Neutrinos in Cosmology and Astrophysics

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Talk Summary

- Lecture 1: Introduction to Cosmology
- Lecture 2: Neutrino Cosmology
- Lecture 3: Neutrino Astrophysics

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- **Lecture 1: Introduction to Cosmology**
- Lecture 2: Neutrino Cosmology
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Einstein Equations

$$R_{\mu\nu} - \frac{1}{2} g_{\mu\nu} R = 8\pi G T_{\mu\nu}$$

$$g_{\mu\nu} : \mathrm{metric}$$

$$R_{\mu\nu} = \Gamma^{\alpha}_{\mu\nu,\alpha} - \Gamma^{\alpha}_{\mu\alpha,\nu} + \Gamma^{\alpha}_{\beta\alpha}\Gamma^{\beta}_{\mu\nu} + \Gamma^{\alpha}_{\beta\nu}\Gamma^{\beta}_{\mu\alpha}$$

$$R = g^{\mu\nu} R_{\mu\nu}$$

Isotropic, uniform, expanding universe:

$$g_{\mu\nu} = diag(-1, a^2, a^2, a^2)$$

FRW metric

$$d\tau^2 = dt^2 - a^2 |d\vec{r}|^2$$

Important Equations of Universe Evolution

$$R_{00} = (...) = -3\frac{\ddot{a}}{a}$$

$$R_{ij} = (...) = \delta_{ij}(2\dot{a}^2 + a\ddot{a})$$

$$R = -R_{00} + \frac{1}{a^2}R_{ii} = 6\left[\frac{\ddot{a}}{a} + \left(\frac{\dot{a}}{a}\right)^2\right]$$

Left-hand side of Einstein Equation:

$$R_{\mu\nu} - \frac{1}{2}g_{\mu\nu}R = (...) = \left(\frac{\dot{a}}{a}\right)^2$$

Important Equations of Universe Evolution

For the right-hand side, we assume a perfect isotropic fluid:

$$T_{\mu\nu} = diag(\rho, P, P, P)$$

Energy density

Pressure

Putting everything together for the 00 element of Einstein Equation:

$$\left(\frac{\dot{a}}{a}\right)^2 = \frac{8\pi G}{3}\rho$$

Energy content of the Universe dictate its evolution

Important Equations of Universe Evolution

Same equation, new nomenclature:

$$\begin{aligned} H &= \dot{a}/a \\ H_0 &= \dot{a}/a)_{\text{today}} \\ \rho_{cr} &= 3H_0^2/8\pi G \\ \Omega_i &= \rho_i/\rho_{cr} \end{aligned}$$

$$\frac{H^2}{H_0^2} = \sum_i \frac{\rho_i}{\rho_{cr}} = \sum_i \Omega_i$$

Friedmann Equation

and some numbers:

$$H_0 = 67.3 \pm 1.2\,\mathrm{km/s/Mpc} = h \times (100\mathrm{km/s/Mpc})$$

latest Planck results

Important Equations of Universe Evolution

Energy Conservation:

$$T^{\mu}_{\nu,\mu} = \frac{\partial T^{\mu}_{\nu}}{\partial x^{\mu}} + \Gamma^{\mu}_{\alpha\mu} T^{\alpha}_{\nu} - \Gamma^{\alpha}_{\nu\mu} T^{\mu}_{\alpha} = 0$$

For $\nu = 0$:

$$\frac{\partial \rho}{\partial t} + \frac{\dot{a}}{a}\left[3\rho + 3P\right] = 0$$

Important Equations of Universe Evolution

Radiation (in thermal equilibrium): $P = \rho/3$

$$\frac{\partial \rho}{\partial t} + 4\frac{\dot{a}}{a}\rho = \frac{1}{a^4}\frac{\partial}{\partial t}\left[\rho a^4\right] = 0 \rightarrow \rho \sim a^{-4}$$

How does the Universe evolve?

$$\left(\frac{\dot{a}}{a}\right)^2 = \frac{8\pi G}{3}\frac{\rho_0}{a^4} \longrightarrow a(t) = \left(\frac{32\pi G\rho_0}{3}\right)^{1/4} t^{1/2}$$

Important Equations of Universe Evolution

(non-relativistic) matter: $P \sim v^2 << \rho$

$$\frac{\partial \rho}{\partial t} + 3\frac{\dot{a}}{a}\rho = \frac{1}{a^3}\frac{\partial}{\partial t}\left[\rho a^3\right] = 0 \quad \rightarrow \quad \rho \sim a^{-3}$$

How does the Universe evolve?

$$\left(\frac{\dot{a}}{a}\right)^2 = \frac{8\pi G}{3}\frac{\rho_0}{a^3} \quad \rightarrow \quad a(t) = \left(\frac{18\pi G \rho_0}{3}\right)^{1/3} t^{2/3}$$

