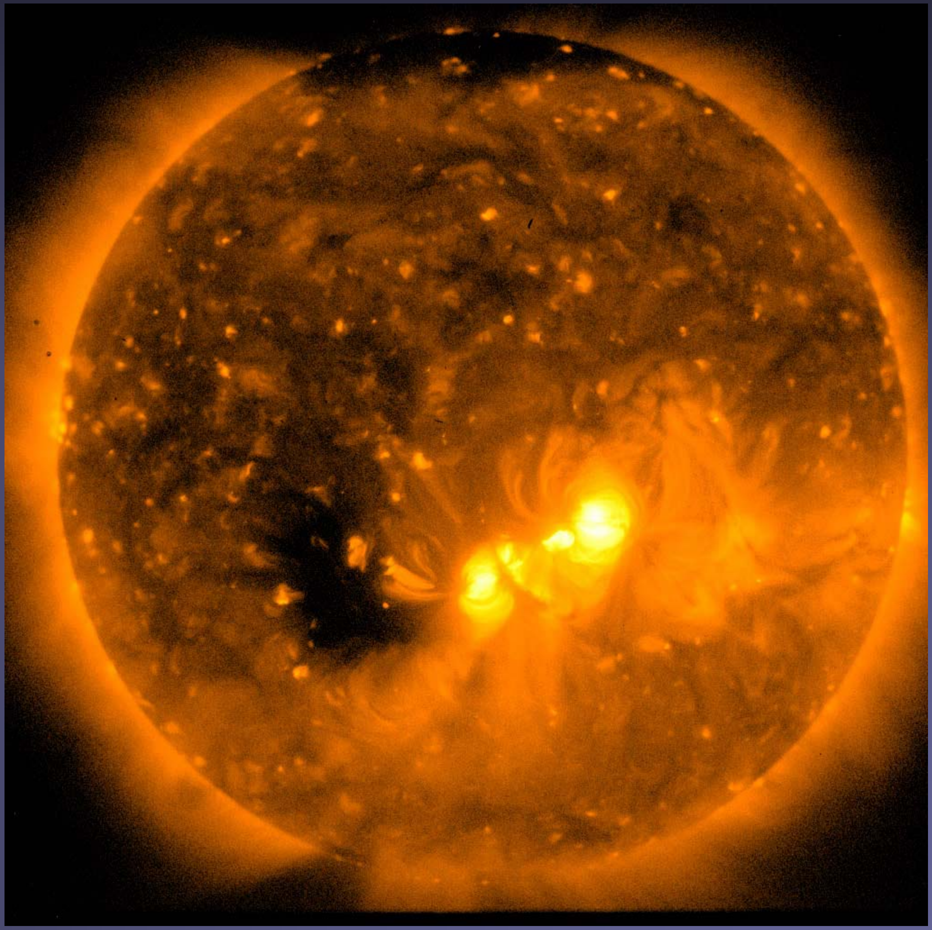
- Intensity vs. distance R from limb (I_0=disk centre intensity)

$$\frac{I}{I_O} = 10^{-6}\left(\frac{0.0532}{R/R_O^{2.5}} + \frac{1.425}{R/R_O^{7}} + \frac{2.565}{R/R_O^{17}}\right)$$

Dominates: far intermediate close

Coronal temperature

- Different temperatures & densities co-exist in the corona
- Range of temps: <1 MK (Coronal hole) to 10 MK (act. region)
- Range of e^- densities (inner corona):
  - Loop: 10^{10}
  - coronal hole: 10^7

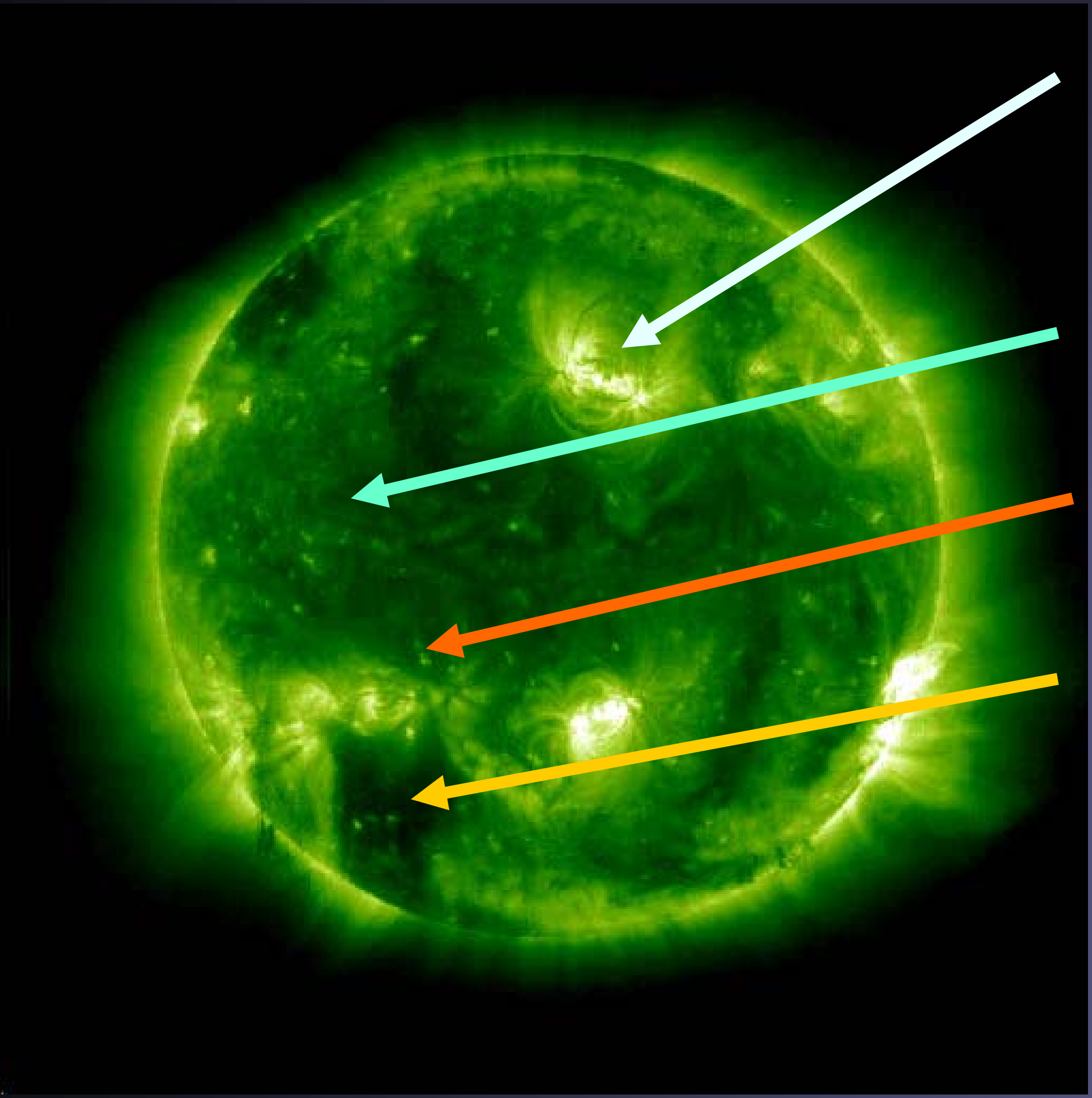


Hinode XRT: 2-5MK gas

Coronal spectrum

- Coronal spectrum: spectrum of the radiation produced in the corona. In contrast to K & F corona, which is radiation coming from the photosphere & only scattered in the corona.
- It is dominated by emission lines of highly ionized species (e.g. Fe IX – Fe XX) with no continuum contribution
- **GET FIGURE !!!!!!!!!!!!!!!!!!!!!!!!!!! X-RAY SPECTRUM**

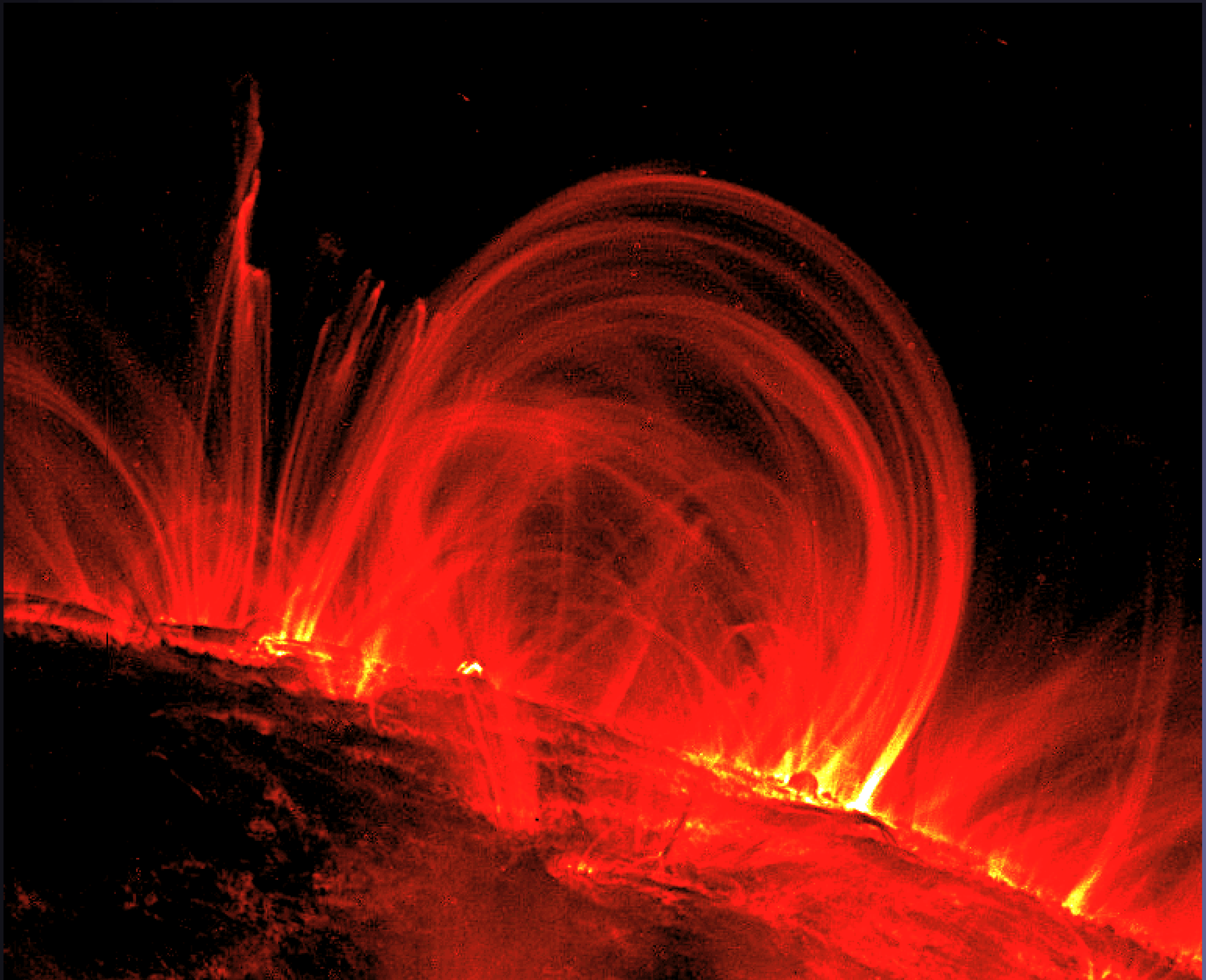
Coronal structures



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

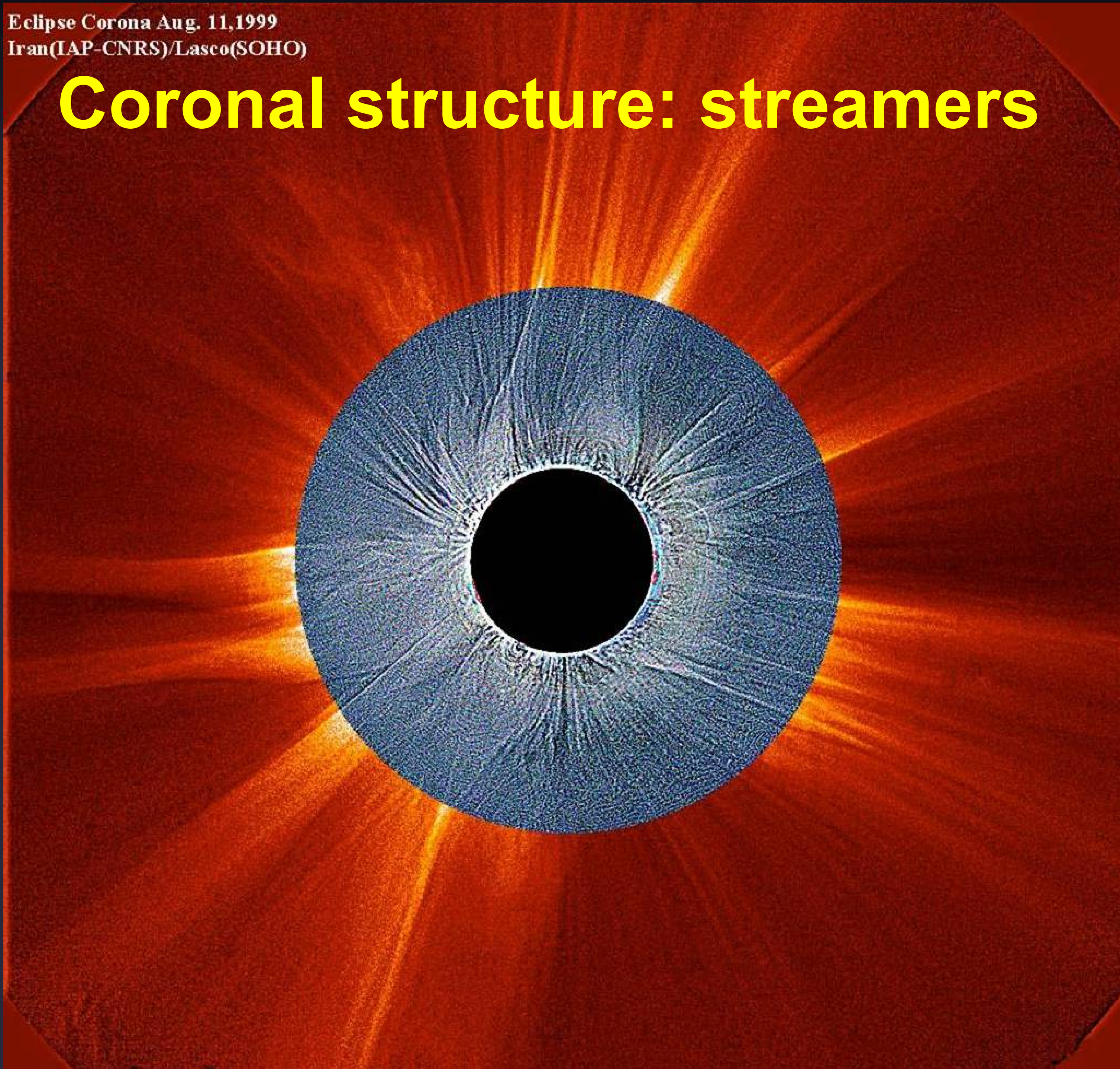
Fe XII 195 Å
(1.500.000 K)
17 May - 8 June 1998

Coronal structure: active region loops

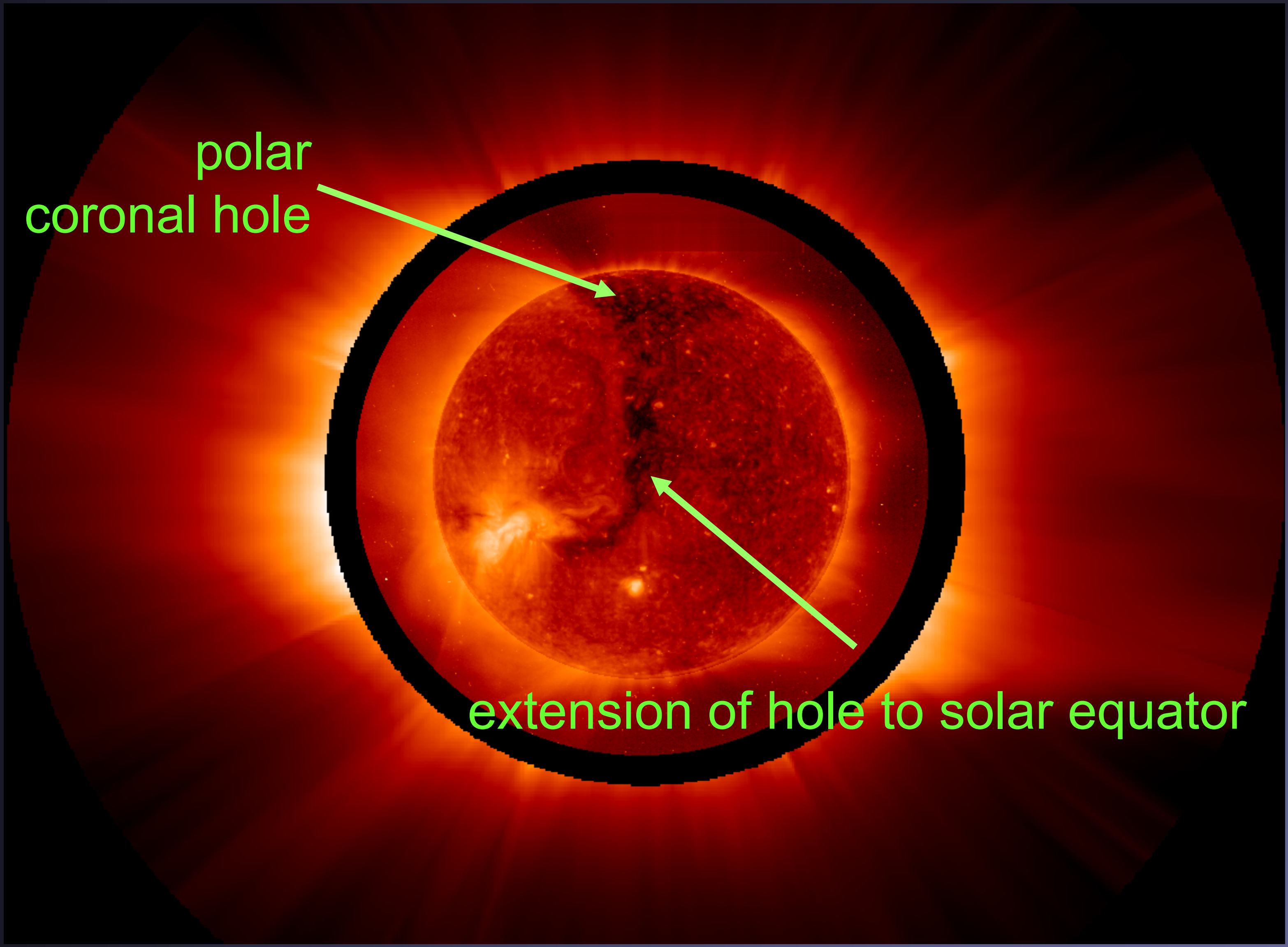


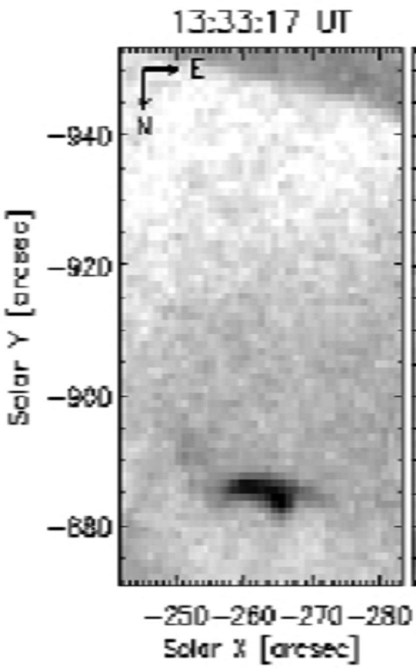
TRACE, 1999

Coronal structure: streamers



Coronal structures: coronal holes



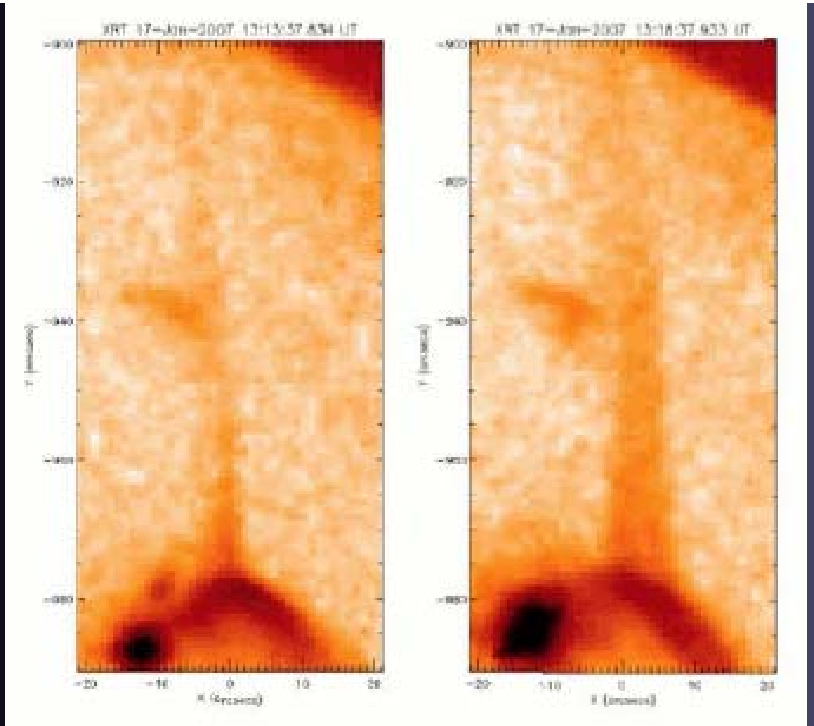


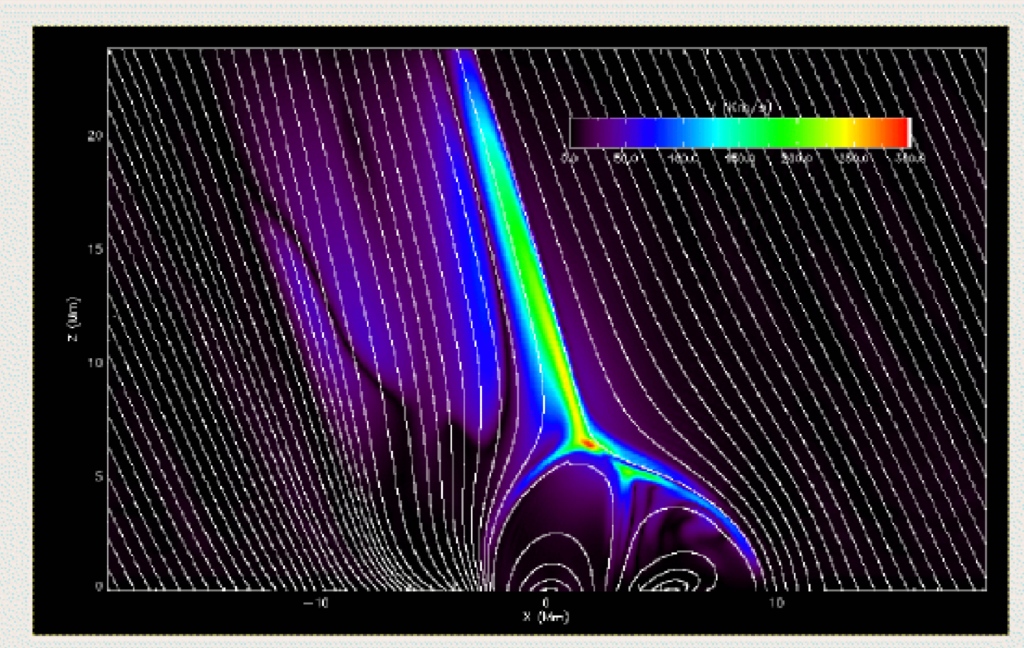
Coronal structures: coronal jets

XRT observations

Model (Moreno Insertis et al.

Velocity map

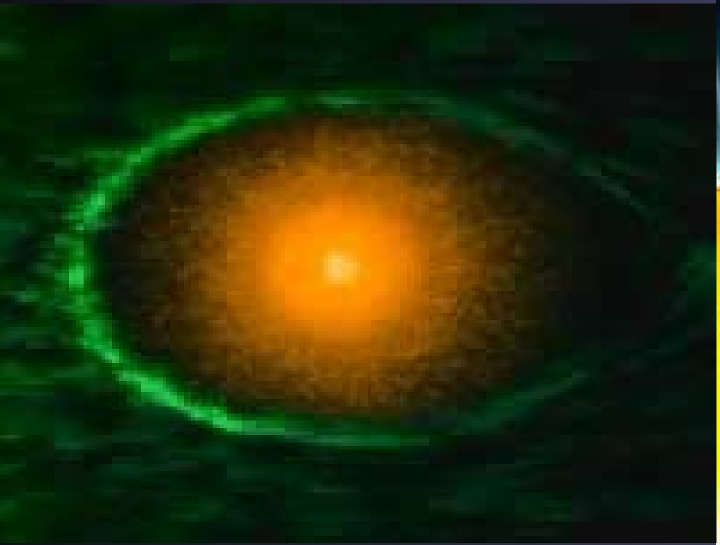


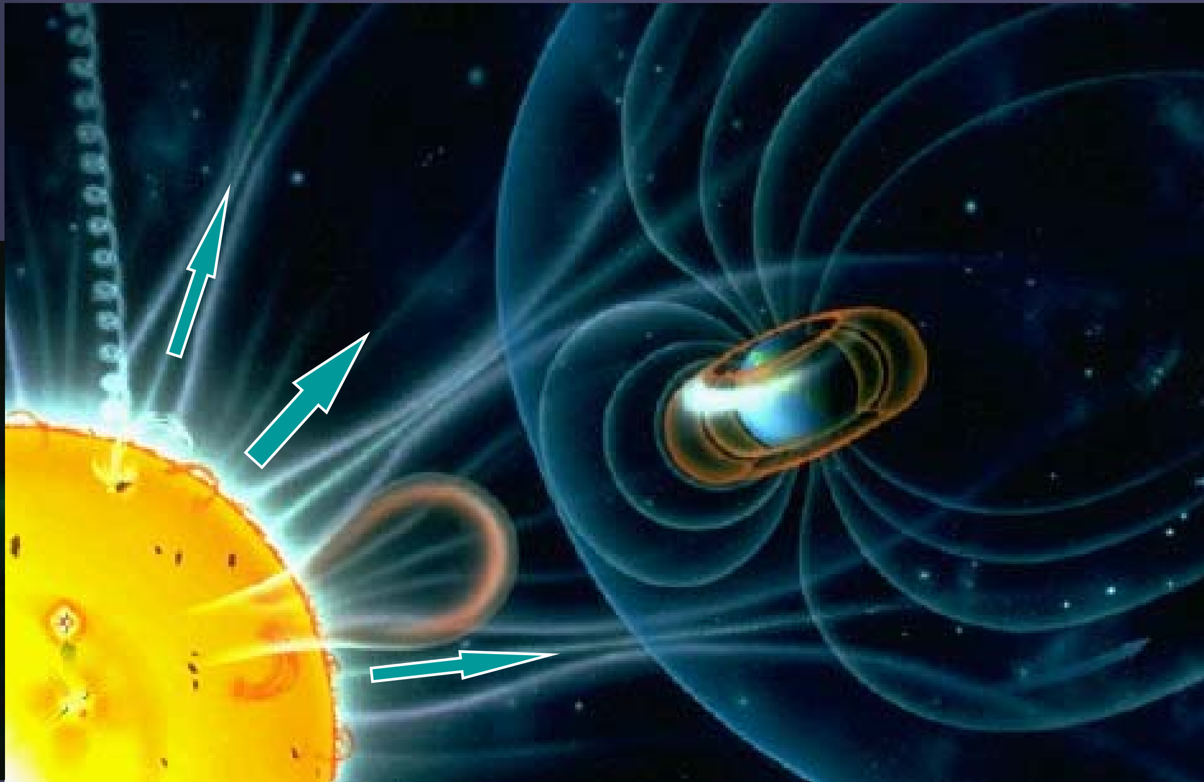


(t=22 min)

The solar wind

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





Discovery of the solar wind

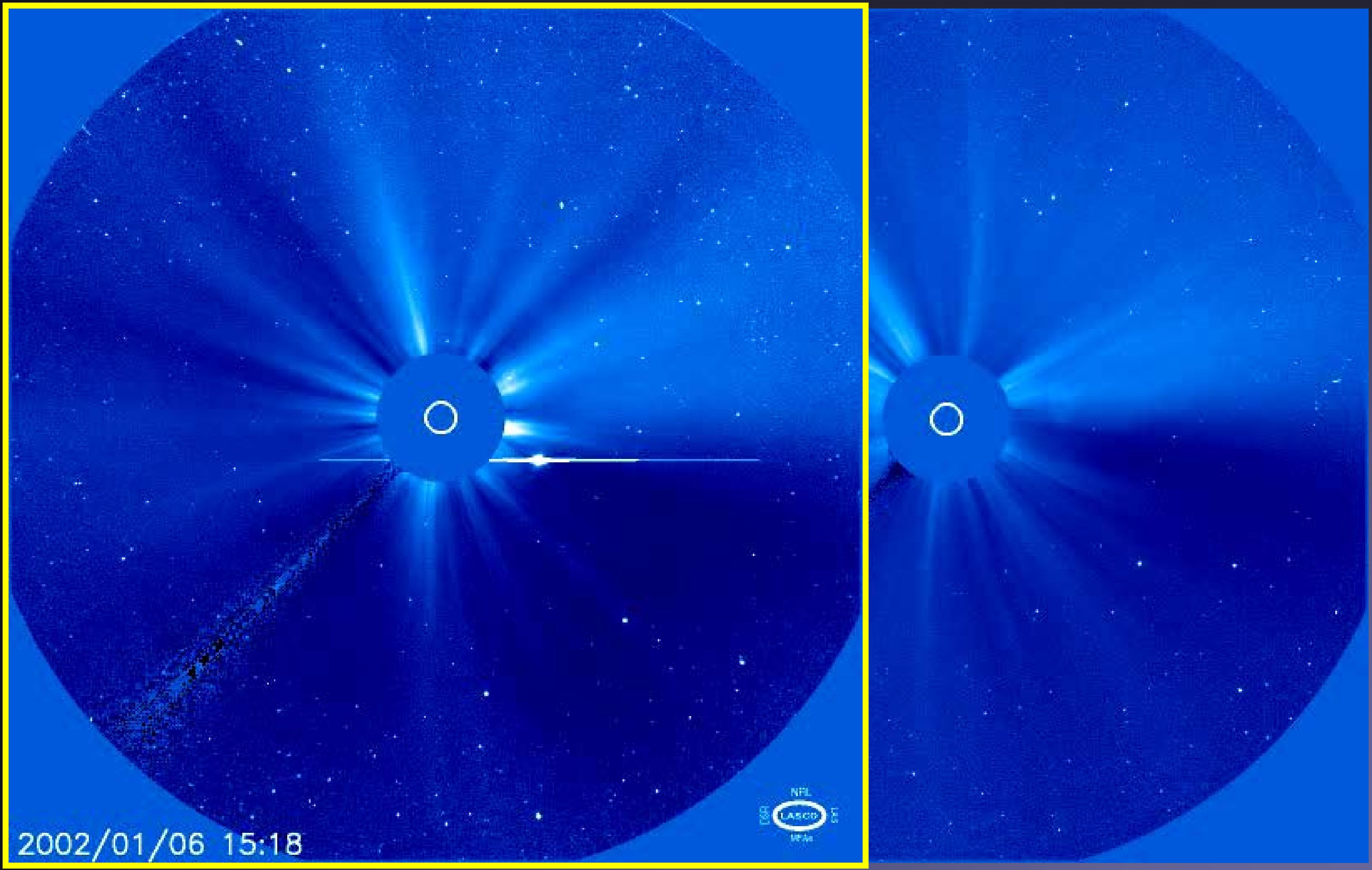
- Ludwig Biermann at MPI für Physik und Astrophysik noticed that the tails of comets always pointed away from the Sun. Solar radiation pressure was insufficient to explain this.

➔ Postulated a solar wind

- Independently, Parker (1958) realized that a hot corona must expand if it was to be in equilibrium with the interstellar medium. Only a supersonic solar wind was compatible with theory and observations.

➔ Supersonic solar wind

Comets and the solar wind



Solar wind characteristics at 1AU

Fast solar wind

- speed > 400 km/s
- $n_p \approx 3\ cm^{-3}$
- homogeneous
- $B \approx 5\ nT = 0.00005G$
- 95% H, 4% He
- Alfvenic fluctuations
- Origin: coronal holes

Slow solar wind

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

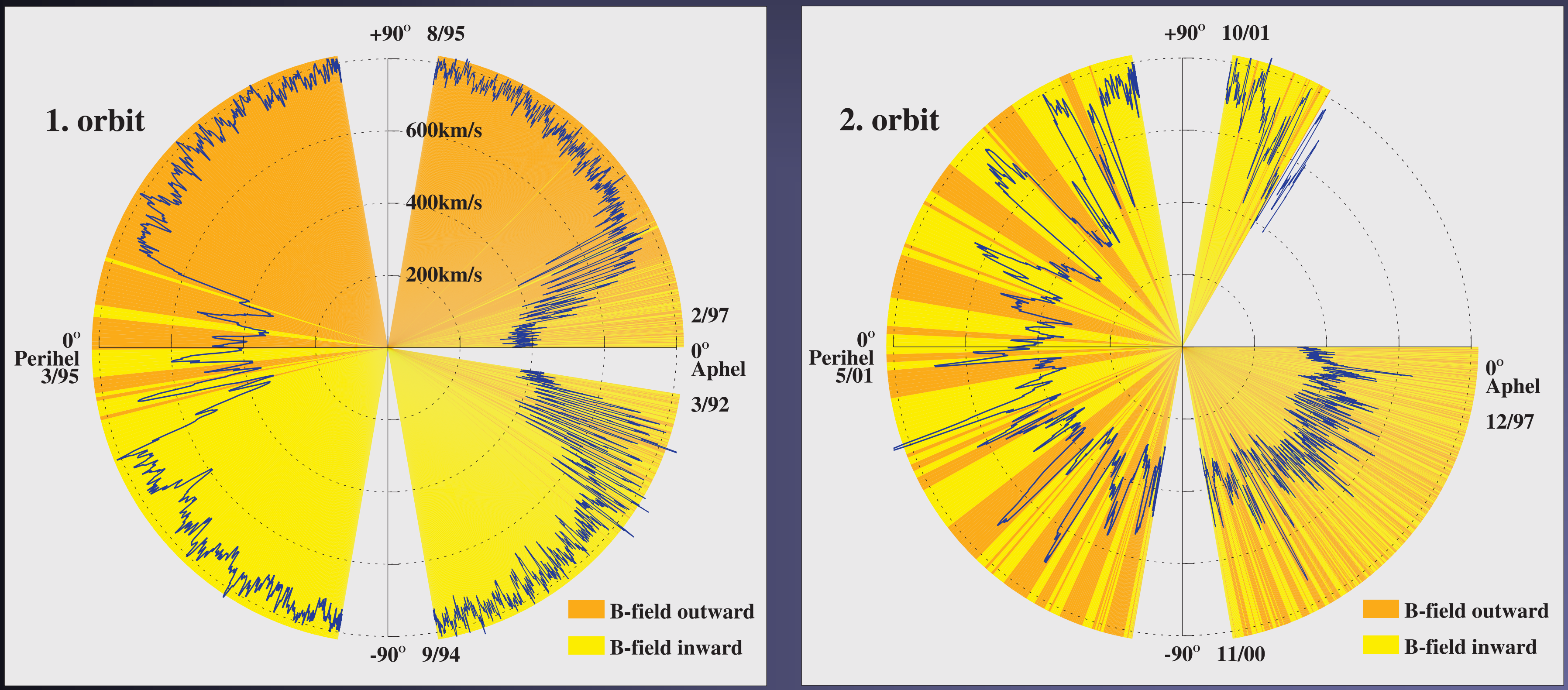
Transient solar wind

- speed from < 300 km/s up to >2000 km/s
- Variable B, with B up to 100 nT (0.001G)
- 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

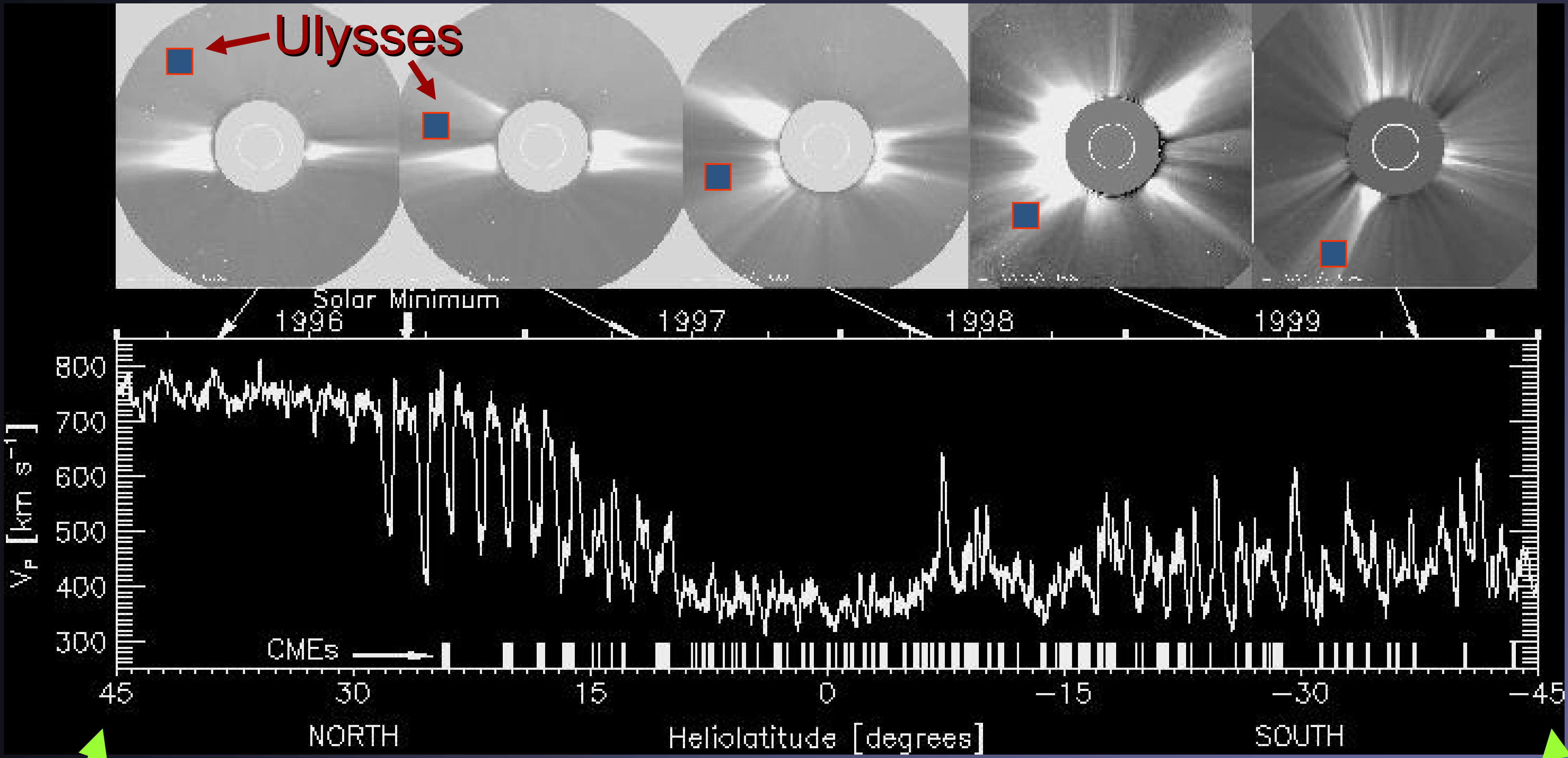
2nd Orbit: 12/1997 - 2/2002
rising / maximum phase



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:

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

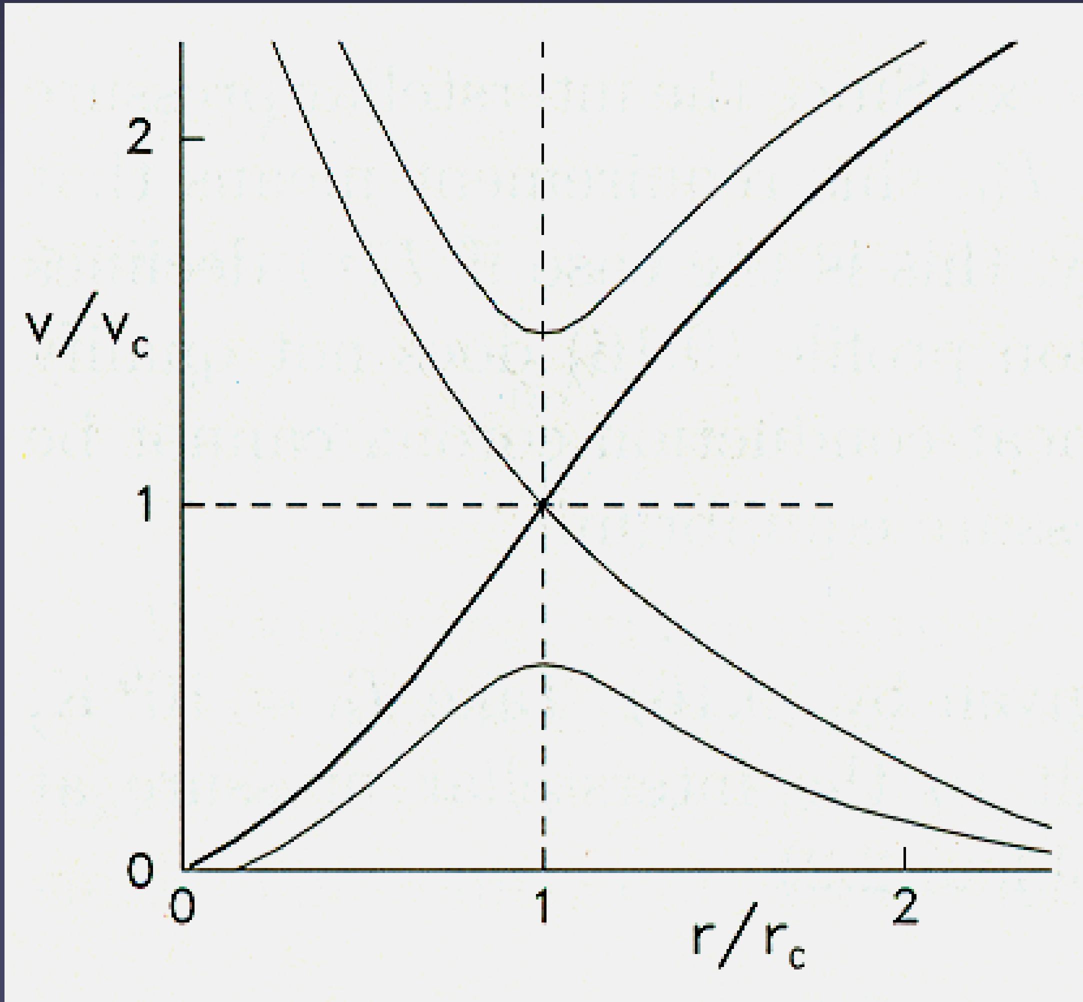
$$v\frac{dv}{dr} = -\frac{1}{\rho}\frac{dP}{dr}\frac{GM}{r^2}$$

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

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

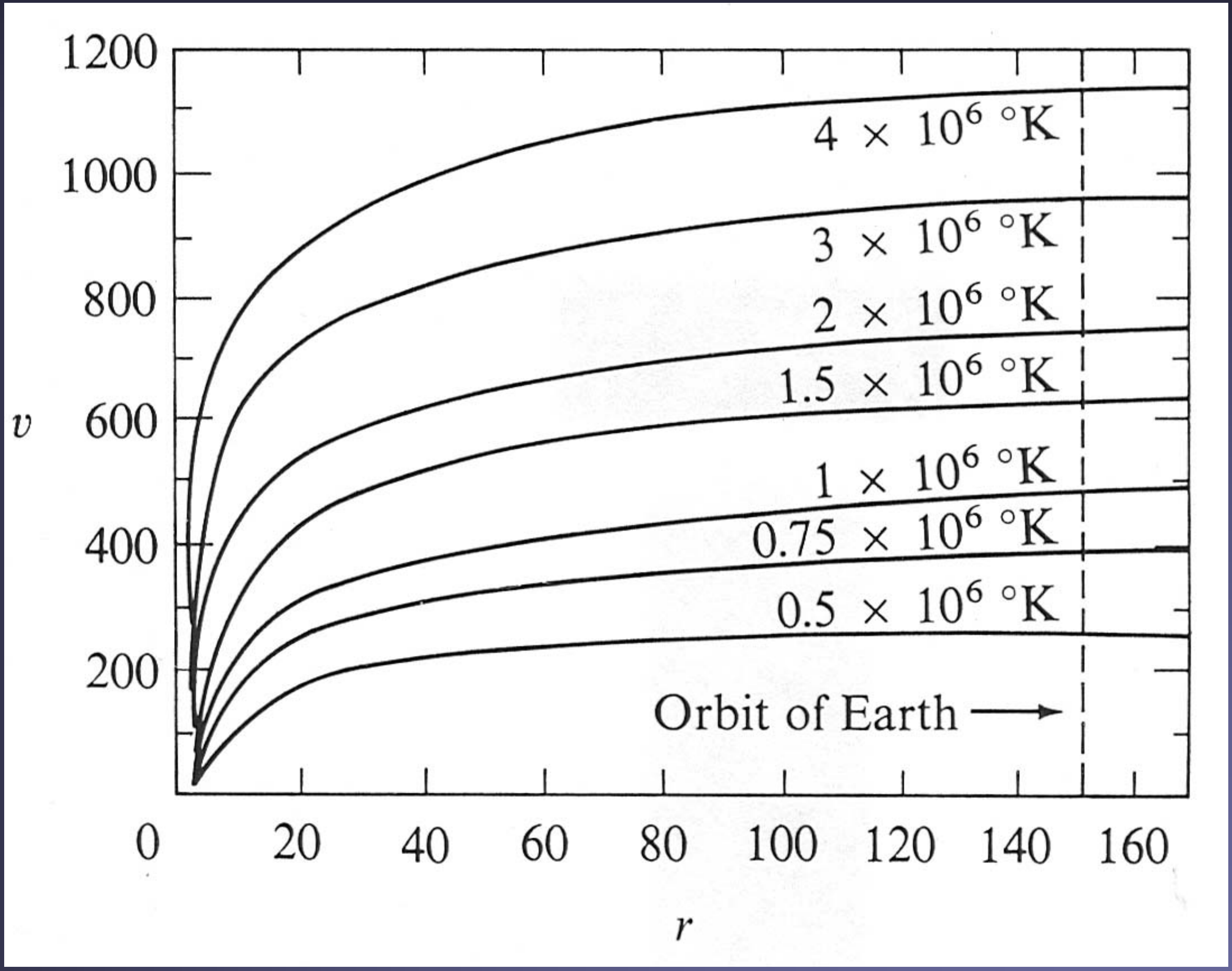
Parker's solar wind solutions

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

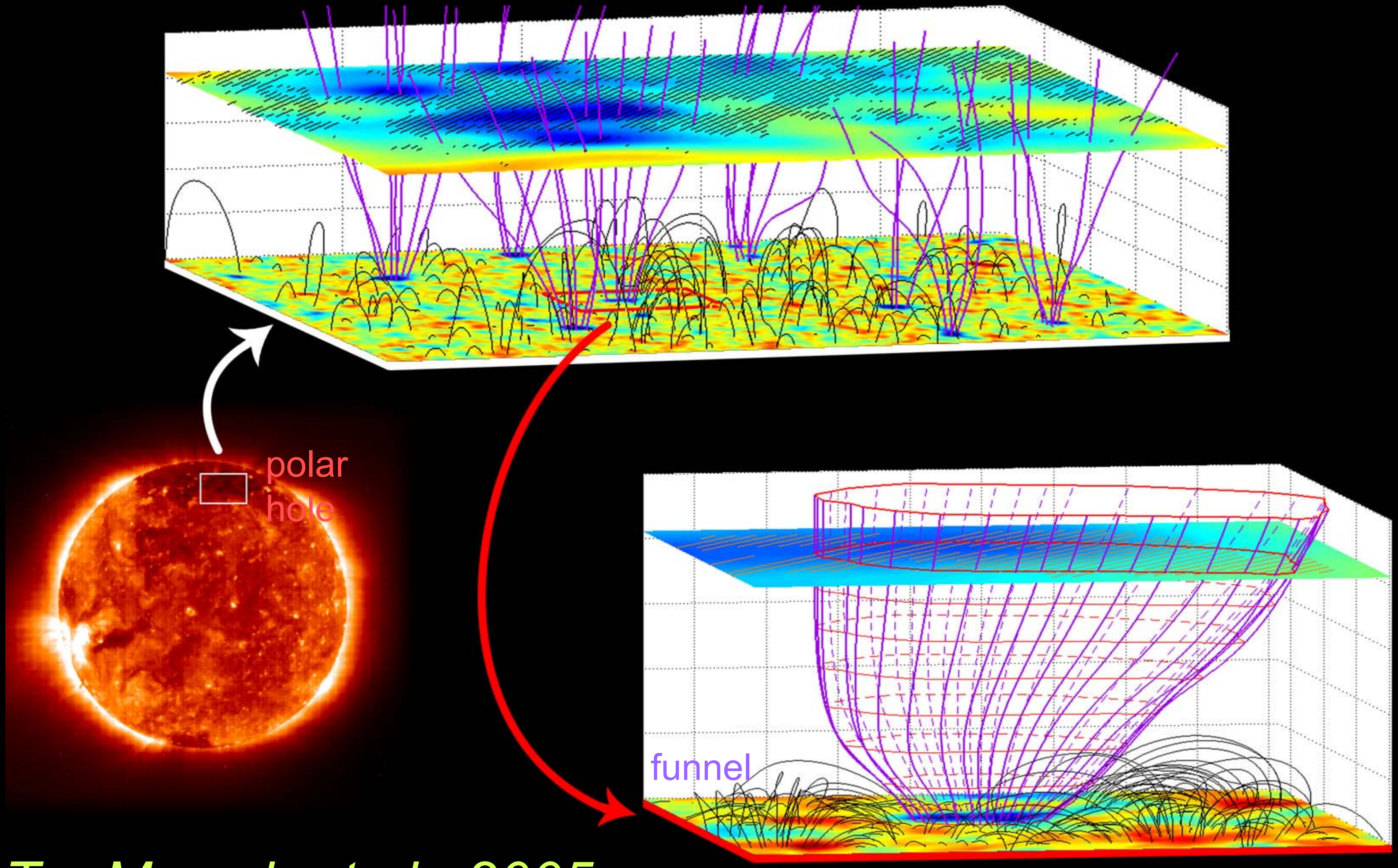


Solar wind speed

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

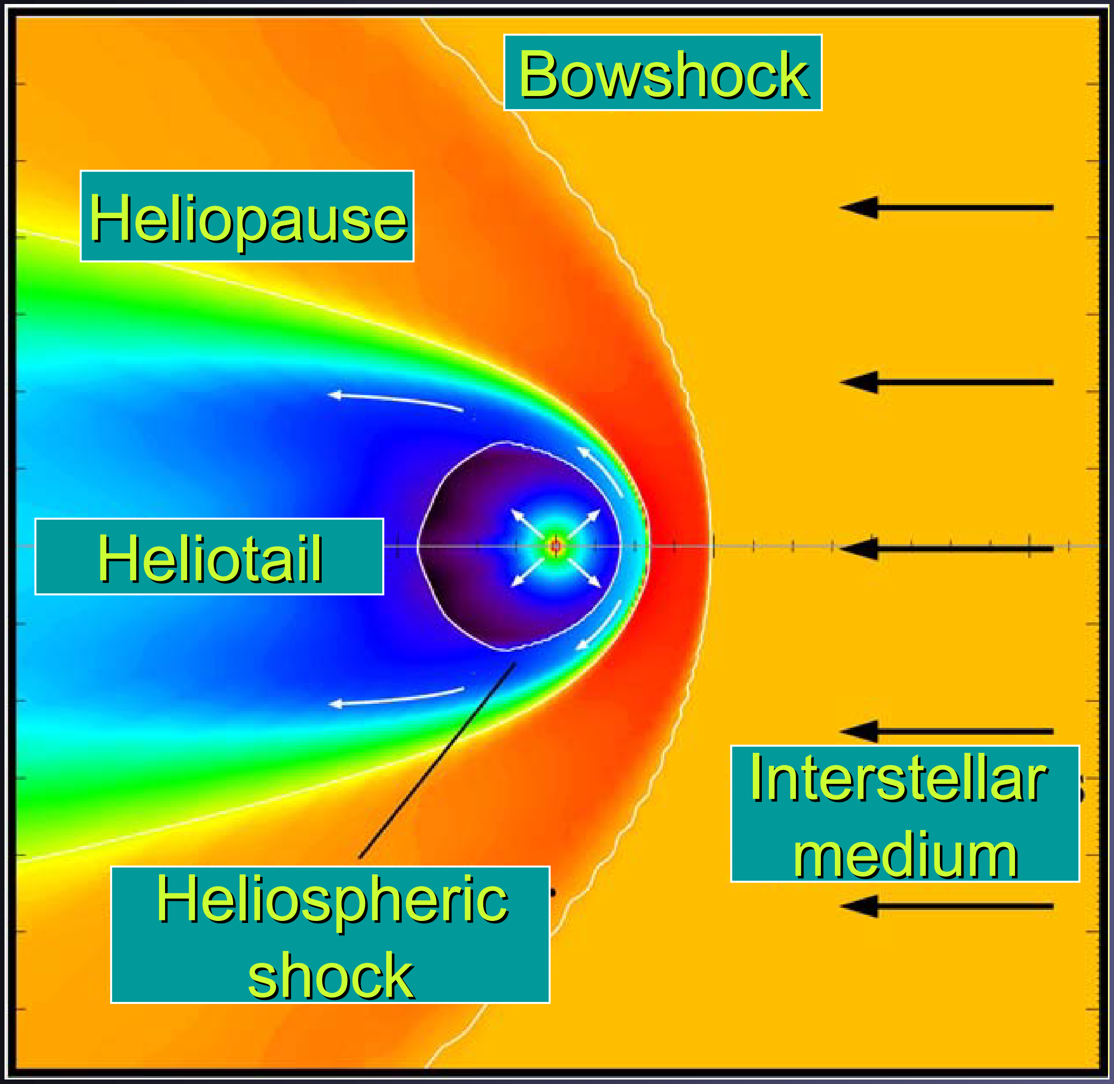


Sources of solar wind: fast wind

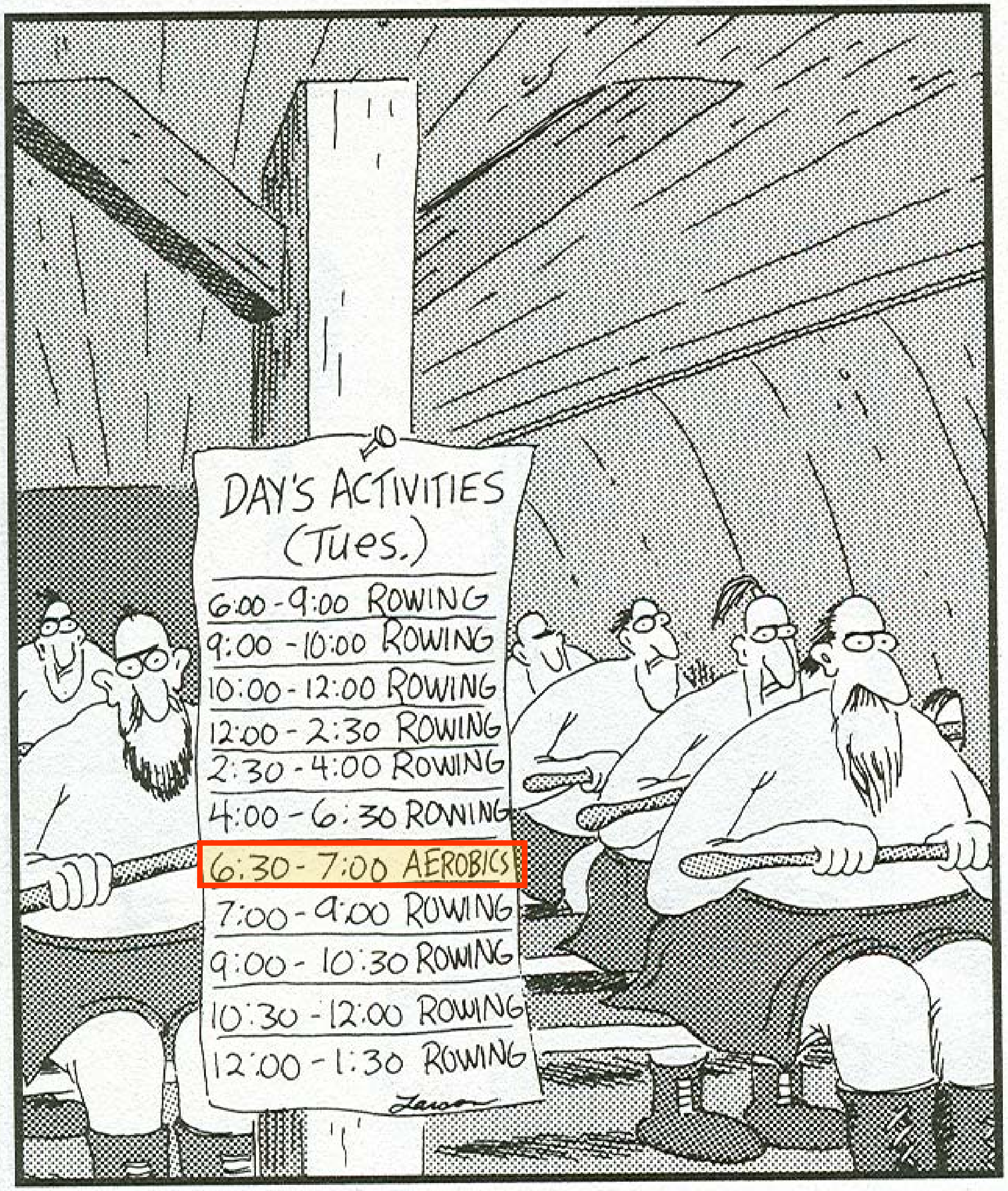


Tu, Marsch et al., 2005

The Heliosphere



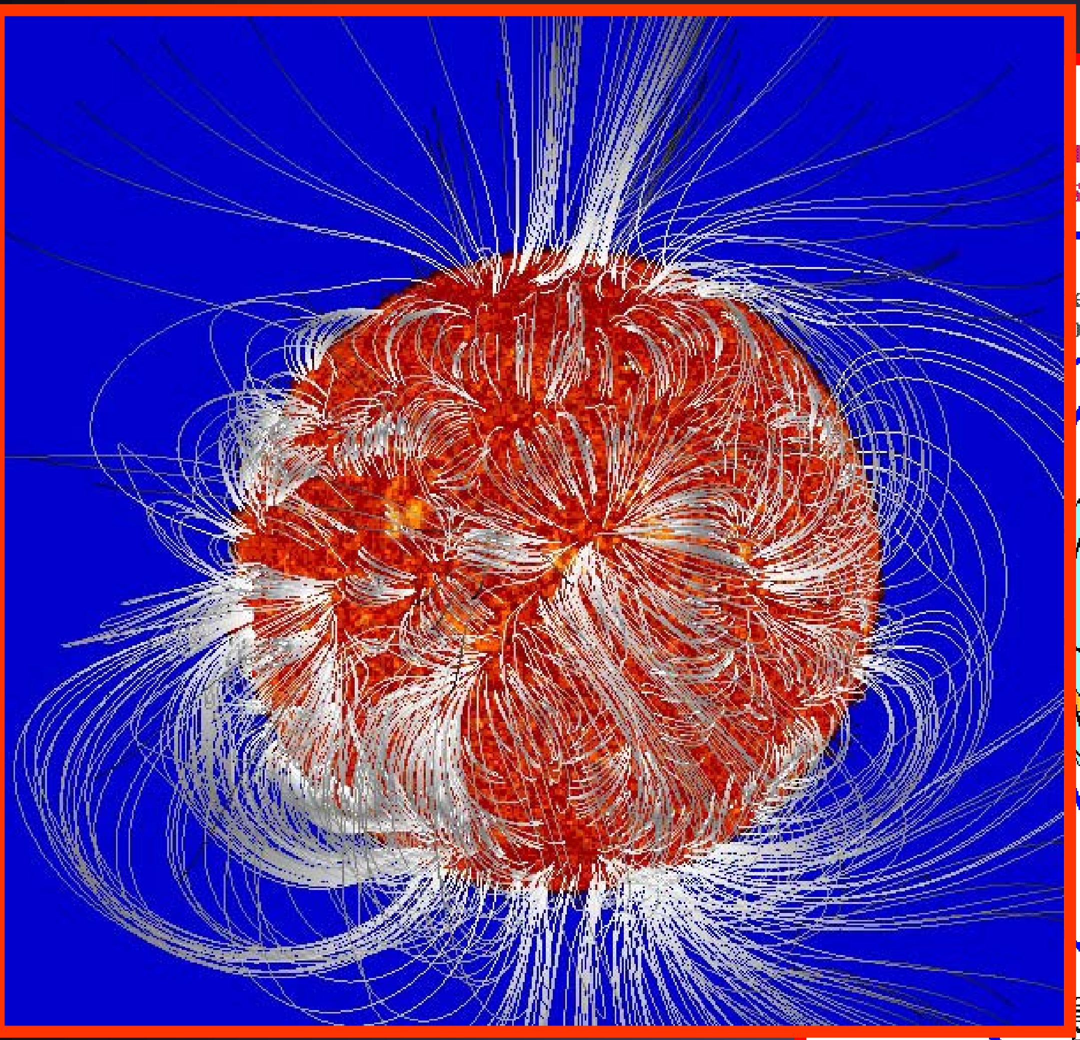
- Heliosphere = region of space in which the solar wind and solar magnetic field dominate over the interstellar medium and the galactic magnetic field.
- Bowshock: where the interstellar medium is slowed relative to the Sun.
- Heliospheric shock: where the solar wind is decelerated relative to Sun
-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

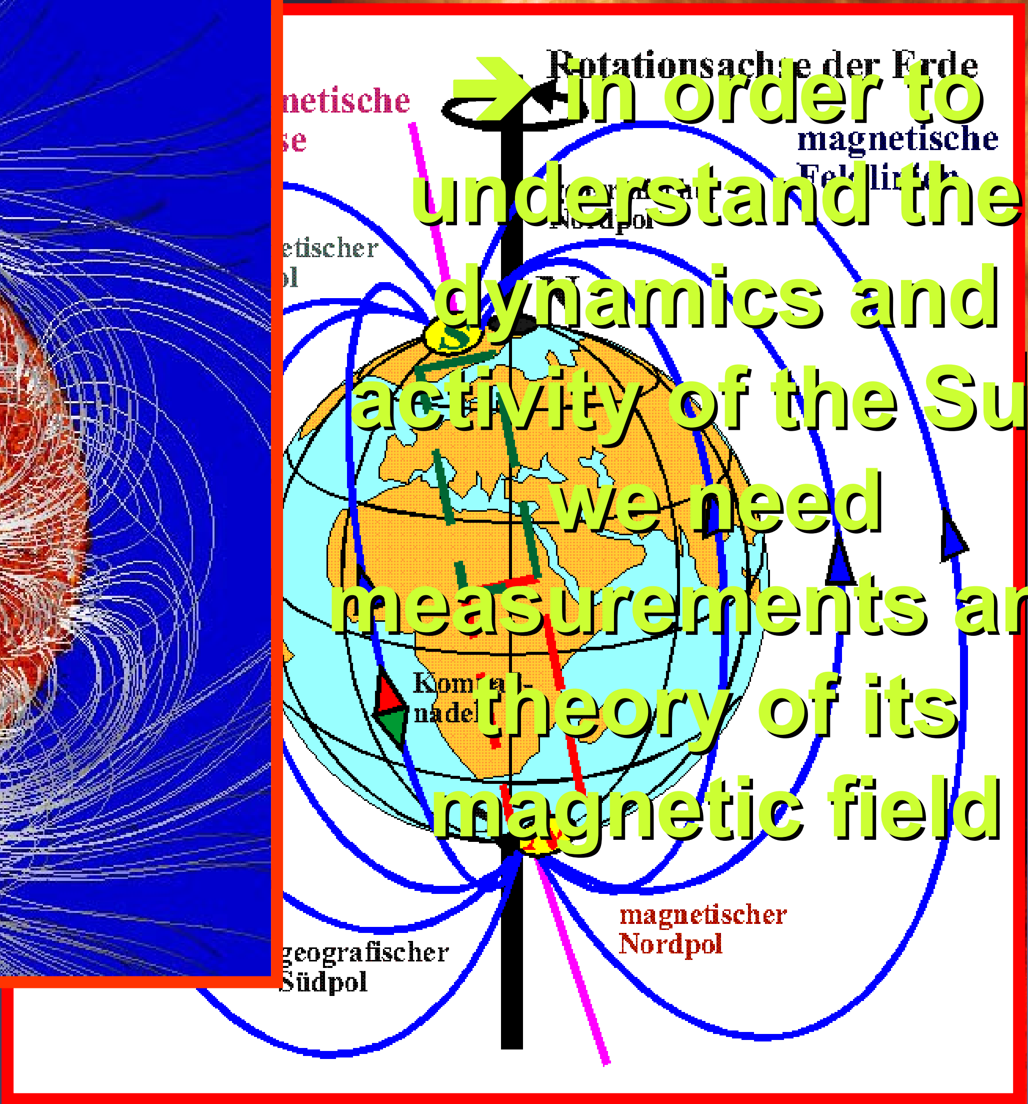


Slave-ship daily schedules

Magnetic Field

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

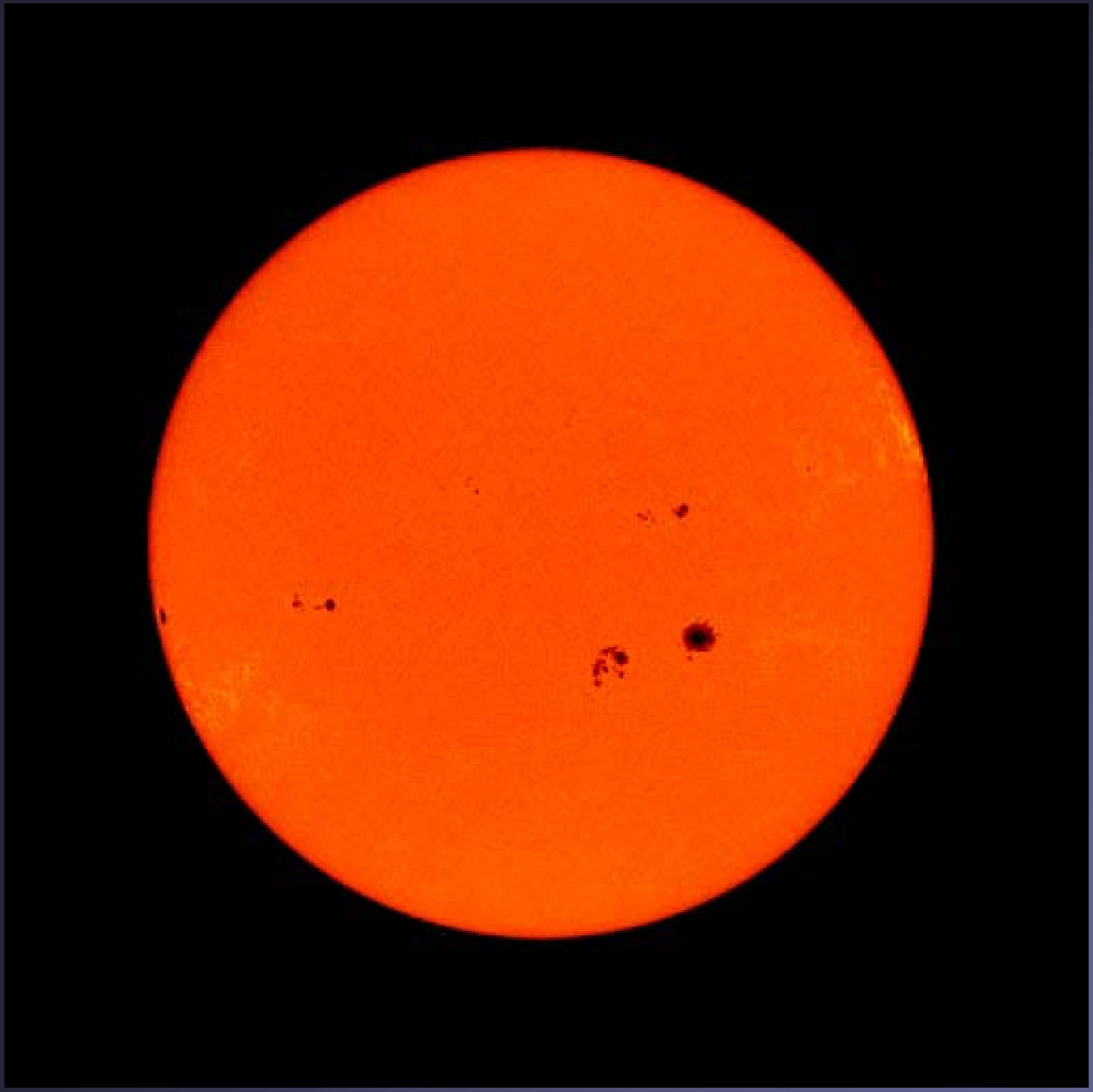




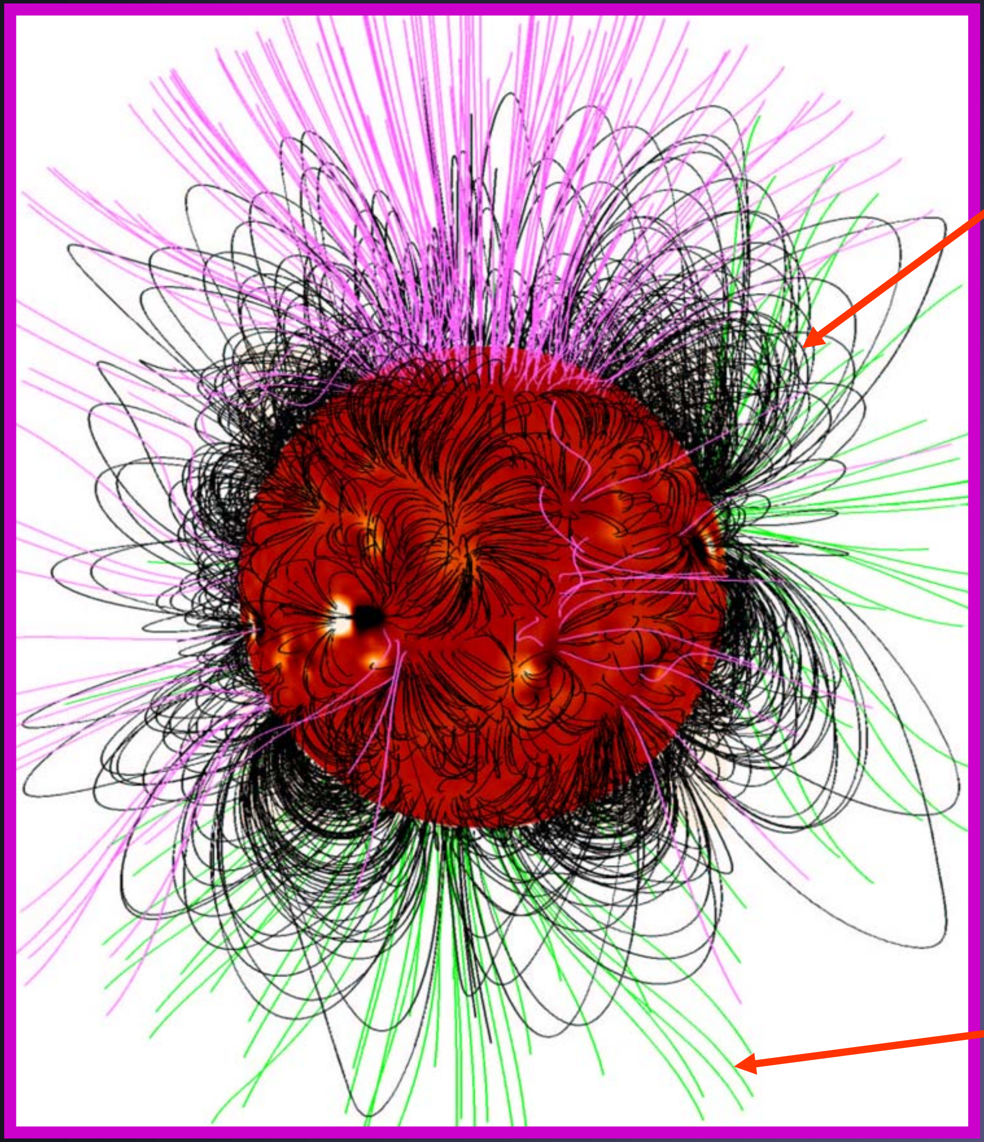
➔ in order to understand the dynamics and activity of the Sun we need measurements and theory of its magnetic field

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)

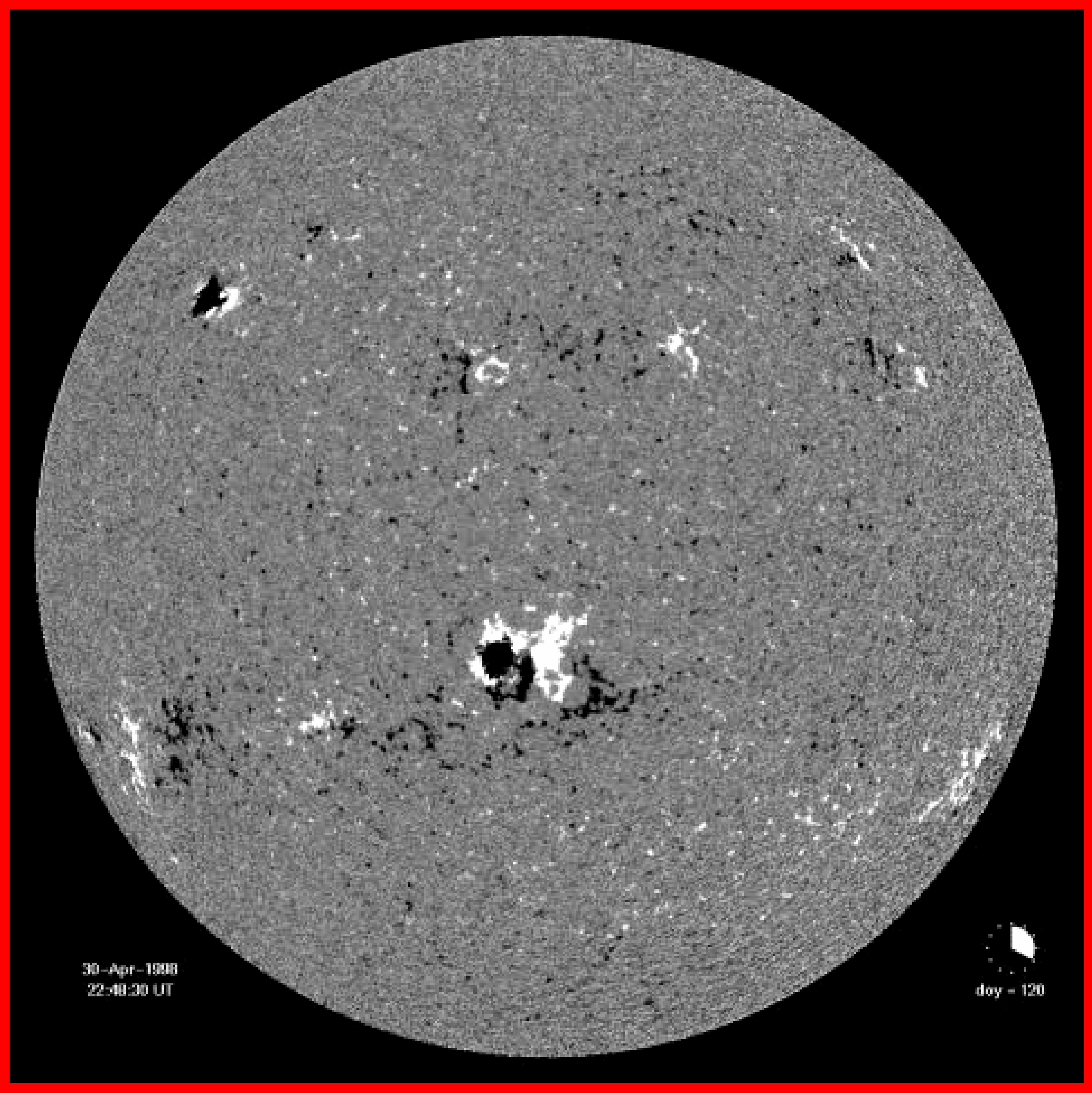
A small part of the total flux through the solar surface connects as open flux to interplanetary space

Open flux: fast solar wind

Measured Magnetic Field at Sun's Surface

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

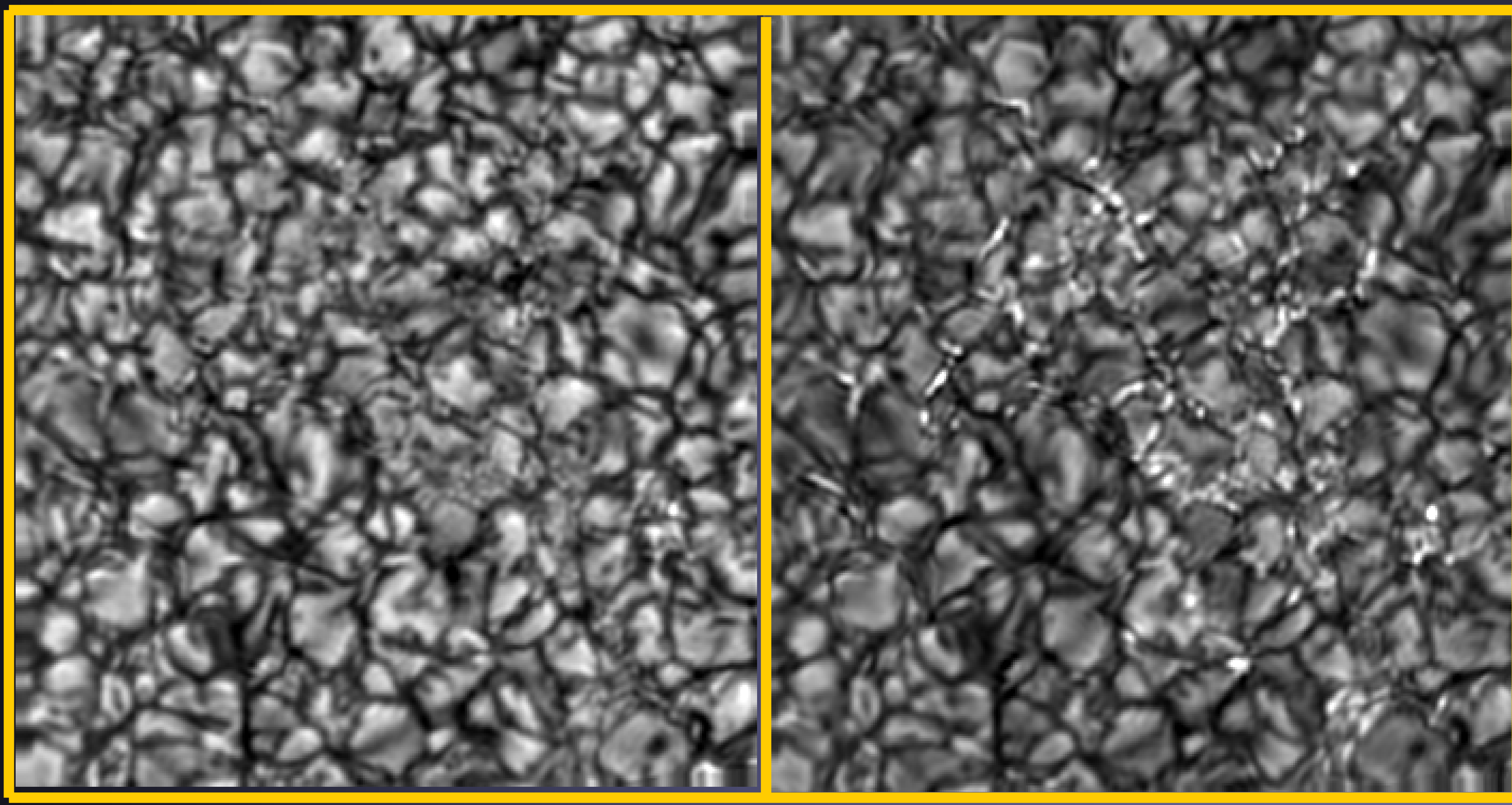
MDI/SOHO
May 1998



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

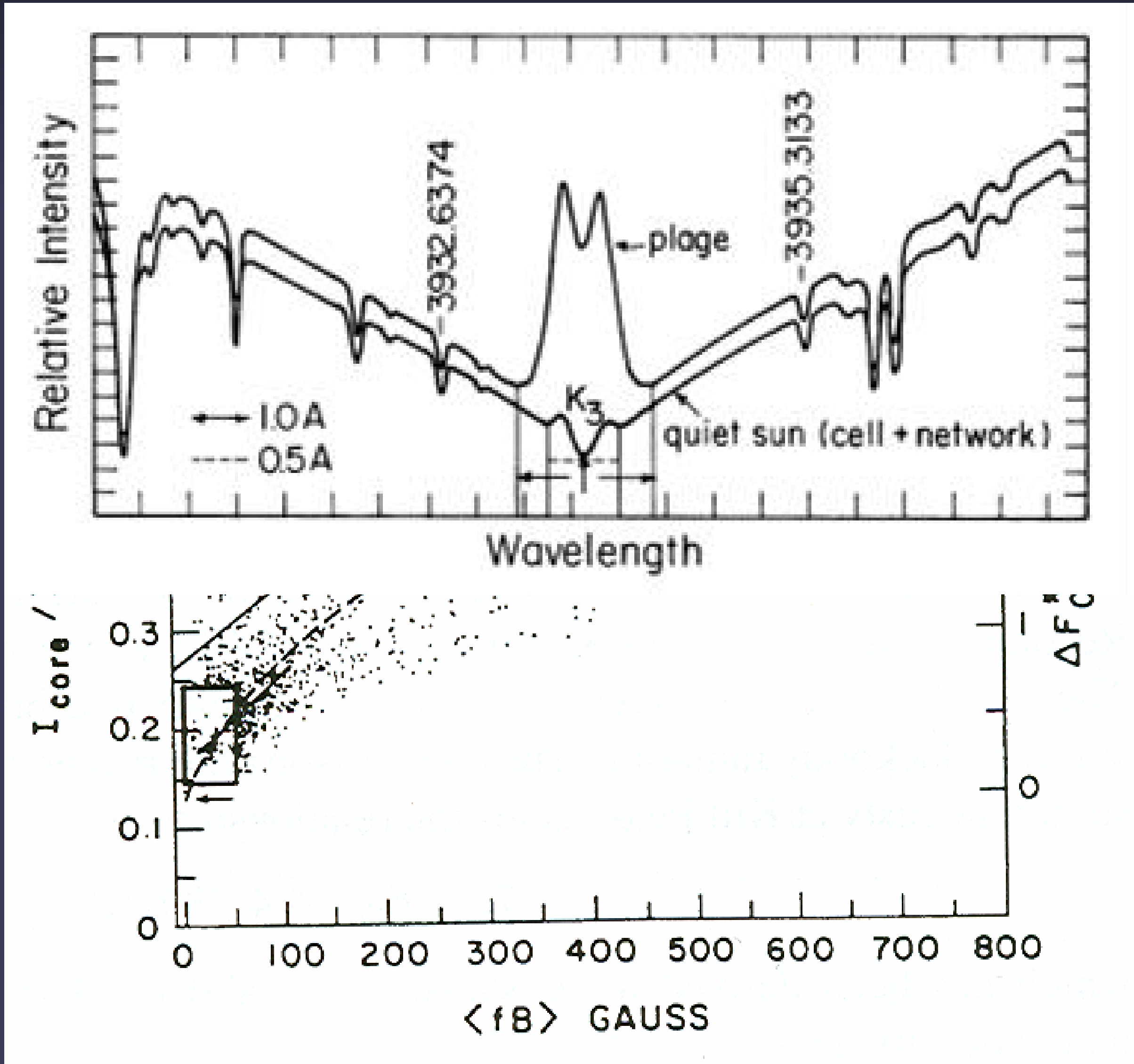
Example of proxies: Continuum vs. G-band