Important Equations of Universe Evolution

Cosmological Constant:

$$\rho = \text{constant} = -P$$

How does the Universe evolve?

$$\left(\frac{\dot{a}}{a}\right)^2 = \frac{8\pi G}{3}\rho_0 \longrightarrow a(t) = exp\left[\left(\frac{8\pi G\rho_0}{3}\right)^{1/2} t\right]$$

Back to Friedmann equation:

$$\left(\frac{\dot{a}}{a}\right)^2 = \frac{8\pi G}{3}\left[\frac{\rho_0^{rad}}{a^4} + \frac{\rho_0^{nr}}{a^3} + \rho_\Lambda\right]$$

$$\begin{aligned} \rho_0^{rad} &: \quad \text{photons and neutrinos} \\ \rho_0^{nr} &: \quad \text{atoms and cold dark matter} \\ \rho_\Lambda &: \quad \text{cosmological constant} \end{aligned}$$

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1978 Nobel, Penzias and Wilson
2006 Nobel, Mather and Smoth

2011 Nobel, Perlmutter, Schmidt and Riess

Our concern here

Back to Friedmann equation:

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more numbers, again from Planck:

$$\begin{aligned}
\Omega_\gamma h^2 &= 2.47 \times 10^{-5} \\
\Omega_c h^2 &= 0.1199 \pm 0.0027 \\
\Omega_b h^2 &= 0.02205 \pm 0.00028 \\
\Omega_\Lambda &= 0.685^{+0.018}_{-0.016}
\end{aligned}$$

Cosmological Neutrinos

- Hom many neutrinos?

  - thermal equilibrium (with Fermi-Dirac distribution) with plasma for high enough temperature:

$$
\begin{aligned}
n + \nu_e &\leftrightarrow p^+ + e^- \\
n + e^+ &\leftrightarrow p^+ + \bar{\nu}_e
\end{aligned}
$$

  - neutrinos and photons (through scattering with electrons) share the same temperature, i.e., the same number density for each neutrino flavor, apart from a 7/8 numerical factor distinguishing boson and fermions distributions.

$$
n_\nu = 3 \times \frac{7}{8} n_\gamma \sim 1180\,\mathrm{cm}^{-3} \left(\times \frac{1}{a^3} \right)
$$

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If in thermal equilibrium

Cosmological Neutrinos

- Hom many neutrinos?

- outside thermal equilibrium: boltzman equations for *1+2* ↔ *3+4* reaction in expanding universe:

$$a^{-3}\frac{d(n_1 a^3)}{dt} = n_1^{(0)} n_2^{(0)} < \sigma v > \left\{ \frac{n_3 n_4}{n_3^{(0)} n_4^{(0)}} - \frac{n_1 n_2}{n_1^{(0)} n_2^{(0)}} \right\}$$

Roughfly expansion rate

Reaction rate

<< equilibrium
>> decoupling

For neutrinos: T ~ 1MeV, but still same number density of photons
Let's calculate this number?

Cosmological Neutrinos

- Hom many neutrinos?

  - photons reheat after neutrino decoupling due to e+e- anihilation:

$$\frac{T_\gamma}{T_\nu} = \left(\frac{11}{4}\right)^{1/3}$$

Calculated through entropy density arguments

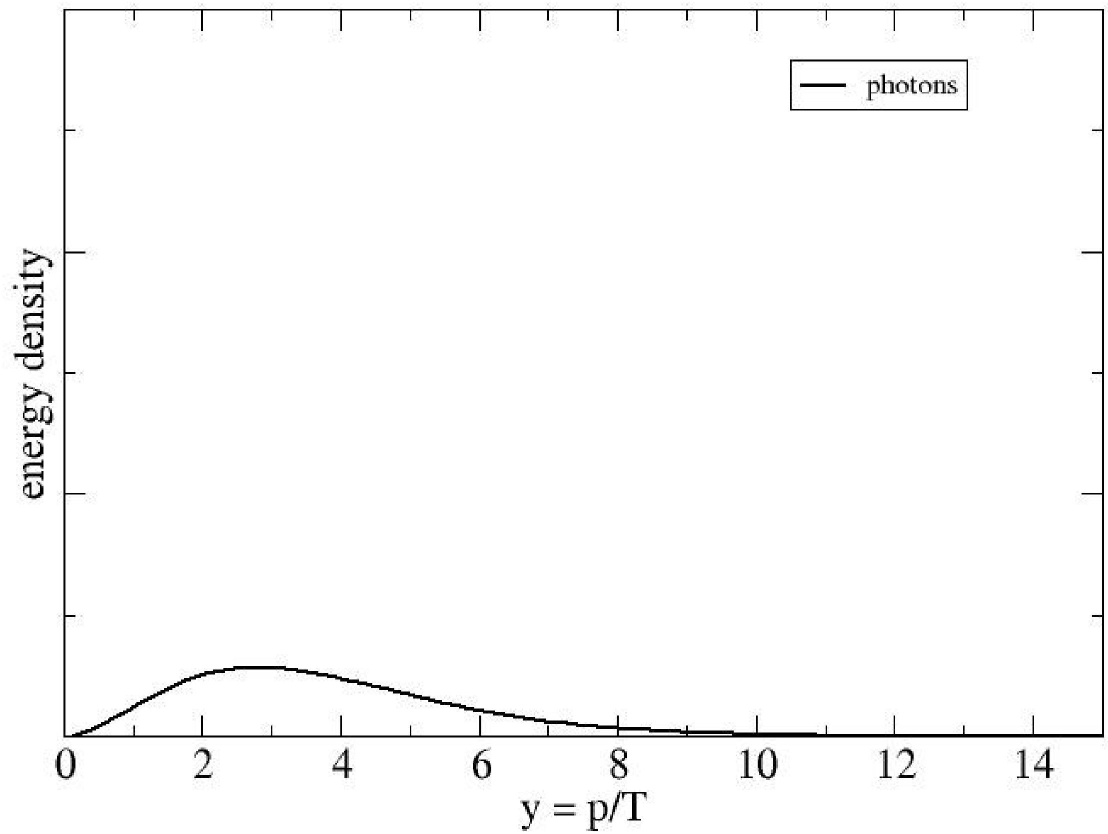
  - leaving fewer neutrinos than photons:

$$\frac{n_\gamma}{n_\nu} = \frac{3}{11}$$

But still a lot:
113 ν/cm^3 per flavour

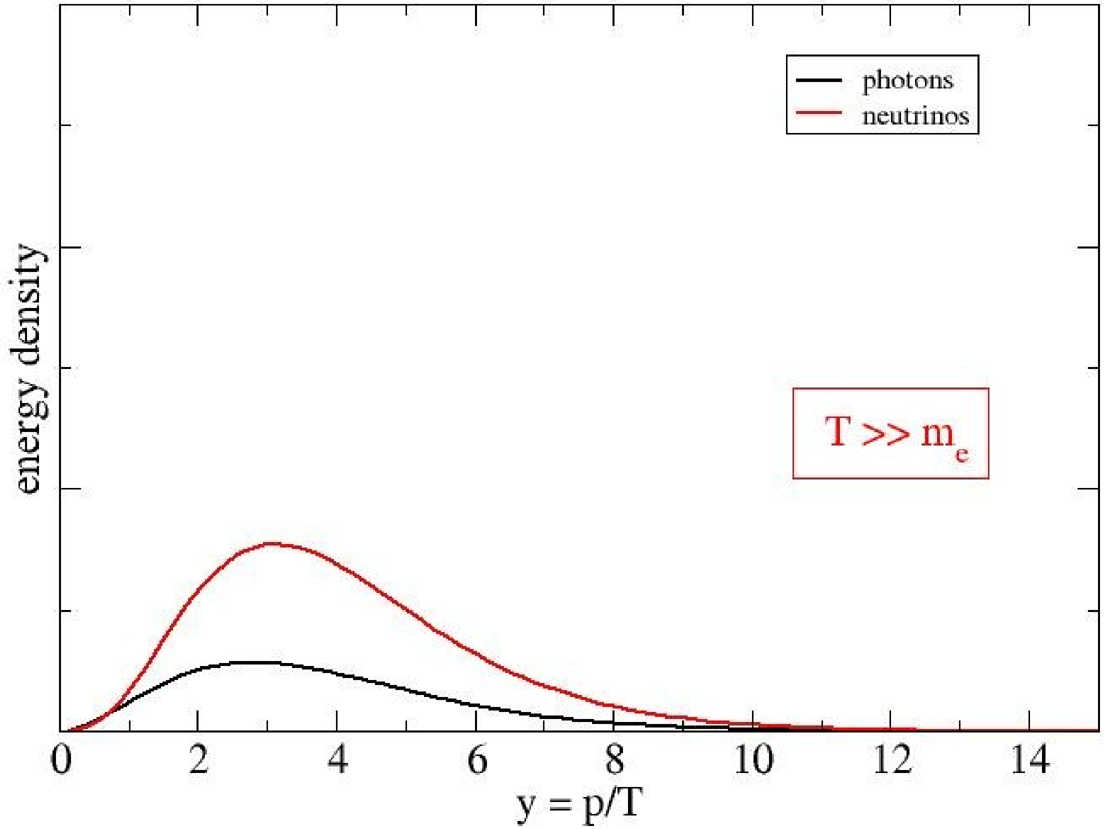
Cosmological Neutrinos

- cosmic microwave background spectrum, CMB (Blackbody radiation)



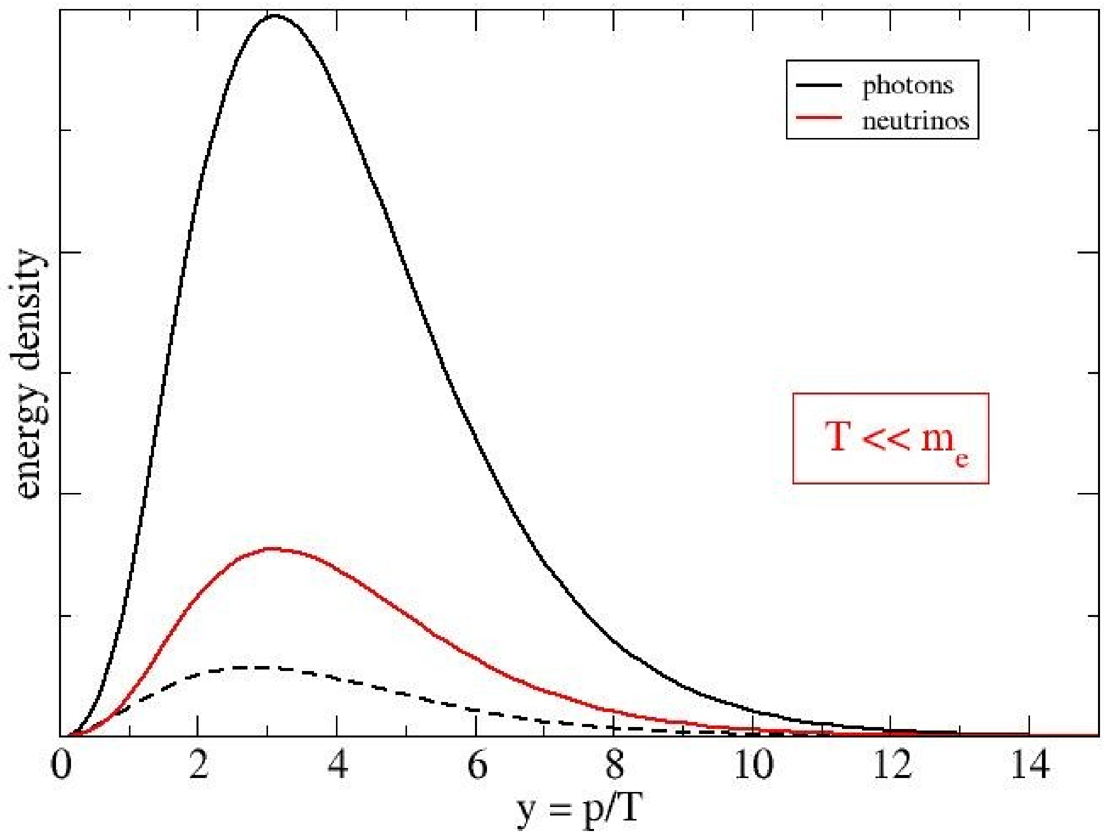
Cosmological Neutrinos

- 3 neutrinos for each photon, sharing the same temperature (but following a Fermi-Dirac distribution). Neutrino decoupling occurs for T ~ few MeV.