Continuum G-band

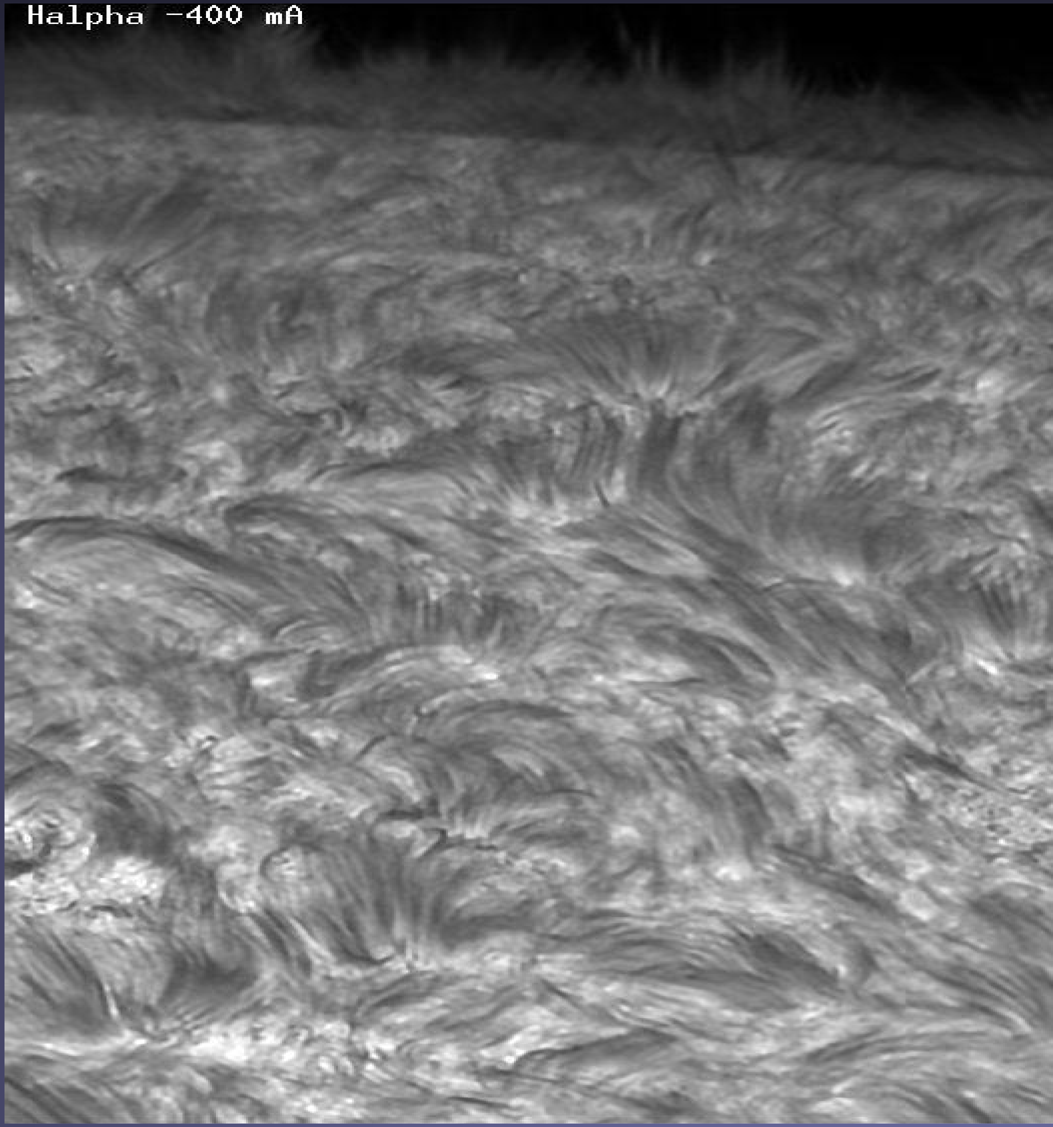
Ca II K as a magnetic field proxy

- Ca II H and K lines, the strongest lines in the visible solar spectrum, become brighter with non-spot magnetic flux.
- $I_{core}/I_{wing} \sim \langle B \rangle^{0.6}$
- Magnetic regions (except sunspots) appear bright in Ca II H+K → Ca plage and network regions



Hα and the chromospheric field

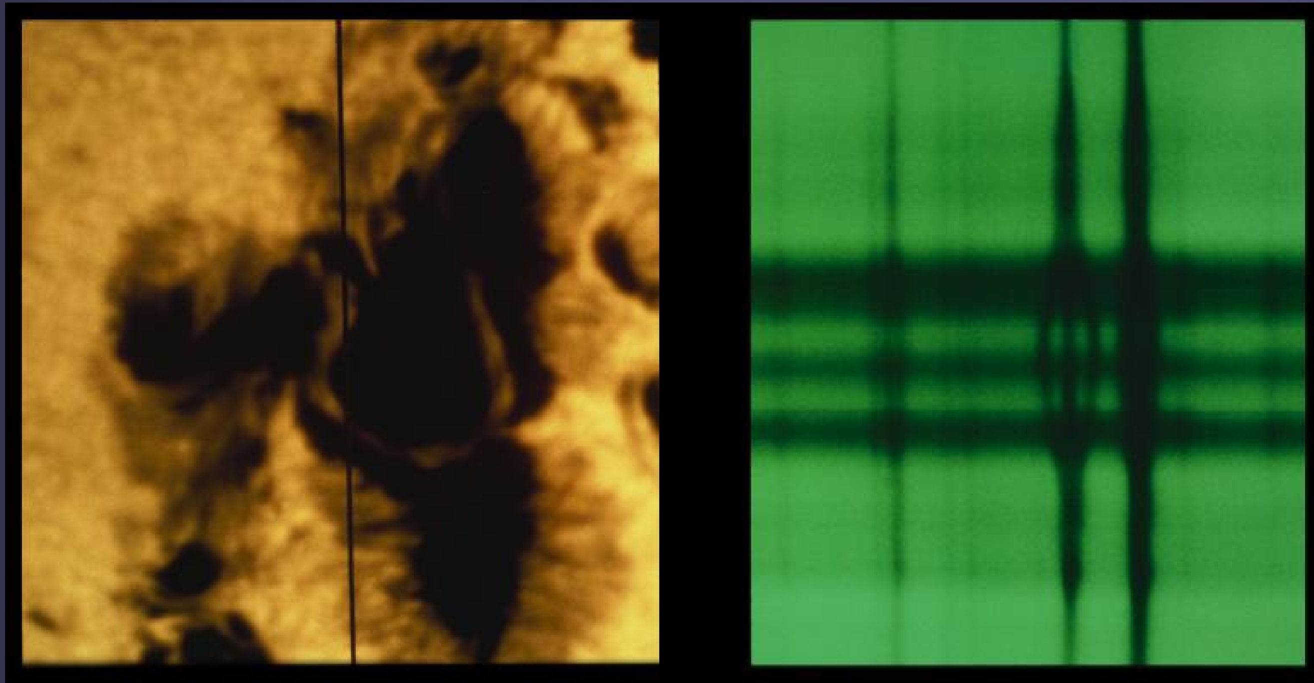
- Hα image of quiet Sun shows many loop-like structures. Do they (roughly) follow the field lines?
- Do they imply a relatively horizontal chromospheric field?



Zeeman diagnostics

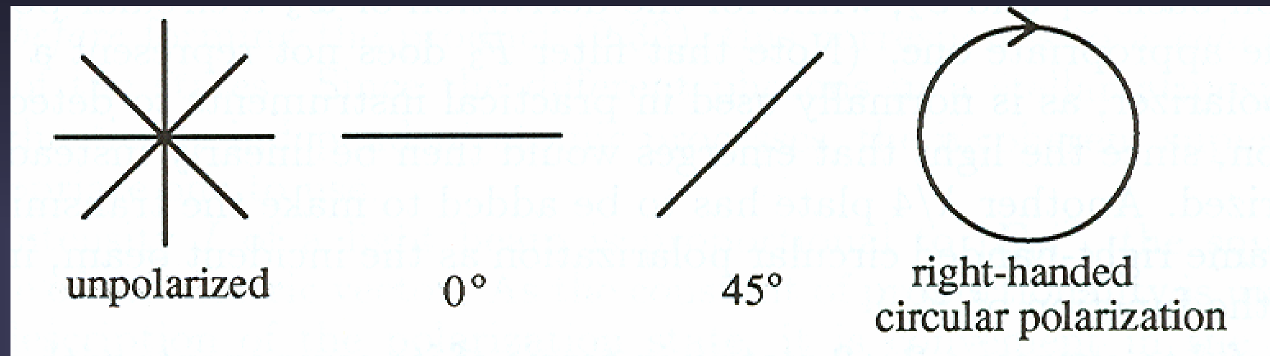
- Direct detection of magnetic field by obs. of magnetically induced splitting and polarisation of spectral lines
- Ob Sun: Zeeman effect changes not just spectral shape of a spectral line (often subtle and difficult to measure), but also introduces a unique polarisation signature

→ 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

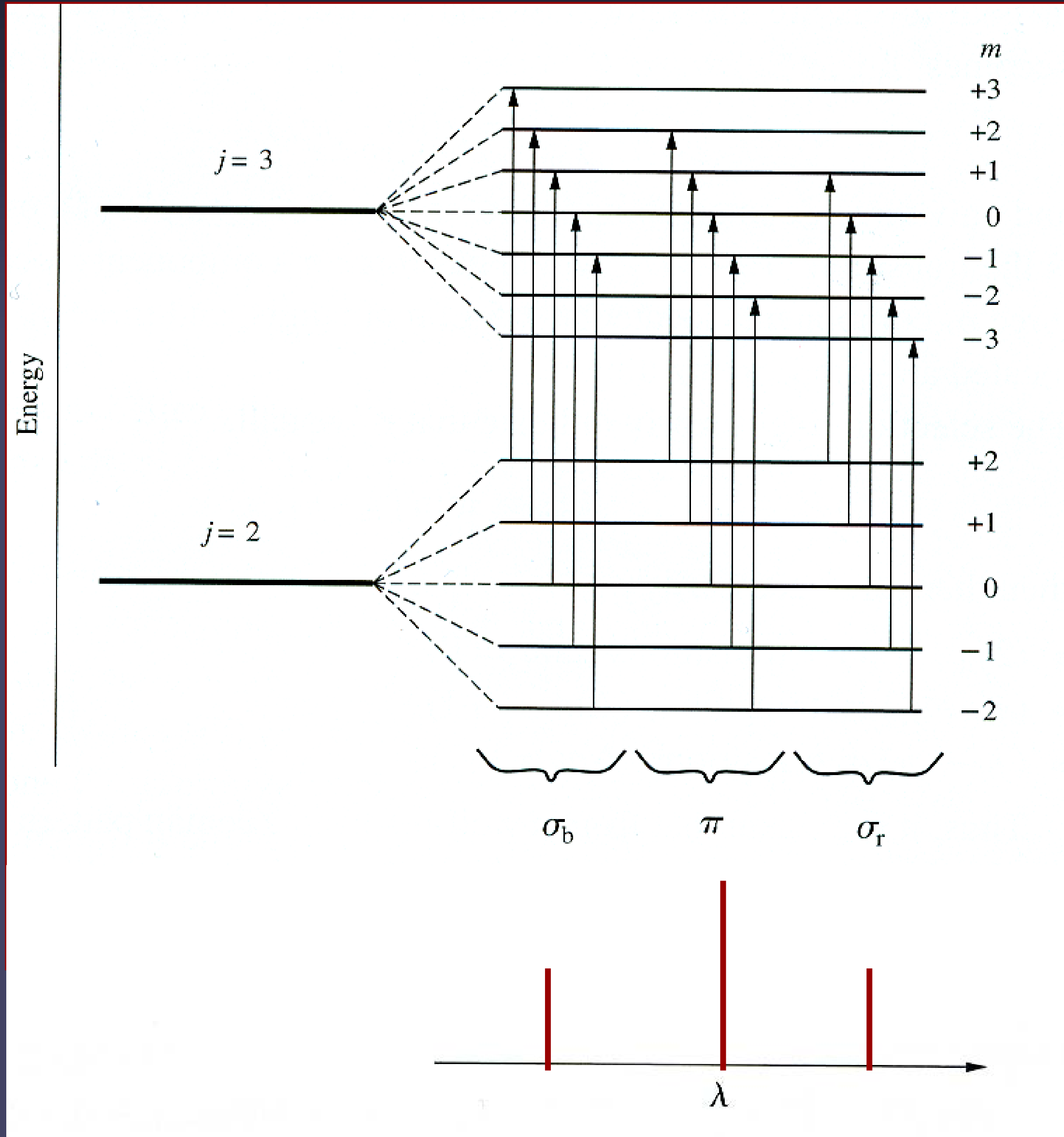


- I = total intensity $= I_{\mathrm{lin}}(0^{\mathrm{o}}) + I_{\mathrm{lin}}(90^{\mathrm{o}}) = I_{\mathrm{lin}}(45^{\mathrm{o}}) + I_{\mathrm{lin}}(135^{\mathrm{o}}) = I_{circ}(\mathrm{right}) + I_{\mathrm{circ}}(\mathrm{left})$
- $Q = I_{\mathrm{lin}}(0^{\mathrm{o}}) - I_{\mathrm{lin}}(90^{\mathrm{o}})$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

Zeeman splitting of atomic levels

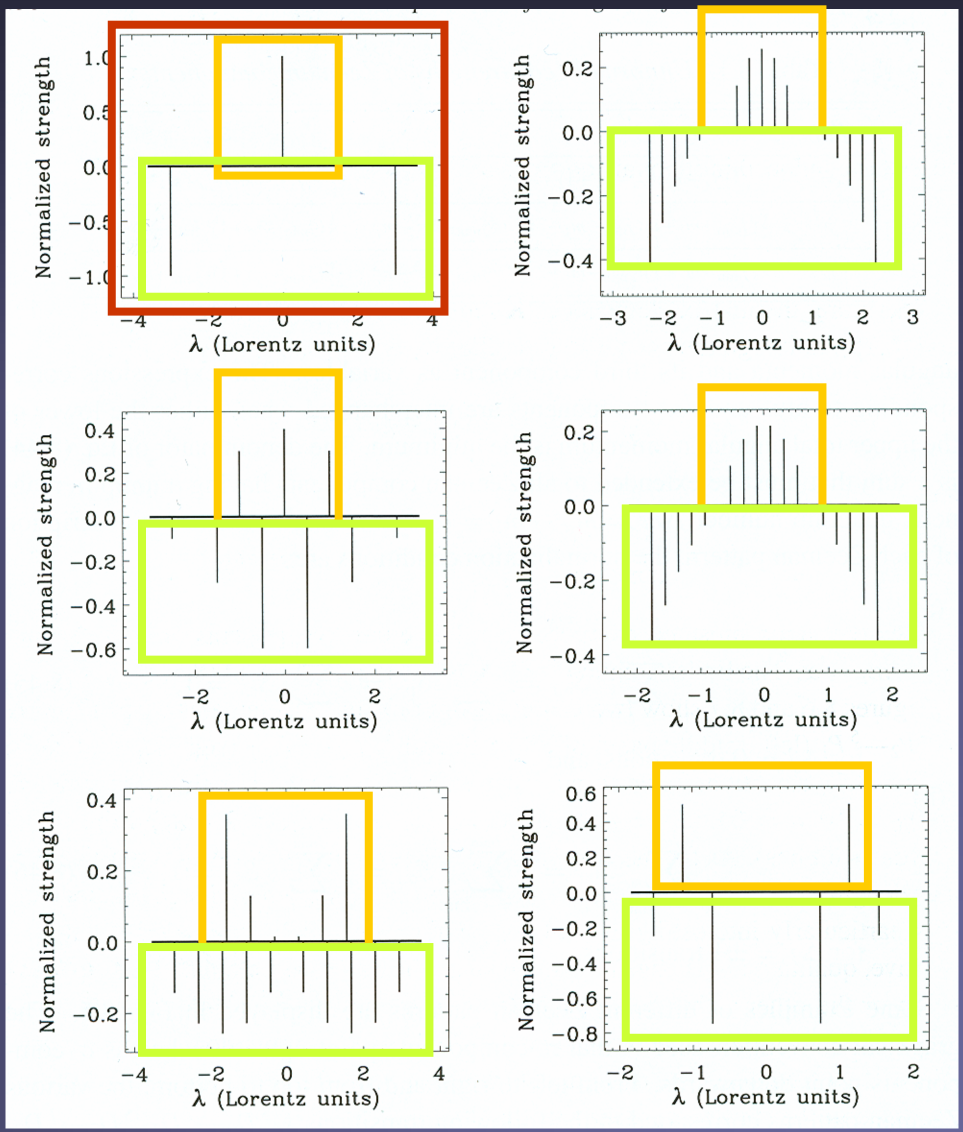
- In a B-field a level with total angular moment J splits into $2J+1$ sub-levels with different M.
- $E_{J,M} = E_J + \mu_0\, g M_J B$
- Transitions are allowed between levels with
  $\Delta J = 0, \pm 1$ &
  $\Delta M = 0$ (π), ± 1 (σ_b, σ_r)
- Splitting is determined by Lande factor g :

  $g(J,L,S) = 1+(J(J+1)+ S(S+1)-L(L+1))/2J(J+1)$

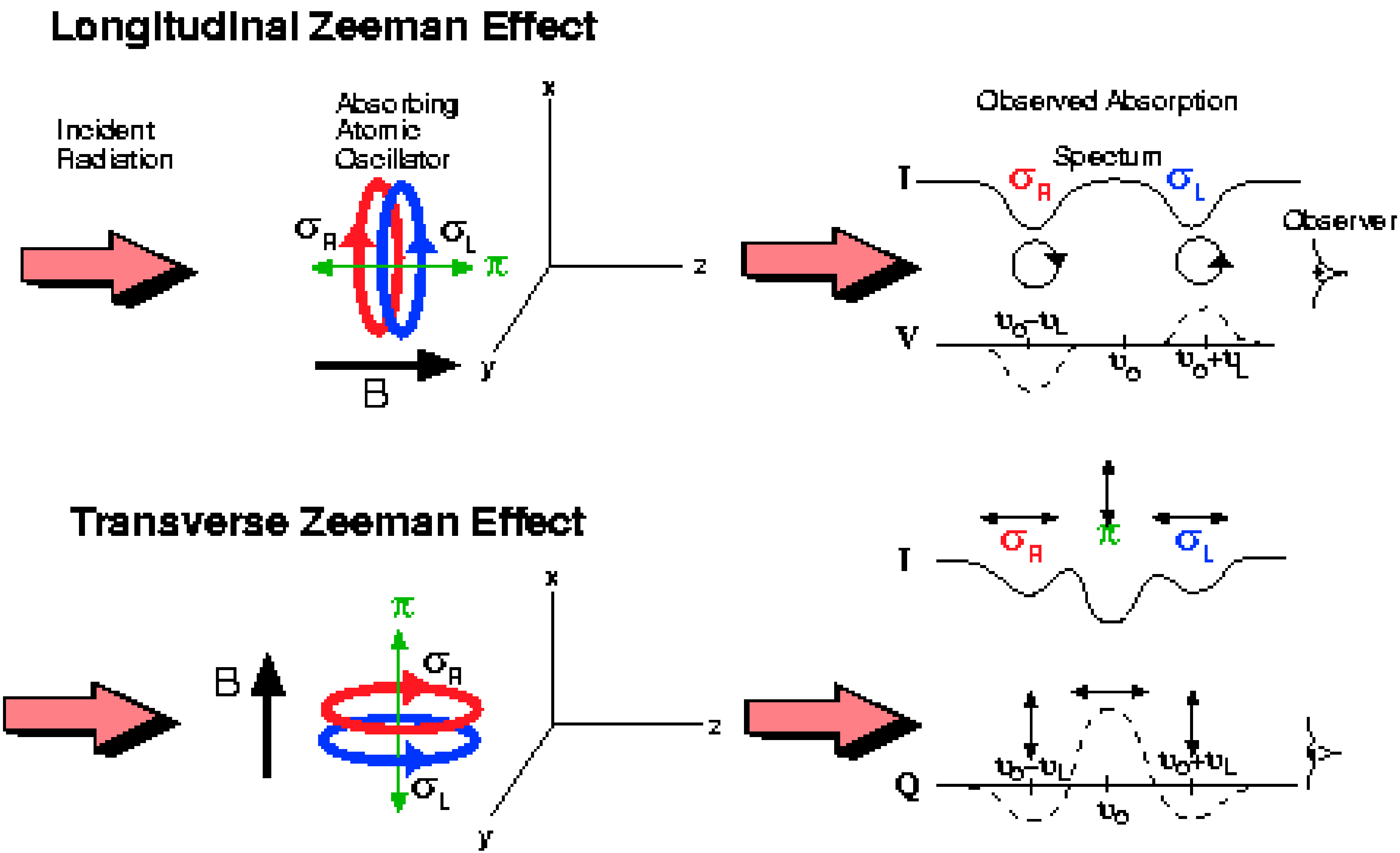


Splitting patterns of lines

- 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: $\Delta M=\pm 1$
- Top left: normal Zeeman effect (rare)
- Rest: anomalous Zeeman effect (usual)



Polarization and Zeeman effect



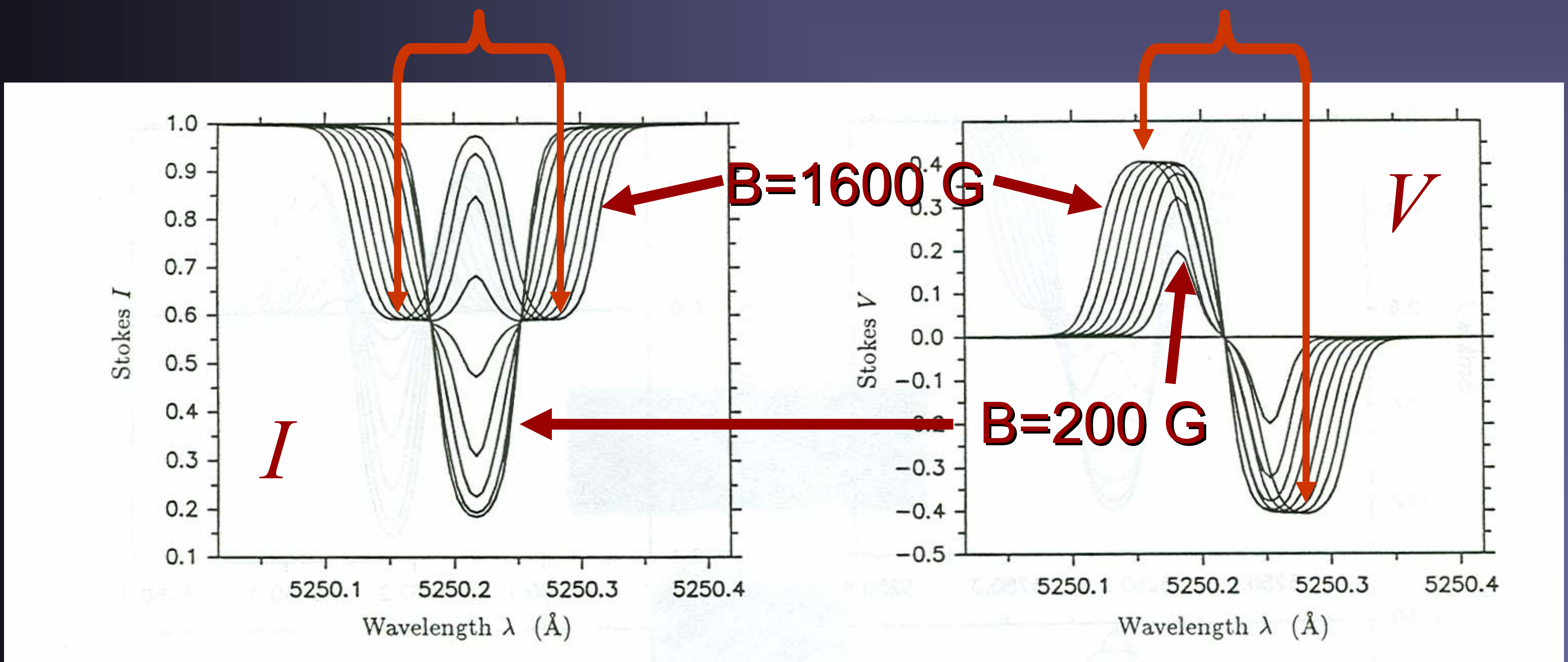
Effect of changing field strength

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

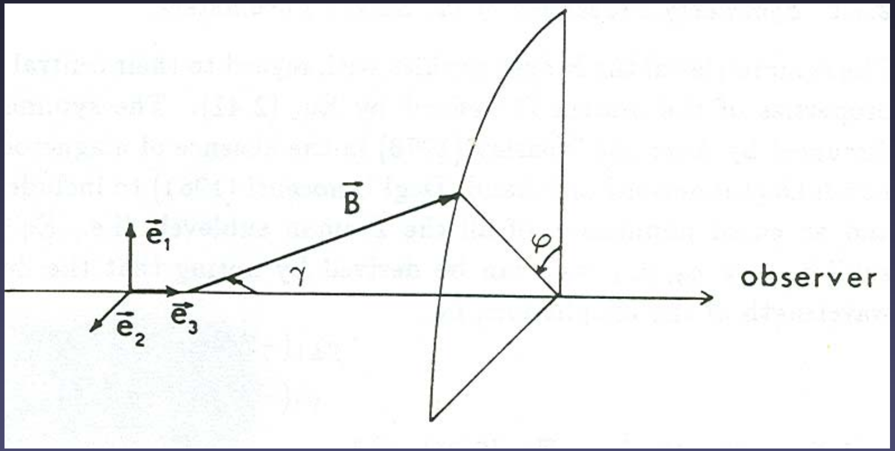
$$\Delta\lambda_H = 4.67\ 10^{-13}\ g_{eff} B \lambda^2 \quad [\text{Å}]$$

g_{eff}=effective Landé factor of line

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



Dependence on B, γ, and φ



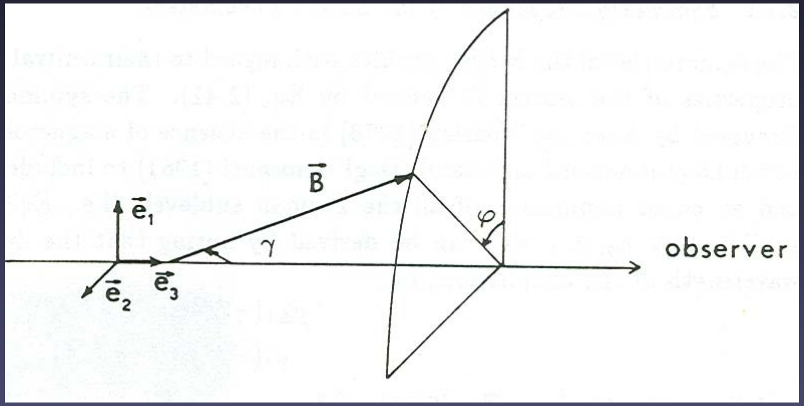
- $I \sim \kappa_\sigma(1+\cos^2\gamma)/4 + \kappa_\pi \sin^2\gamma/2$
- $Q \sim B^2 \sin^2\gamma \cos 2\varphi$
- $U \sim B^2 \sin^2\gamma \sin 2\varphi$
- $V \sim B \cos\gamma$

- V: longitudinal component of B
- Q, U: transverse component of B
- Above formulae for Q,U,V refer to relatively weak fields (e.g. B and B^2 dependence of field)
- Zeeman splitting etc. is hidden in κ_σ and κ_π. For Q, U, V these dependences have not been given for simplicity.

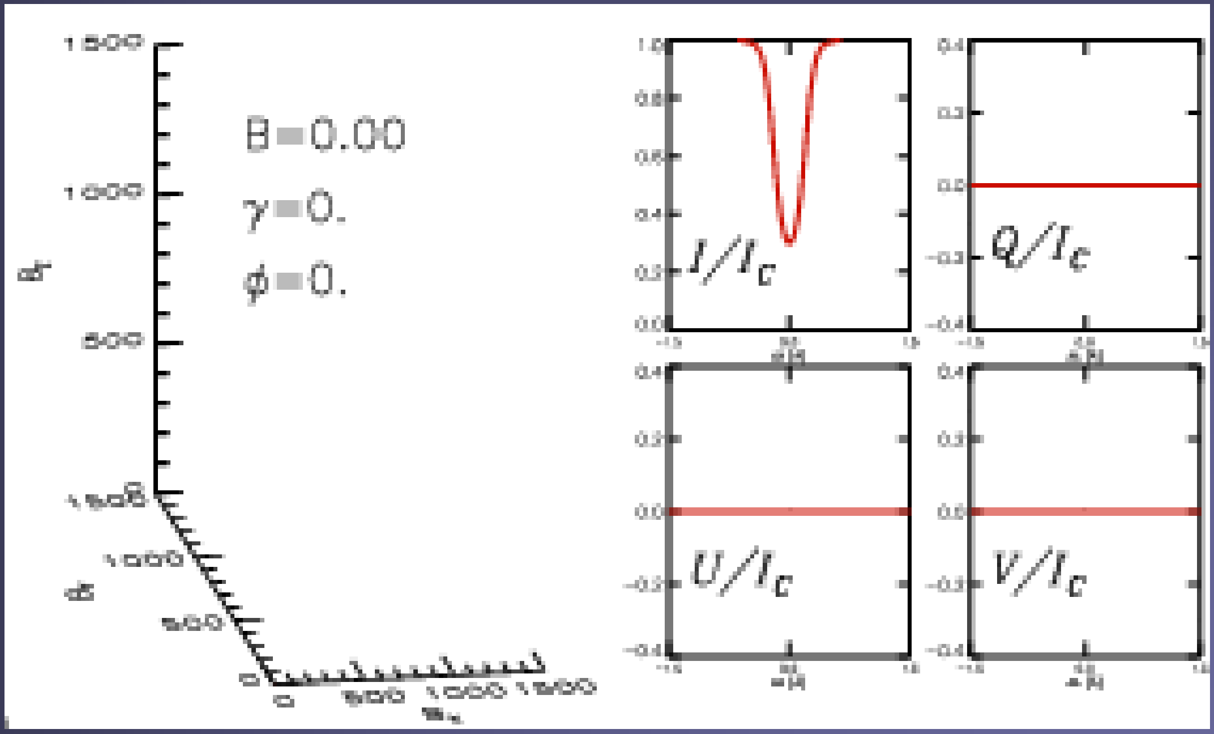
Juanma Borrero

Dependence on B, γ, and 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



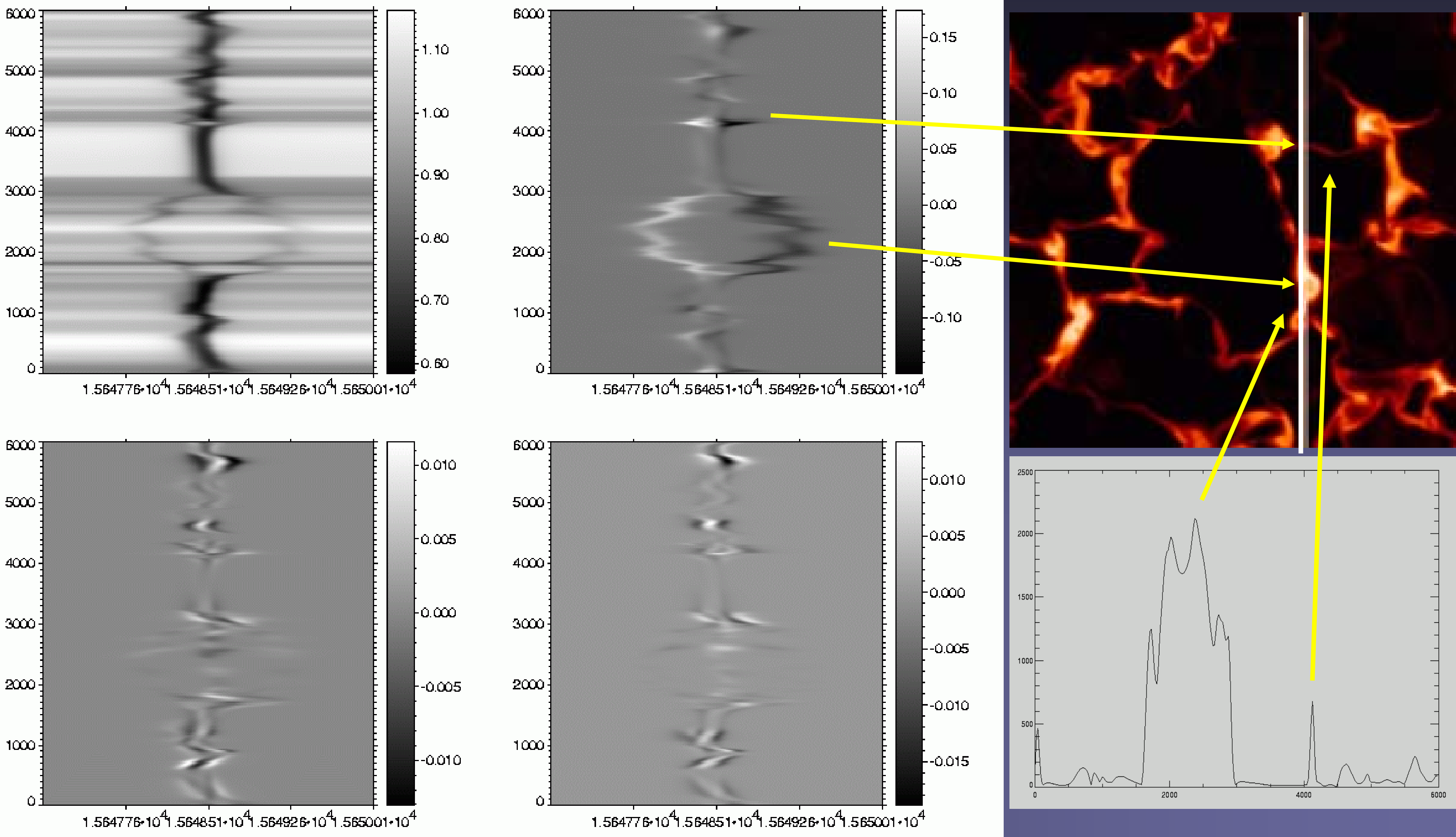
Juanma Borrero

Zeeman polarimetry

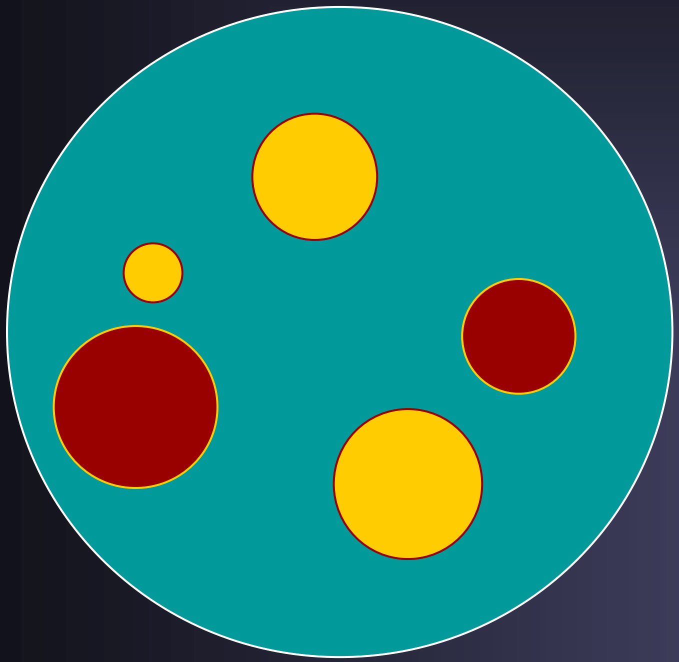
- Most widely used remote sensing technique of astrophysical (and certainly solar) magnetic fields
- Effective measurement of field strength if Zeeman splitting is comparable to Doppler width or more: $B > 200$ G … 1000 G (depending on spectral line) ➔ works best in photosphere and chromosphere
- Splitting scales with λ ➔ works best in IR
- Sensitive to cancellation of opposite magnetic polarities ➔ needs high spatial resolution (but usually small-scale magnetic structure is not resolved)