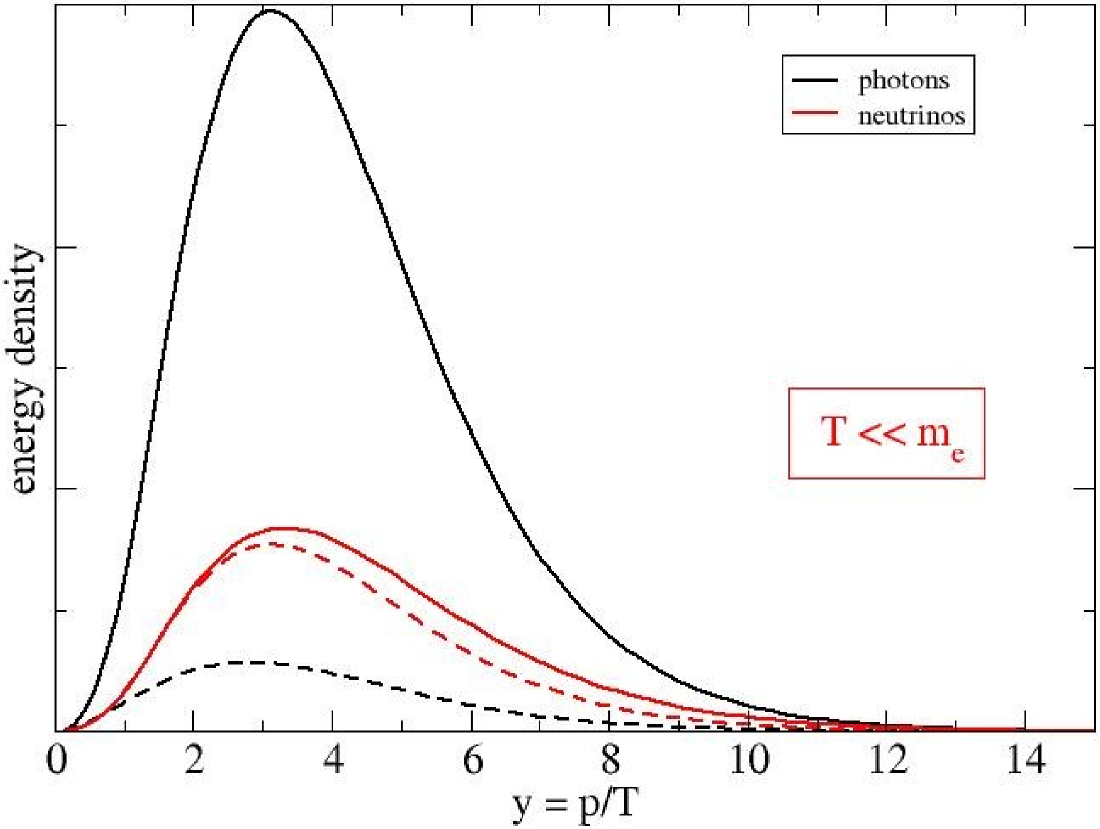
Cosmological Neutrinos

- Photon Reheating through $e^+ + e^-$ annihilation.



Cosmological Neutrinos

- Since higher energy neutrinos decouple last, some are still in contact with plasma during $e^+ + e^-$ annihilation → partial heating of neutrinos (x10 in plot).



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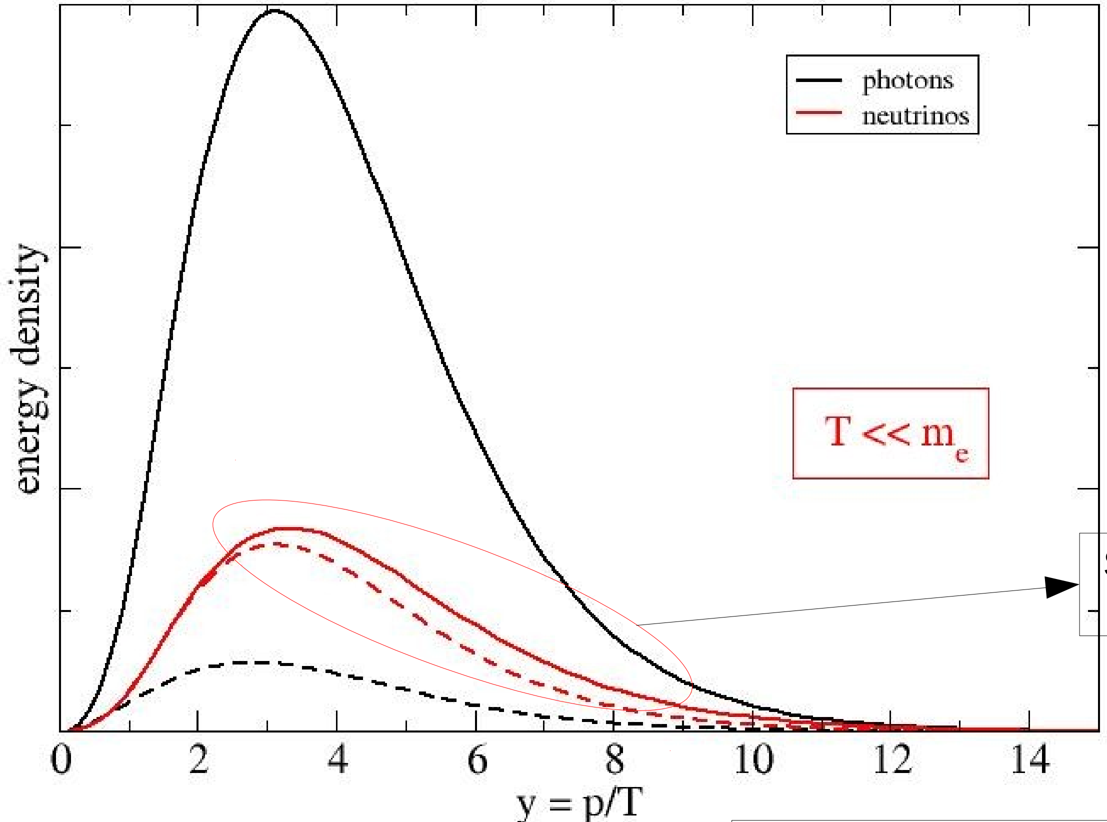

Slightly non-thermal distribution

Neff=3.046

Multiplicative factor which would provide the same energy density with perfect thermic spectrum

Cosmological Neutrinos

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  - Relativistic neutrinos:

$$\rho_\nu(m_\nu << T_\nu) = \frac{7\pi^2}{120}\left(\frac{4}{11}\right)^{4/3} T_\gamma^4 \rightarrow \Omega_\nu h^2 = 1.68 \times 10^{-5}$$

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$$\Omega_\nu h^2 > 5 \times 10^{-4}$$

Other important moments in Universe history

- At first, radiation dominated the energy content of the Universe (scales as a^{-4})

- As the universe cools down, matter grows in importance (scales as a^{-3})

Equality happens when:

$$\frac{\rho_0^m}{a^3} = \frac{\rho_0^{rad}}{a^4} \longrightarrow a = \frac{\rho_0^{rad}}{\rho_0^m} \sim 10^{-4}$$

Transition between pressured to pressureless Universe

Structures start to grow!

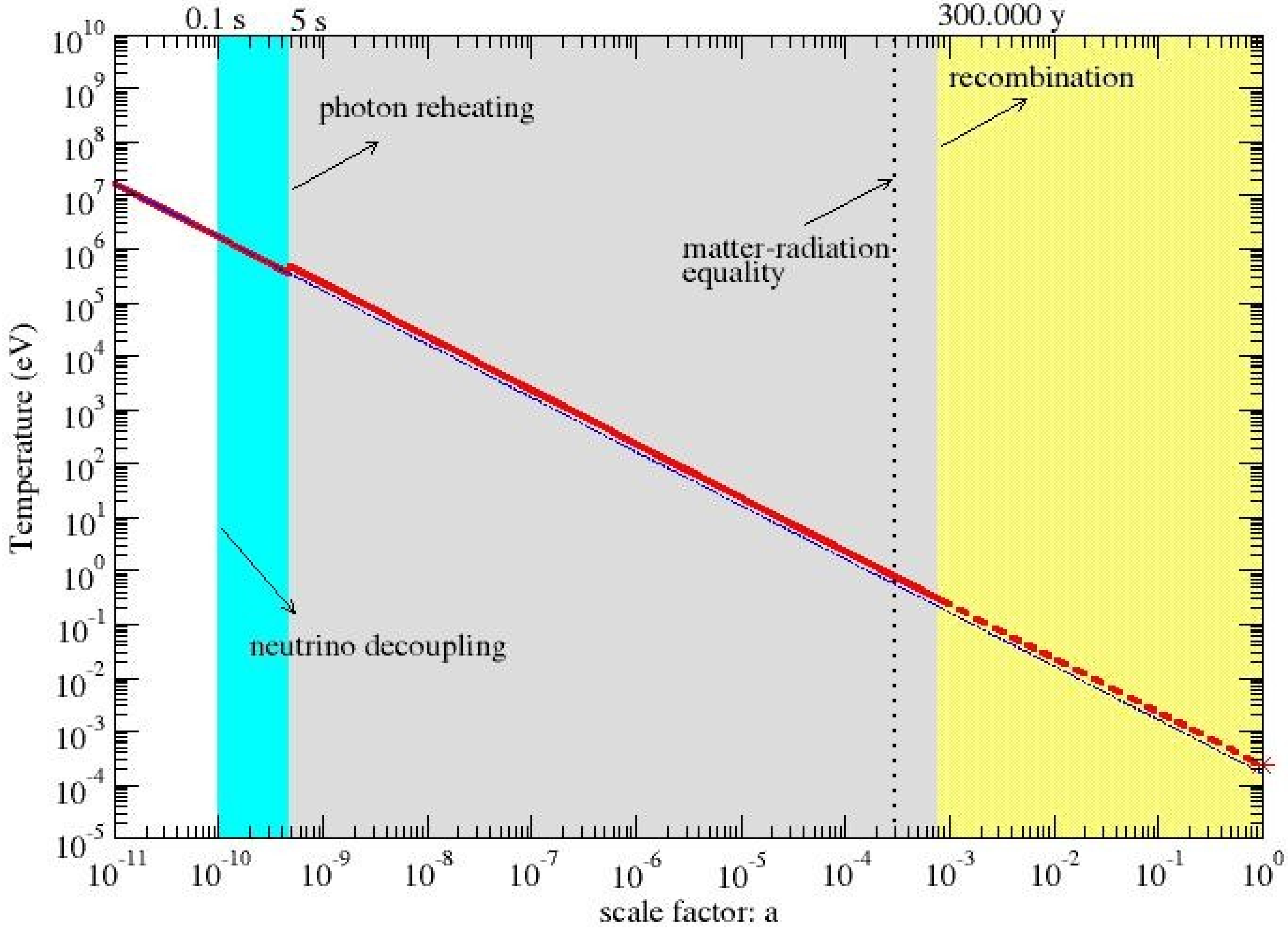
Other important moments in Universe history

- CMB decoupling: when universe cools down enough even the most energetic photons are not energetic or numerous enough to prevent atoms to form.

- Electrons are captured by protons, forming Hydrogen atoms.

- Photons scattering cross-section drops, and photons travel freely since.