Zeeman splitting ~ λ^2

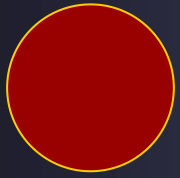
Fe I 1564.28nm



Cancellation of magnetic polarity

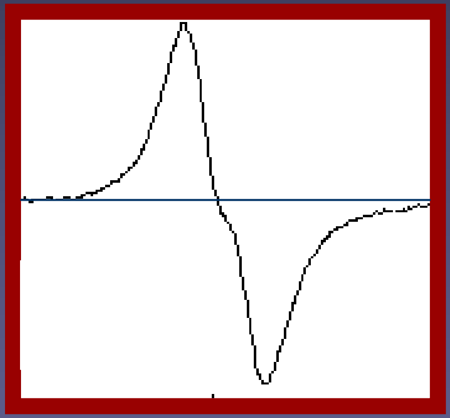


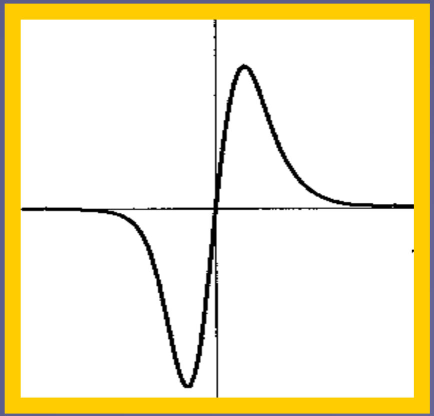
Spatial resolution element



= positive polarity magnetic field

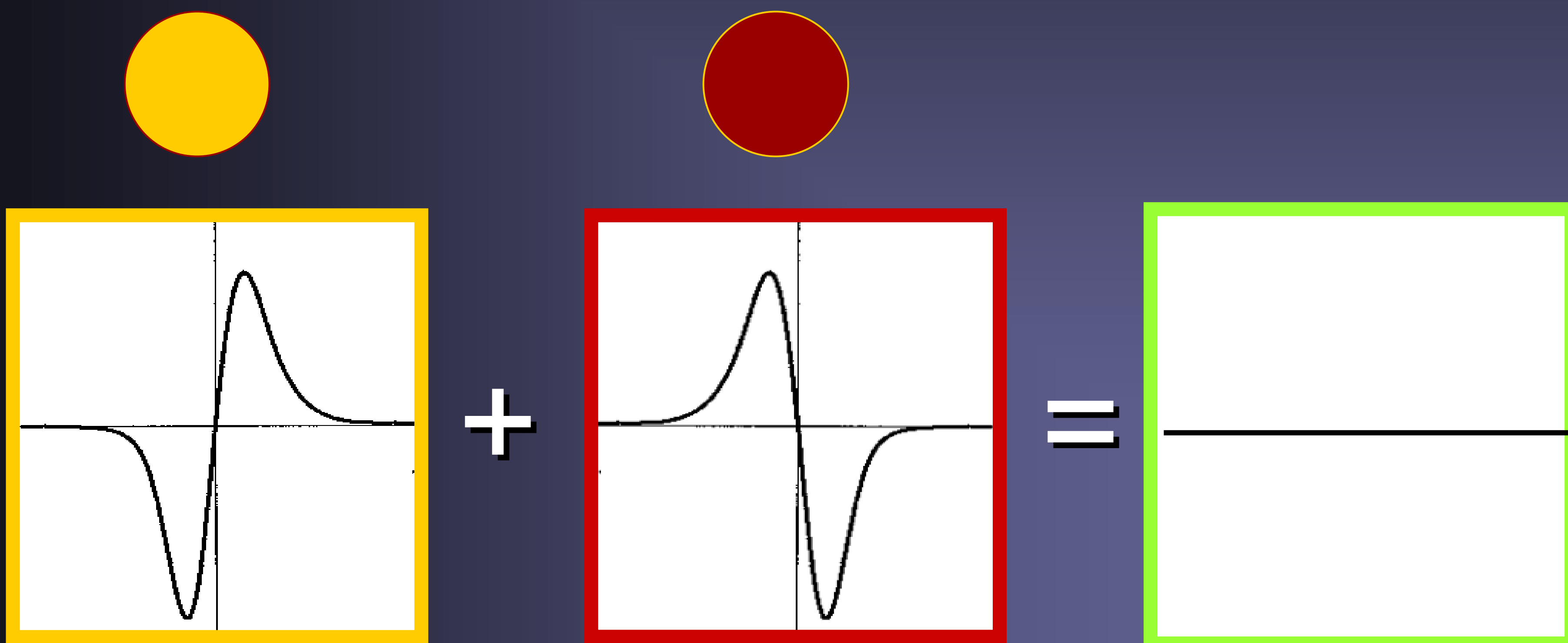
= negative polarity magnetic field





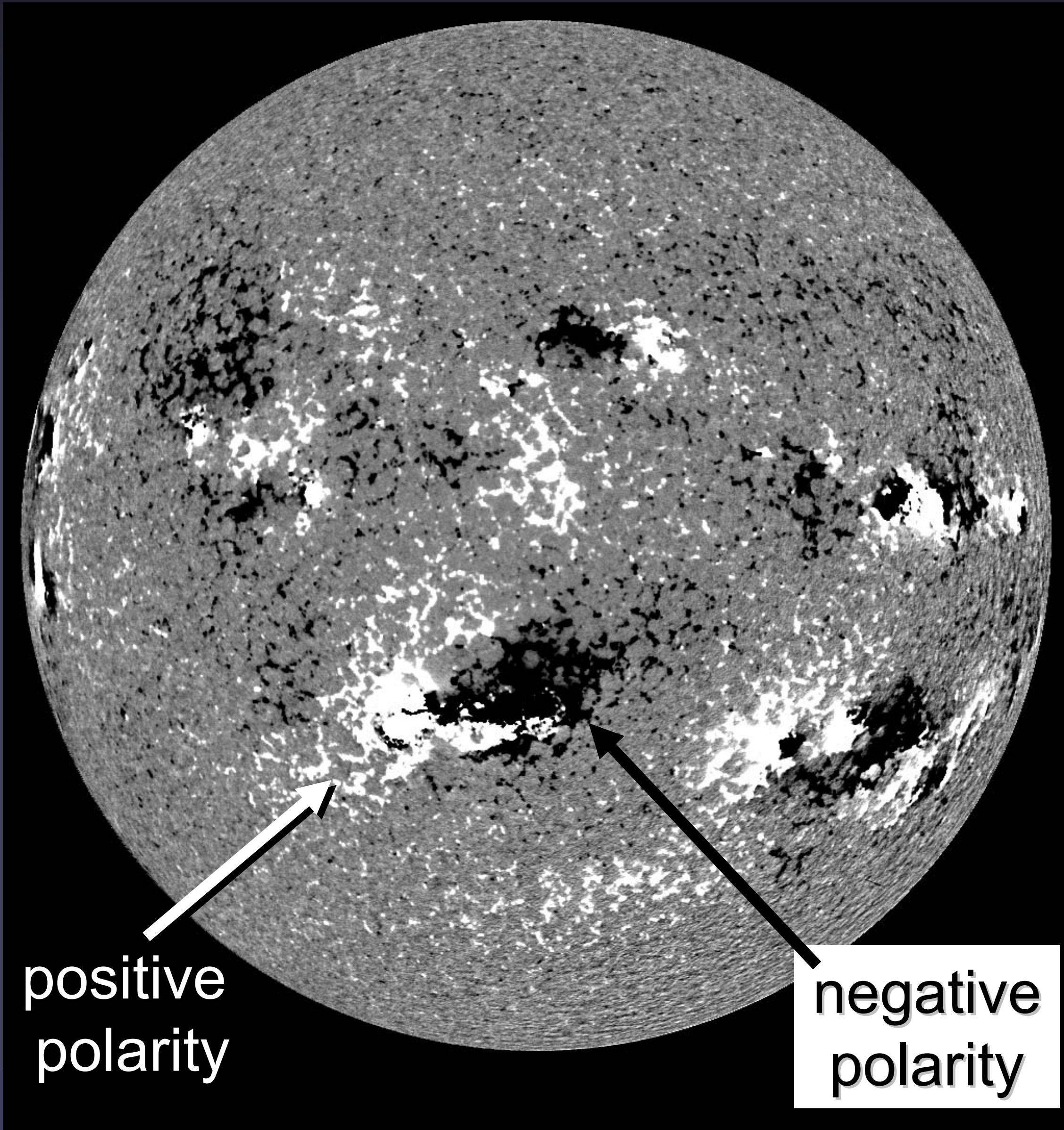
Stokes V signal cancellation

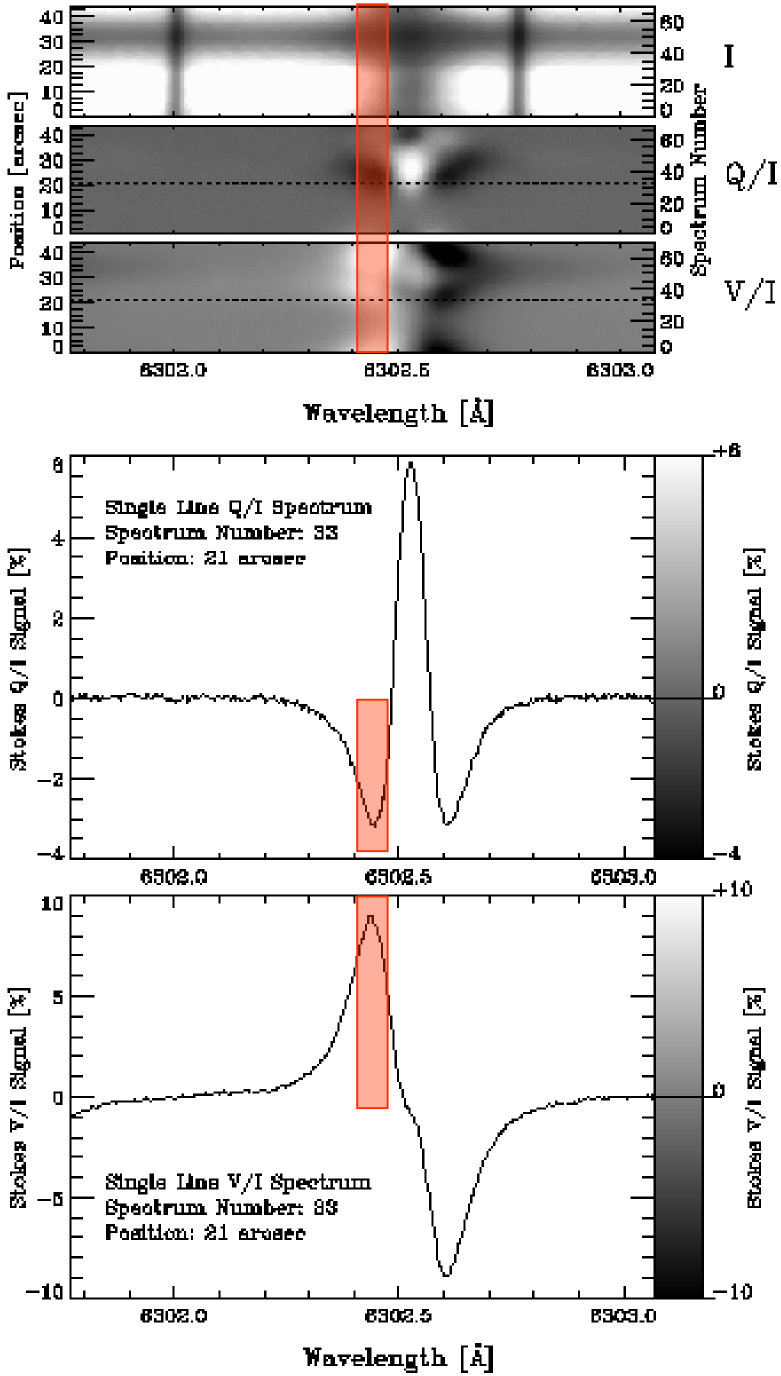
negative polarity positive polarity



Magnetograms

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

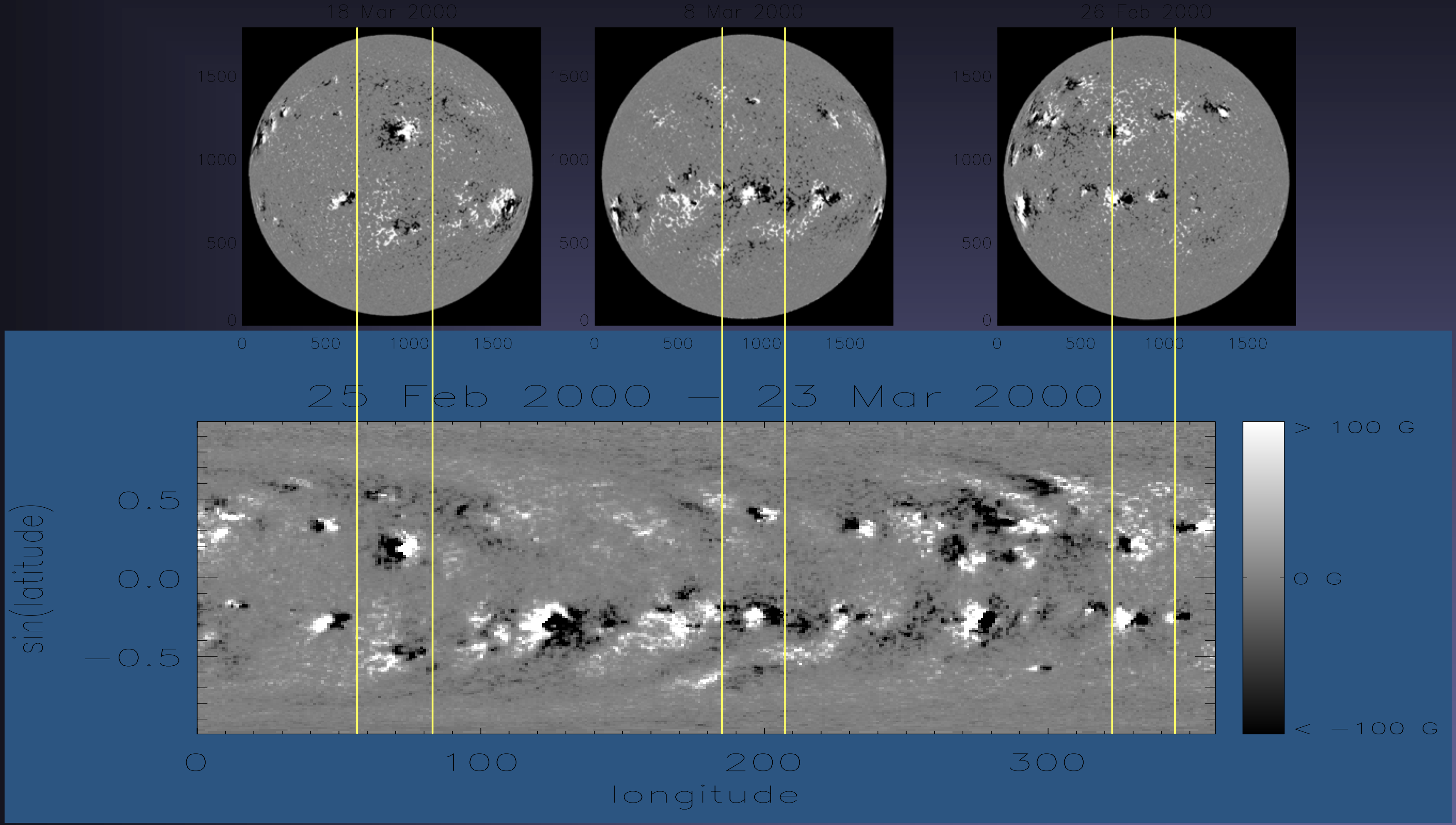




What does a magnetogram show?

- 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

- If light is polarized than a separ. RTE is required for each of 4 Stokes parameters: Written as differential equation for Stokes vector $\mathbf{I}_\nu = (I_\nu, Q_\nu, U_\nu, V_\nu)$
- 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

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

$\eta_I, \eta_Q, \eta_U, \eta_V$ are the line-to-continuum absorption ratios for each of the Stokes parameters, respectively ($\eta_I = \kappa_{IL}/\kappa_C$)

ρ_Q, ρ_U, ρ_V are magneto-optical coefficients for *Q, U* and *V*

Polarized radiative transfer III

- The Zeeman effect only enters through $\mathbf{\Omega}_\nu$
- $\mathbf{\Omega}_\nu$ contains
  - absorption due to Zeeman-split line $(\eta_I, \eta_Q, \eta_U, \eta_V)$. I.e. influence of B on absorption coefficient κ of medium
  - magnetooptical effects, (ρ_Q, ρ_U, ρ_V) . I.e. influence of B on refractive index n of medium
- Example magnetooptical effect: Faraday rotation, i.e. 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.: Unno solution

- General solution best done numerically (even formal solution is non-trivial: needs 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 .

MHD = Magnetohydrodynamics

- Here only a very few basics. See also special courses on MHD within the IMPRS curriculum
- MHD is an extension of hydrodynamics to a partly ionized gas (plasma) including the influence of a magnetic field. It is also a particular simplification of the particle-based (statistical mechanics) description of plasmas.
- The MHD equations are obtained by combining Maxwell's equations with equations of fluid mechanis (hydrodynamics).
- MHD is a good approximation in most of the Sun

MHD equations I

- Relevant Maxwell's equations, Ohm's law:

$$c\nabla\times\mathbf{B}=4\pi\mathbf{j},\qquad c\nabla\times\mathbf{E}=-\frac{\partial\mathbf{B}}{\partial t},\qquad \nabla\mathbf{B}=0$$

$$\mathbf{j}=\sigma(\mathbf{E}+\frac{1}{c}\mathbf{v}\times\mathbf{B}),\qquad \sigma=\text{electrical conductivity}$$

- Mass conservation:

$$\frac{\partial\rho}{\partial t}+\nabla\cdot(\rho\mathbf{v})=0$$

- j=current density, $\mathbf{E}$=electric field, $\mathbf{B}$=magnetic field, c=speed of light, $\mathbf{v}$=velocity, ρ= mass density

MHD equations II

- Force balance = equation of motion = momentum conservation

$$\rho \frac{\partial \mathbf{v}}{\partial t} + \rho(\mathbf{v}\nabla)\mathbf{v} = -\nabla P + \frac{1}{c}\mathbf{j}\times\mathbf{B} - \rho g + \rho \upsilon \Delta \mathbf{v}$$

- Equation of state (e.g. ideal gas law)

$$P = f(\rho, T)$$

- Energy equation (which takes on different forms in different parts of the Sun)

Terms of force balance equation

$\rho \dfrac{\partial \mathbf{v}}{\partial t} + \rho(\mathbf{v}\nabla)\mathbf{v}$	total derivative of impulse-dens
∇p	pressure gradient
$\dfrac{1}{c}\mathbf{j}\times\mathbf{B}$	magnetic forces
ρg	gravity
$\rho \upsilon \Delta \mathbf{v}$	viscous friction

Some simplifications

- Ideal MHD: no viscosity: $\rho\upsilon\Delta\mathbf{v} = 0$
- Stationary state: $\partial/\partial t = 0$
- Magnetohydrostatics: $\mathbf{v} = 0$
- No magnetic tension: $\frac{1}{4\pi}(\mathbf{B}\cdot\nabla)\mathbf{B} = 0$
- Force-free field: $\mathbf{j}\times\mathbf{B} = (\nabla\times\mathbf{B})\times\mathbf{B} = \mathbf{0}$
- Potential field: $\nabla\times\mathbf{B} = \mathbf{0}$
- Force-free field: valid when magnetic terms in force balance equation dominate over all the others.
- Potential field:valid when additionally there are no currents

MHD: 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, solar wind
  - $\beta < 1$ in atmosphere, particularly in corona $\beta << 1$

Plasma β vs. height in solar atmosphere

Field dominates Gas dominates

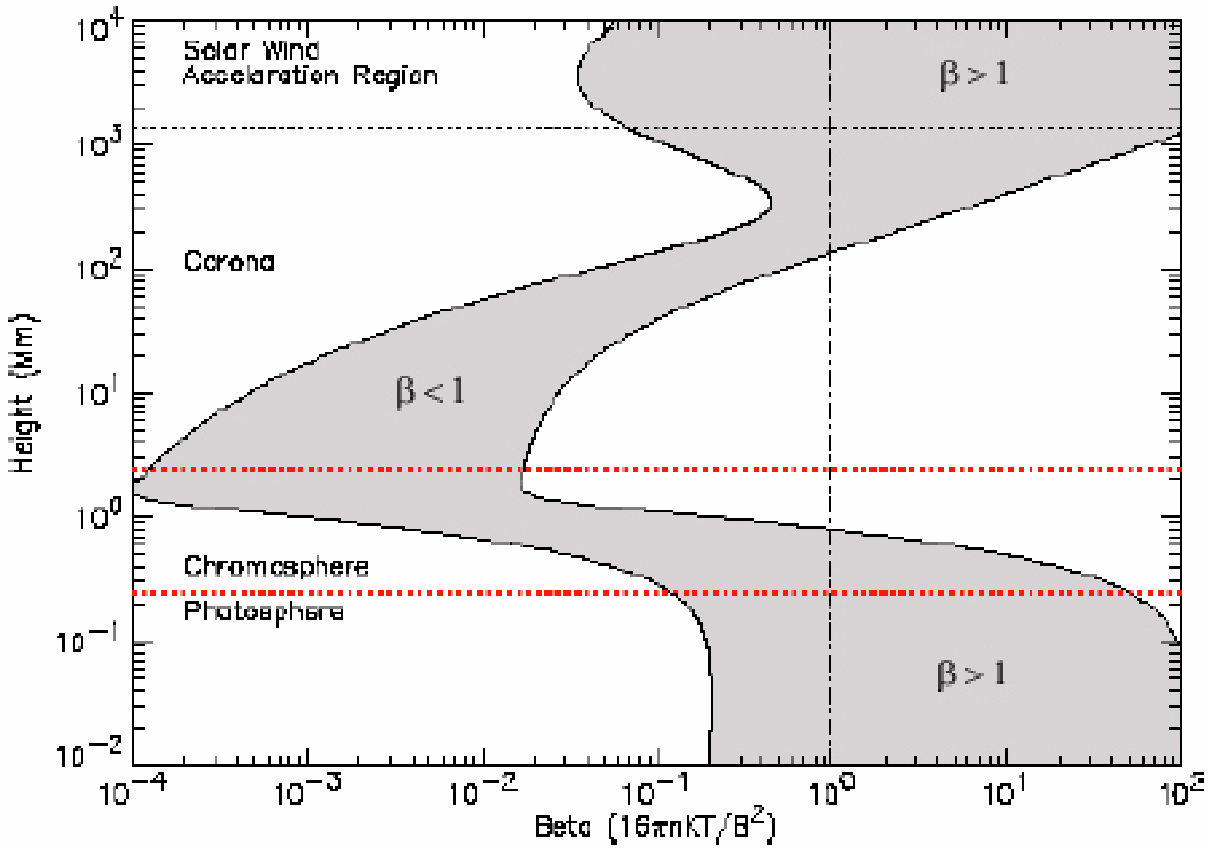


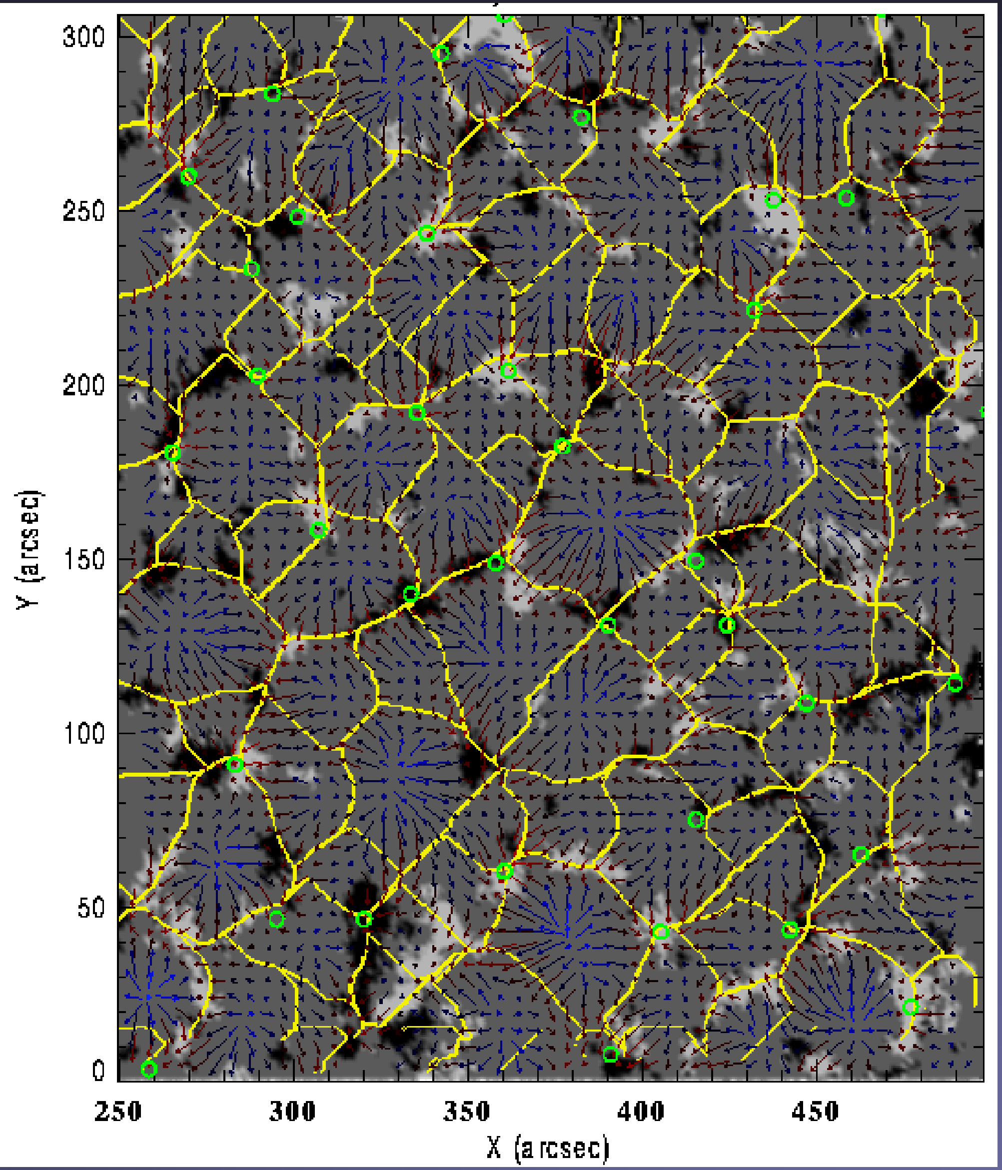
Figure 1. Ratio of gas pressure to magnetic pressure (β) as a function of height[1].

(Gary 2000)

Supergranules and magnetic field

- Magnetogram: black and white (oppos. polarities)
- 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 of supergranules



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)
- Valid even in photosphere of sunspots (only 10^{-4} of all particles are ionized), due to common collisions

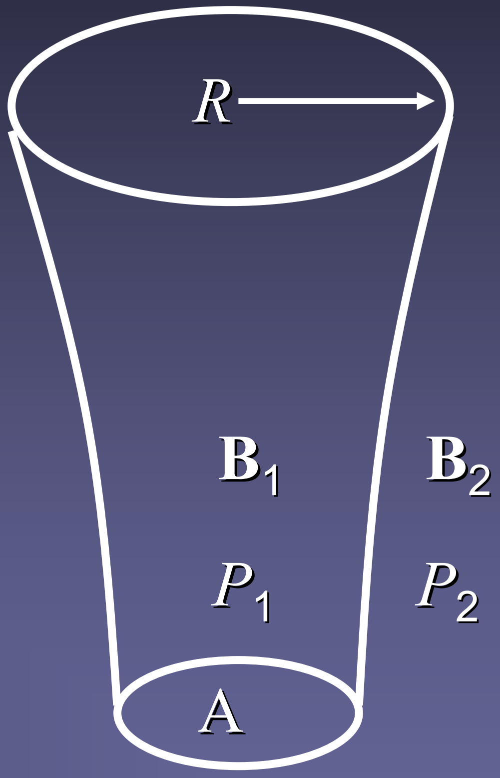
→ If plasma moves perpendicularly to B-field,

  - $\beta > 1$: it drags the field with it
  - $\beta < 1$: it is stopped by the field

- Flows parallel to the field are unaffected.

Magnetic flux tubes

- In convection zone and in photosphere most of magnetic energy is in concentrated magnetic flux tubesbundles of magnetic field (bounded by topologically simple surface)
- The flux tube has a current sheet at its boundary
- Consider a thin flux tube ($R<H_P$) that is homogeneous inside (no variation of B and P across cross-section)



Rump of a flux tube

Simplified force balance: pressure balance

- Consider a static, vertical, thin magnetic flux tube, FT (typical in lower solar atmosphere):
  - Interior of FT 1: field $\mathbf{B}_i$, pressure p_i, density ρ_i
  - Exterior of FT: field $\mathbf{B}_e$, pressure p_e, density ρ_e

➔ Vertical force balance is independent in the individual components.

  - For a vertical field (for each comp):

$$\nabla p - \rho g = 0$$

- Horizontal force balance between the components is then reduced to pressure balance.

Pressure balance

- pressure balance between thin flux tube interior i and exterior e

$$\frac{B_i^2}{8\pi} + P_i = P_e + \frac{B_e^2}{8\pi}$$

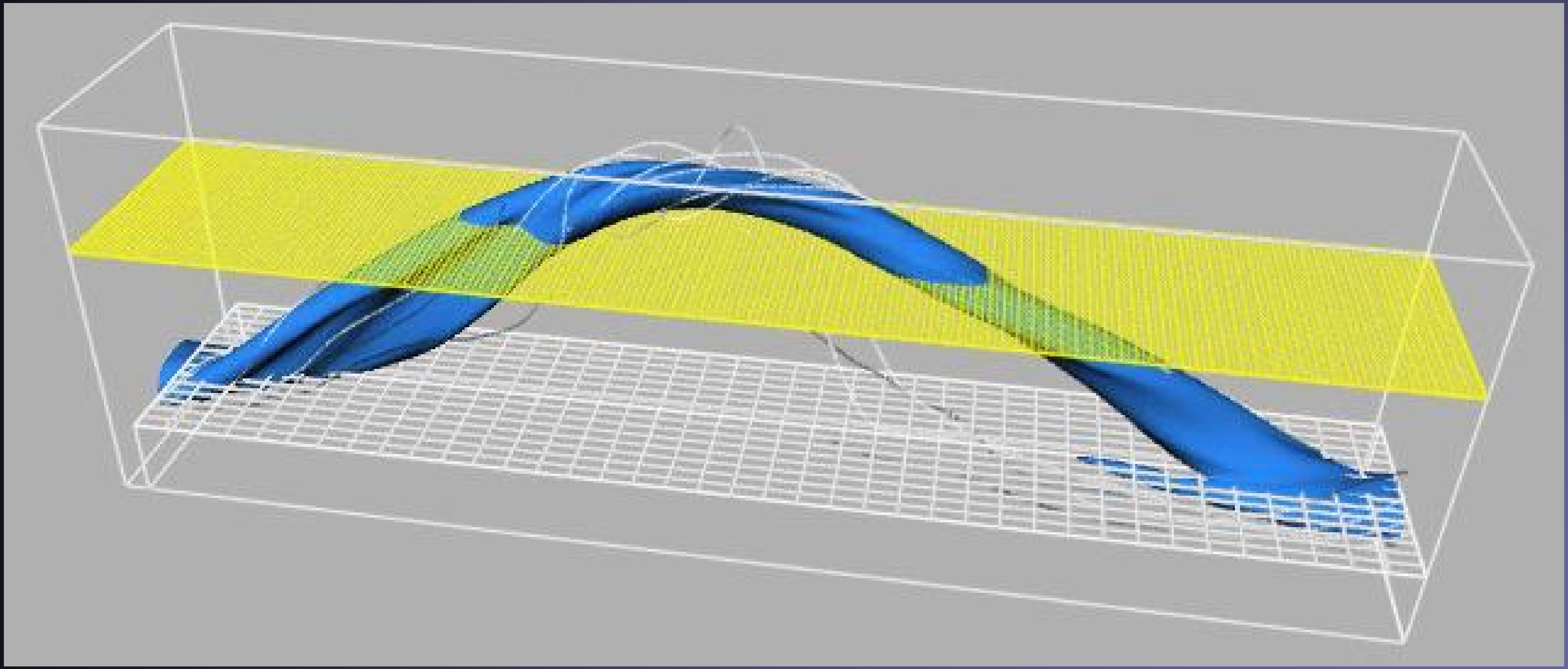
- If, e.g. $B_\mathrm{e} = 0$, then $P_\mathrm{i} < P_\mathrm{e}$ and it follows:

➔ Magnetic features are evacuated compared to surroundings.

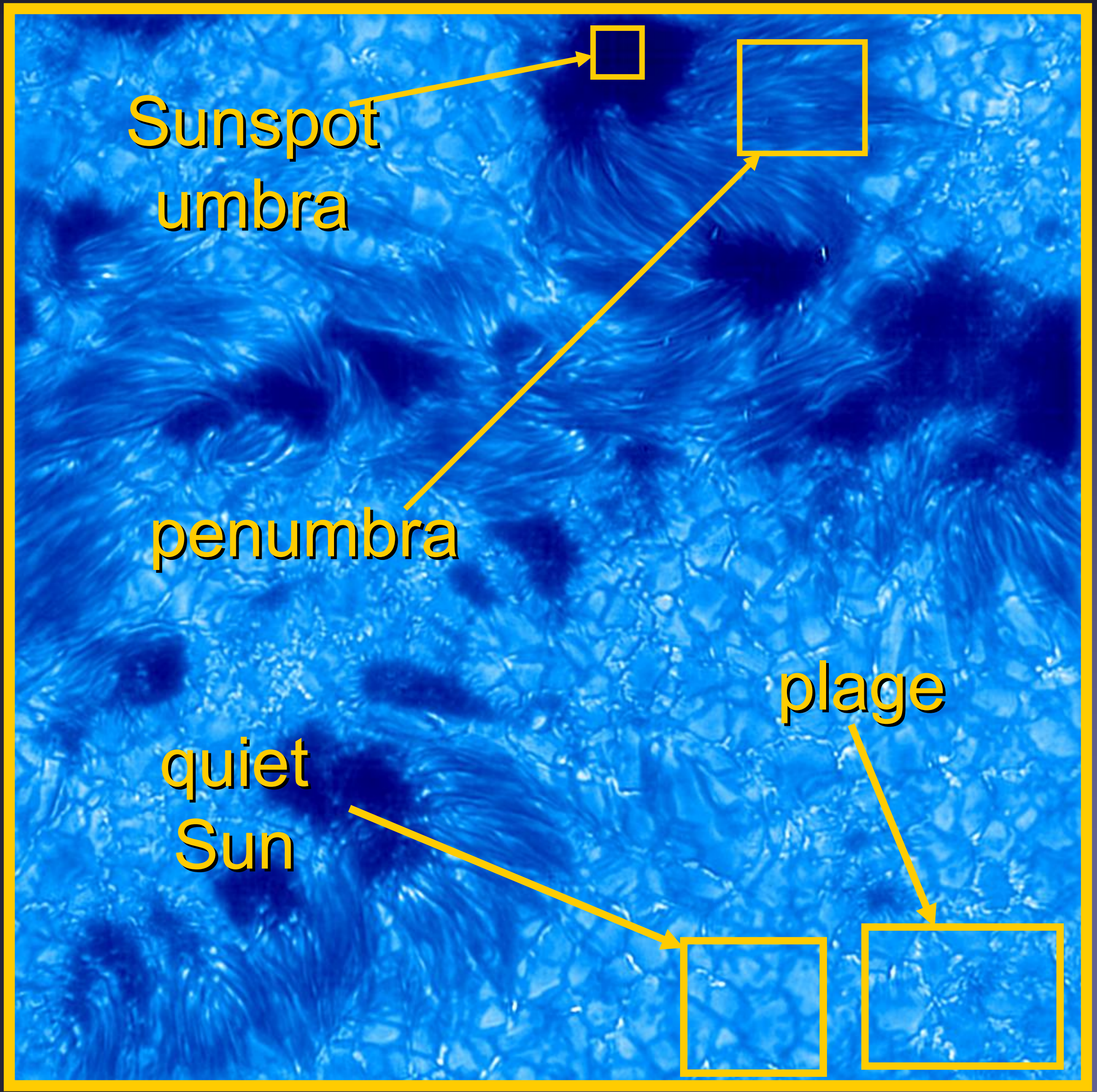
- If $B_\mathrm{r} = 0$ and $T_i = T_e$, then also $\rho_\mathrm{i} < \rho_\mathrm{e}$, so that the magnetic features are buoyant compared to the surrounding gas.
- In convection zone this buoyancy means that rising magnetic flux tubes keep rising (unless stopped by another force, e.g. magnetic curvature force) ➔ field cannot be stably stored in convection zone

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 next section) a magnetic field will rise towards the solar surface. At the solar surface it will produce a bipolar active region.



Photosphere: magnetoconvection



Sunspots

Umbra

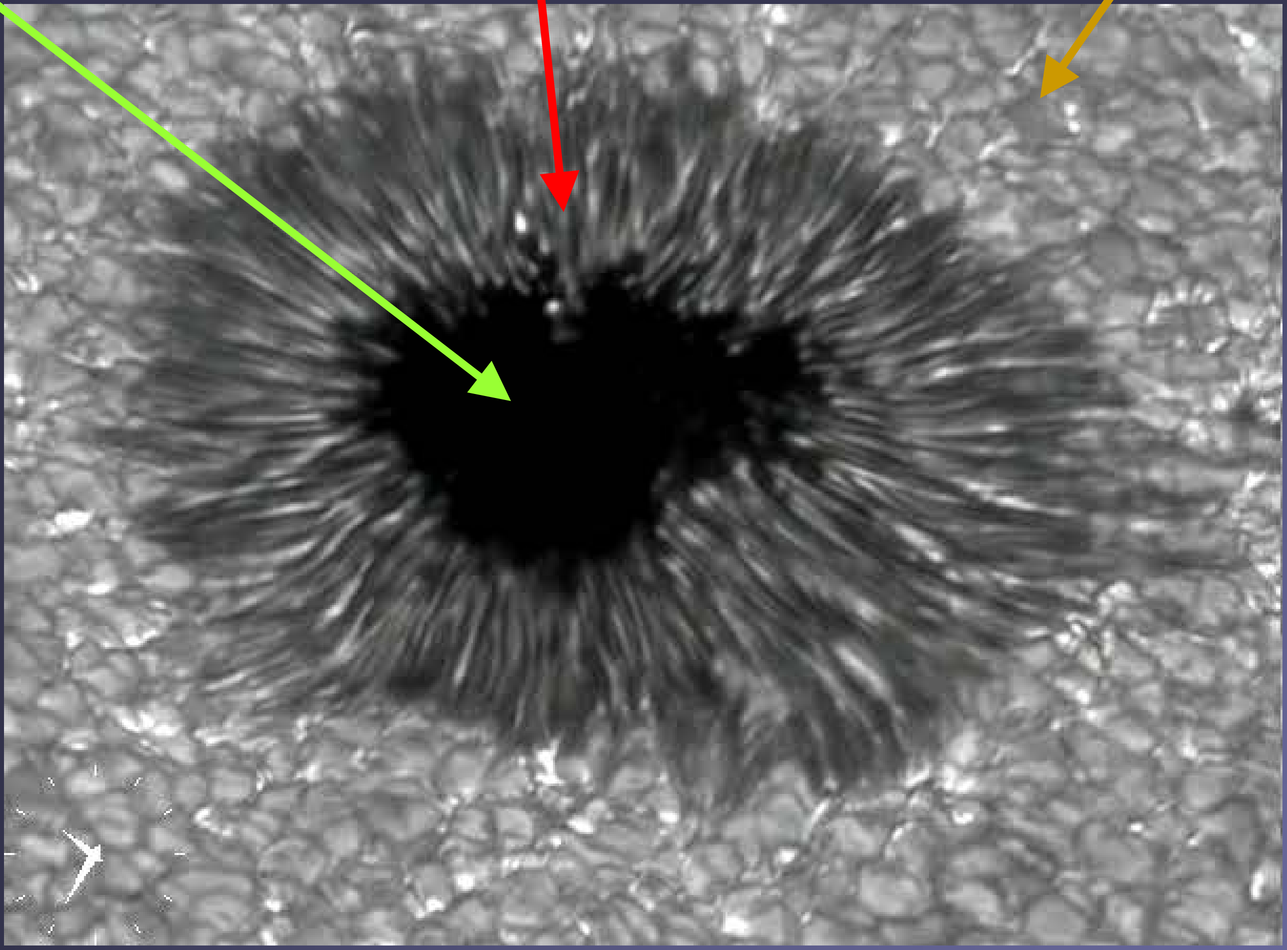
Penumbra

Granule

$T_{\mathrm{eff}} \approx 4500$ K

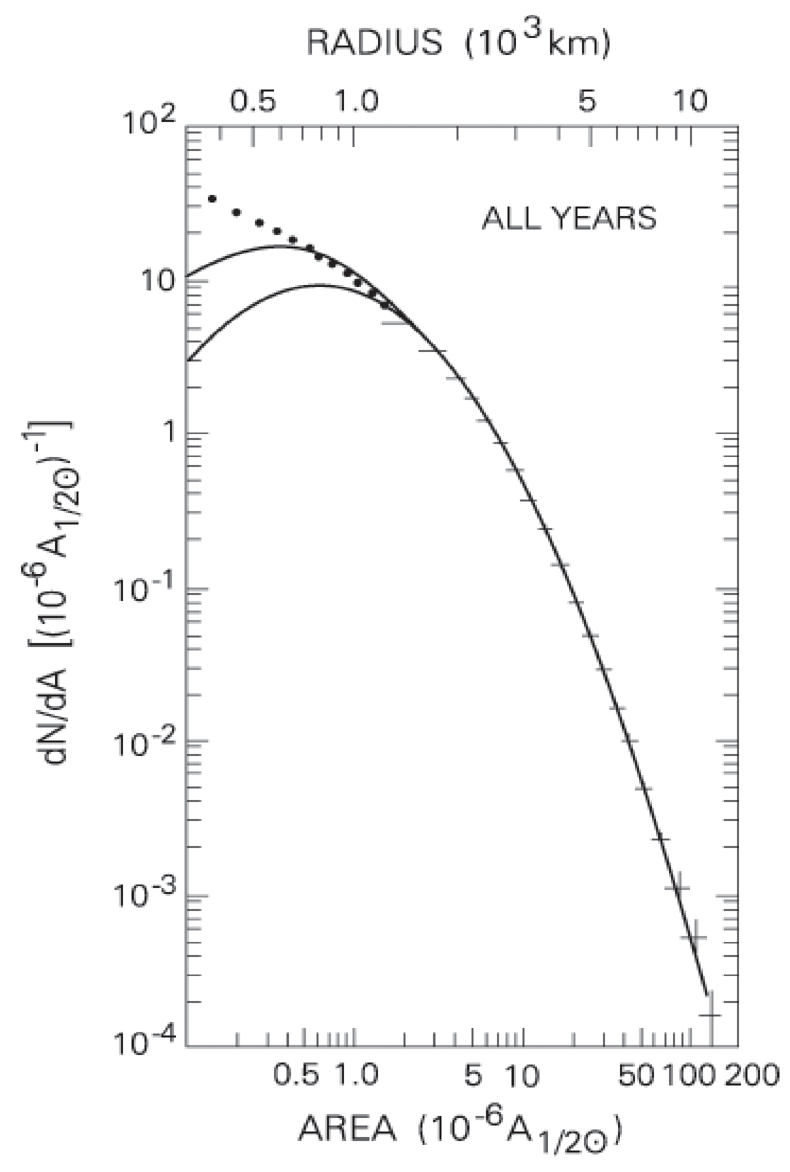
$T_{\mathrm{eff}} \approx 5500$ K

$T_{\mathrm{eff}} \approx 5800$ 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.