RECOMBINATION



0.1 s
5 s
300.000 y
photon reheating
recombination
matter-radiation equality
neutrino decoupling
Temperature (eV)
scale factor: a

Cosmological Observables

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BBN: Big-Bang Nucleosynthesis:

What is the proton/neutron ratio in our Universe?

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What is the proton/neutron ratio in our Universe? (tip: take Sun as ”typical”)

Cosmological Observables

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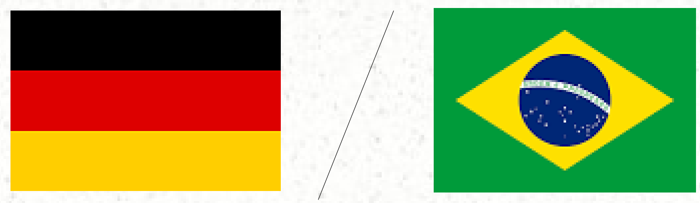
Answer: 

Cosmological Observables

BBN: Big-Bang Nucleosynthesis:

What is the proton/neutron ratio in our Universe?

Answer: 1 neutron for each ~7 protons

Why?

- neutrons outside an atom are unstable → no neutrons today
- neutrons and protons were onde in thermal equilibrium → same number

Cosmological Observables

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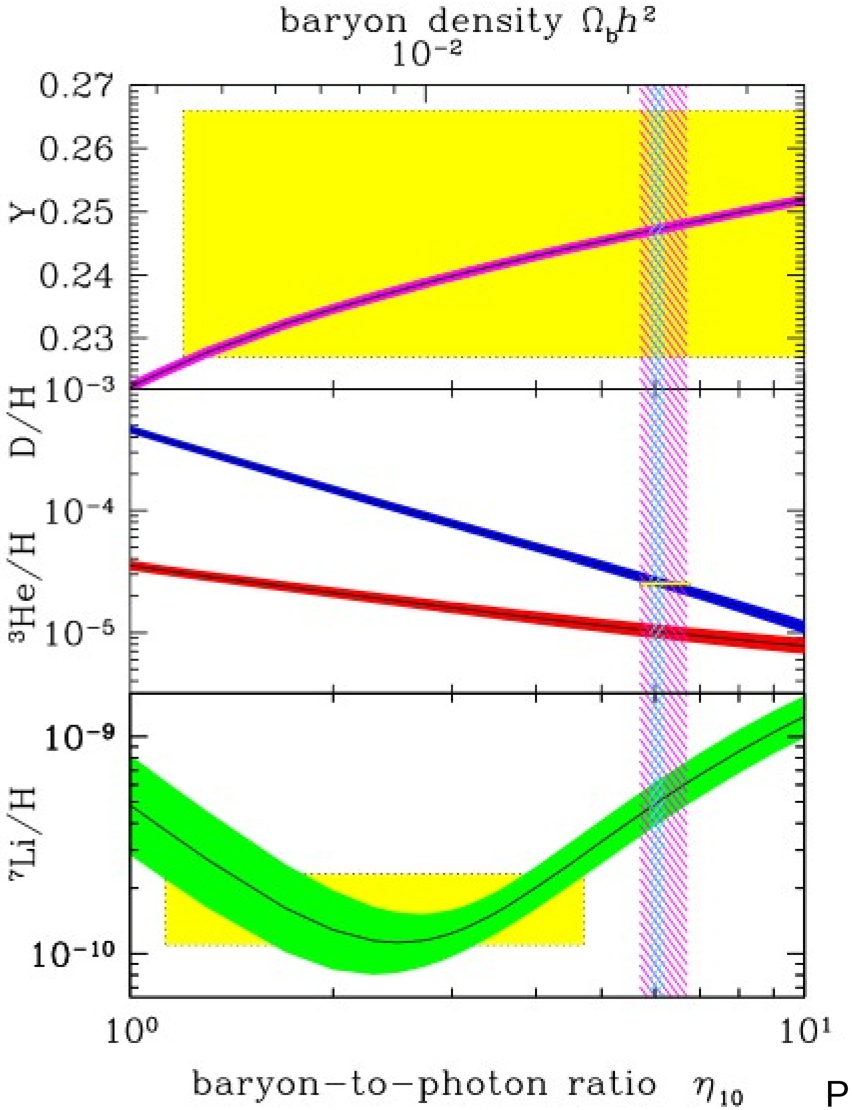
Answer: neutrinos

$$\begin{aligned} n + \nu_e &\leftrightarrow p^+ + e^- \\ n + e^+ &\leftrightarrow p^+ + \bar{\nu}_e \end{aligned}$$

- guarantees equilibrium when $T \gg m_{n,p}$.
- enhance neutron conversion when $T < m_{n,p}$.
- decouple at the right moment to produce 7:1.

Cosmological Observables

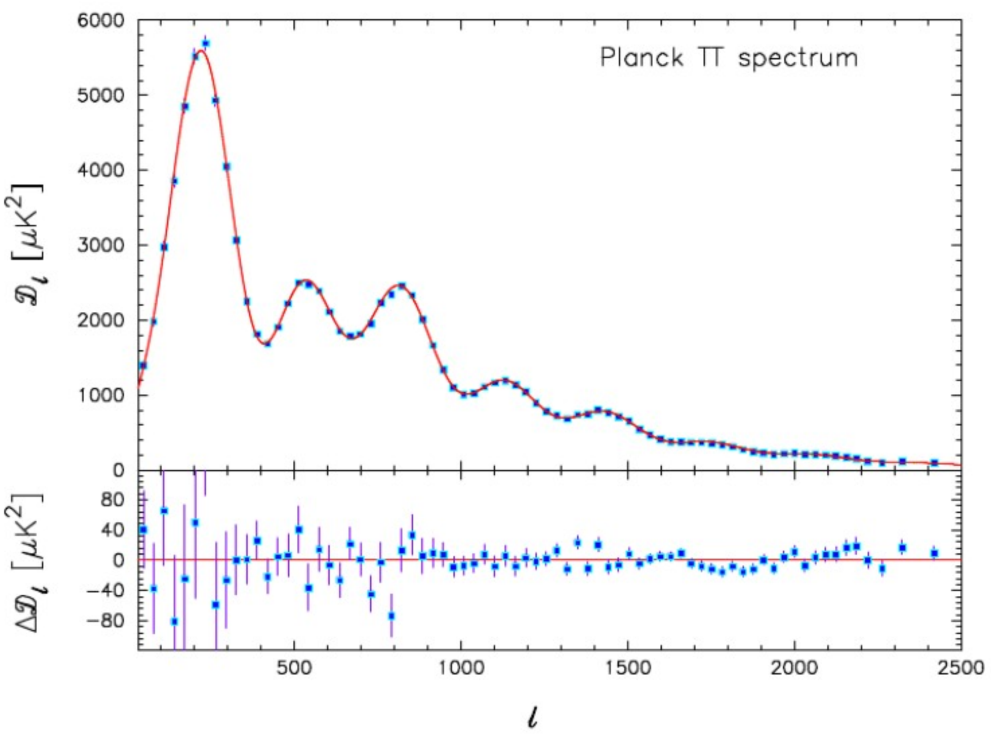


PDG

Cosmological Observables

CMB anisotropies:

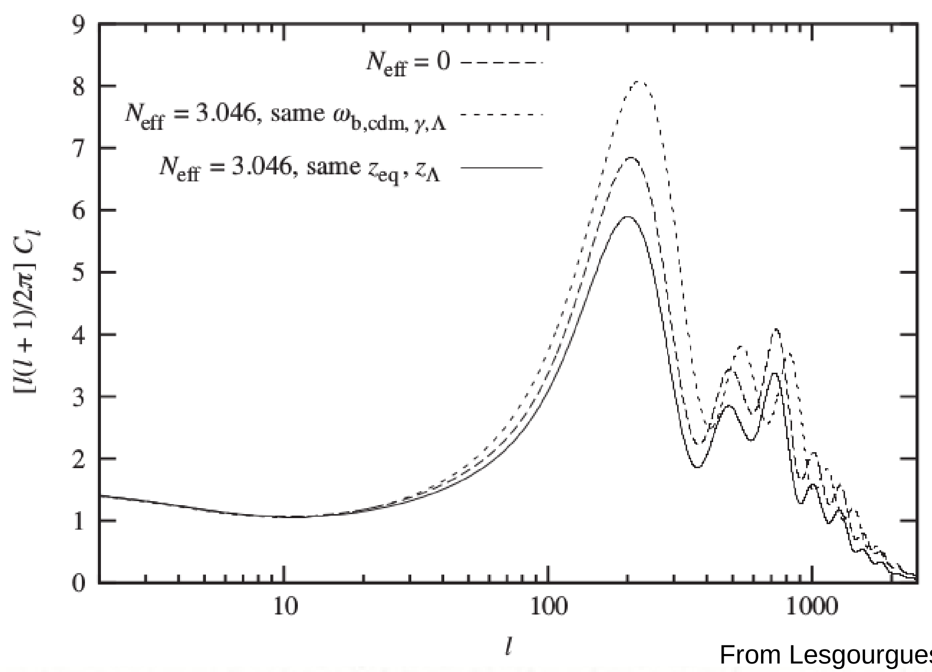
Early universe was almos isotropic, and small unisotropy is imprinted in CMB fluctuations.



Cosmological Observables

CMB anisotropies:

(Massless) Neutrinos affect anisotropies through:
- matter-radiation energy density equality time
- perturbations



From Lesgourgues et al, ”Neutrino Cosmology”

How we know they are there

Structure Formation:

- Left out because choices must be made (and because we would be talking about non-linear effects, which is hard)...

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Enought for today. Tomorrow we put mass no neutrinos!