Lognormal distributions

- Normal distribution = Gaussian distribution
- Log-normal distribution = Gaussian distribution if plotted on the logarithm of the x-axis (params: σ, μ)

$$f(x) = \frac{1}{x\sigma\sqrt{2\pi}} \exp\left(-\frac{1}{2\sigma^2}\left(\log(x) - \mu^2\right)\right)$$

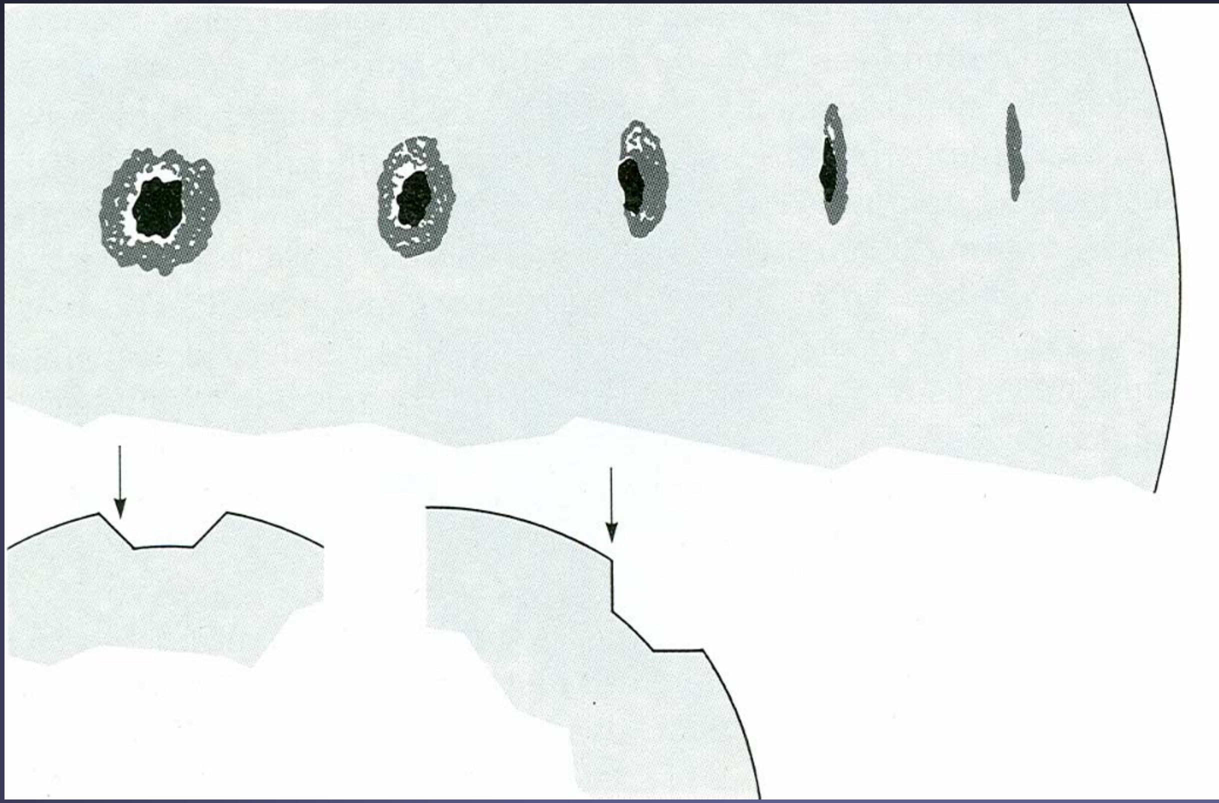
- Lognormal distributions are found in many natural and man-made systems, particularly when values cannot be negative, means, μ, are low and variances, σ, large.
- E.g. lengths of latent periods of infectious diseases, file sizes on disks, concentration of chemicals in honey samples, sunspot areas, solar UV radiances, ….
- Gaussians are due to additive processes, while lognormals are often produced by multiplicative processes, which are additive on a log-scale.

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

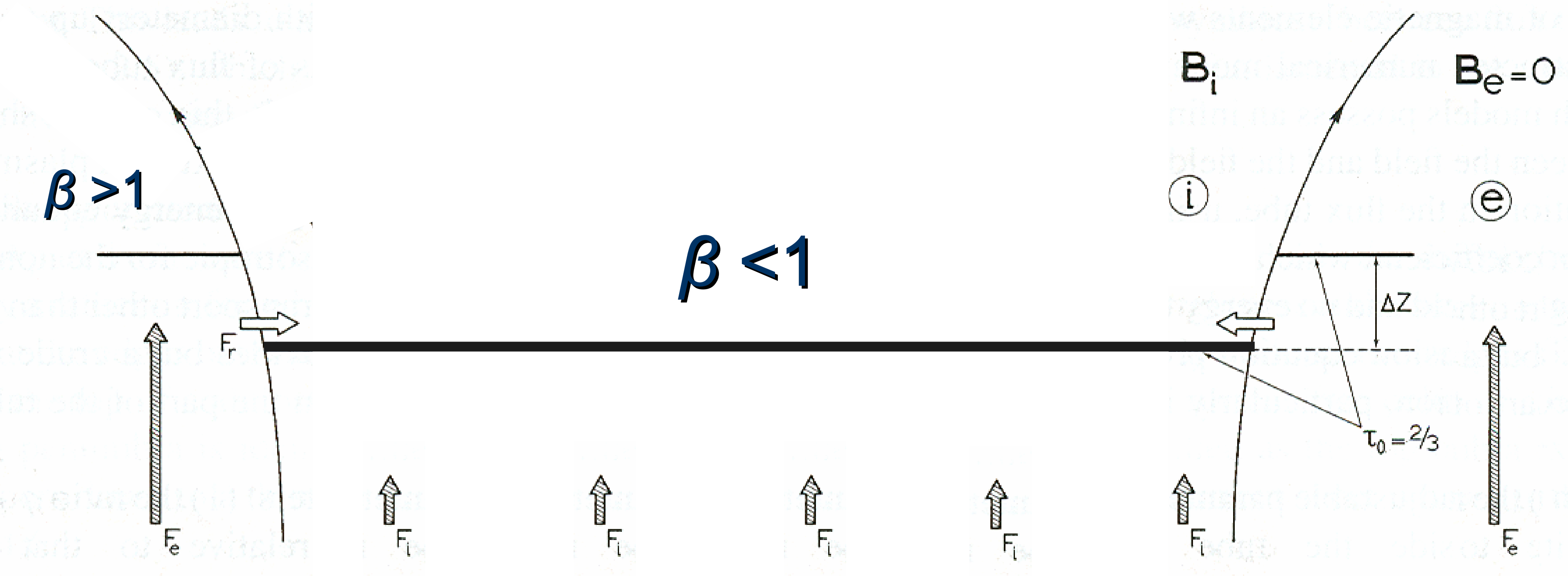
- Correct interpretation by Wilson (18^{th} century).



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.

Why are sunspots dark?

- Basically the strong nearlly 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 convection zone (CZ) gas is very large

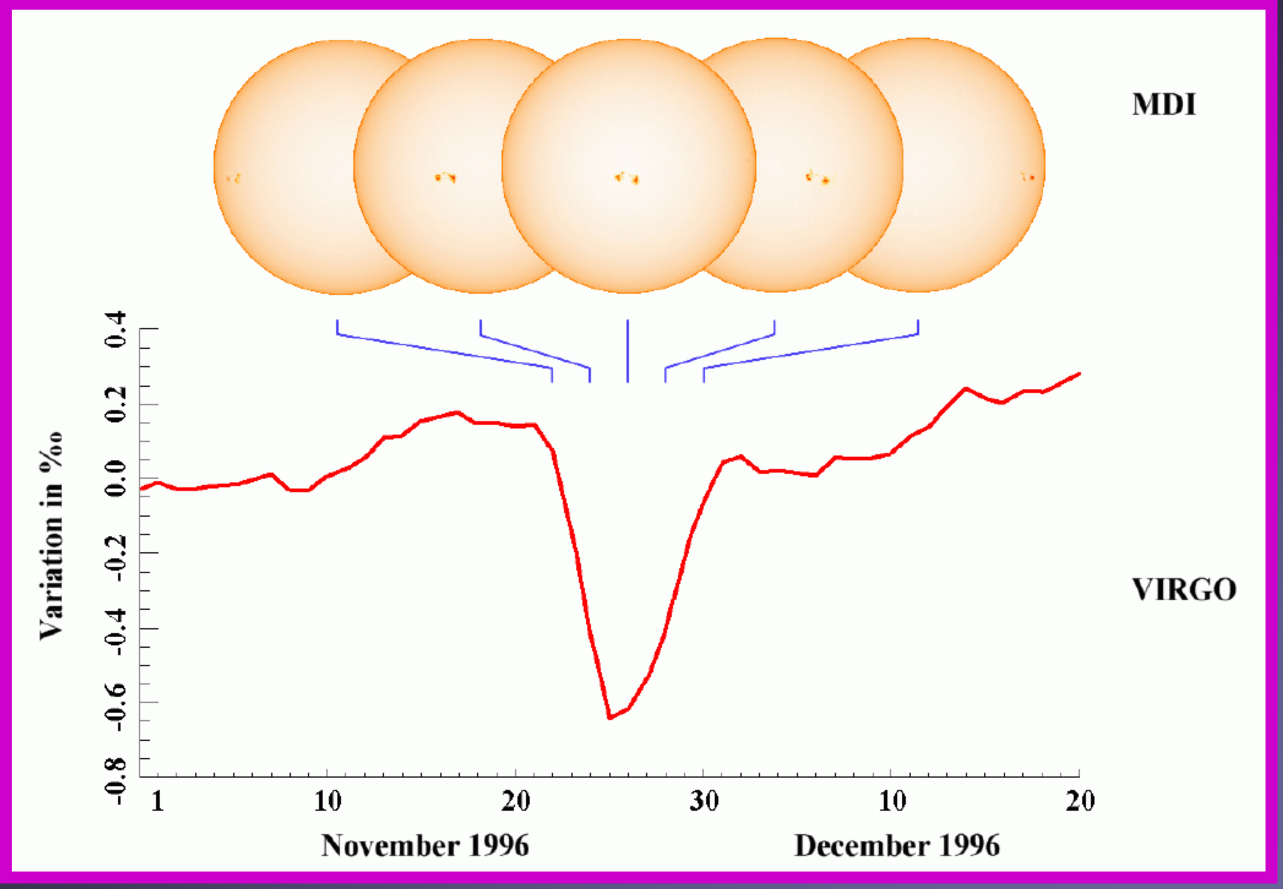
➔ High thermal conductivity: blocked heat is redistributed throughout CZ (no or only very weak bright rings around sunspots)

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.

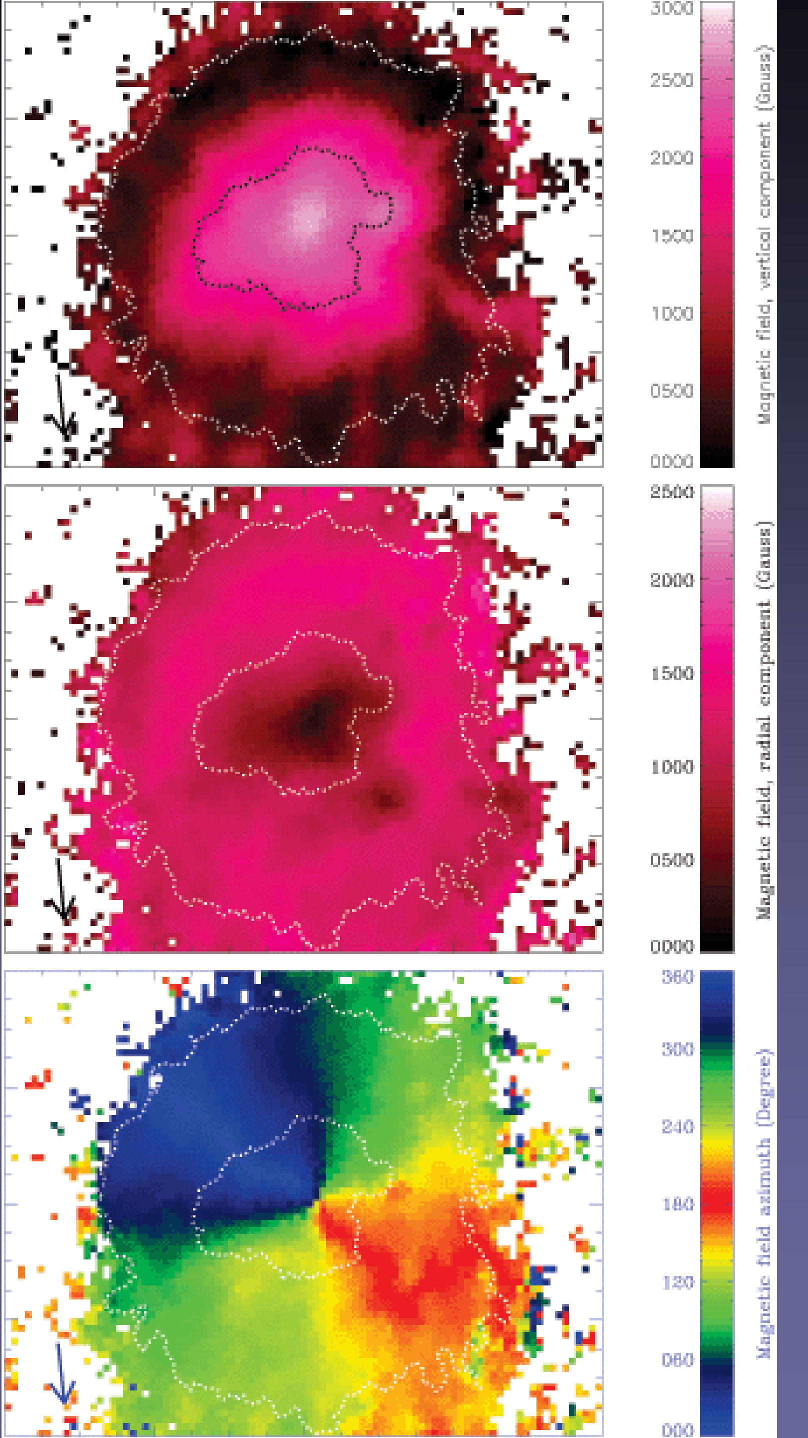
Solar irradiance during passage of a sunspot group

- The Sun as a whole darkens when spots move across its disc
- I.e. the blocked heat does not reappear somewhere else
- I.e. no strong bright rings around spots



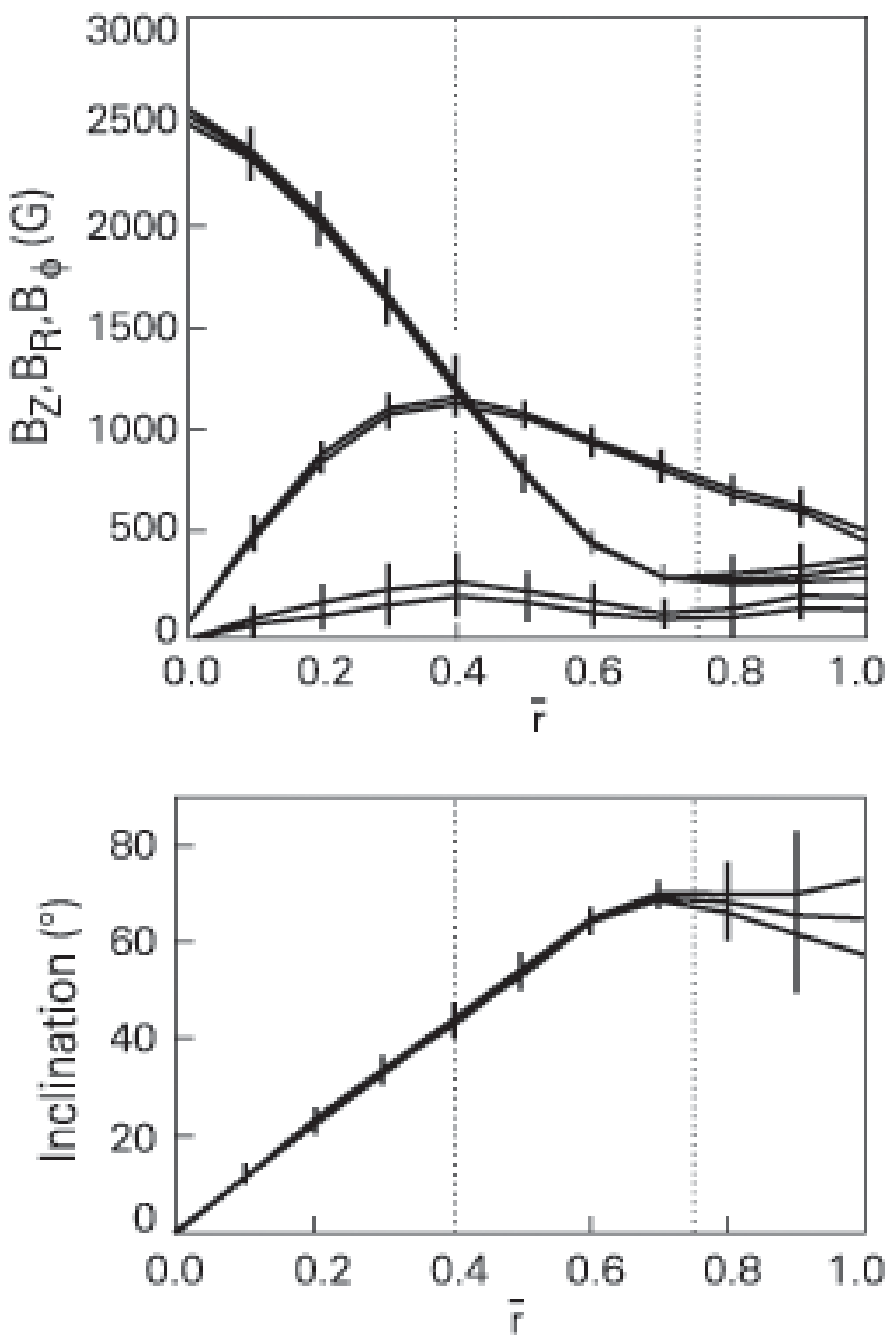
Magnetic structure of sunspots

- 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. It becomes almost horizontal near R_{spot} .
- Regular spots have a field structure similar to a buried dipole



Magnetic structure of sunspots II

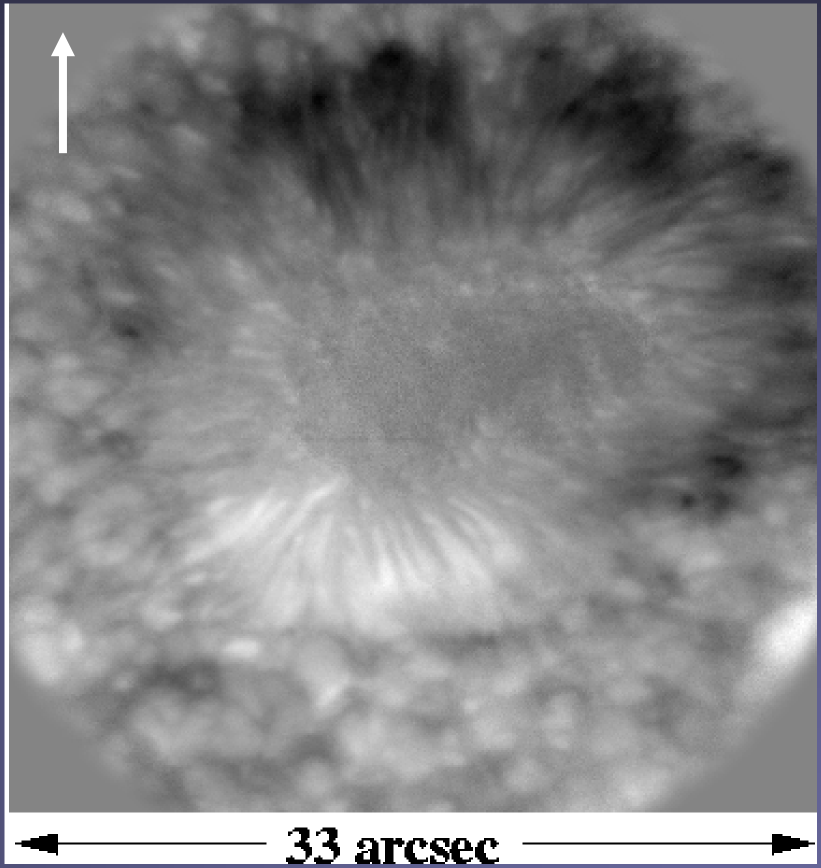
Azimuthal averages of the various magnetic field components in a sample of regular (near-circular) medium-sized sunspots.



Evershed effect

- Observation: Penumbra seen at $\mu<1$ shows
  - on limb side: Doppler red shift
  - on disc side: Doppler blue shift
- Interpretation: horizontal OUTflow of material from inner penumbra to outer

Limb

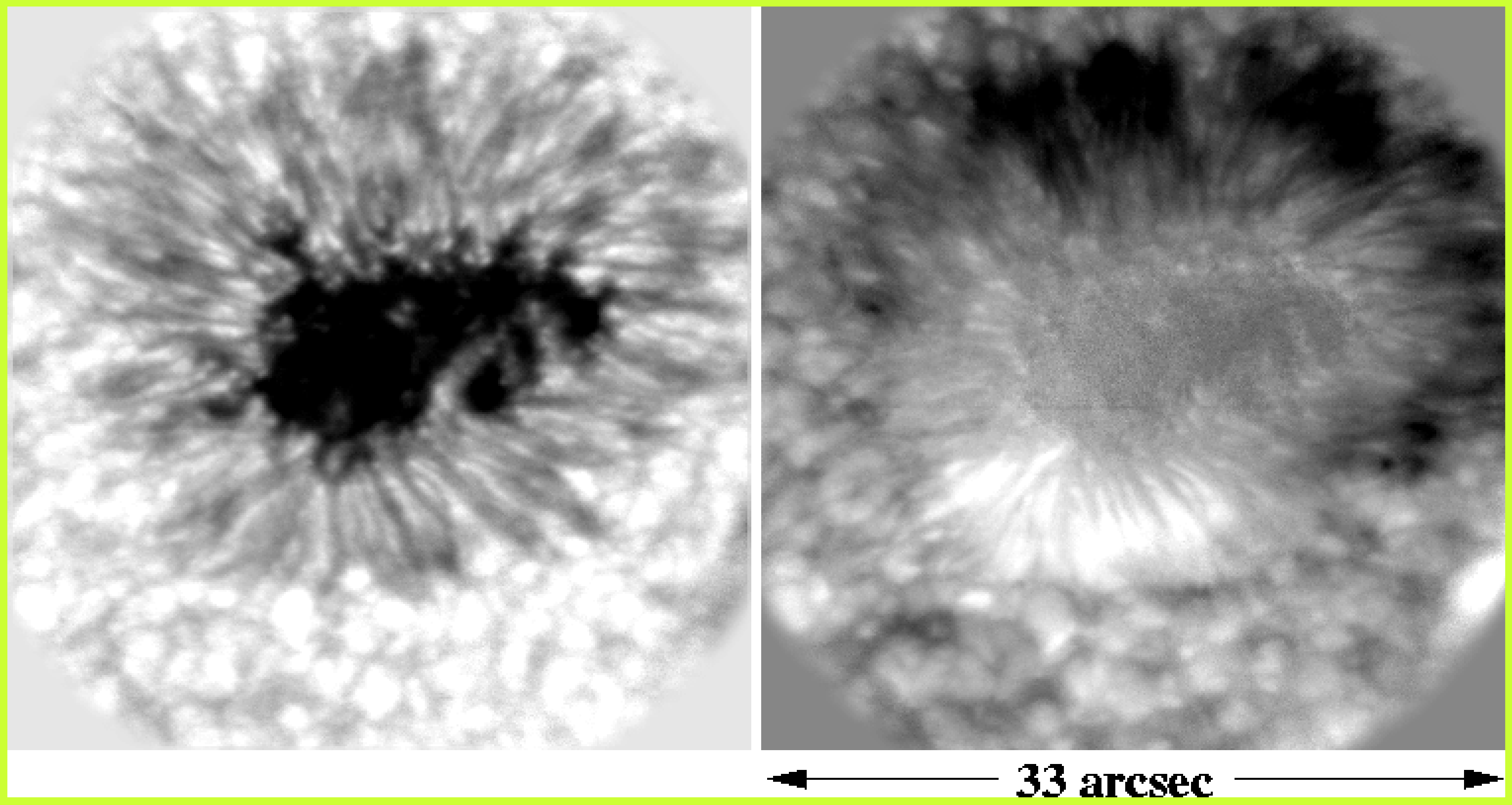


dark: redshift, light: blueshift

Evershed effect

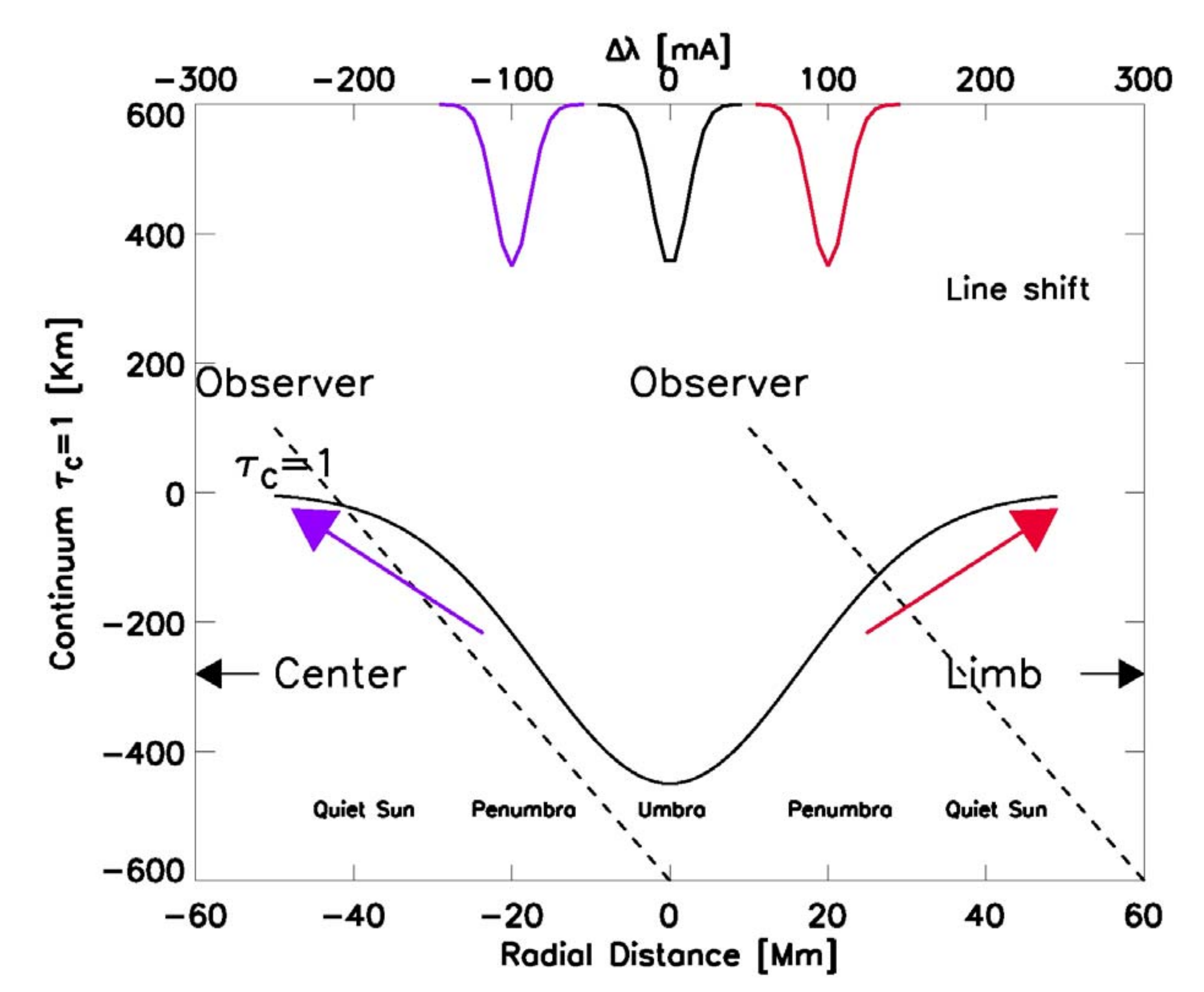
In photospheric layers penumbra shows nearly horizontal outward flows of 1-2 km/s on average (peak flows: supersonic)
Umbra remains at rest. In chromosphere: inward directed flow

Brightness Doppler shift



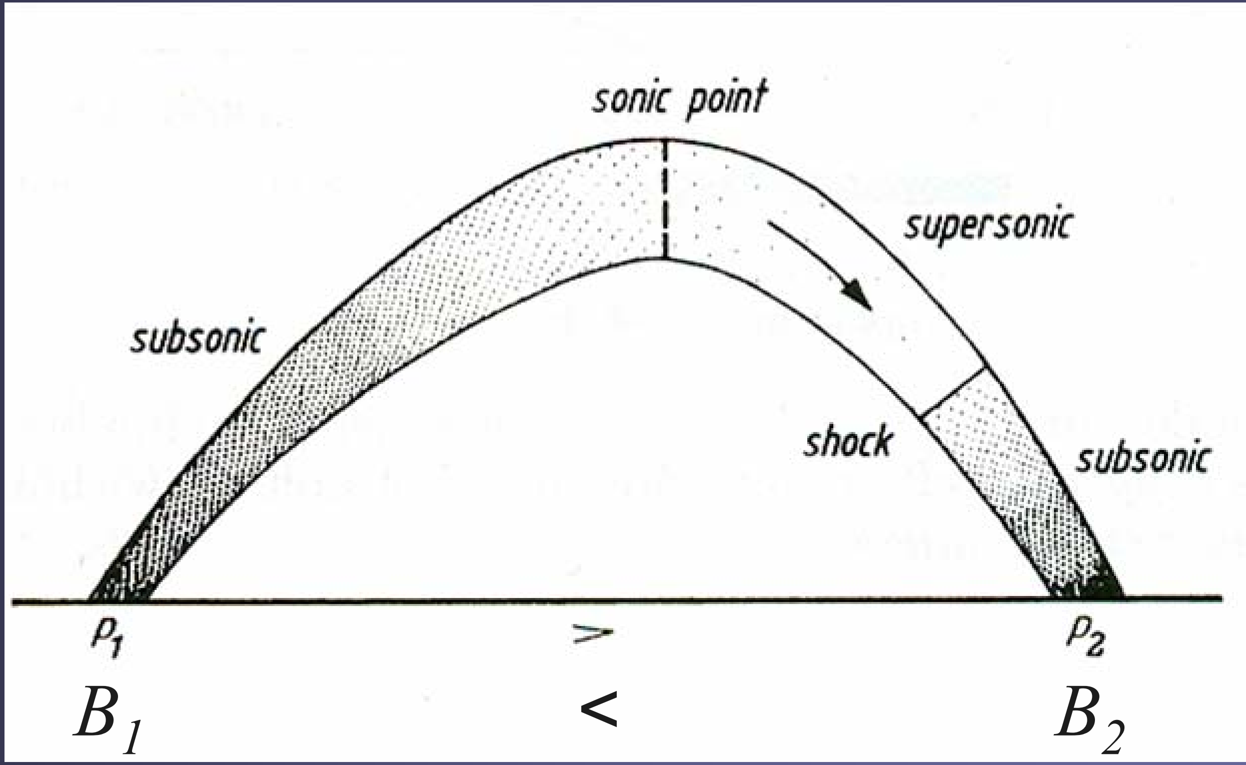
Evershed effect: illustration

- Almost horizontal outflow of matter seen in the penumbra of a sunspot.
- Thought to be driven by either
  - siphon flow
  - convection

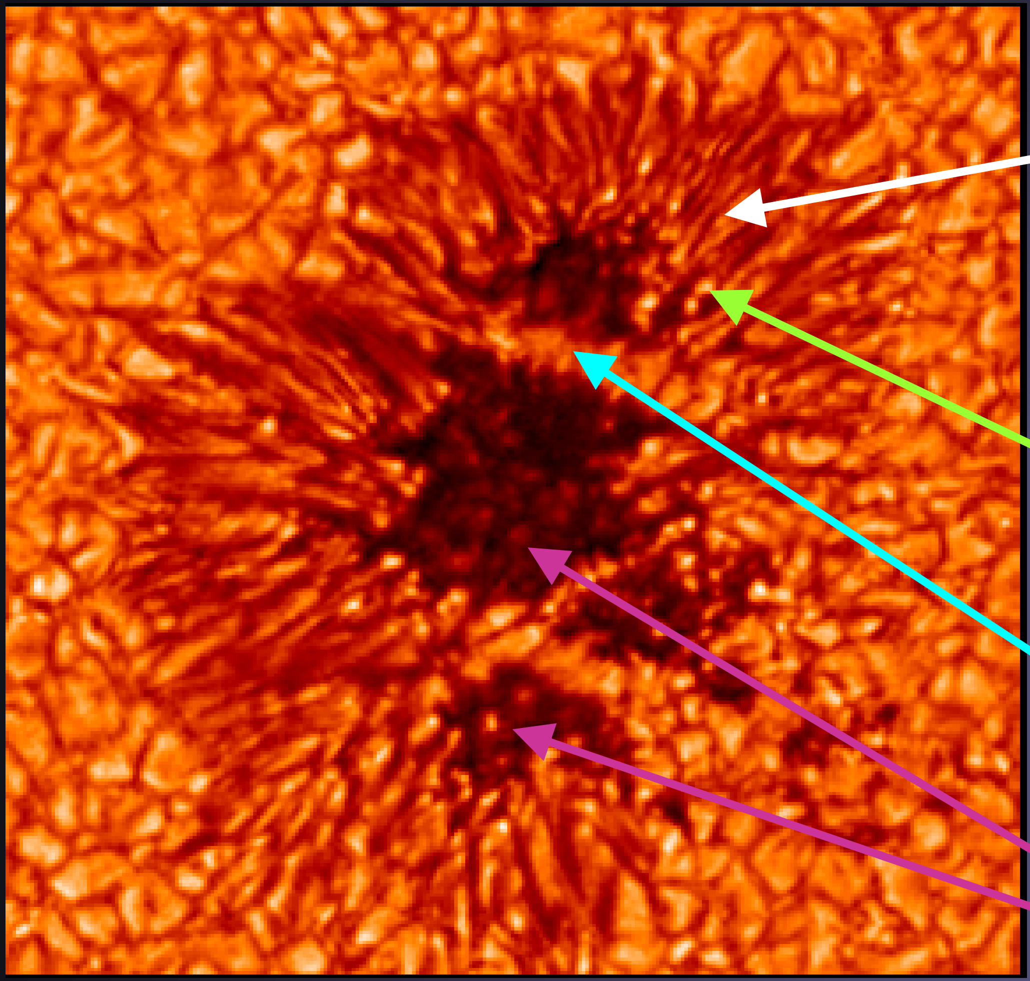


Siphon flow model of Evershed effect

- 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.



Sunspot fine structure



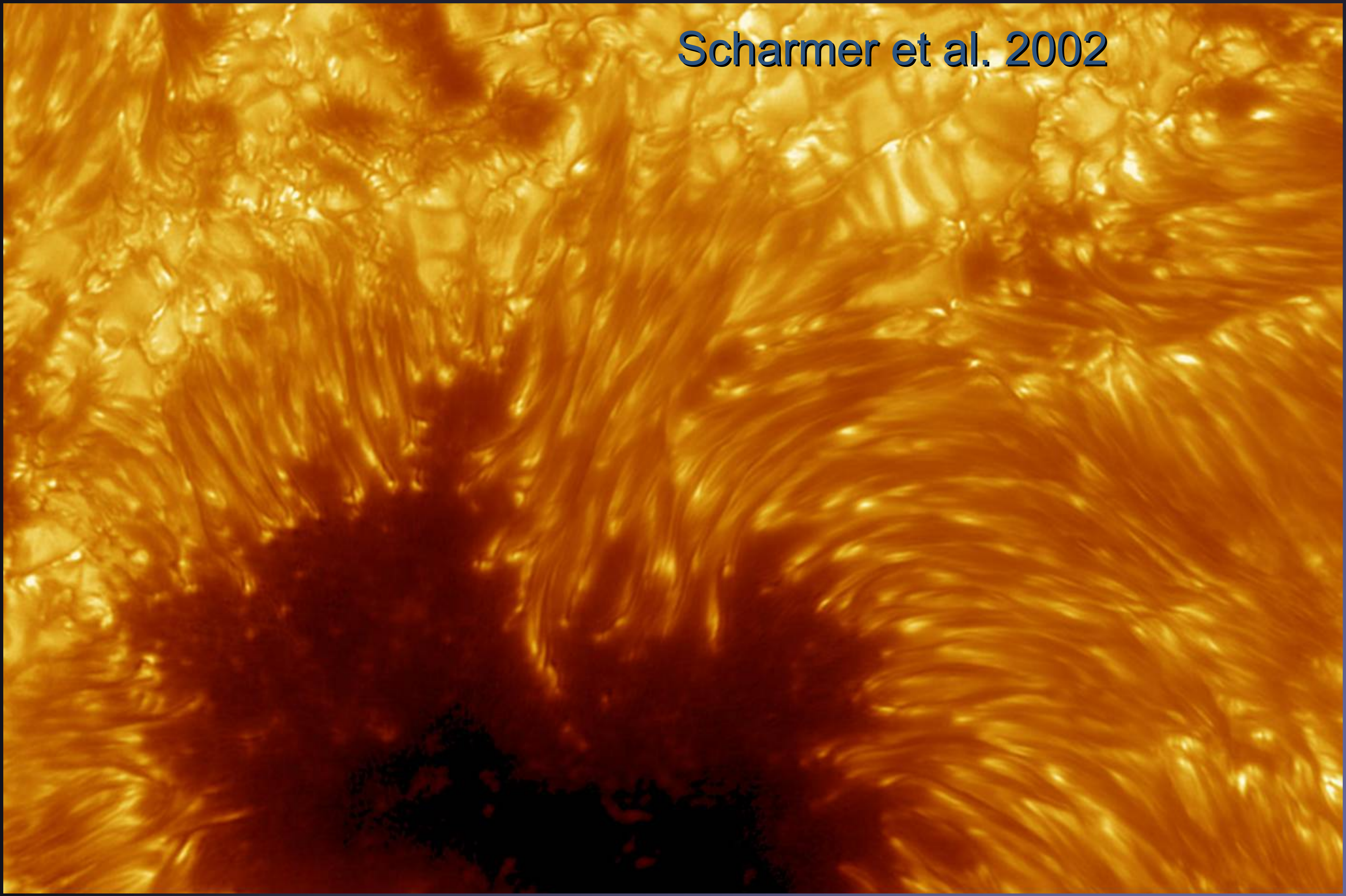
Penumbral filaments (bright and dark)

Penumbral grain (seen to move inward)

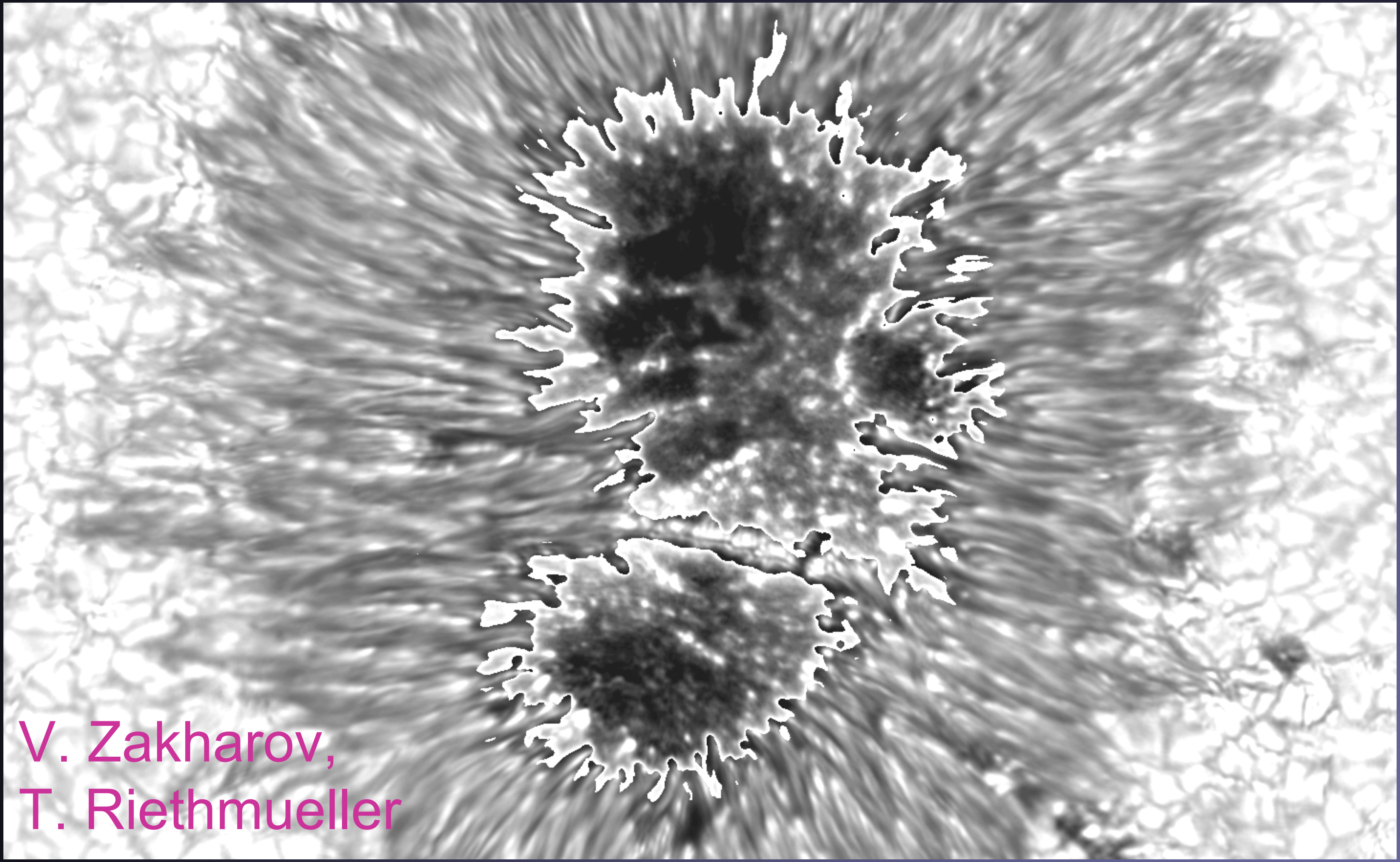
Light bridge

Umbral dot

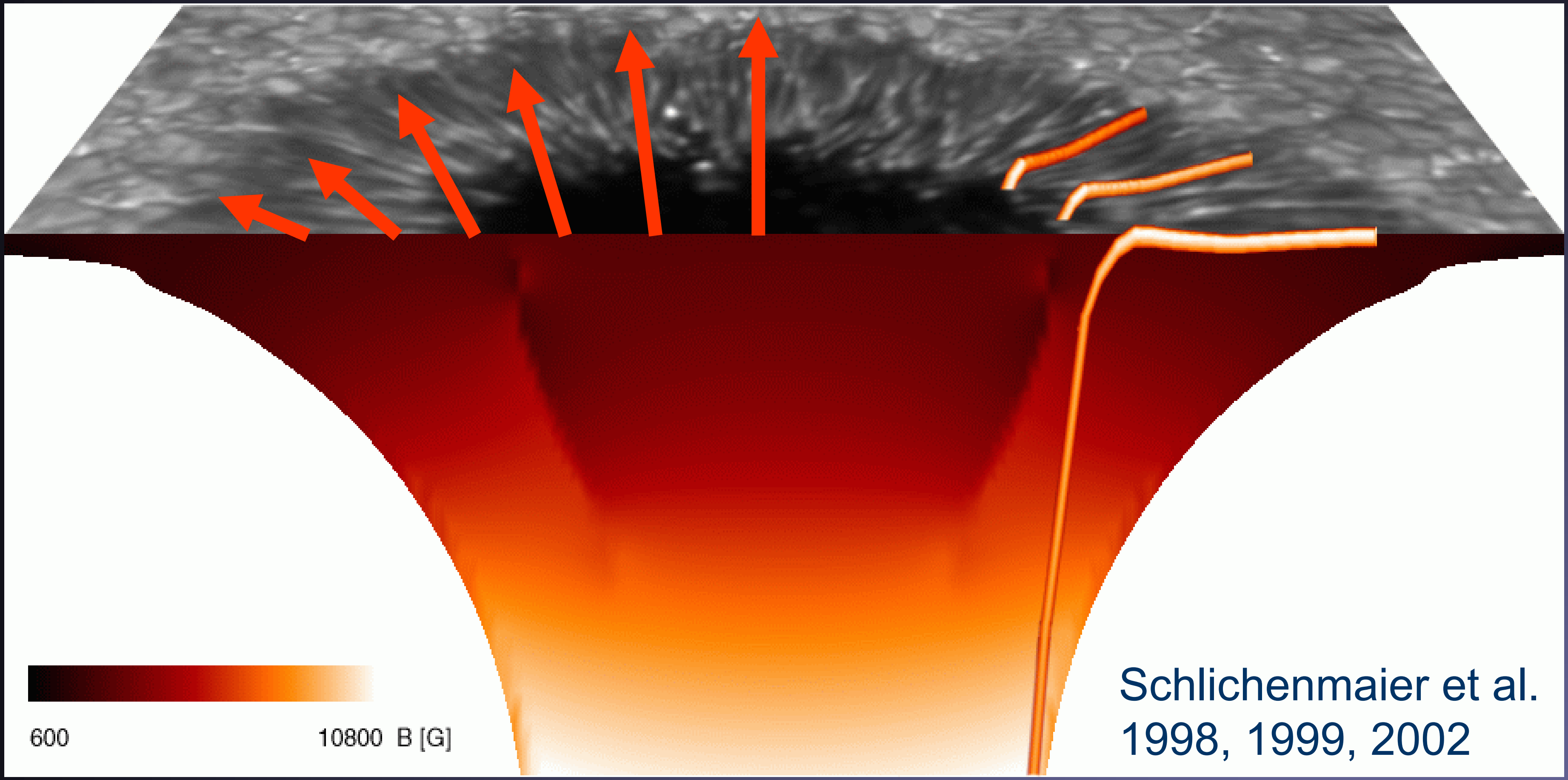
Highest resolution



Umbral dots seen in TiO

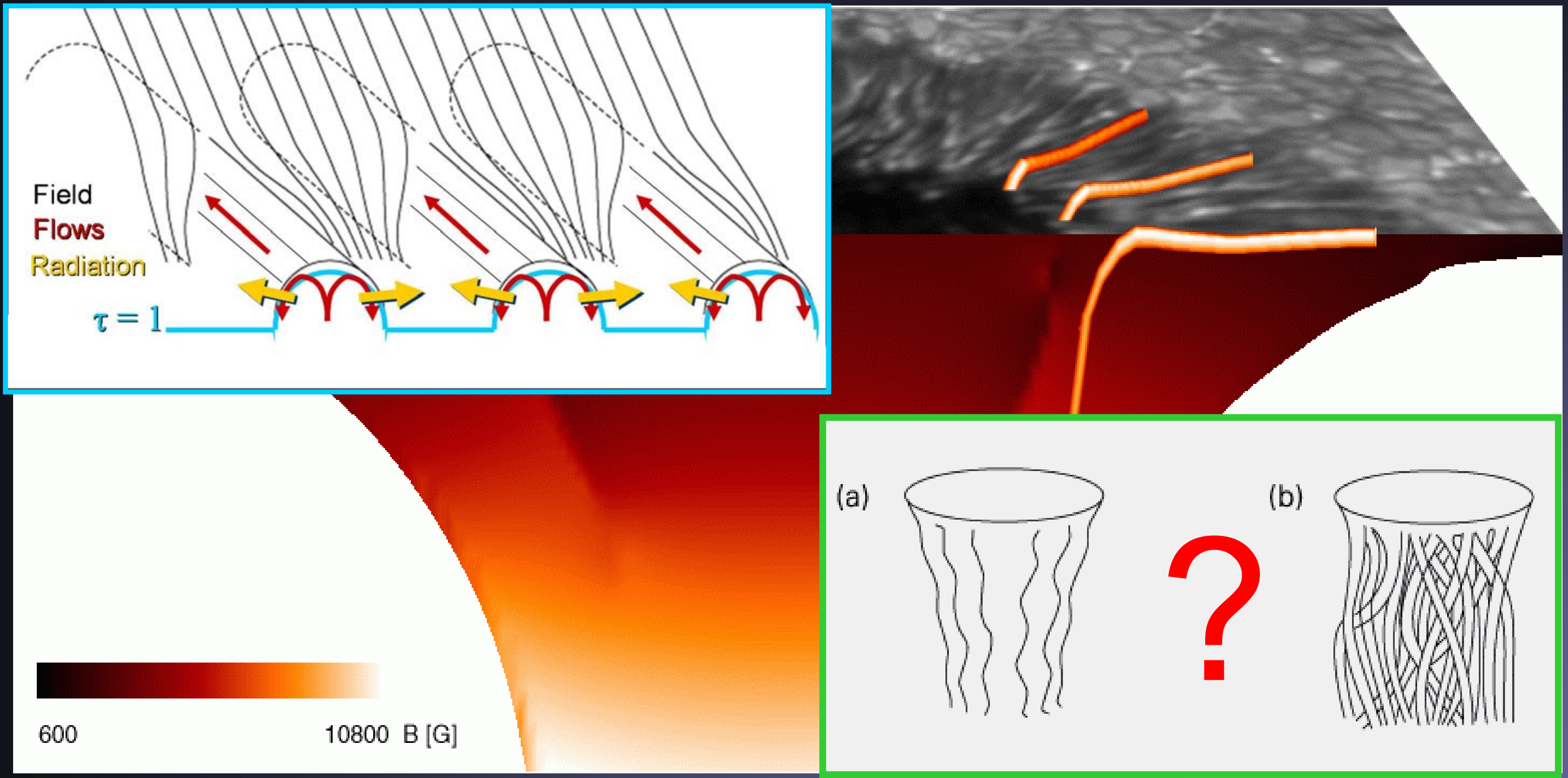


Magnetic structure of sunspots



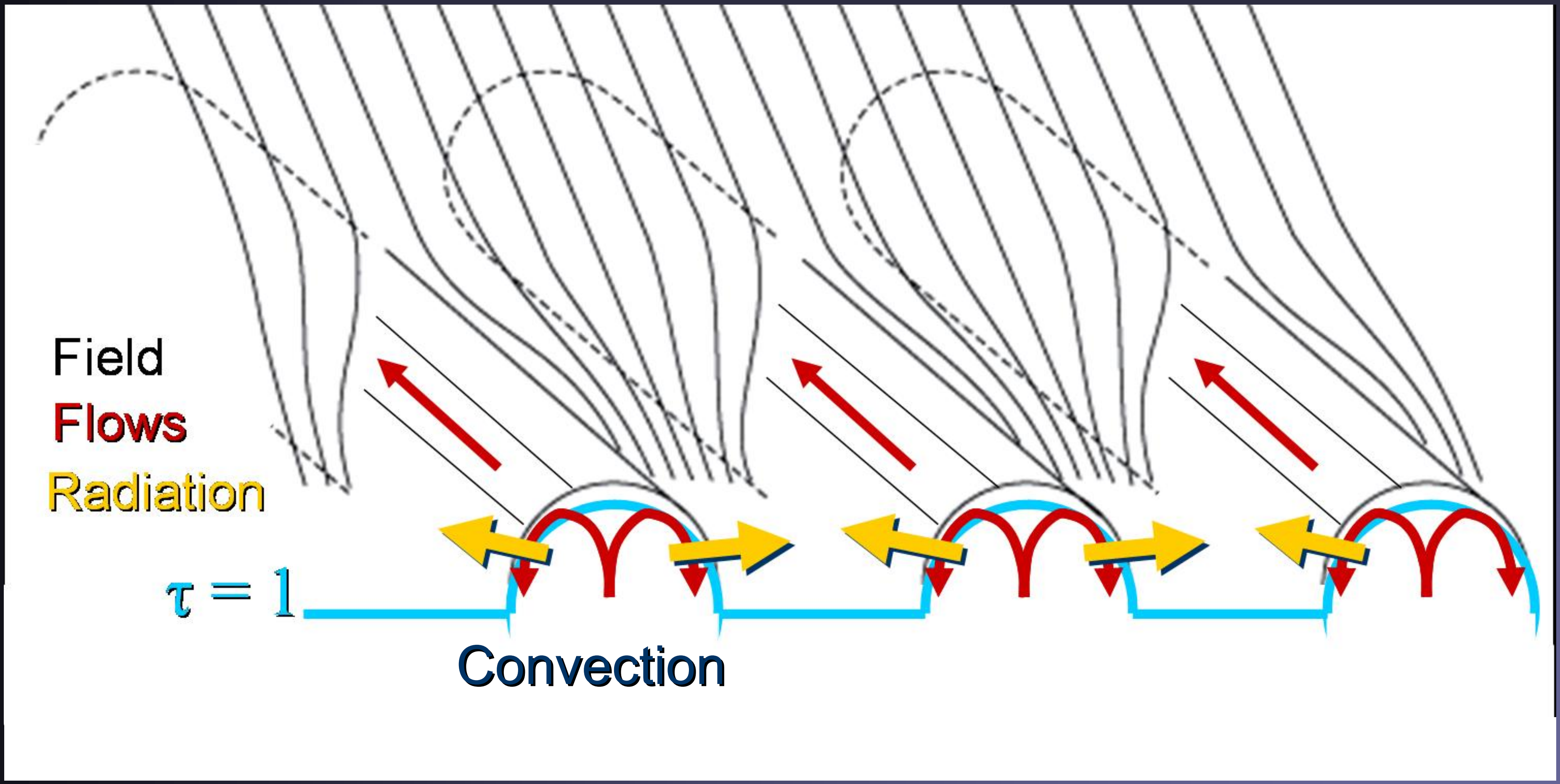
- Regular on large scales (≈ dipole, B_{max} ≈ 2500 G, for simple spots)
- Extremely complex on small scales (penumbra, subsurface)

Magnetic structure of sunspots



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

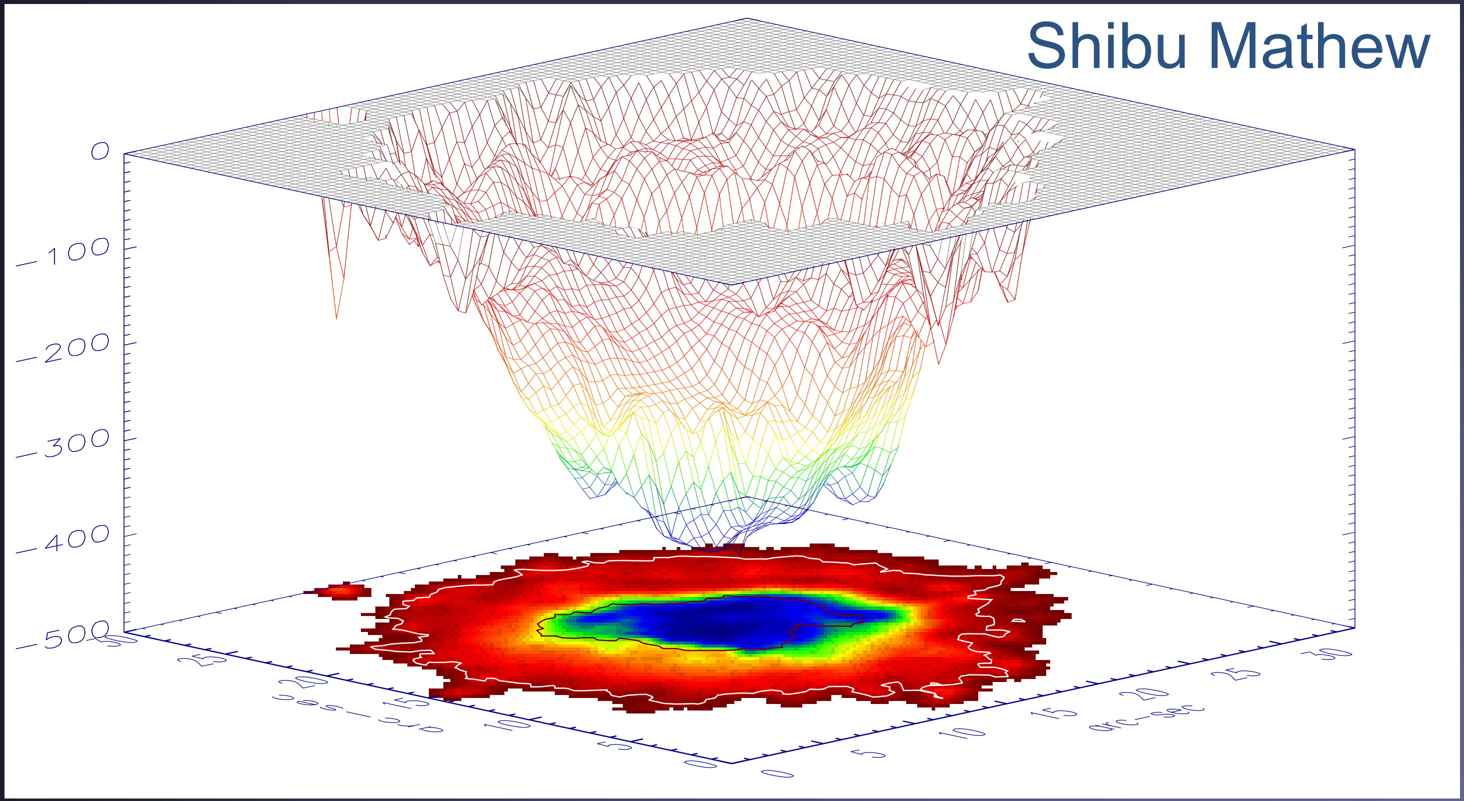
Current view of fine-structure of penumbra



Zakharov et al. 2008, Rempel et al. 2008

Sunspot Wilson depression

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

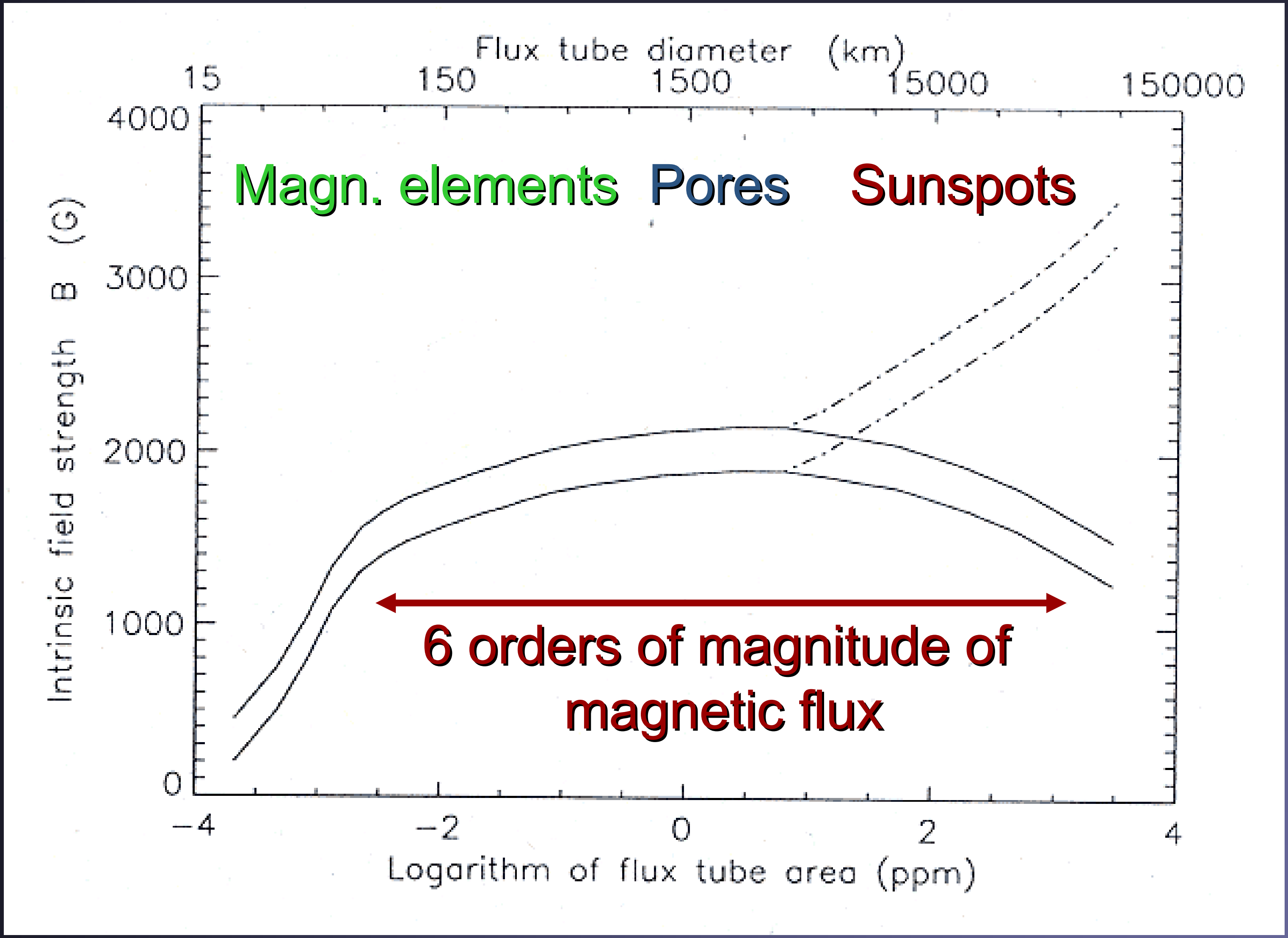


$B^2/8\pi + p = \mathrm{const.}$ ➔ p lower in spot than outside ➔ density also lower ➔ opacity also lower ➔ we see deeper into spot

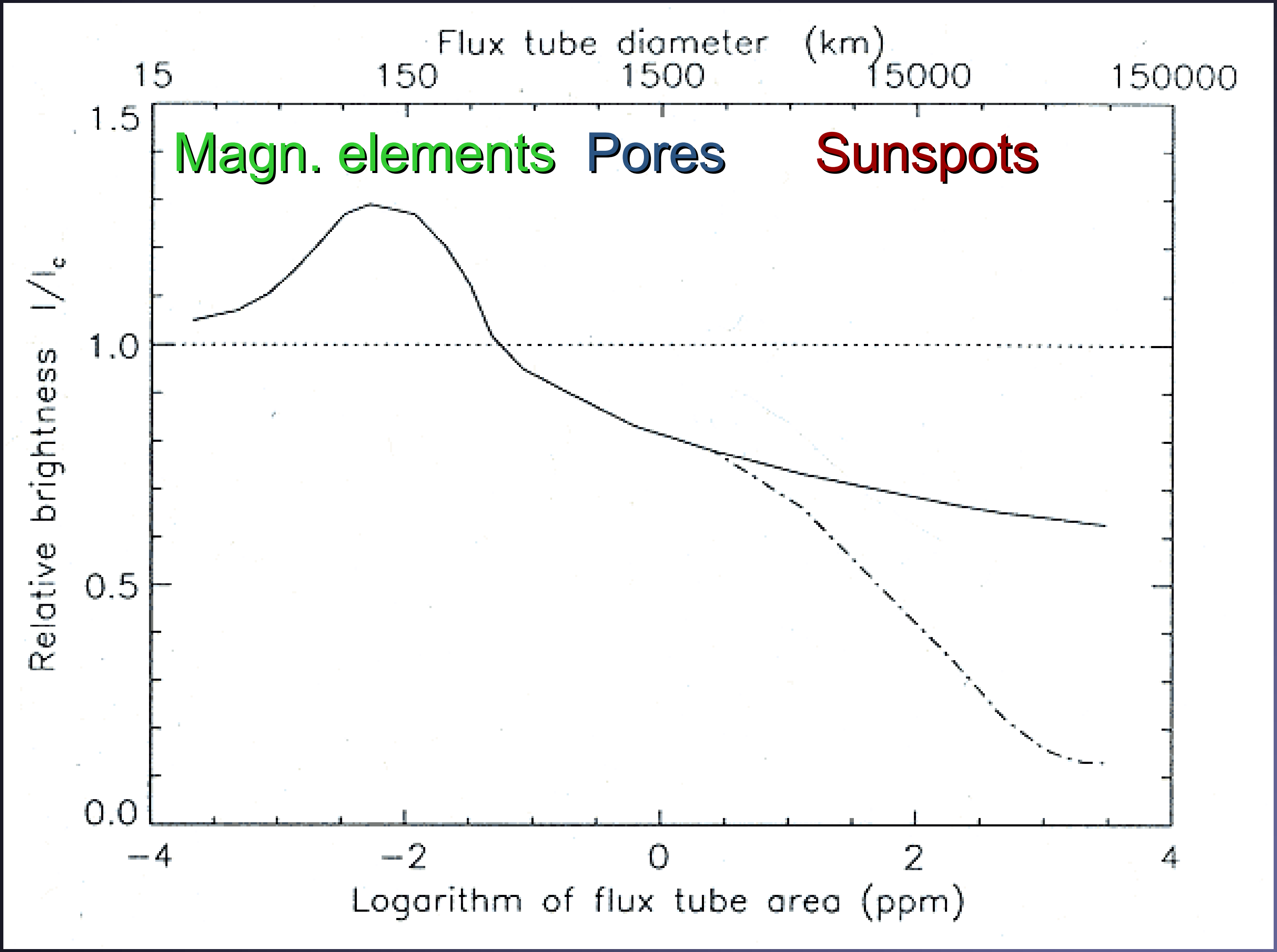
Magnetic elements

- 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 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, expanding with height), like smaller, simpler versions of sunspots.

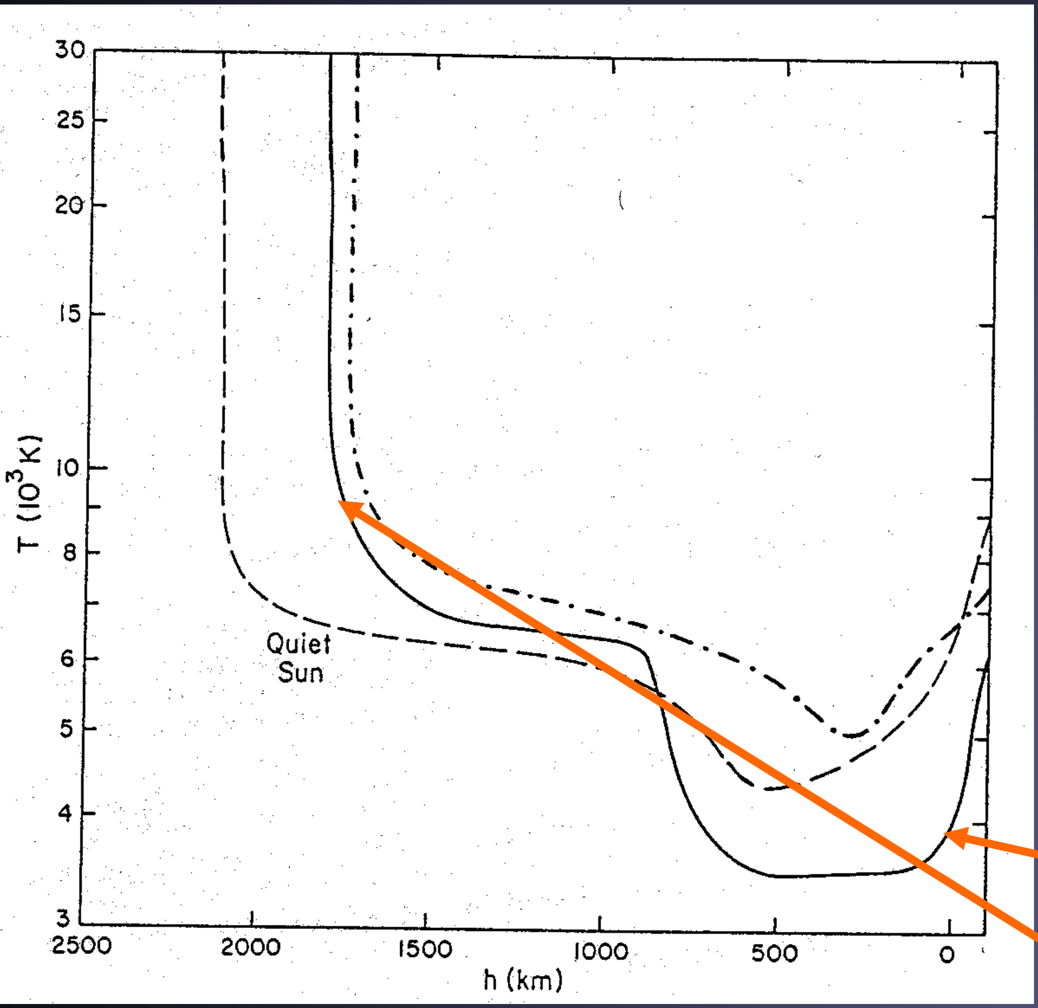
Surprisingly constant field strength



Temperature contrast vs. size



Temperature stratifications of quiet Sun, sunspot, magnetic element

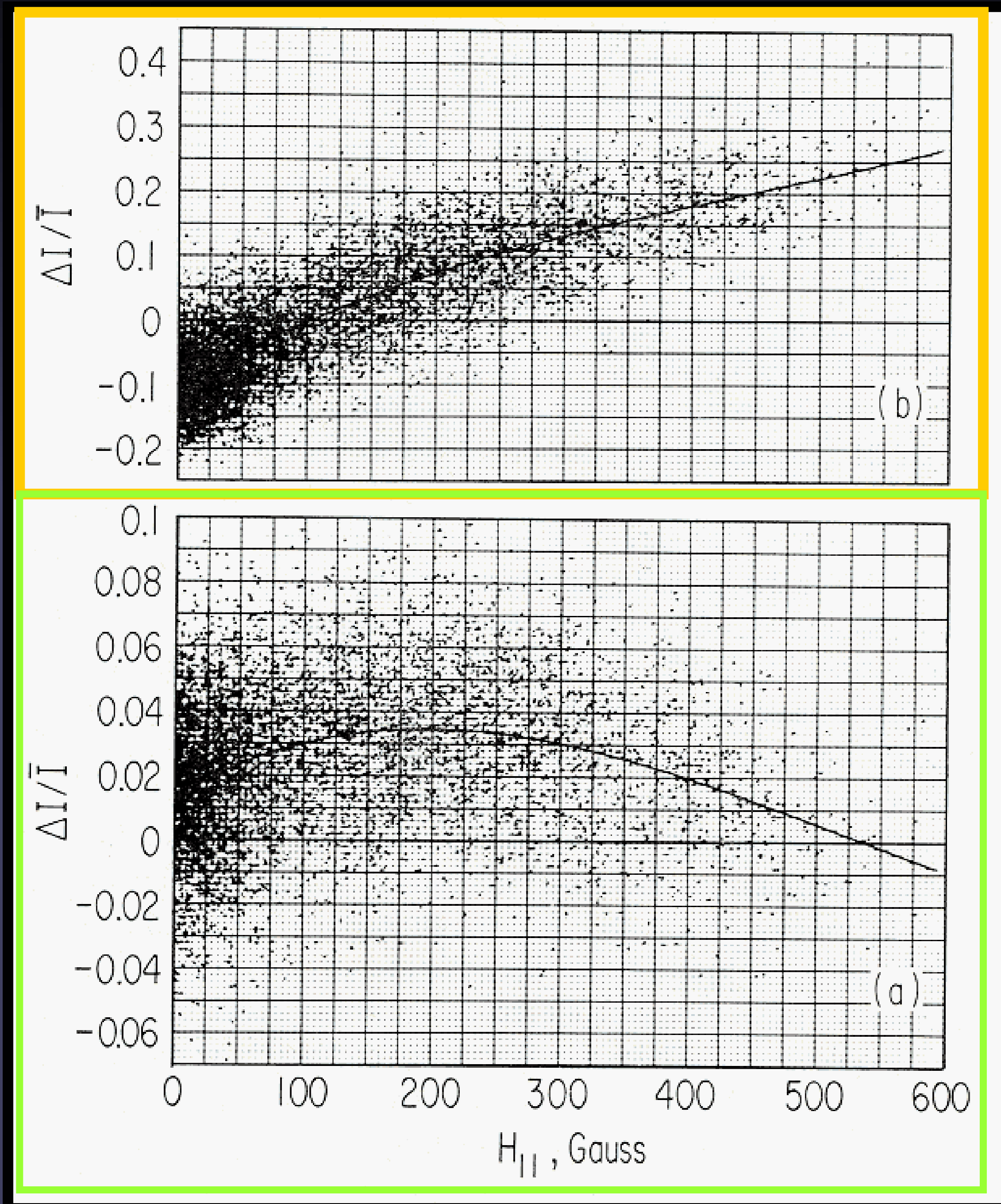


- Dashed: Quiet Sun atmosphere
- Solid: sunspot atmosphere
- Dot-dashed: active region plage, magnetic element
- Plage is hottest everywhere in atmosphere
- Sunspot cold (dark) in photosphere, but hot (bright) in chromosphere & transition region

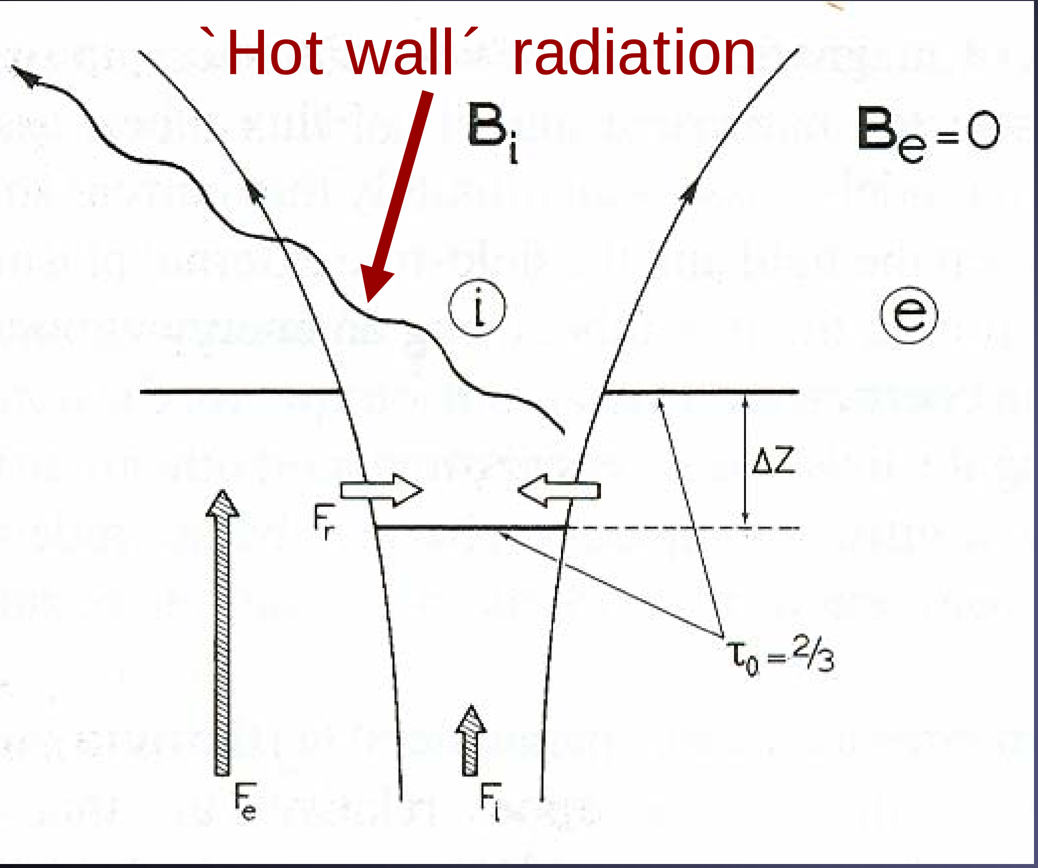
Contrast of magnetic elements

Lower image: continuum contrast (i.e. brightness of magnetic elements relative to QS brightness) vs. mean B in a pixel.

Upper image: contrast in core of a line



Why are magnetic elements bright?

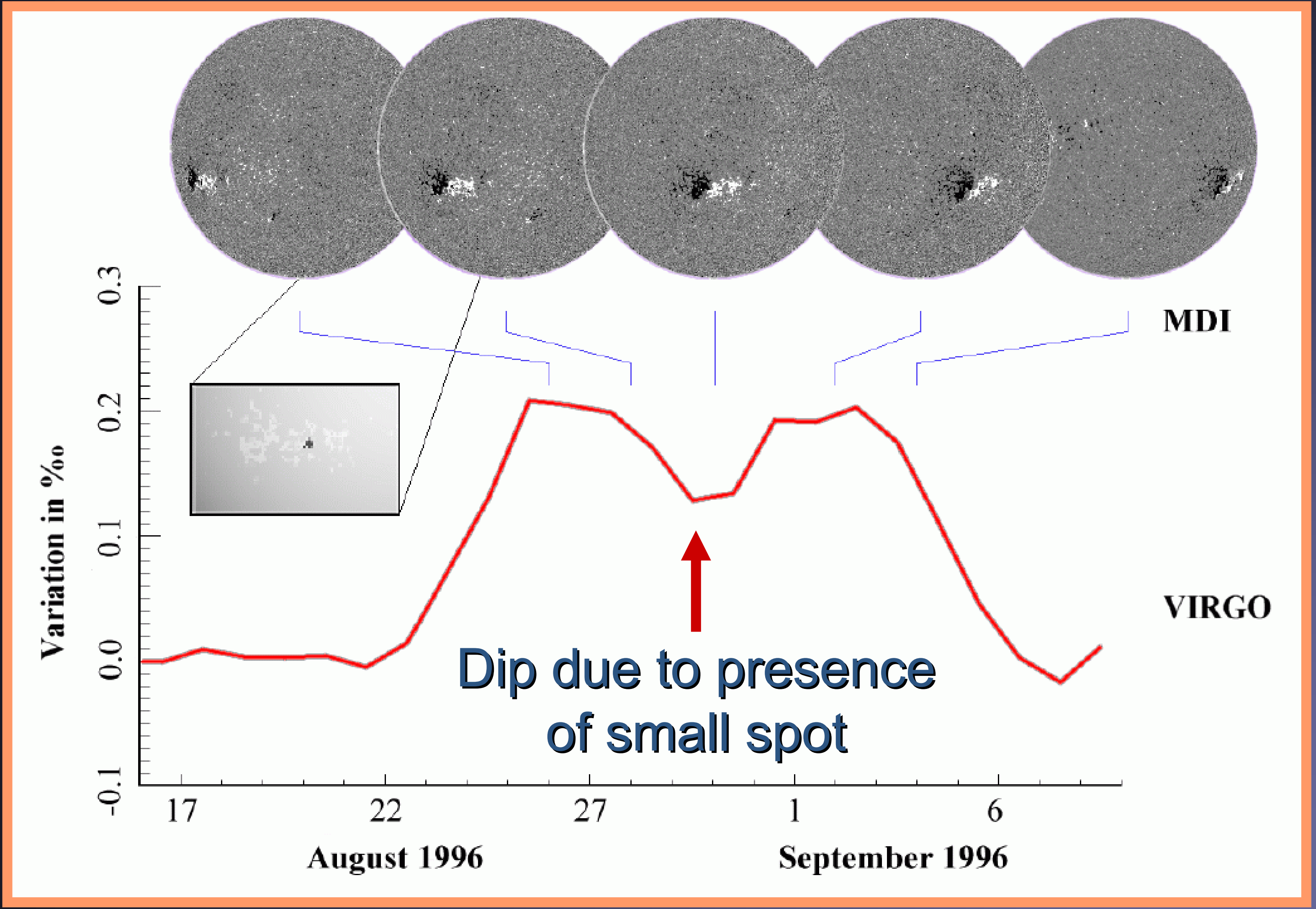


- Quenching of convection
- Partial evacuation
  - → enhanced transparency
  - → heating by `hot walls´
  - → *local flux excess*
- Inflow of radiation wins because the flux tubes are narrow (diameter ~ Wilson depression).
- High heat conductivity
  - → flux disturbance partly propagates into the deep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

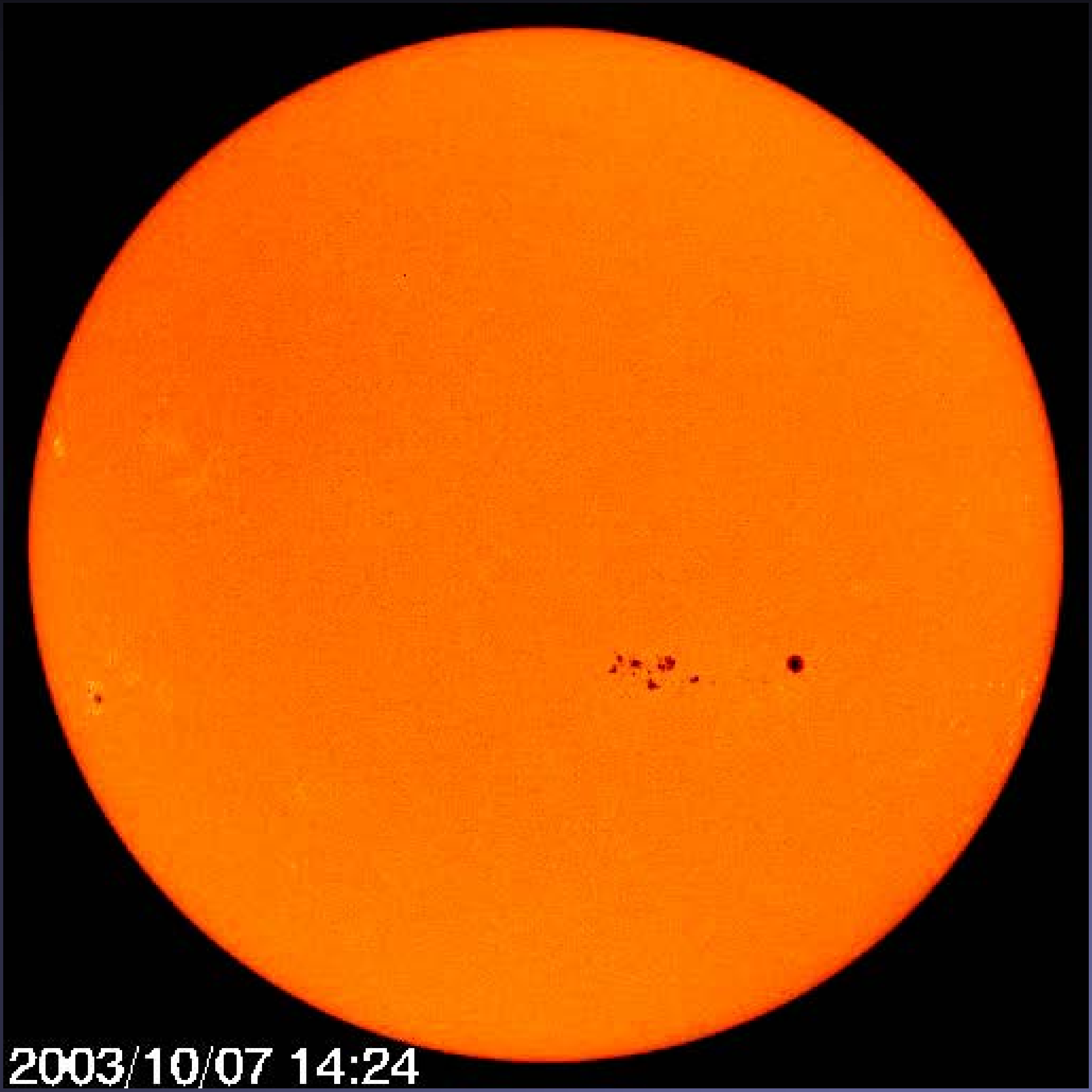
Why this magnetic field structure?

- B-field structure of magn. elements driven largely by
  - horizontal pressure balance: $B^2/8\pi + p = \mathrm{const}$
  - hydrostatic stratification (vertical pressure balance). For an isothermal atmosphere: $p = p_0 \exp(-z/H)$
- Since gas pressure drops exponentially with height, (for T=const), so must the field strength: $B = B_0 \exp(-z/2H)$
- Flux conservation: $\mathrm{div}\, B = 0$, or: $\iint B(x,y,z)\,dx\,dy = \mathrm{const}$
- B decreases exponentially with z ➔ area A of sunspot (=magnetic flux tube) increases exponentially: $A = A_0 \exp(z/H)$

Faculae lead to brightening of the whole Sun



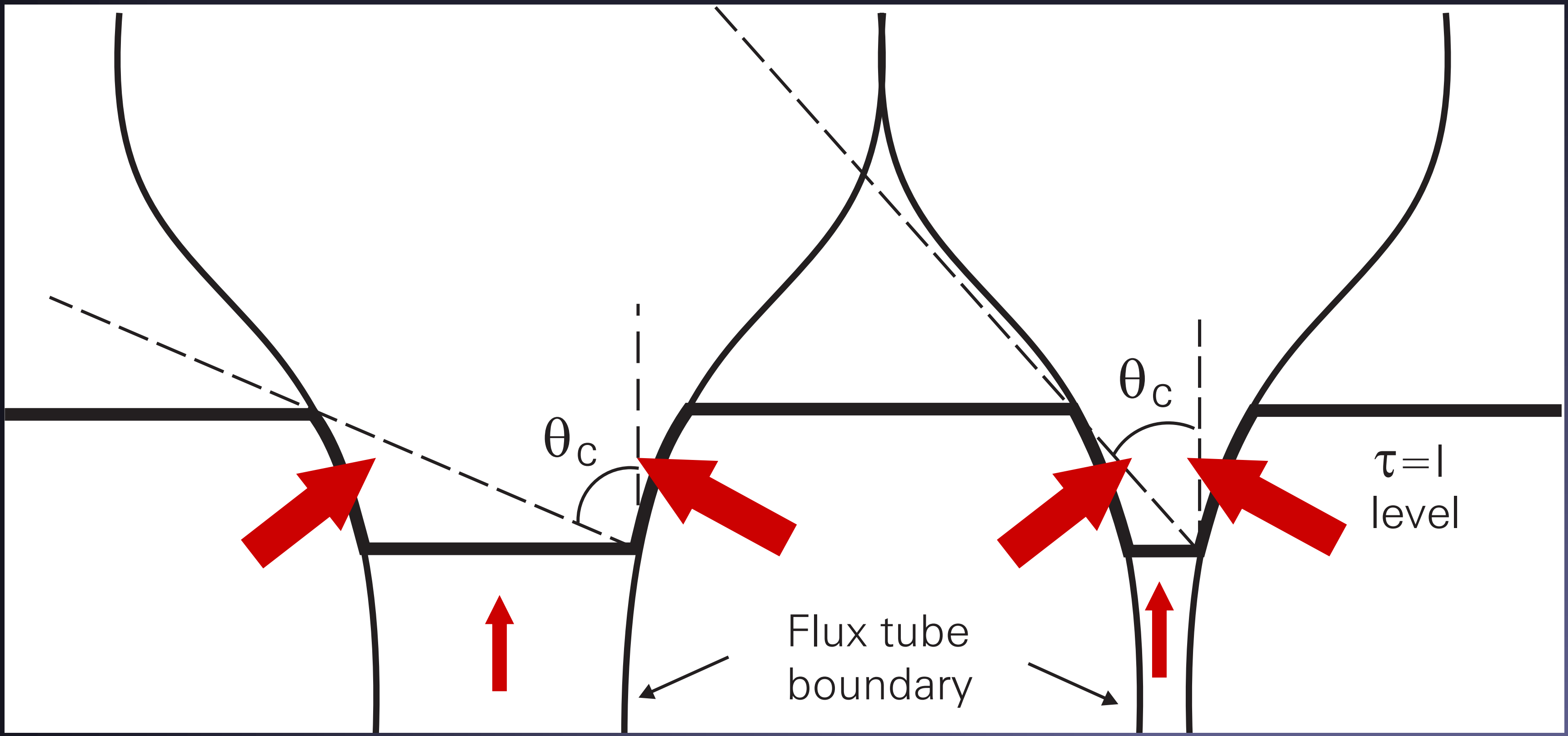
Why are faculae best seen near limb?



The Sun in White Light, with limb darkening removed

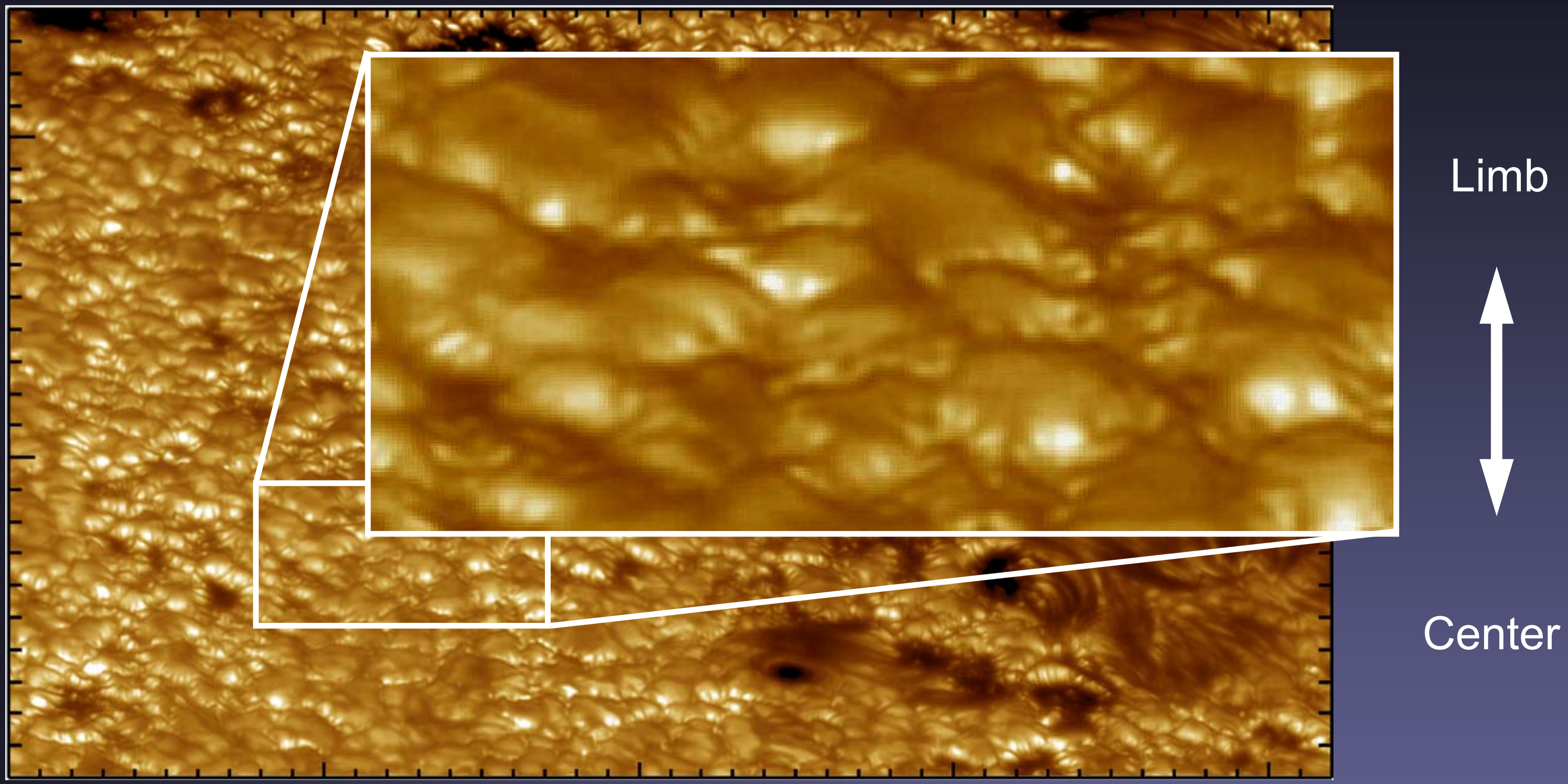
MDI on SOHO

Flux-tube brightening near limb



- The flux tubes expand with height (pressure balance
- Most energy radiates into them through walls, which are hot.
- They appear brightest when hot walls are well seen, i.e. near limb (closer to limb for larger tubes)

Facular brightening



(continuum image: SST, La Palma $\theta=60°$ $\lambda=488nm$)

Recent observations reveal:
(Lites et al. 2004)

- 3D appearance of faculae
- extension up to 0.5"
- narrow dark lanes centerward of faculae

3-D radiation MHD simulations

- Similar to hydrodynamic simulations describing granulation, except that now the full MHD equations need to be solved ➔ Additional complication.
- Include realistic radiative transport of energy. Include solar surface to allow comparison with observations
- Suffer from same shortcomings as the HD simulations (too low Reynolds number) etc.

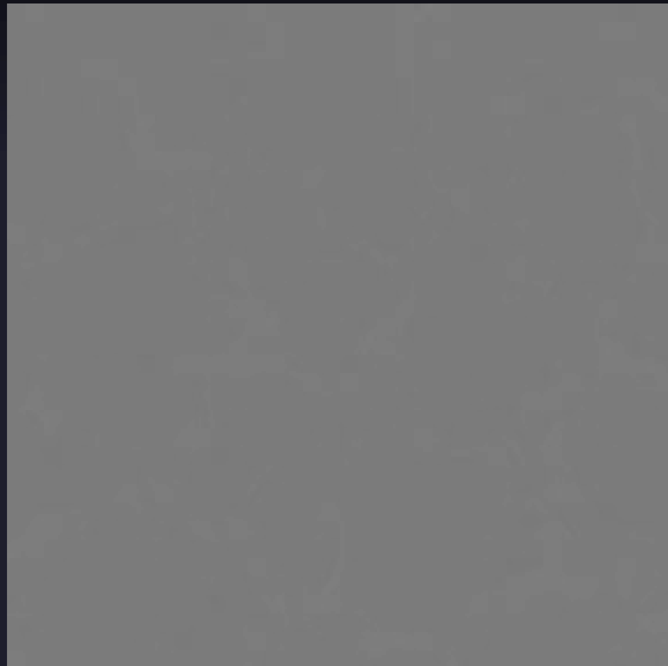
B_z

(Z=0)

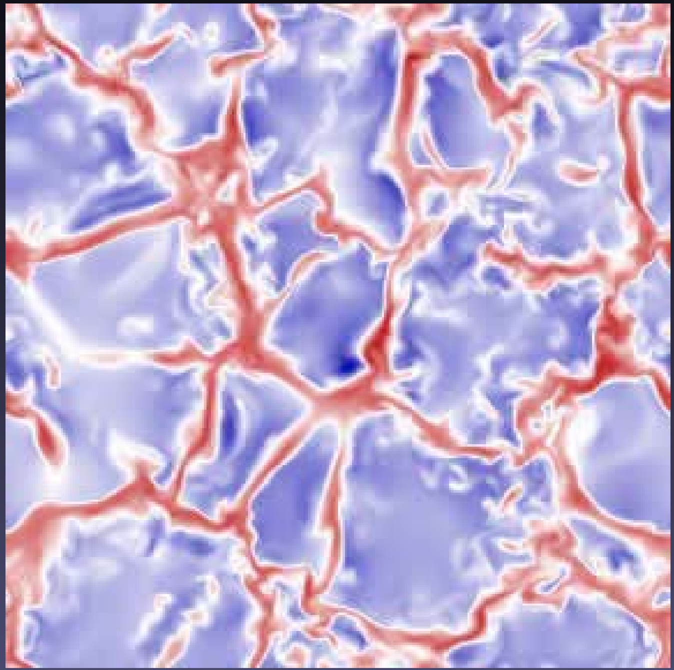
>500G

>1000G

>1500G



6 Mm



v_z

(Z=0)

3-D compressible radiation-MHD simulations

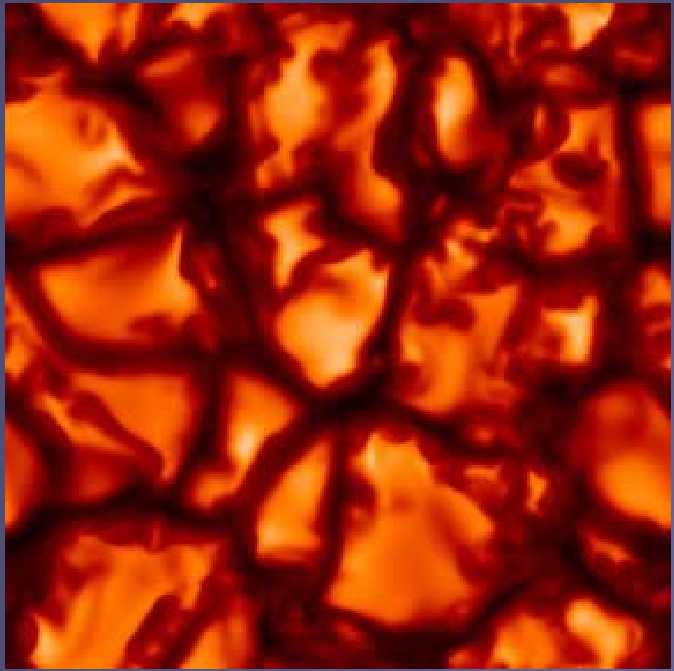
Plage: $B_Z(t=0) = 200$ G

Grid Size: 288 x 288 x 100

Vertical extent: 1.4 Mm

Horizontal extent: 6 Mm

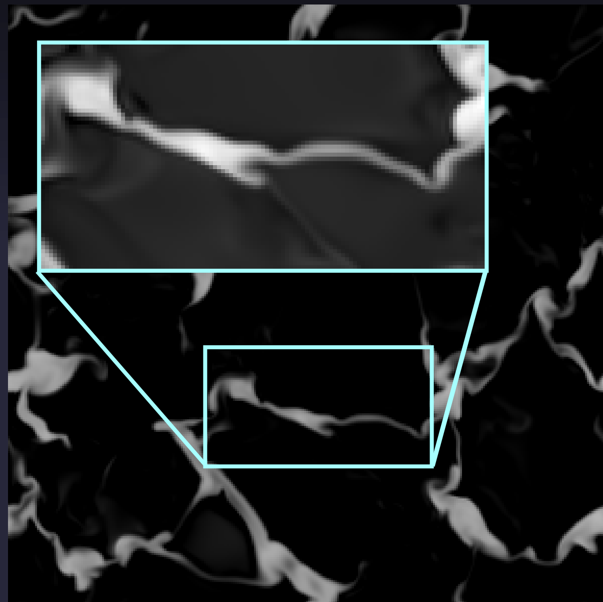
Alexander Vögler et al.

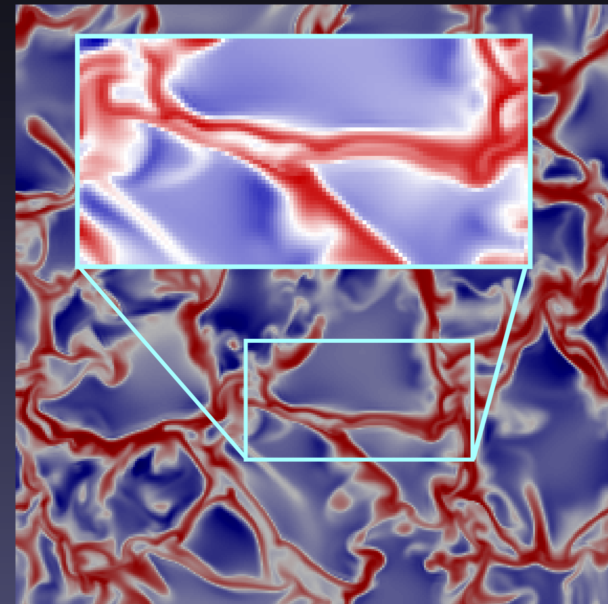


I_c

(Z=0)

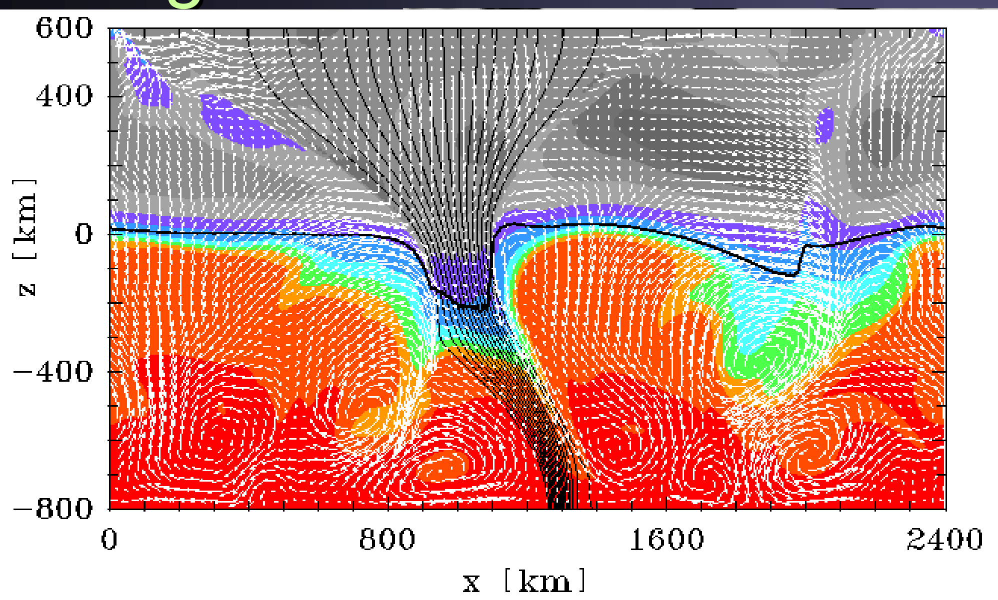
B_z

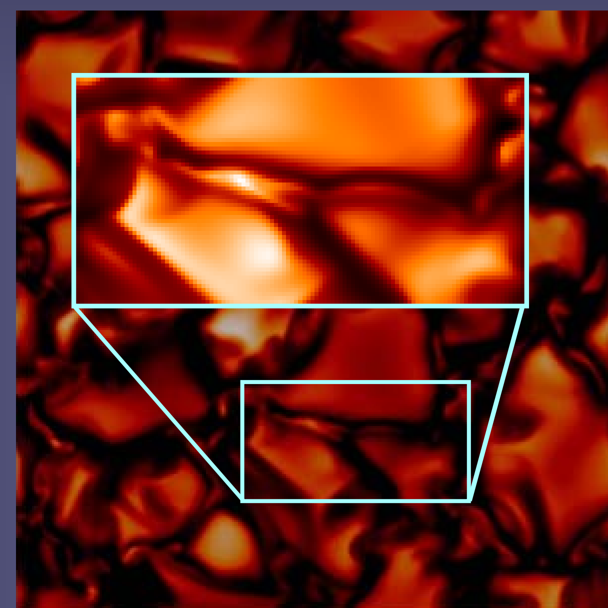




v_z

Details of thin magnetic





I_C

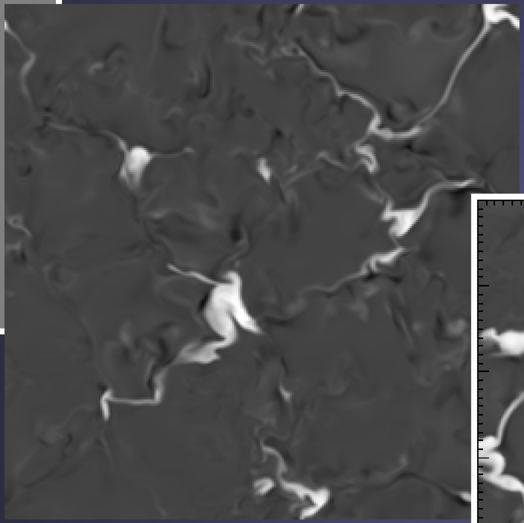
Horizontal cuts near surface level

Vögler et al. 2005

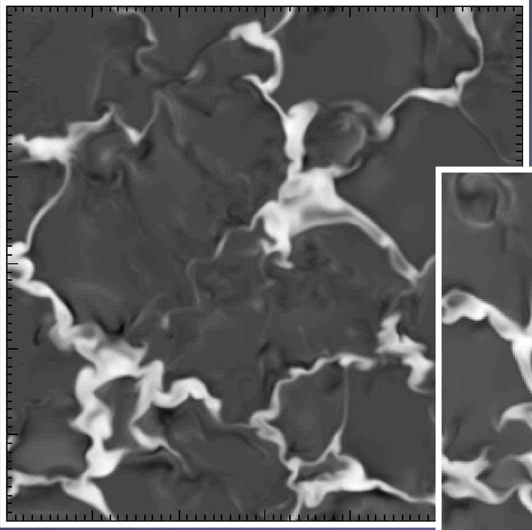
MHD simulations: from quiet Sun to strong plage

0 G

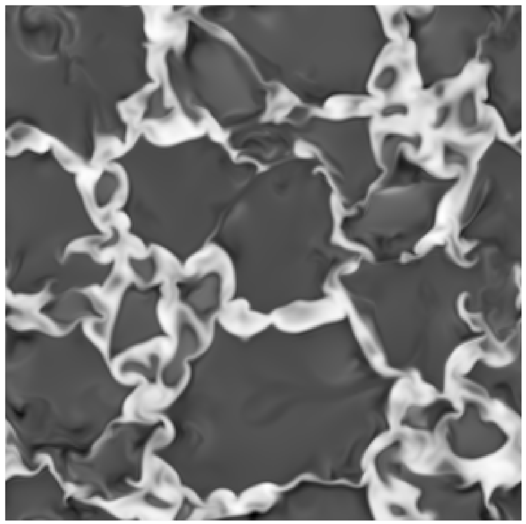
50 G



200 G



400 G

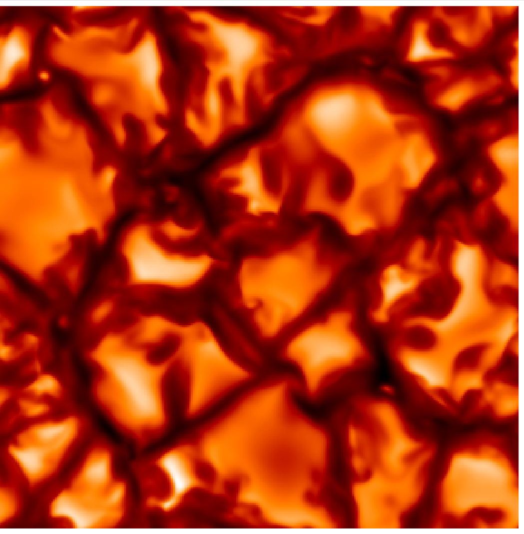


Magnetic field

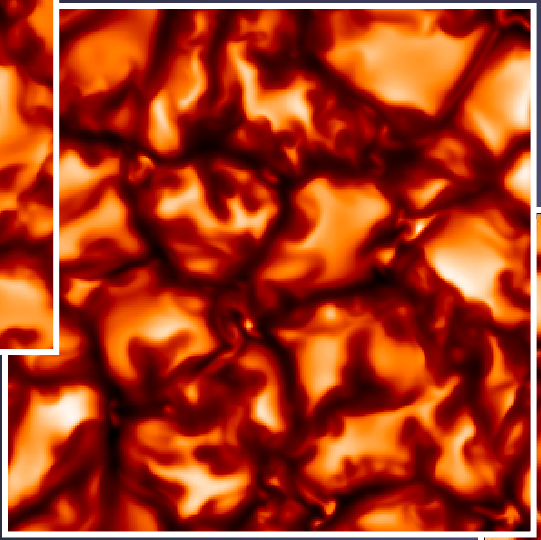
Radiation MHD simulations of solar surface layers. Open lower boundary with fixed value of entropy for bottom inflow (i.e. assume irradiance changes in surface layers)

MHD simulations: from quiet Sun to strong plage

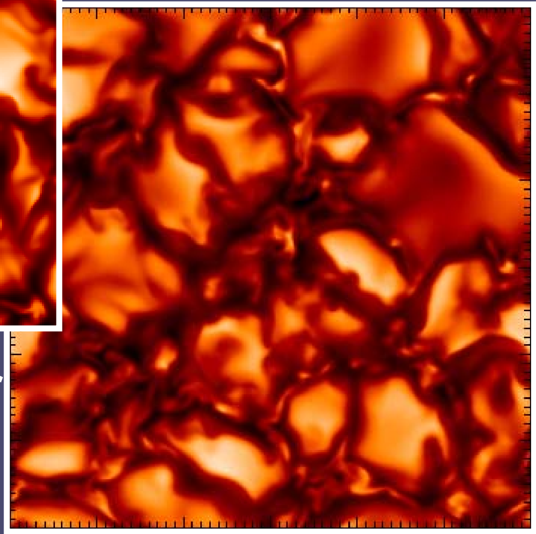
0 G



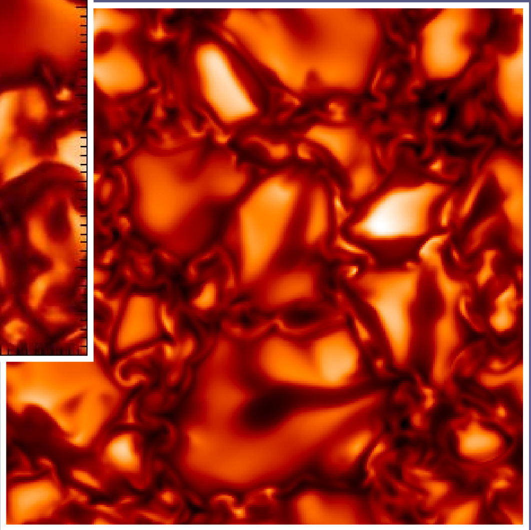
50 G



200 G



400 G

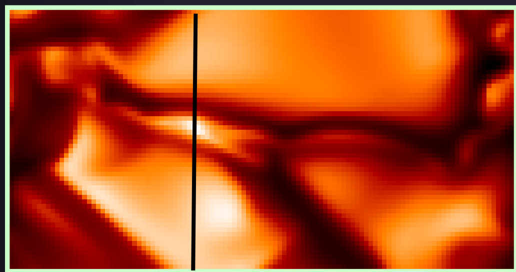


Vögler et al. 2005

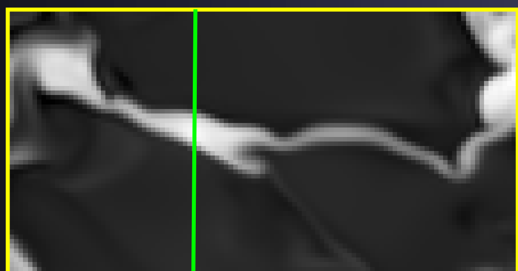
6000x6000x1400 km box, 20km grid

Radiation MHD simulations of solar surface layers. Open lower boundary with fixed value of entropy for bottom inflow (i.e. assume irradiance changes in surface layers)

Vertical cut through sheet-like structure



I

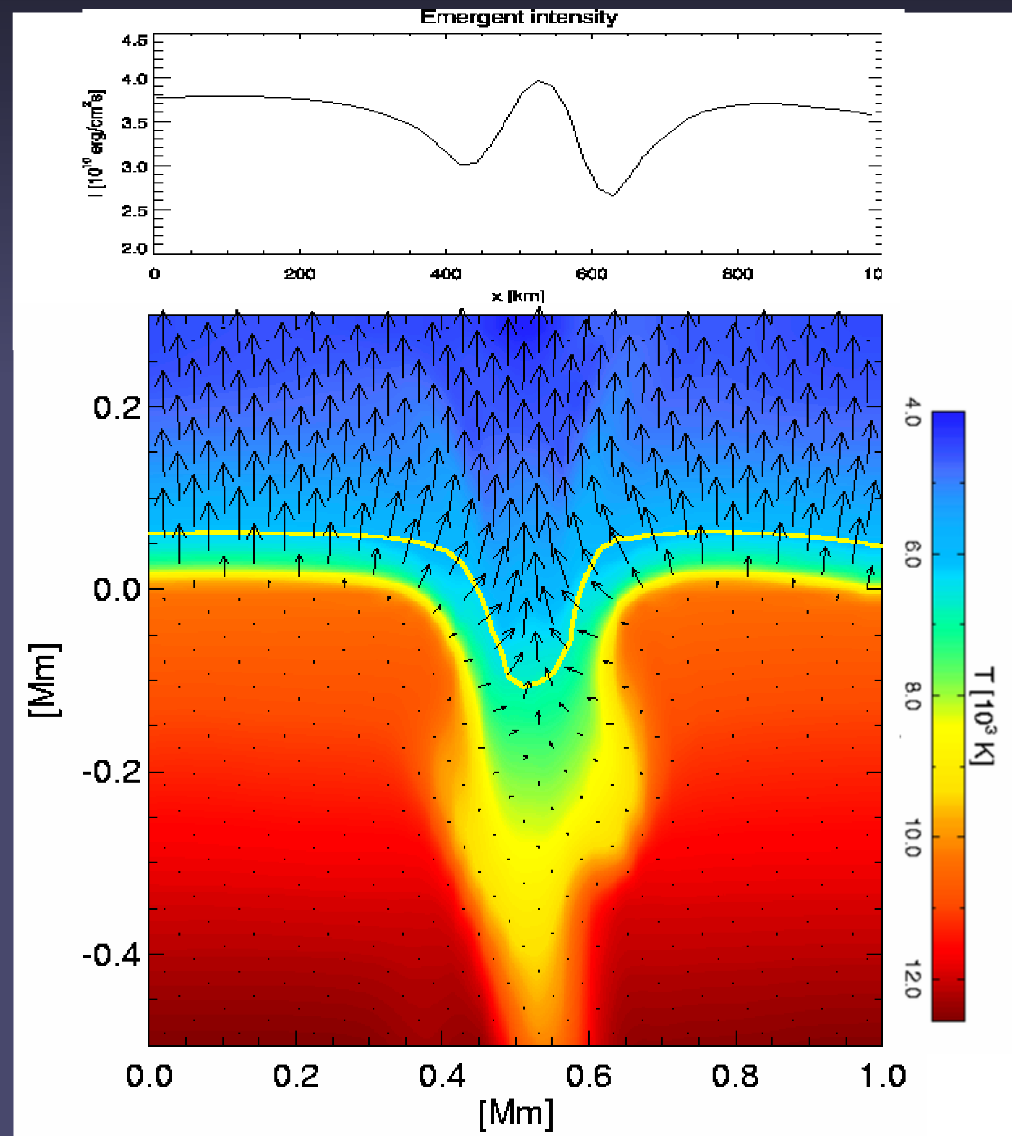


B_z

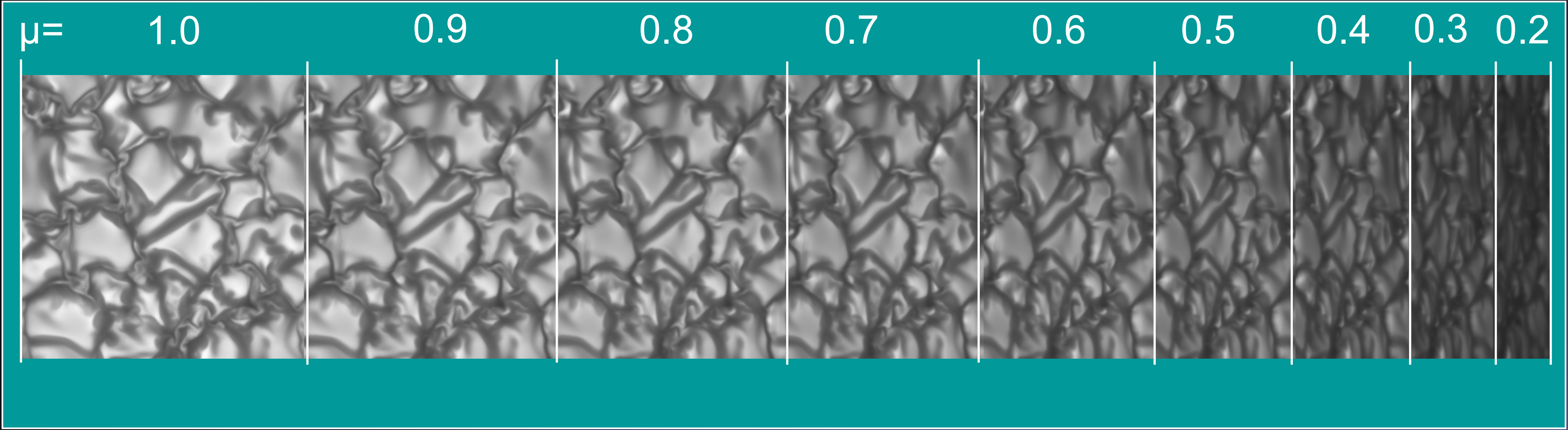
- partial evacuation leads to a depression of the τ=1 level
- lateral heating from hot walls (Spruit 1976)

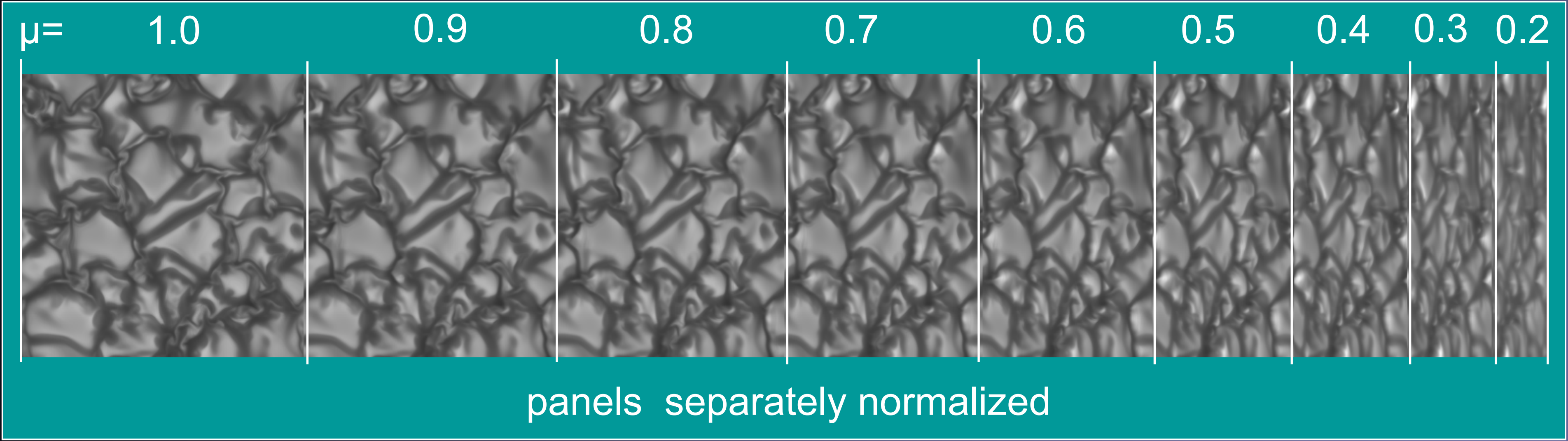
➔ Brightness enhancement of small structures

Radiation flux vectors &

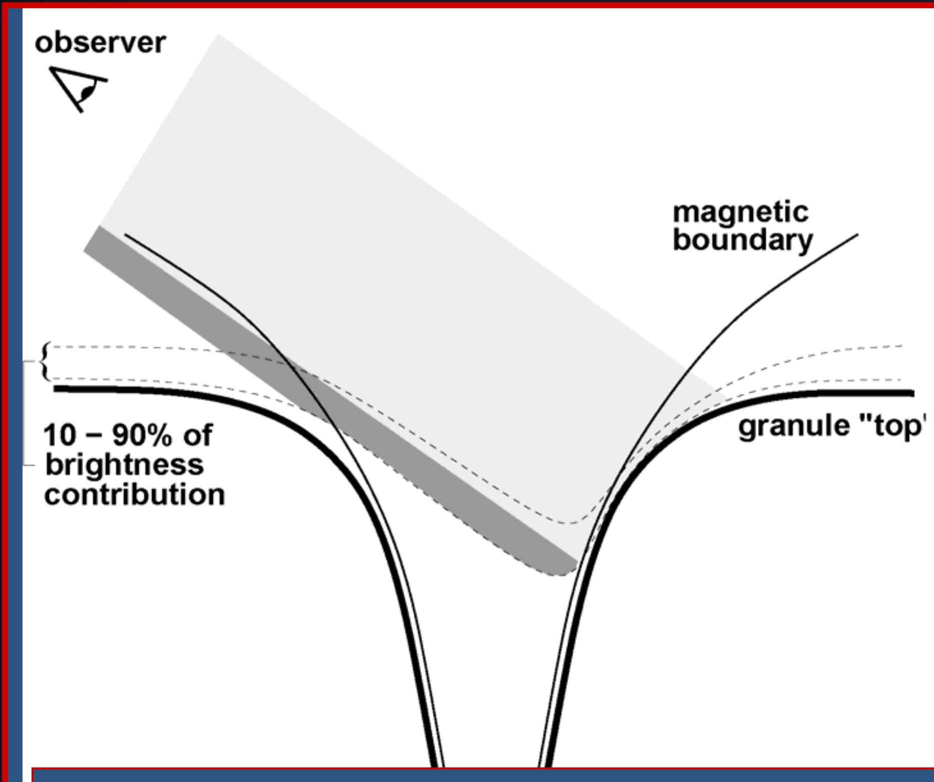


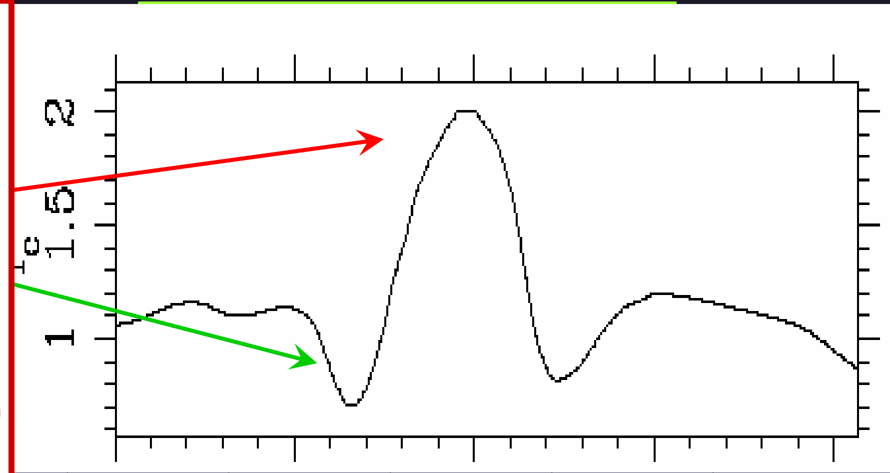
B_0=200 G: CLV of wavelength-integrated brightness

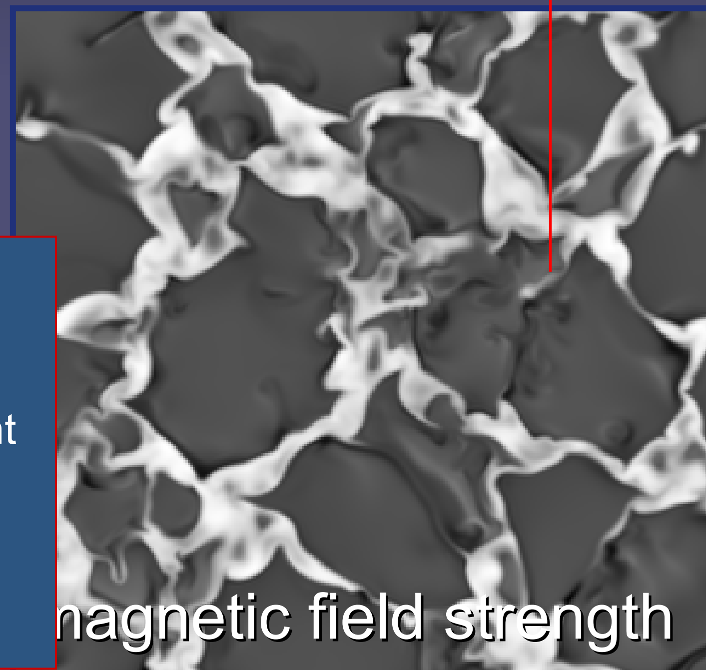




Facular brightening







Facula: narrow layer of hot material on side and top of adjacent granule

Dark lane: - cool & tenuous material in adjacent flux concentration
- cool & dense material above neighbouring granule

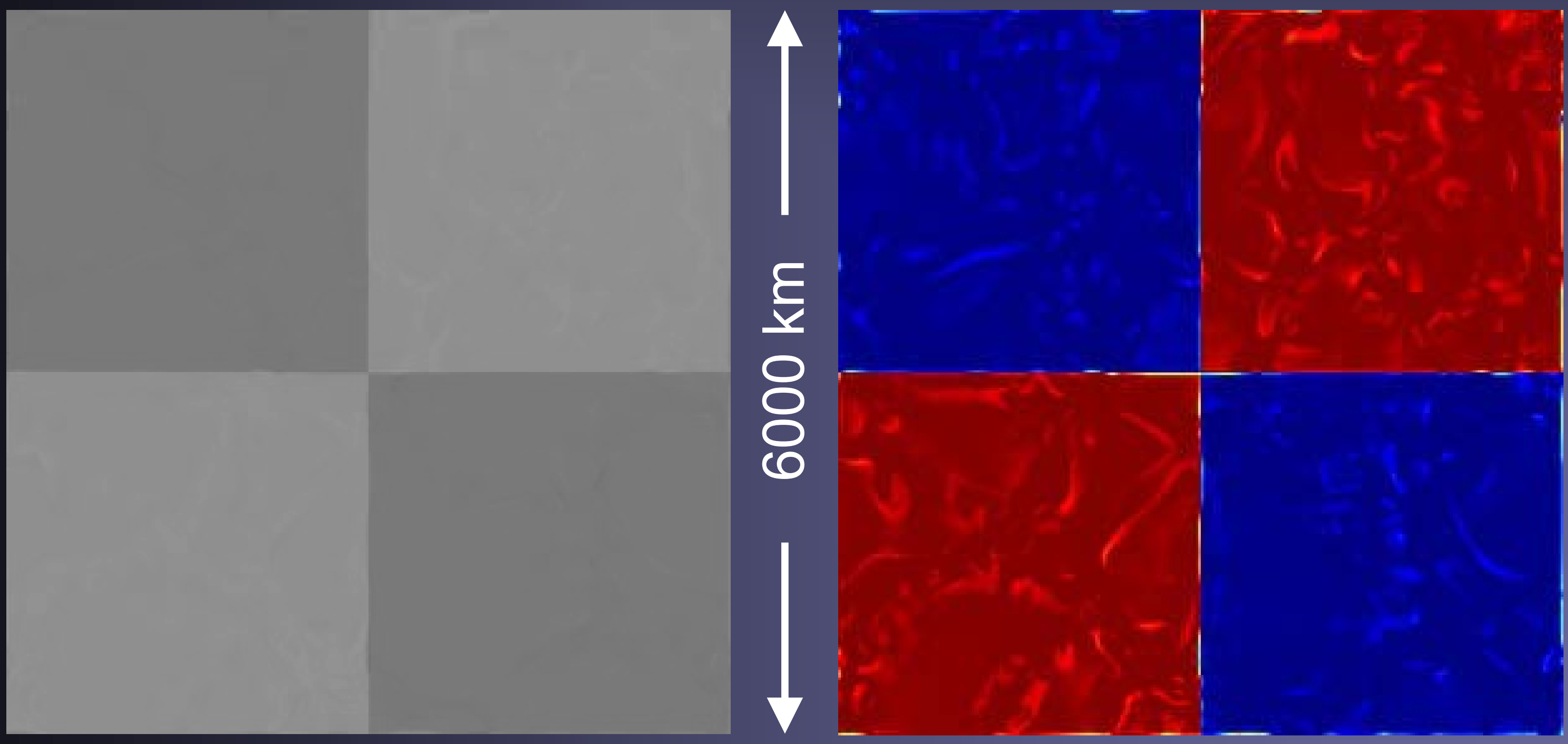
(Keller et al. 2004)

3-D simulations of mixed-polarity fields

Alexander Vögler, Robert Cameron, Manfred Schüssler

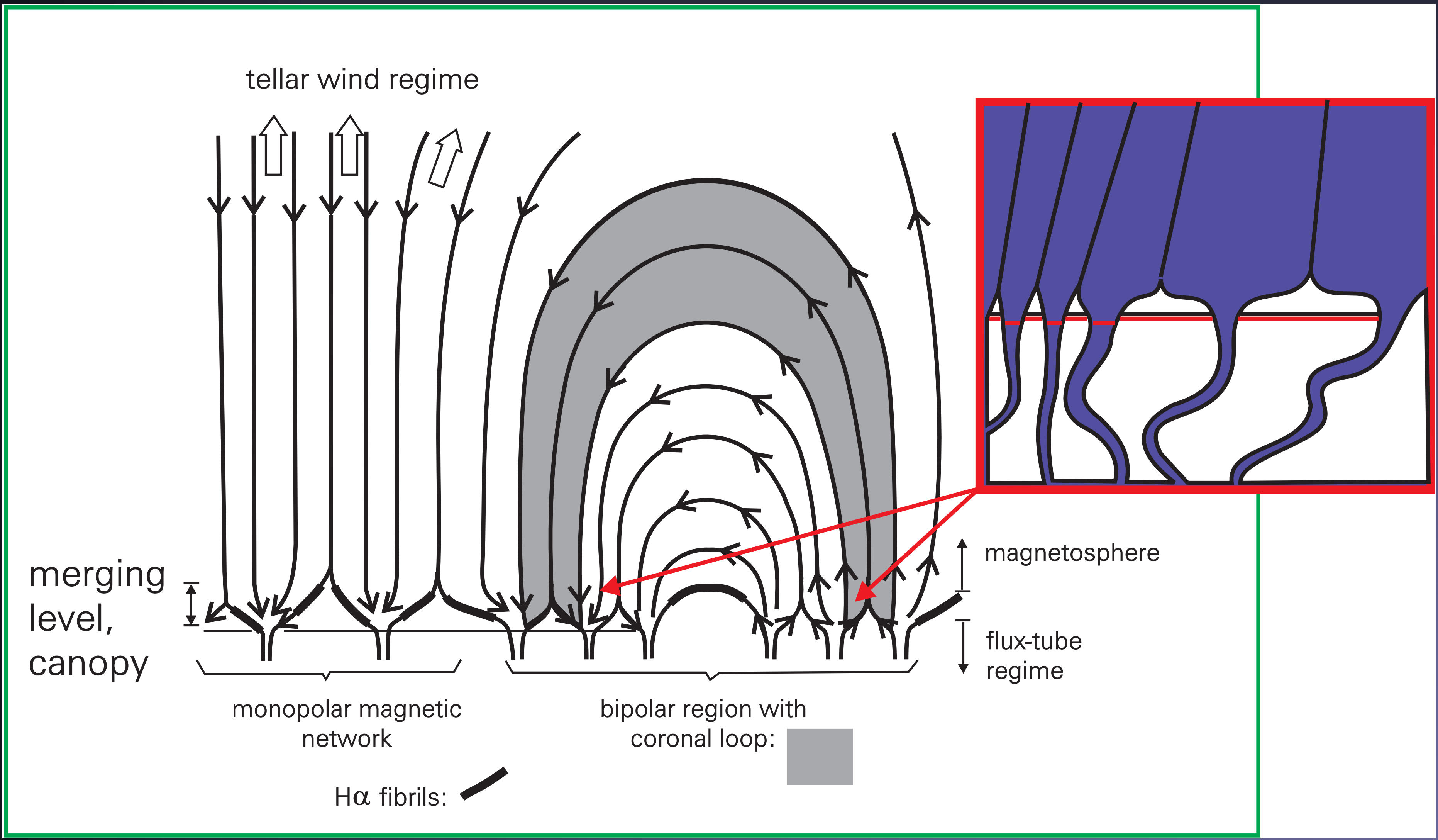
Mixed polarity simulations: diffusion & cancellation of opposite polarities (20 km resolution): $<B_{initial}>$=200G.

B_z $\qquad$ $B_z < 200$ G



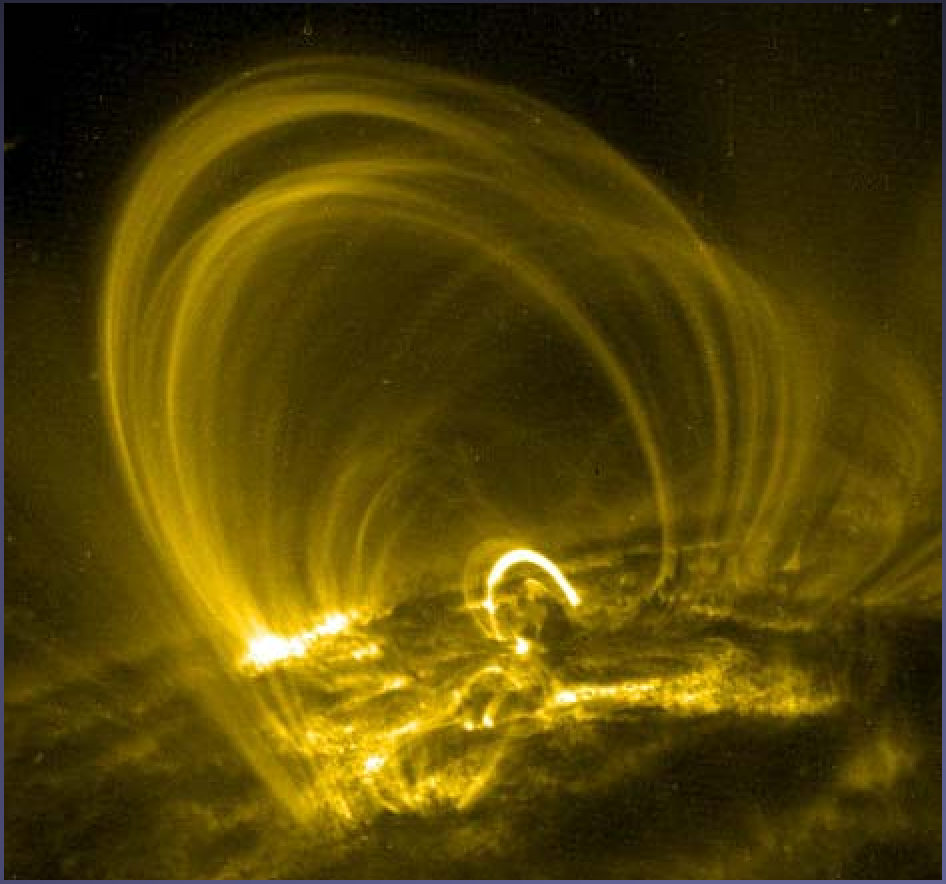
Simulations require computers...

Flux Tubes, Canopies, Loops and Funnels



Coronal loops

- Coronal loops are closed field lines in the corona. However, closed flux must be *filled with plasma* before it can be called a *coronal loop*.
- Loop temps range from below 0.1MK to 10MK
- Often a given observations (in a given spectral band) samples radiation in a narrow range of temps. It sees only a small fraction of all loops.



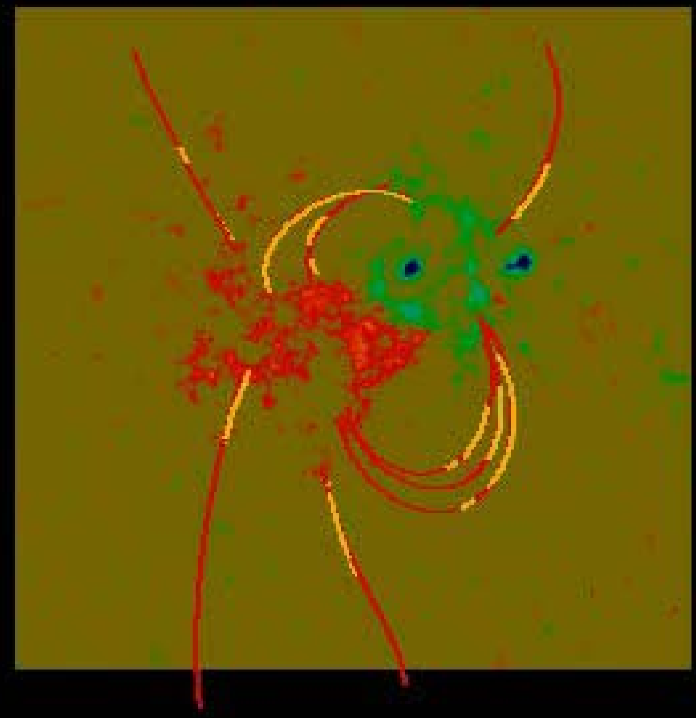
Loops at 0.9MK (TRACE Fe IX 171Å)

Coronal loops in 3-D

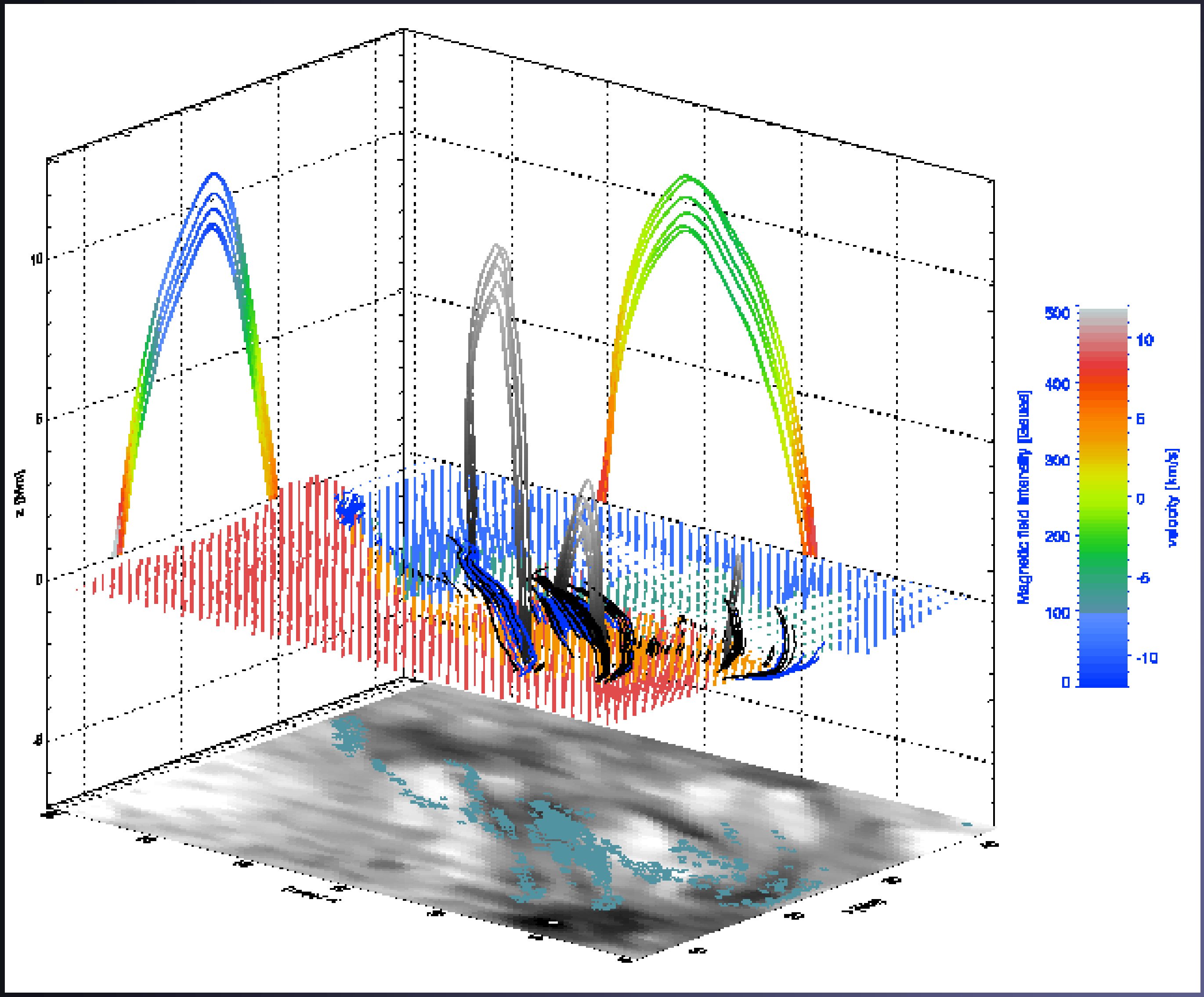
Yellow lines: First stereoscopic reconstruction of coronal loops observed by the two STEREO spacecraft looking at the Sun from different directions.

Red lines: magnetic field extrapolations starting from magnetogram on solar surface

Feng et al. 2007



Structure of Cool 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

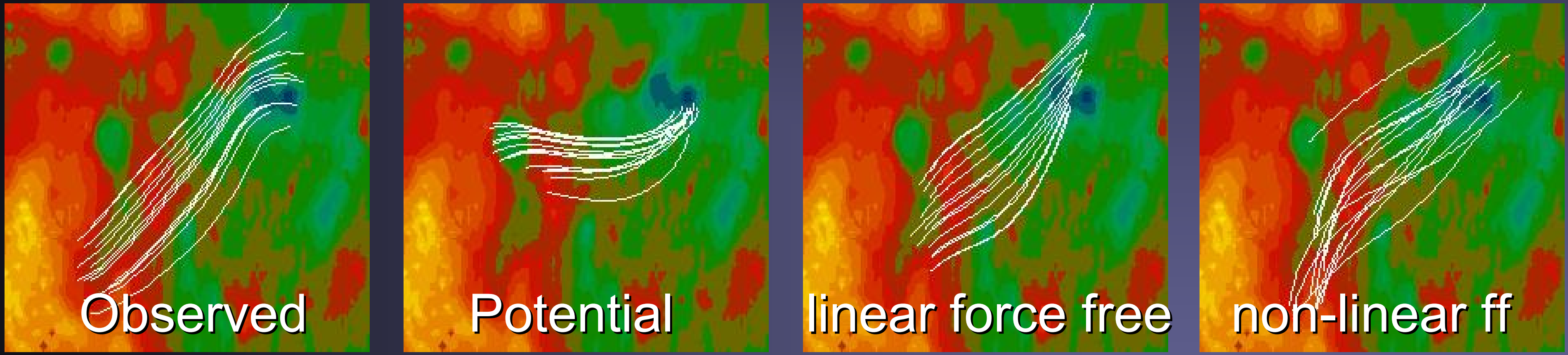
- General problem in solar physics: Magnetic field is measured mainly in the photosphere, but it makes music mainly in the corona.

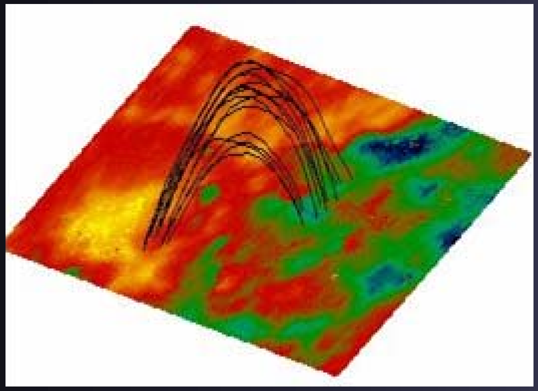
→ Either improve coronal field measurements (currently difficult) or extrapolate from photospheric measurements into the corona.

- 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).

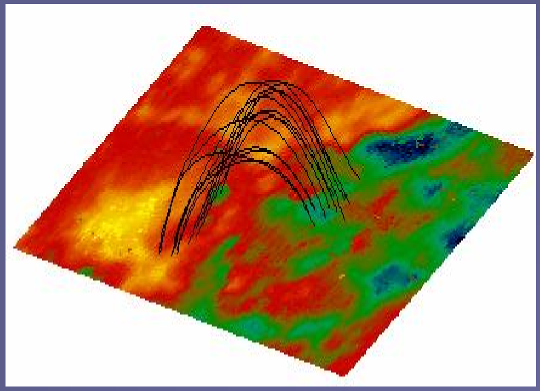
Testing Magnetic Extrapolations

- 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.

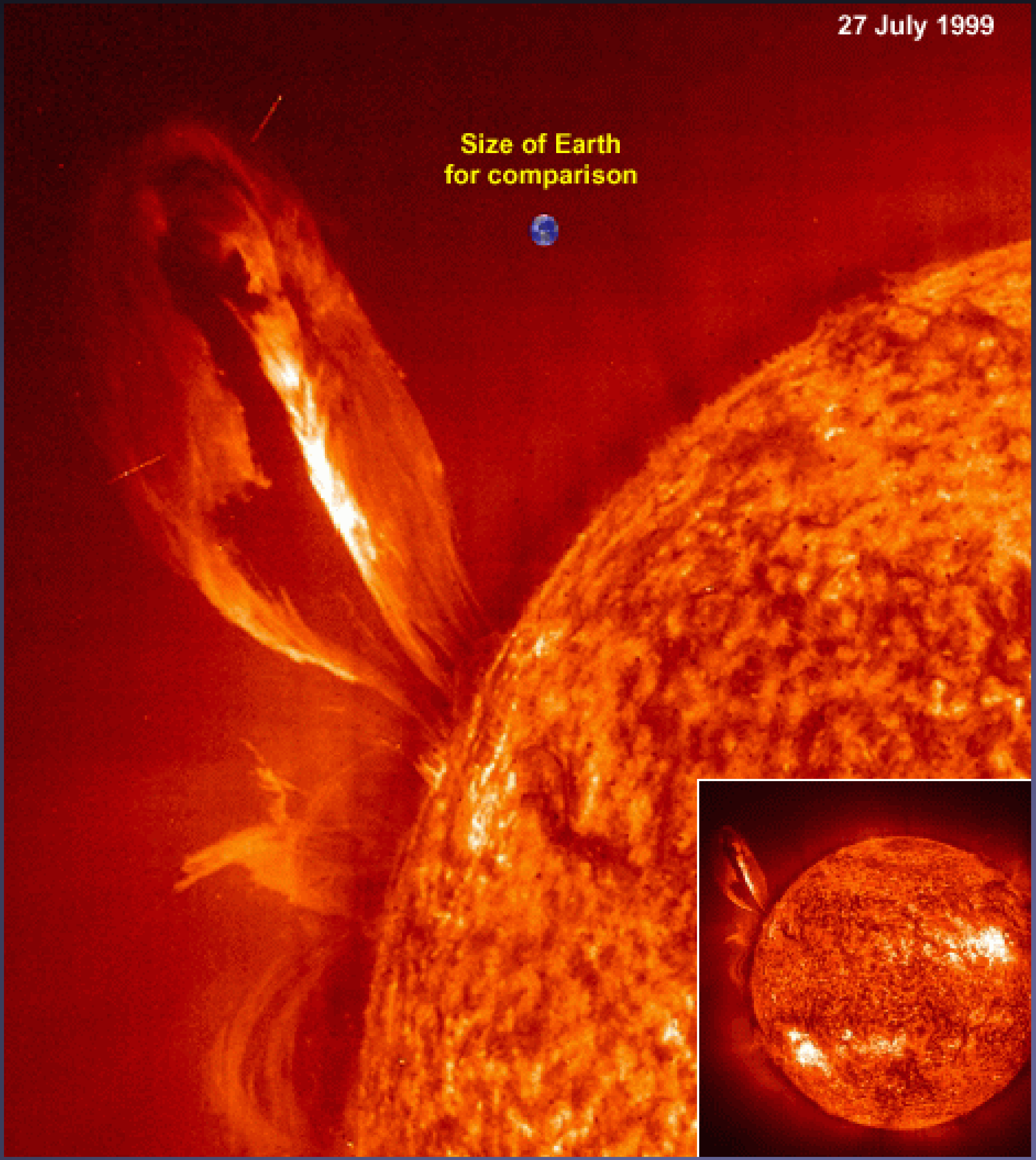




Wiegelmann et al. 2004

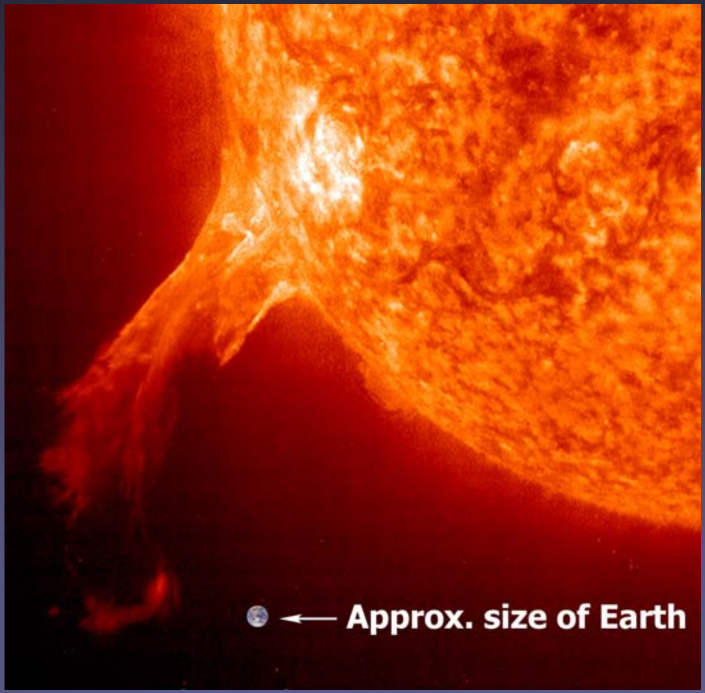


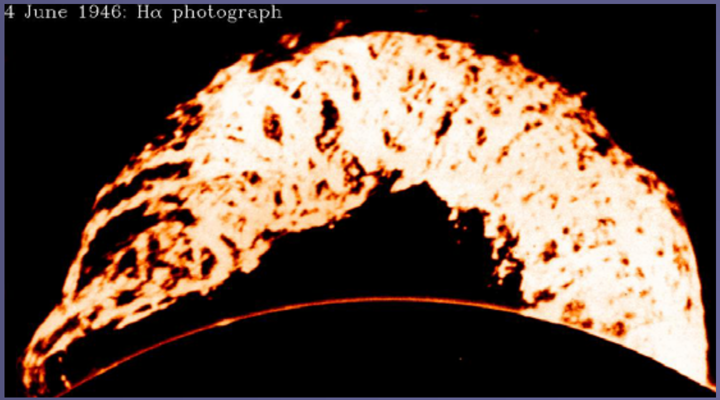
Prominences



Prominence material supported by magnetic field

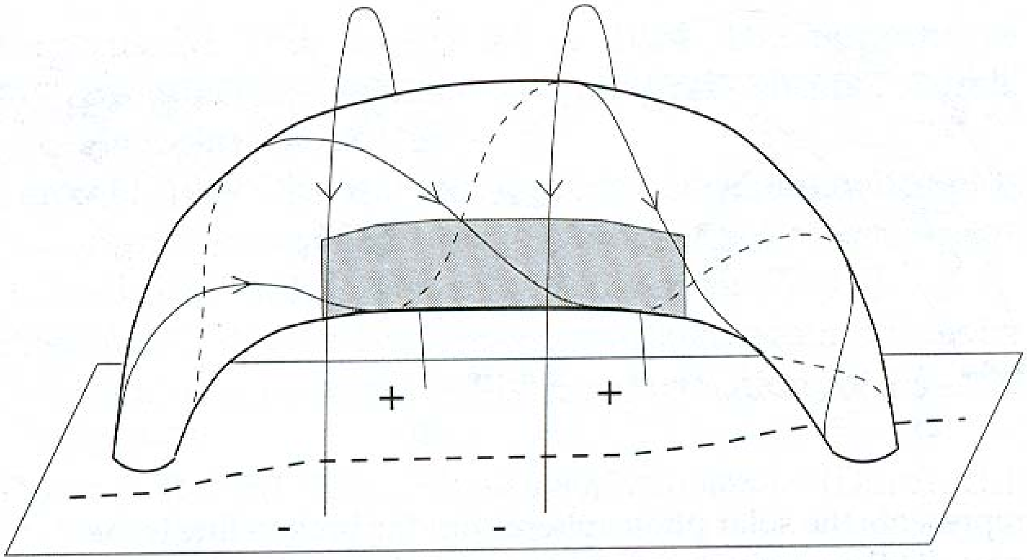
- Density of prominence material is orders of magnitude higher than of surrounding corona.
- Gas has to be supported against gravity.
- Magnetic field can provide this support, since ionized gas can only flow along field lines.

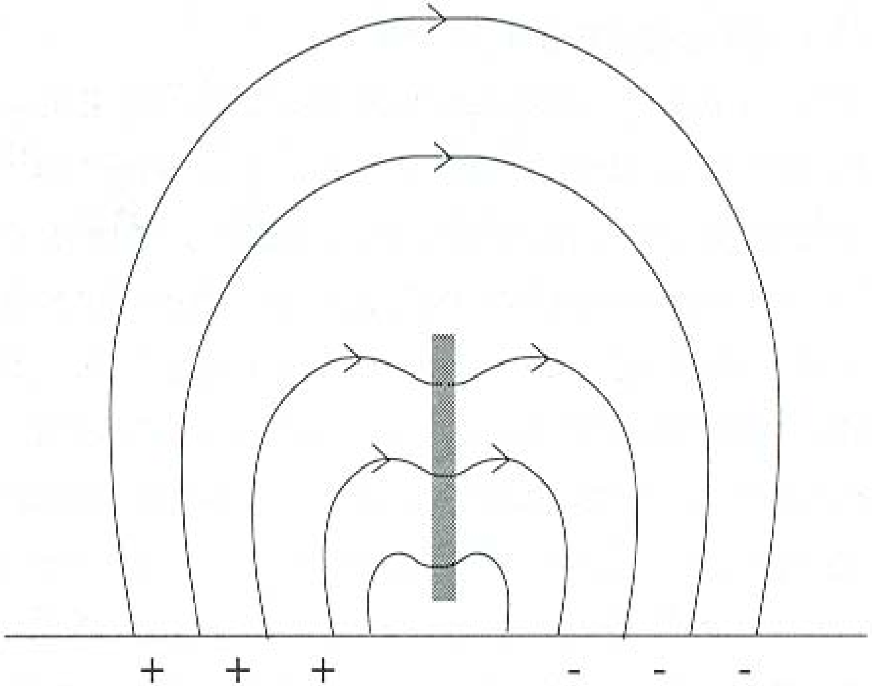


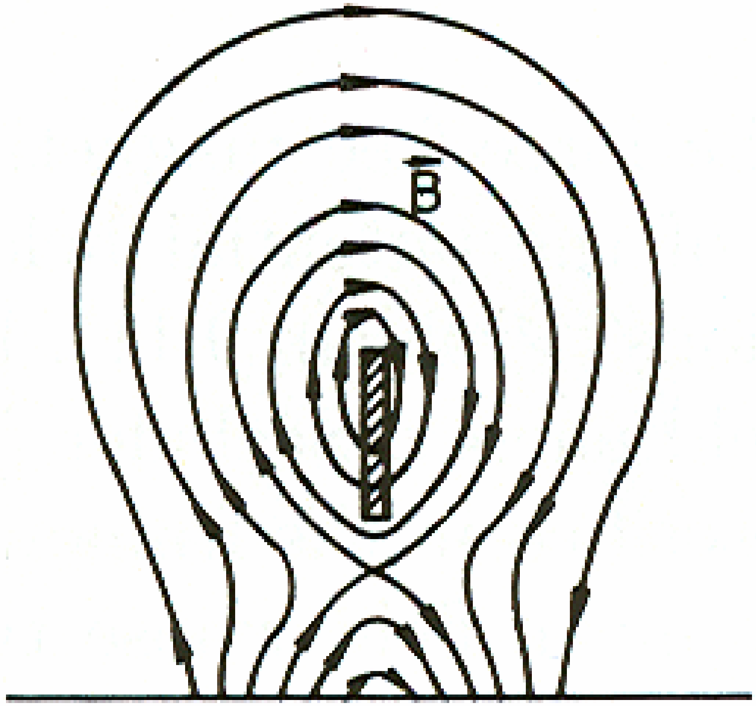


Prominence models

Kippenhahn-Schlüter (below), Kuperus-Raadu (below right) and flux tube (right; 3-D Kuperus-R.)

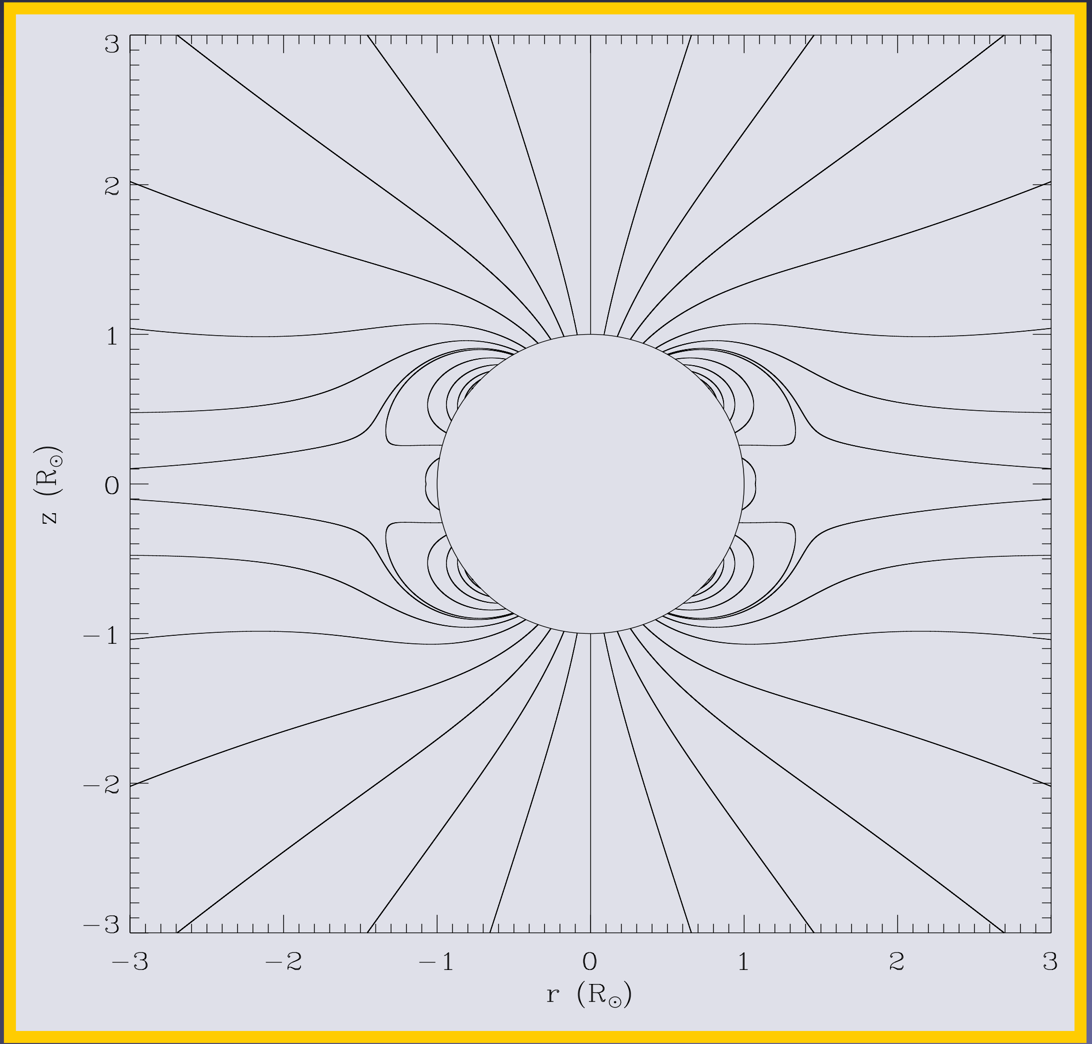






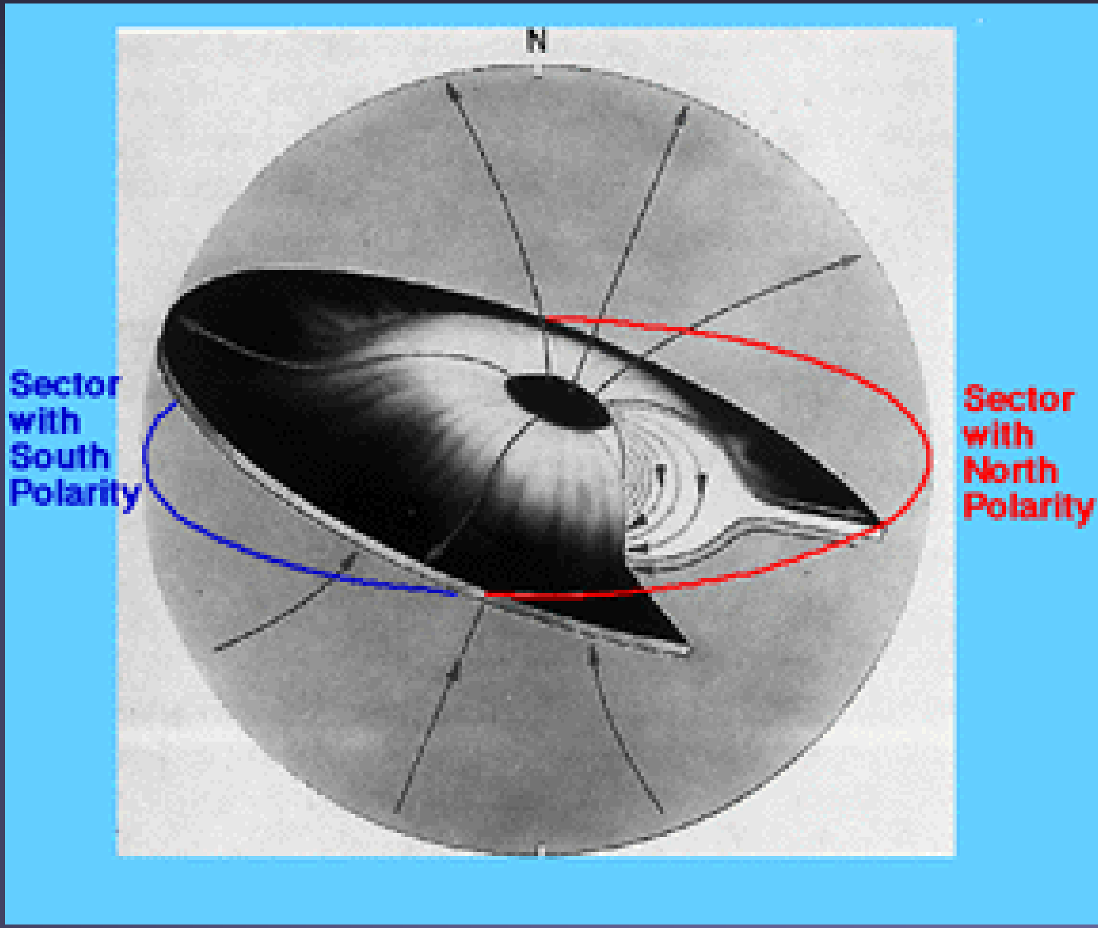
Large scale magnetic structure of the quiet Sun

- At large scales only dipolar component of 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.